Observing the Galactic Plane with the Balloon-borne Large-Aperture Submillimeter Telescope

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Abstract

Stars form from collapsing massive clouds of gas and dust. The UV and optical light emitted by a forming or recently-formed star is absorbed by the surrounding cloud and is re-radiated thermally at infrared and submillimetre wavelengths. Observations in the submillimetre spectrum are uniquely sensitive to star formation in the early Universe, as the peak of the thermal emission is redshifted to submillimetre wavelengths. The coolest objects in star forming regions in our own Galaxy, including heavily-obscured proto-stars and starless gravitationally-bound clumps, are also uniquely bright in the submillimetre spectrum. The Earth’s atmosphere is mostly opaque at these wavelengths, however, limiting the spectral coverage and sensitivity achievable from ground-based observatories.

The Balloon-borne Large Aperture Submillimeter Telescope (BLAST) observes the sky from an altitude of 40 km, above 99.5% of the atmosphere, using a long-duration scientific balloon platform. BLAST observes at 3 broad-band wavelengths spanning 250–500 μm, taking advantage of detector technology developed for the space-based instrument SPIRE, scheduled for launch in 2008. The greatly-enhanced atmospheric transmission at float altitudes, increased detector sensitivity and large number of detector elements allow BLAST to survey much larger fields in a much smaller time than can be accomplished with ground-based instruments. It is expected that in a single 10-day flight, BLAST will detect \( \sim 10000 \) extragalactic sources, \( \sim 100 \times \) the number detected in 10 years of ground-based observations, and 1000s of Galactic star-forming sources, a large fraction of which are not seen by infrared telescopes.

The instrument has performed 2 scientific flights, in the summer of 2005 and winter of 2006, for a total of 16 days of observing time. This thesis discusses the design of the instrument, performance of the flights, and presents the analysis of 2 of the fields observed during the first flight. A failure in the optical system during the first flight precluded sensitive extragalactic observations, so the majority of the flight was spent observing Galactic targets. We anticipate exciting extragalactic and Galactic results from the 2006 data.
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Glossary

bolometer A thermal detector sensitive to a wide range of radiation frequencies.

cross-linking The process of scanning a field along different directions to provide map-making constraints on large spatial scales.

CSBF (Columbia Scientific Balloon Facility) A NASA facility that provides launch, communication, and recovery services for high altitude scientific balloons.

FIR (far infrared) A region of the electromagnetic spectrum defined roughly as wavelengths spanning 8–1000 μm.

flux density The brightness of a point source per unit frequency, usually denoted $S_\nu$ and measured in Janskies (Jy).

FWHM (full-width at half-maximum) A characteristic width of a single-peaked profile defined as the distance between the points on the curve at half of the peak value. For a Gaussian profile of width $\sigma$, the FWHM = $2\sqrt{2\ln 2} \sigma$.

HMPO (high-mass proto-stellar object) A pre-stellar object, with no associated radio emission, that may eventually form a massive star.

IRAS (Infrared Astronomy Satellite) A NASA space telescope designed to observe the entire sky at infrared wavelengths. It observed at 12, 25, 60 and 100 μm and was launched in January 1983.

JCMT (James Clerk Maxwell Telescope) A 15-m submillimetre telescope located at the peak of Mauna Kea in Hawaii.

Jy A unit of flux density named after astronomer Carl Jansky, defined as $10^{-26}$ W m$^{-2}$ Hz$^{-1}$.
**ΛCDM (Lambda-Cold Dark Matter)**  The simplest cosmological model that explains current observations, including the cosmic microwave background, large-scale structure and supernovae. It states that the mass content of the Universe is dominated by a cosmological constant $\Lambda$ and cold dark matter.

**modified blackbody**  A blackbody spectrum multiplied by emissivity $(\nu/\nu_0)^{\beta}$.

**NEFD (noise equivalent flux density)**  The flux density required to produce a signal equal to the RMS noise in a detector measured in 1s.

**NEP (noise equivalent power)**  The power absorbed by a detector required to create a signal equal to the RMS noise in a 1-Hz bandwidth.

**pointing solution**  The position and orientation of the focal plane on the sky as a function of time.

**PSF (point spread function)**  A telescope’s beam pattern.

**redshift**  Since the Universe is expanding, distant objects are moving away from us, and thus emitted light is redshifted to longer frequencies. Redshift, denoted $z$, is frequently used to describe an object’s distance from us, as the inferred distance is model-dependent, but the redshift is directly observed.

**responsivity**  The change in voltage in a bolometer due to a change in incident power.

**SCUBA (Submillimetre Common-User Bolometer Array)**  A submillimetre camera operating from the JCMT.

**SED (spectral energy distribution)**  The flux density of a source as a function of frequency or wavelength.

**ULIRG (ultraluminous infrared galaxy)**  A class of infrared-bright galaxies discovered by IRAS in the mid 1980s.

**ultra-compact (UC) H II region**  A later phase of HMPO in which a star has begun to form and exhibits radio emission.
Preface

BLAST is a complex experiment and has required the expertise of a large team of scientists. Approximately 15–20 people, professors and students, from 9 institutions have spent considerable amounts of time developing, building and running the telescope, as well as analysing the resulting data. This thesis describes the complete instrument design and data analysis procedures, including work not done by the author.

Throughout the thesis, work where the author was particularly involved is described in greater detail. This includes: the design and construction of the gondola frame (Sec. 2.4), active pointing system (Sec. 2.5.1) and other miscellanea (Secs. 2.7 and 2.8); the pointing solution (Sec. 4.3); the preliminary stages of analysis of the Galactic Plane field (Secs. 5.2–5.6.2); and the deconvolution and profile-fitting of a nearby spiral galaxy (Sec. 6.1 and 6.2). Additionally, the author was an integral part of the field team, helping to assemble, test and monitor the telescope in each of the three flights to date (Secs. 3.4–3.6). During this time, the author was at least peripherally involved with nearly all subsystems and tasks described in the thesis. These points are re-iterated in the introductions to Chapters 2 and 4.

It is also noted that all of the text in this thesis is written by the author. Certain sections, in particular the descriptions of the gondola frame, active pointing system and Galactic Plane analysis are very similar to the text in the related journal articles, but this is because these were also written by the author. All plots, unless marked otherwise, were created by the author. Figs. 5.3, 5.10, 5.15 and 5.16, although marked as appearing in journal articles, were also created by the author. Photos and diagrams are, for the most part, modified versions of the originals made by others.
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Chapter 1

Introduction

Our understanding of structure formation throughout the history of the Universe is becoming increasingly detailed. In the early part of the 20th century, the astronomer Georges Lemaître proposed a primitive version of the Big Bang theory, that the Universe began in a state much denser than seen today. This theory was born out of the discovery of the expanding Universe solution to Albert Einstein’s theory of general relativity by Alexander Friedmann in 1922. The theory was supported by observations of Cepheid variable stars in distant galaxies, pulsating stars with a well-defined period-luminosity relationship, by Edwin Hubble in 1929 [Hubble 1929]. Measurements of the period and apparent luminosity of these objects, along with recession velocities measured by Vesto Slipher, led Hubble to conclude that the recession velocity of distant objects is directly proportional to their distance, easily explained if the Universe is in fact expanding. Development of the Big Bang theory continued with a primitive form of Big Bang Nucleosynthesis, the prediction of the relative proportion of the light elements created in the early Universe, by Ralph Alpher and George Gamow in 1948 [Alpher et al. 1948]. Later in 1948, Alpher and Robert Herman predicted the Cosmic Microwave Background (CMB), a background radiation left over from the early hot phase of the Universe [Alpher and Herman 1948]. Robert Dicke independently predicted the existence of the CMB in the early 1960s [Dicke et al. 1965]. Dicke’s colleagues David Wilkinson and Todd Roll built a microwave receiver in order to detect the CMB, which was in fact discovered serendipitously by Arno Penzias and Robert Wilson in 1965 [Penzias and Wilson 1965]. The identification of this “excess antenna temperature” as a cosmic signal proved to be very strong support for the Big Bang theory. Since then, more detailed cosmological observations have shown that the CMB spectrum is very well-fit by the Planck function [COBRA, Gush et al. 1990] and that the temperature of the CMB across the sky is remarkably uniform [COBE DMR, Smoot et al. 1992]. These and other observations have continued to support and refine the Big Bang theory, from which the ΛCDM (“cold
dark matter” with cosmological constant \( \Lambda \)) concordance model (e.g. Tegmark et al. 2001) has been developed. Modern observations of the CMB with the Wilkinson Microwave Anisotropy Probe (WMAP) (Spergel et al. 2007), along with a wide range of other astronomical data, including galaxy clustering (e.g. Tegmark et al. 2004; Cole et al. 2005), Type 1a supernovae (e.g. Riess et al. 2004) and weak gravitational lensing (e.g. Hoekstra et al. 2006), have constrained the parameters of \( \Lambda \)CDM to great precision, including the surprising fact that the baryon contribution to the mass in the Universe is only 4%, with roughly 22% being non-baryonic dark matter and the remaining 74% made up of some sort of dark energy with negative pressure.

The \( \Lambda \)CDM model accurately describes the structure of the Universe on large scales. It also states that the Universe began approximately 13.7 Gyr ago in a very hot and dense phase. There may well have been an early inflationary epoch, when the Universe underwent exponential expansion; inflation is needed to explain the coherence of structure on scales larger than the light travel time in the age of the Universe. Approximately 380,000 years after the Big Bang, the Universe had cooled so much that protons and electrons recombined to form neutral hydrogen and helium, at which time photons decoupled from matter and travelled freely towards us to form the CMB we see today. These dark ages proceeded until roughly 100 Myr later, when stars first started to form by gravitational collapse, in a period known as reionization. Analysis of the WMAP data implies that reionization occurred at \( t \approx 400 \) Myr (\( z \approx 10 \)), but little is known about this era and the subsequent era of galaxy formation, as the relevant non-linear physical processes are difficult to model. The history of galaxy formation has been studied extensively with the use of N-body simulations (e.g. Navarro et al. 1997; De Lucia et al. 2006) based on \( \Lambda \)CDM. The initial conditions for the simulations are set by \( \Lambda \)CDM, and parameters governing structure formation are tuned in order to reproduce the large-scale structure seen by astrophysical and cosmological observations, making use of the entire electromagnetic spectrum, from radio up to X- and \( \gamma \)-rays. Progress in understanding the details of galaxy formation and evolution is driven by the availability of deep extragalactic data over a wide range of frequencies.

Extragalactic astronomy really only began with observations in the optical spectrum in the early 20th century, when it was realized that many objects in the sky are exterior to the Milky Way. Radio astronomy progressed through the first half of the century and flourished with technological developments made during World War II and with the invention of radio interferometry in the late 1940s, providing sensitive measurements and improvements in resolution
from degrees to arcminutes. The technology for detecting submillimetre radiation is relatively new; the first useful arrays of submillimetre detectors were developed in the 1990s, with the Submillimetre Common-User Bolometer Array (SCUBA) \cite{Holland1999} on the 15-m James Clerk Maxwell Telescope (JCMT) on Mauna Kea in Hawaii, and the Max-Planck Millimetre Bolometer Array (MAMBO) \cite{Kreysa1998} on the 30-m Institut de Radio Astronomie Millimétrique (IRAM) telescope on Pico Veleta, Spain. Observations in the submillimetre have proven to be very useful for the study of structure formation in the Universe.

1.1 Cosmic Infrared Background

Infrared astronomy, defined roughly as the study of radiation from $\lambda \sim 1$–200 $\mu$m, began in earnest with the Infrared Astronomical Satellite (IRAS), launched in 1983. IRAS mapped the full sky at 12, 25, 60 and 100 $\mu$m, eventually producing a catalogue of more than 15,000 extragalactic objects with known redshifts \cite{Saunders2000}. The vast majority of these galaxies are late-type spirals, but a new population of ultra-luminous infrared galaxies (ULIRGs) was found \cite{Soifer1987, Sanders1996}. These galaxies are thought to play an important part in star formation in the higher redshift Universe. It is believed that they harbour large amounts of star formation producing strong UV and optical light which is absorbed by interstellar dust, which then thermally re-radiates in the mid- to far-infrared (FIR).

The Cosmic Background Explorer (COBE) satellite, launched in 1989, featured 3 instruments. Two of these instruments, the Differential Microwave Radiometer (DMR) and the Far-InfraRed Absolute Spectrophotometer (FIRAS) were developed to study the CMB. The third instrument, the Diffuse Infrared Background Experiment (DIRBE), was designed to detect the hypothesized cosmic infrared background (CIB), a diffuse background due to large amounts of unresolved ULIRGs in the early Universe. The CIB was tentatively detected by FIRAS \cite{Puget1996}; it was later firmly detected by DIRBE at 140 and 240 $\mu$m \cite{Hauser1998} and at 60 and 100 $\mu$m \cite{Finkbeiner2000}. The intensity (power per collection area per solid angle, W m$^{-2}$ sr$^{-1}$) found in the CIB exceeds the energy found by integrating the optical light from galaxies in the Hubble Deep Field \cite{Pozzetti1998} by a factor of $\sim 2.5$, and the integrated IRAS counts by a factor of 5–10, indicating that the FIR indeed plays an important part in the history of star formation, and that ULIRGs were much more important
in the past.

### 1.2 Star Formation History

Cosmological observations tell us that baryons form only a small fraction of the total mass in the Universe. Baryons are attracted by the gravitational potentials of large-scale over-densities in dark matter, however, and thus the stellar light tends to trace large-scale structure. The bulk of stellar mass in the present-day Universe is found in elliptical galaxies (Fukugita and Peebles 2004), massive objects typically found in galaxy clusters and mostly devoid of interstellar dust.

It is found that the stars in elliptical galaxies are very old, with little to no star formation since a redshift of \( z \sim 1 \) (Oke 1971; Stanford et al. 1998). Observations of high-redshift galaxies in the UV through infrared by the Canada-France Redshift Survey (CFRS) showed that star formation space density increases strongly out to \( z \sim 1 \) (Lilly et al. 1996). At the same time, a large population of galaxies at \( z > 3 \) was found using the Lyman break technique, whereby candidate high-redshift galaxies are identified in 3-colour UV/optical observations of blank fields by the drop-out of emission in the short-wavelength band due to Lyman limit absorption by the intervening intergalactic medium (Steidel et al. 1996). The spectra of these galaxies are similar to local-Universe star forming galaxies, indicating that these objects might be the progenitors of modern-day ellipticals, or perhaps the spheroids of more common galaxies. It is noted, however, that there does not appear to be enough star formation activity, as traced by optically-detected galaxies, to account for the number of stars seen in local ellipticals (Madau et al. 1996).

The solution to the problem of where these stars formed seems to lie in the ULIRGs discovered by IRAS. If ULIRGs existed in much higher numbers in the early Universe, they could be the progenitors of elliptical galaxies. A single-temperature modified blackbody spectrum is generally used to describe the spectral energy distribution (SED) of dust emission from ULIRGs,

\[
S_\nu \propto \left( \frac{\nu}{\nu_0} \right)^\beta B_\nu(T),
\]

with \( B_\nu(T) \) the Planck function, \( T \) the dust temperature and \( \beta \) the emissivity index. The exact shape of the emissivity function is determined by dust composition and geometry, neither of which are well known. Additionally, since star-forming sources are typically not well resolved in the submillimetre, the emissivity term is a volume average over the emitting region. It is found observationally that the emissivity term is well described by a power law (e.g. Schwartz...
as shown above. ULIRGs observed in the local Universe are well-fit by this model, with
dust temperatures of $T \sim 40$ K and emissivity indices $\beta \sim 1.5$, resulting in peak emission at
$\lambda \sim 100$ $\mu$m (e.g. Eales and Edmunds 1996; Dunne et al. 2000). For similar objects at redshift
$z = 1–5$, we therefore expect peak emission at $\lambda \sim 200–600$ $\mu$m, and hence observations in the
submillimetre should be sensitive to the history of dust-enshrouded star formation out to very
large redshifts.

1.3 Extragalactic Submillimetre Astronomy

Extragalactic submillimetre surveys, starting with SCUBA and MAMBO and continuing with
Bolocam (Glenn et al. 1998) and SHARC II (Dowell et al. 2003) on the Caltech Submillimeter
Observatory (CSO) on Mauna Kea in Hawaii, have now detected hundreds of high-redshift
FIR-bright galaxies, the majority of which are presumed to be strongly star-forming. The
most ambitious submillimetre project to date, the SCUBA Half-Degree Extragalactic Survey
(SHADES; Coppin et al. 2006), finds 120 point sources brighter than 2 mJy in a survey area
of 720 arcmin$^2$, for a total of $10^4$ mJy deg$^{-2}$, resolving 20–30% of the submillimetre part of the
CIB. In the following, we describe what can be inferred about the star-formation history of
the Universe given a large sample of submillimetre-detected objects, after first discussing the
negative K-correction.

1.3.1 Negative K-correction

An interesting feature of extragalactic observations in the submillimetre is what is known
as the negative K-correction (e.g. Peebles 1993). The term comes from optical astronomy
(Humason et al. 1956) and refers to the fact that when a very distant object is observed, a
different part of its spectrum is sampled than if it was nearby, due to the Doppler shift that
results from the expansion of the Universe. The Doppler shift is known as the redshift of the
object, as the shift is towards longer, or redder, wavelengths. Observing an object at high
redshift with a bandpass filter with a given central wavelength $\lambda_0$ samples a shorter rest-frame
wavelength $\lambda' = \lambda_0/(1 + z)$. Galaxies tend to be fainter at shorter wavelengths in the optical,
so, for two identical objects, one nearby and one distant, the distant one will be fainter due to
the normal $1/R^2$ fall-off, but will be additionally fainter due to the fact that an intrinsically
Figure 1.1: Flux density for a modified blackbody with $T = 40$ K and $\beta = 1.5$ as a function of wavelength at a variety of redshifts (left panel) and as a function of redshift for a variety of wavelengths (right panel). Note that at longer wavelengths, flux density increases with redshift.

fainter part of its spectrum is sampled. In the submillimetre, however, this effect is reversed. A ULIRG typically has dust temperatures of $T \sim 40$ K, and the submillimetre passbands ($\sim 250$–1000 $\mu$m) fall on the Rayleigh-Jeans portion of the spectrum, for which the brightness increases for decreasing wavelength. To demonstrate this effect, we plot a range of flux density curves for objects at a range of redshifts (Fig. 1.1). We choose an SED, the specific luminosity as a function of wavelength, described by a modified blackbody,

$$L_\nu \propto \nu^\beta B_\nu(T),$$  \hspace{1cm} (1.2)

where $B_\nu(T)$ is the Planck function for a blackbody with temperature $T$ and $\beta$ is the emissivity index. The measured flux density $S_\nu$ for an object at redshift $z$ is then

$$\nu \, dv \propto \frac{L_\nu \, dv'}{4\pi D_L^2}, \hspace{1cm} (1.3)$$

where $\nu' = (1 + z) \nu$ is the emitted frequency corresponding to the observed frequency $\nu$ and $D_L$ is the model-dependent luminosity distance corresponding to the redshift $z$. This can be
written in terms of the specific luminosity at the rest frame frequency $\nu$, 

$$S_\nu \, d\nu \propto \left[ (1 + z) \frac{L_{\nu'}}{L_\nu} \right] \frac{L_\nu \, d\nu}{4\pi D_L^2}, \quad (1.4)$$

where the term in square brackets is the “K-correction” and the factor of $1 + z$ is due to the bandwidth $\Delta\nu' = (1 + z) \Delta\nu$. The curves in Fig. 1.1 are for an SED with $\beta = 1.5$ and $T = 40$ K. For observed wavelengths greater than $\sim 600 \mu$m, the K-correction factor exceeds the distance factor for $z \gtrsim 1$. This presents a great advantage for observations of high-redshift star-forming systems at submillimetre wavelengths. The spectrum of star-forming objects decreases with increasing wavelength in the UV, optical, near-IR and radio, thus the submillimetre is unique in its capability to select objects at high redshifts.

### 1.3.2 Number Counts

The “differential number counts” $dN/dS$ is defined as the number of objects with flux density between $S$ and $S + dS$. It depends on: (i) the evolving luminosity function $\phi(L, z)$, the number of objects with luminosity between $L$ and $L + dL$ per unit volume at a redshift $z$; (ii) luminosity as a function of flux $S$ and redshift $z$, $L(S, z)$; and (iii) the luminosity distance to the object $D_L(z)$, which depends on cosmology,

$$\frac{dN}{dS} = f \left( \phi(L, z), L(S, z), D_L(z) \right). \quad (1.5)$$

With a model of the evolving luminosity function $\phi(L, z)$ and a particular cosmological model, the number counts $dN/dS$ can be derived analytically. A measure of $dN/dS$ therefore allows one to constrain the evolution of the luminosity function.

Differential number counts have been fit with the empirically-motivated broken power law,

$$\frac{dN}{dS} = \begin{cases} \frac{N'}{S'} \left( \frac{S}{S'} \right)^{-\beta}, & S > S' \\ \frac{N'}{S'} \left( \frac{S}{S'} \right)^{-\alpha}, & S < S' \end{cases}, \quad (1.6)$$

where $N'$, $S'$, $\alpha$ and $\beta$ are free parameters constrained by the data (e.g. Borys et al. 2003). The Schechter function (Schechter 1976),

$$\frac{dN}{dS} = \frac{N'}{S'} \left( \frac{S}{S'} \right)^{\alpha} \exp(-S/S'), \quad (1.7)$$

with $N'$, $S'$, $\alpha$ free parameters, also fits the data well (see Coppin et al. 2006). The number counts found in submillimetre surveys far exceed that expected if the locally-observed $\phi(L)$
holds at all redshifts (Saunders et al. 1990), indicating strong evolution with densities of the brightest ULIRGs \( \sim 400 \) times greater at \( z \sim 2 \) than at \( z = 0 \) (Blain et al. 2002). The differential number counts function can be integrated to provide an estimate of the total flux density of the resolved sources, to compare data from different surveys, and also provides an observable that can be used to constrain numerical simulations of models.

### 1.3.3 Luminosity Evolution

Given a model of a source’s SED, flux density measurements can be converted to an integrated far-infrared flux. If the distance to the submillimetre-detected sources can be determined, one can then calculate the FIR luminosity. From this, one can estimate star formation rates and constrain the evolution. Current results show that star formation peaked around \( z = 2–3 \) (e.g. Wall et al. 2007), but the specifics are still unclear, given continuing difficulties in finding redshifts of submillimetre galaxies.

### 1.3.4 Clustering

The clustering properties of high-redshift submillimetre galaxies provides an additional observable with which one can further constrain models. Ellipticals are found in the densest regions in the local Universe and are much more strongly clustered than spiral galaxies (e.g. Guzzo et al. 1997). Strong clustering has been measured for Lyman break galaxies at \( z \sim 3 \) (Giavalisco and Dickinson 2001), and the class of galaxy known as extremely red objects (EROs) is also strongly clustered (Daddi et al. 2000). If submillimetre-detected galaxies are indeed the progenitors of modern-day ellipticals, the clustering of both populations should be similar. One can measure the 2-dimensional (without distances) and 3-dimensional (with distances) clustering signals, given sufficiently large catalogues. The 2-dimensional clustering signal is parametrized by the angular correlation function \( w(\theta) \), the excess probability of finding two sources within a radius \( \theta \) compared to that found for a uniformly random field,

\[
dP = N^2 [1 + w(\theta)] d\Omega_1 d\Omega_2,
\]  

where \( N \) is the source surface density and \( d\Omega_1 \) and \( d\Omega_2 \) are elements of solid angle. Attempts to measure the angular correlation function in SCUBA data have been attempted, but so far only
weak detections have been reported (e.g. Blain et al. 2004; Scott et al. 2006), as the number of known sources is low and the observed areas are small.

1.3.5 The Confusion Limit

Source confusion is the contribution to noise due to multiple faint sources falling within a single telescope beam (Condon 1974). Confusion provides the limiting noise level for extracting individual point sources, since observing for longer times no longer increases their detectability. The confusion limit is typically defined as the flux density at which the surface density of sources exceeds 0.03 beam$^{-1}$. At 850 $\mu$m (SCUBA), the confusion limit is $\sim 2$ mJy. Depending on the slope of the number counts, the confusion limit can be a strong function of the telescope beam size (Takeuchi and Ishii 2004).

1.4 Galactic Submillimetre Emission

Observations of star-forming regions in our own Galaxy allow for the study of the evolutionary stages of star formation in detail. There is currently no predictive theory of star formation, but observations are beginning to build the framework for the development of such a theory (e.g. Evans 1999). Current evidence shows that massive stars are born deep within dense clouds of gas and dust. These early-stage objects are known as “high-mass protostellar objects” (HMPOs) and are characterized by cold dust emission, high luminosities and no associated radio emission (e.g. Zhang et al. 2007). Eventually, the massive stars formed inside HMPOs ionize the surrounding gas and can be seen in H$\text{II}$ radio emission. These objects are then known as ultra compact (UC) H$\text{II}$ regions. Submillimetre measurements are sensitive to both of these types of object, but it is the former that are more interesting, since mid-IR instruments such as IRAS and Spitzer MIPS are not sensitive to the cold dust emission. A submillimetre telescope with the capability to make large, sensitive maps has the potential to discover thousands of these objects and thus constrain models for the early stages of star formation. By measuring the relative fraction of pre-stellar cores to stars in the later stages of evolution, the time scales for accretion and collapse can be determined.

Measurements of Galactic star forming regions by SCUBA at 850 $\mu$m (e.g. Johnstone et al. 2000) have already revealed large numbers of such sources, but with only one band are unable to
constrain the temperature of emission and thus cannot classify a particular source in terms of its stage of evolution. An instrument with detectors at several submillimetre wavelengths spanning the peak of thermal emission can strongly constrain the dust temperature and thus allow for the identification of the coldest sources. An accurate constraint on the dust temperature is required to accurately infer clump mass, and thus an instrument sensitive to the peak of the thermal dust emission provides a direct measure of the dust mass.

Several models of the formation of stars from molecular clouds have been proposed. One theory suggests that molecular cores fragment during collapse (Klessen et al. 1998; Myers 1998; Padoan and Nordlund 2002). In such a scenario, the mass function of the molecular clouds, the number of objects found with a given mass, should be similar to the stellar initial mass function (IMF), the mass function of stars at the time of formation (Salpeter 1955). Another suggested process is that the masses of stars are determined by powerful stellar outflows (Adams and Fatuzzo 1996). Adams and Fatuzzo (1996) predict a transfer function between the molecular cloud mass function and the stellar IMF based on this model. Bonnell et al. (1997) propose a “competitive” accretion model whereby the mass of a star is strongly dependent on its ability to accrete. Accurate measurements of the molecular cloud mass function, together with knowledge of the stellar IMF (e.g. Massey 2003), can help constrain models of star formation. Observations of a Galactic star formation region finding large numbers of molecular clouds with accurate determination of masses can help to constrain the transfer function from molecular clouds to stars.

In addition to finding new objects, submillimetre measurements can also help constrain the physical parameters of warmer objects detectable in the IR. Measurements of the Rayleigh-Jeans portion of the thermal emission allows for determination of the emissivity index $\beta$, and thus more precisely constrains physical parameters of the dust clouds, such as mass and luminosity.

Submillimetre instrumentation is also sensitive to dust present in clouds of H I known as cirrus clouds (Gibson et al. 2000), a relatively recently discovered phase of the interstellar medium. Observations of cirrus can help clarify the evolutionary importance of this phase.

Measurements of star formation in nearby galaxies are also useful. Nearby resolved galaxies allow for probing the details of star formation; substructure within the galaxy can be resolved and the details of where star formation occurs in relation to the distribution of inter-stellar dust can be studied. This provides details at an intermediate scale between the small-scale...
measurements of star formation in the Galaxy and that occurring in massive, unresolved galaxies in the local Universe and at high redshift.

### 1.5 A Simple Model

Given the current knowledge of submillimetre-detected high-redshift galaxy number counts, we pose the following question: what should a new submillimetre telescope be capable of detecting in order to significantly advance the field? There are currently \( \sim 300 \) sources detected by SCUBA at \( 850 \mu m \), using non-optimal (and different) observing strategies, and the number counts appear to disagree by factors of \( \sim 3 \). A new experiment could make a strong impact by detecting a uniform sample of \( \sim 300 \) sources. The new sample would allow us to test for systematic errors in the old sample, and by combining the data, would significantly reduce poisson errors. In order to estimate redshifts (see Sec. 1.6.2), one would like to sample near the peak of star-formation emission (in intensity units, \( \nu S_\nu \)), at \( \lambda = \sim 200 \mu m \) for \( z = 1-3 \). These wavelengths are largely inaccessible from the ground due to absorption of water vapour in the atmosphere (Fig. 1.2); thus we are driven to observe from space or an environment with space-like conditions.

We present here a simple source model in order to estimate whether a telescope can be built that can detect 300 sources at \( 250 \mu m \) in a reasonable amount of time. We make the naive assumption that the \( 250 \mu m \) number counts are made up of only the \( 4 \) mJy sources found at \( 850 \mu m \) and that they are all described by modified blackbodies with \( T = 40 \) K and \( \beta = 1.5 \). We assume 300 sources per degree, as measured by Coppin et al. (2006). In order to not be confusion limited, we require a beam size such that we have 30 beams per source. With 300 sources per \( \text{deg}^2 \), the beam size must be less than \( \sim 40'' \), and thus the diameter of the primary mirror must be \( D = 1.22 \lambda / \theta \sim 2 \) m. We now calculate how long it would take to detect the \( S_{850} = 4 \) mJy sources with a 2-m primary. Assuming the sources all lie at \( z = 1 \), the flux density at \( 250 \mu m \) is \( S_{250} \approx 160 \) mJy. We use the point source sensitivity for the telescope derived later in Sec. 1.6.1, finding that we map 1\( \text{deg}^2 \) to a 5\( \sigma \) limit of 32 mJy in 1 hr. If we additionally require that the sources are detected to 5\( \sigma \) at \( 500 \mu m \), we find observing times of \( \sim 75 \) hr. We emphasize that this is a very naive extrapolation of existing data, but that it does suggest that building a submillimetre experiment with a 2 m primary mirror and existing bolometer
Figure 1.2: Atmospheric transmission plots at far-infrared and submillimetre wavelengths at a variety of altitudes. The lower panel is the transmission at an altitude of $\sim 4$ km, at the location of the JCMT and the CSO, the summit of Mauna Kea in Hawaii. We note that the sky is almost completely opaque at 250 $\mu$m. The upper panels show transmission at a variety of altitudes, including 40 km, the nominal long-duration balloon altitude. Due to the low atmospheric column density ($\sim 0.5\%$ of that from the ground), the atmosphere is mostly transparent.
technology has significant potential for deep cosmological surveys.

1.6 BLAST

Current submillimetre instrumentation is hampered by slow mapping speeds, due primarily to low atmospheric transmission and the small number of detector elements in a given instrument. Future instrumentation will have vastly improved mapping speeds due to much larger numbers of detector elements (e.g. SCUBA-2, Holland et al. 2006) and increased transmission (e.g. the space-based SPIRE, Griffin et al. 2006, one of the instruments aboard the Herschel satellite).

The Balloon-borne Large-Aperture Submillimeter Telescope (BLAST) takes advantage of new large-format detector arrays developed for SPIRE, as well as a 2-m carbon-fibre primary mirror built as a test for the 3-m Herschel primary. The high-altitude balloon platform provides a method for observing with the new detector technology at near-space observational conditions on a much shorter time scale than for the satellite platform. BLAST will thus produce the first sensitive large-area surveys in the submillimetre, yielding catalogues of thousands of submillimetre-detected sources. This will represent a substantial advance in the field, as current dedicated ground-based surveys, most notably SHADES, have detected a total of \( \sim 300 \) sources in the 10-year history of extragalactic blank-field surveys at submillimetre wavelengths.

1.6.1 Sensitivities and Predicted Number Counts

The BLAST instrument is described in detail in Chapter 2, but we present detector and telescope parameters here in order to estimate the instrument’s sensitivity. The BLAST focal plane consists of 3 arrays of detectors with contiguous 30% bandpasses centred at 250, 350 and 500 \( \mu \text{m} \). The number of detector elements and the diffraction-limited “full-width at half-maximum” (FWHM) beam sizes are listed in Table 1.1. Detector noise is usually quoted as the “noise equivalent power” (NEP), the power absorbed by the detector required to create a signal equal to the root mean square (RMS) noise in a 1-Hz bandwidth (e.g. Griffin et al. 2002). NEP scales with the square root of detector bandwidth, assuming independent Poisson noise, so it is quoted in W Hz\(^{-1/2}\). Based on estimates of optical efficiencies, atmospheric transparency and detector characteristics, we estimate background loading \( Q \) and the resulting NEP. Given the NEP, and the relevant telescope parameters, we can calculate the resulting “noise equivalent
Table 1.1: Instrumental Sensitivities

<table>
<thead>
<tr>
<th>$\lambda_c$ (µm)</th>
<th>$N_{\text{det}}$</th>
<th>FWHM (&quot;)</th>
<th>$\Delta \lambda/\lambda_c$</th>
<th>$\eta_{\text{opt}}$</th>
<th>$\eta_{\text{det}}$</th>
<th>$\eta_{\text{tel}}$</th>
<th>$\eta_{\text{sky}}$</th>
<th>$Q$ (pW)</th>
<th>NEP (W Hz$^{-1/2}$)</th>
<th>NEFD (mJy s$^{1/2}$)</th>
</tr>
</thead>
<tbody>
<tr>
<td>250</td>
<td>149</td>
<td>30</td>
<td>0.3</td>
<td>0.4</td>
<td>0.8</td>
<td>0.15</td>
<td>0.051</td>
<td>149</td>
<td>$60.4 \times 10^{-17}$</td>
<td>181</td>
</tr>
<tr>
<td>350</td>
<td>88</td>
<td>41</td>
<td>0.3</td>
<td>0.4</td>
<td>0.8</td>
<td>0.15</td>
<td>0.034</td>
<td>102</td>
<td>$44.2 \times 10^{-17}$</td>
<td>182</td>
</tr>
<tr>
<td>500</td>
<td>43</td>
<td>59</td>
<td>0.3</td>
<td>0.4</td>
<td>0.8</td>
<td>0.15</td>
<td>0.017</td>
<td>67</td>
<td>$31.9 \times 10^{-17}$</td>
<td>184</td>
</tr>
</tbody>
</table>

* Instrument parameters leading to sensitivity calculation (Eqn. 1.9). See Sec. 1.6.1.

flux density” (NEFD), the incident flux density required to produce a signal equal to the RMS noise in the detector measured over 1 s, and is given in Jy s$^{1/2}$. NEFD is related to NEP by

$$\text{NEFD} = \frac{\text{NEP}}{\sqrt{2A_{\text{tel}} \eta_{\text{opt}} \eta_{\text{det}} (1 - \eta_{\text{tel}})(1 - \eta_{\text{sky}})\Delta \nu}},$$

where $A_{\text{tel}}$ is the effective area of the telescope, $\eta_{\text{opt}}$ is the optical efficiency of the full optical system, $\eta_{\text{det}}$ is the quantum efficiency of the detector, $\eta_{\text{tel}}$ and $\eta_{\text{sky}}$ are the emissivities of the telescope and sky, respectively, and $\Delta \nu$ is the optical bandwidth. The factor of $\sqrt{2}$ comes from the fact that NEP is given per frequency bandwidth and NEFD is given per integration time; by the Nyquist theorem, $\Delta f = (2\Delta t)^{-1}$. $A_{\text{tel}} = \frac{1}{4}\pi D_{\text{eff}}^2$ is the area of the telescope that is used for light collection, correcting for the blockage from the secondary and struts. For BLAST, $D_{\text{eff}} \approx 1.8$ m. The resulting NEPs and NEFDs are listed in Table 1.1. Note that, despite significantly different background loading at each wavelength, the larger (frequency) bandwidths at shorter wavelengths conspire to produce very similar NEFDs.

Given the estimated NEFDs, we can calculate the point source sensitivity of a given survey strategy. If we observe a region of sky with area $\Omega_m$ for a time $\tau$, the resulting map, subdivided into pixels of size $\theta_p$, will have noise per pixel $\xi_p$ of

$$\xi_p = \text{NEFD} \sqrt{\frac{\Omega_m}{\theta_p^2 \tau N_{\text{det}}}},$$

where $N_{\text{det}}$ is the number of detectors at the given wavelength. If a point source of flux density $S$ is observed with a telescope beam of Gaussian width $\theta_b = \text{FWHM}/2.35$, the error on the measurement $\hat{S}$ is

$$\xi_S = \frac{1}{\sqrt{\pi(\theta_b/\theta_p)}}\xi_p = \frac{\text{NEFD}}{\sqrt{\pi\theta_b}} \sqrt{\frac{\Omega_m}{\tau N_{\text{det}}}}.$$
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Table 1.2: Point Source Sensitivity (per $\sqrt{\text{deg}^2 \cdot \text{hr}}$)

<table>
<thead>
<tr>
<th>$\lambda_e$ (µm)</th>
<th>FWHM</th>
<th>$\theta_b$ (&quot;)</th>
<th>$N_{\text{det}}$</th>
<th>NEFD (mJy s$^{1/2}$)</th>
<th>$\xi_s$ (mJy)</th>
</tr>
</thead>
<tbody>
<tr>
<td>250</td>
<td>30</td>
<td>12.8</td>
<td>149</td>
<td>181</td>
<td>39.2</td>
</tr>
<tr>
<td>350</td>
<td>41</td>
<td>17.4</td>
<td>88</td>
<td>182</td>
<td>37.7</td>
</tr>
<tr>
<td>500</td>
<td>59</td>
<td>25.1</td>
<td>43</td>
<td>184</td>
<td>37.8</td>
</tr>
</tbody>
</table>

Table 1.3: Predicted Source Counts in a Variety of 50-hr Surveys.

<table>
<thead>
<tr>
<th>Area (deg$^2$)</th>
<th>1σ Depth (mJy beam$^{-1}$)</th>
<th>$N_s$ (250 µm)</th>
<th>$N_s$ (350 µm)</th>
<th>$N_s$ (500 µm)</th>
<th>$N_s &gt; 5\sigma$ (z &gt; 3)</th>
<th>$N_s &gt; 5\sigma$ (z &gt; 3)</th>
<th>$N_s &gt; 5\sigma$ (z &gt; 3)</th>
</tr>
</thead>
<tbody>
<tr>
<td>1.0</td>
<td>5</td>
<td>1601</td>
<td>429</td>
<td>1276</td>
<td>487</td>
<td>584</td>
<td>279</td>
</tr>
<tr>
<td>2.0</td>
<td>7</td>
<td>1916</td>
<td>518</td>
<td>1513</td>
<td>600</td>
<td>752</td>
<td>341</td>
</tr>
<tr>
<td>4.0</td>
<td>10</td>
<td>1985</td>
<td>542</td>
<td>1554</td>
<td>620</td>
<td>623</td>
<td>319</td>
</tr>
<tr>
<td>9.0</td>
<td>15</td>
<td>2142</td>
<td>597</td>
<td>1664</td>
<td>689</td>
<td>619</td>
<td>321</td>
</tr>
<tr>
<td>36.0</td>
<td>30</td>
<td>2069</td>
<td>594</td>
<td>1592</td>
<td>687</td>
<td>521</td>
<td>284</td>
</tr>
</tbody>
</table>

* This table is taken from Chapin (2004).

A formal derivation is given in Sec. 5.4.2. The point source sensitivity for a 1-hr observation of a 1 deg$^2$ field of view are listed in Table 1.2. Note that the beam size and number of detectors roughly cancel out to give very similar sensitivities. This is not a coincidence, as the detector separation on the array is proportional to the observing wavelength, and so for the same physical area in the focal plane $N_{\text{det}} \theta_b^2 \approx \text{constant}$.

Finally, given a model of the luminosity function and evolution of high-redshift sources at the BLAST wavelengths, we can estimate the number of sources BLAST will detect for a given survey strategy. Chapin (2004) has developed such a model, using a luminosity function based on $\text{IRAS}$ counts, a single SED template and luminosity evolution from $0 < z < 2.2$ (Chapin 2004). The total source counts and number of objects at $z > 3$ detected in a 50-hr survey over a variety of survey areas are listed in Table 1.3.

$^1$The result presented here is different from the one in Sec. 5.4.2 by a factor of $\pi \theta_b^2$, the area of the beam. This is due to a difference in the definition of the calibration of the maps considered.
1.6.2 Photometric Redshifts

If the SED of a high-redshift galaxy is known precisely, then 3-colour submillimetre measurements on or near the emission peak can constrain the redshift of the galaxy \cite{Hughes2002}. The SEDs of the submillimetre-detected galaxies are of course not precisely known, but an estimate of the redshift can still be made given reasonable ranges for $T$ and $\beta$. Based on SED templates of nearby ULIRGs, we estimate that we can measure the redshifts of BLAST-detected galaxies photometrically from $z \sim 1$–3 to an accuracy of $\Delta z \sim 0.5$ \cite{Hughes2002}. Complementary measurements at other wavelengths, in particular longer wavelengths available from SCUBA, can increase the accuracy and extend the reach to $z \sim 5$.

1.6.3 Science Goals

We list the extragalactic and Galactic science cases for flying BLAST. All of these goals should be achievable in a single 10-day long duration balloon flight.

1.6.3.1 Extragalactic

- identify $\sim 10^4$ individual high-redshift galaxies, increasing the number of known sources by nearly 2 orders of magnitude, and thus produce statistically significant number counts
- measure photometric redshifts of the sources detected in all 3 bands, and determine bolometric rest-frame FIR luminosities, star formation rates and the evolution of star formation for 1000s of dusty star forming galaxies
- measure the large-scale clustering of star-bursting galaxies up to scales of 600 Mpc
- measure the confusion limit at the BLAST wavelengths
- test the detectors and filters and influence survey design for SPIRE

1.6.3.2 Galactic

- identify potential pre-stellar cores in star-forming regions and estimate their lifetimes by determining the relative number fraction of cores to later stages of evolution
- estimate column densities, masses and luminosities of the surrounding regions by disentangling the degeneracy between $T$ and $\beta$
Chapter 1. Introduction

- search for high-mass protostellar objects in the vicinity of ultracompact H II regions
- measure Galactic cirrus, molecular clouds and cold dust not associated with star formation

1.7 Instrument Performance

During the 2005 flight, the optical system suffered a severe anomaly, possibly due to delamination of the primary mirror by the freezing of water collected during a rain storm before launch. The point spread function was significantly non-Gaussian, showing asymmetric structure on scales of 3'. The resulting resolution and point source sensitivity were decreased from nominal by a factor of \( \sim 7 \). It was decided that the extragalactic science case could not be fulfilled and instead the observing time was concentrated on Galactic fields. Several steps were taken to avoid this failure in the 2006 flight, however, and we anticipate highly sensitive maps of both Galactic and extragalactic fields.

1.8 Outline

This thesis is organized as follows. Chapter 2 describes the BLAST instrument in detail. The 1-day test flight from North America in 2003, the first long-duration science flight from Sweden in 2005 and a 11-day run from Antarctica at the end of 2006 are described in Chapter 3. Instrument performance and observations made are discussed. The steps to create maps from the raw data are detailed in Chapter 4 and analysis of two of the Galactic fields from the 2005 flight are described in Chapters 5 and 6.
Chapter 2

The BLAST Instrument

In order to produce the most sensitive submillimetre measurements possible, BLAST uses the best current detector technology and pushes the scientific limits of what is possible on a balloon platform. We make use of detectors developed for SPIRE, one of three instruments on board the European Space Agency’s Herschel satellite, scheduled for launch in 2008.

The balloon platform offers many advantages over both ground- and space-based instruments. In particular, atmospheric water vapour is opaque in the short wavelength submillimetre at all but a few narrow-band regions. Observations from a balloon platform at an altitude of 40 km, above 99% of the atmosphere, are much more sensitive, allowing for quicker mapping speeds and access to the shorter wavelengths ($\sim 250 \mu m$) not visible from the ground. Observing from space would clearly be favourable, giving 100% transmission and much longer integration times, but quick development time and low cost (compared to a space-based experiment) are clear advantages of the long-duration balloon platform. Disadvantages compared to ground-based telescopes include small telescope aperture (BLAST has a 2-metre aperture, compared to the 15-metre JCMT and the 30-metre IRAM) and an unstable pointing platform.

The design of the instrument is driven by the primary science goals, to detect large numbers of high-redshift star-forming galaxies and to constrain their redshifts by sampling the peak of the FIR dust emission profile. At redshifts of $z \sim 1-3$, the FIR emission from $T \sim 40$ K dust peaks at observed wavelengths of 200–400 $\mu m$. These wavelengths are mostly inaccessible from the ground, even at the best observing sites, such as at Mauna Kea in Hawaii, at an elevation of 4 km, due to absorption by water in the atmosphere (see lower panel of Fig. 1.2). It is thus necessary to observe from space or environments with near-space conditions. As can be seen from Fig. 1.2, the atmospheric transmission at long-duration balloon altitudes of $\sim 40$ km is nearly 100%, with only very narrow absorption lines across the full submillimetre spectrum. Atmospheric emission is also significant at lower altitudes, even in the best weather at Mauna
Kea, limiting instrumental sensitivity, further driving the need to observe from near-space atmospheric conditions. In order to properly sample the dust peak, we use detectors centred at 250, 350 and 500 µm with nearly contiguous 30% bandpasses (see Fig. 2.7). We have shown in Chapter 1 that a 2 m primary mirror provides sufficient resolution and sensitivity (Sec. 1.5).

In order to map regions of 1 deg² and larger, the detector arrays must be scanned across the sky. The detectors are characterized by 1/f noise which dominates at frequencies less than \( f_0 \sim 0.03 \), and thus drifts in signal baseline will dominate on time scales longer than \( \tau \sim 1/f_0 \sim 30 \text{ s} \). In order to constrain these drifts, a detector should return to the same point on the sky on a timescale shorter than \( \tau \). Mapping speeds are discussed in more detail in Sec. 3.1.

To produce accurate maps after having scanned a region of sky, the pointing of the telescope must be well known. To avoid significant smearing of the telescope beam in the final maps, we wish to know the pointing to better than \( \sim 1/5^{\text{th}} \) of a beam. The beamsize at 250 µm is \( \sim 30'' \); our target accuracy for post-flight pointing reconstruction is thus \( \sim 5'' \). In-flight pointing reconstruction is not as critical, however, and we are satisfied with in-flight pointing errors of \( \sim 30'' \). To achieve this, the feedback rate of the control system should be \( \sim 10 \text{ Hz} \), and the telescope structure must therefore be stiff at frequencies lower than 10 Hz.

The physical characteristics of the gondola frame are constrained by the balloon platform. The strongest constraint is total weight, limited by the maximum lift of the balloon to \( \sim 2000 \text{ kg} \). Weight was a consideration in the design of all components of the instrument. The gondola frame is made mostly of aluminum, with the exception of the motor housings, which are hardened steel. A cable-suspended platform was chosen, as the necessary stiffness is achieved with much less weight than with a rigid structure. The physical dimensions of the frame are limited by the size of the building in which it is built and by the geometry of the launch vehicle.

The BLAST instrument is a complicated system, requiring the expertise and time of a large team of scientists. While the entire instrument is described here, we note that the author was particularly involved with the design and construction of the gondola frame (Sec. 2.4), active pointing system (Sec. 2.5.1), power system (Sec. 2.7), and various miscellaneous subsystems such as the inner frame lock motor, cooling system and thermometry (Sec. 2.8). The bolometer bias diagnostic tool described in Sec. 3.3.1 was also written by the author. Additionally, the author was an integral part of the field team, helping to assemble, test and run the instrument in each of the 3 flights (Secs. 3.4–3.6).
The instrument is described in Pascale et al. (2007), but in this chapter we further discuss each of the components of the telescope assembly, starting with the detectors and working out to the frame.

2.1 Detectors

BLAST uses detectors developed by the Jet Propulsion Laboratory for use on SPIRE, a photometric submillimetre instrument aboard the Herschel satellite. See Bock et al. (1998) for further details.

2.1.1 The Bolometer Device

The detectors, known as bolometers, because they detect thermal radiation over a wide range of frequencies, are semiconductor devices consisting of an absorber and a thermistor (see Fig. 2.1). The absorber is made of silicon nitride and is configured in a finely-patterned spider web mesh, with $\sim 1.5\%$ filling factor, to reduce the heat capacity and cross-section to cosmic rays, while maintaining infrared absorption. The absorber is suspended from the substrate, the support structure from which it is etched, on radial legs of bare silicon nitride. The legs have low thermal conductivity and thus isolate the absorber and at the same time provide rigid mechanical support. The thermistor consists of neutron-transmutation-doped germanium which is bonded to the absorber.

The physics of the bolometer is quite simple: incoming IR photons are absorbed by the bolometer and increase its temperature. The change in temperature is measured via change in resistance in the thermistor. The response time of the bolometer is limited by the thermal conductance of the support legs and heat capacity of the bolometer, thus there is a trade-off between sensitivity and response time. The thermal properties of the BLAST bolometers are chosen so that the electrical time constant is $\tau \sim 2\ms$.

The detectors are fabricated in arrays, one for each of the three bands, 250, 350 and 500 $\mu$m. The physical size of each bolometer is determined by the diffraction limit of the telescope, thus the shorter-wavelength detectors are smaller and more elements will fit into the focal plane. The arrays contain 149, 88 and 43 detectors at 250, 350 and 500 $\mu$m.

The detectors are coupled to the optics with 2$f\lambda$ feed horn arrays which provide maximal
point source sensitivity per detector; this solution is preferred over filling the focal plane when
the number of detectors is limited (Gear and Cunningham 1990).

2.1.2 Thermal Model

We calculate the thermal constant based on a model for the resistance and thermal characteristics of the bolometer. This discussion follows closely the derivation in Rieke (2003, Chapter 9).

The thermal properties of a bolometer are modelled as a thermal mass connected by a weak thermal link, characterized by the thermal conductance $G$ and heat capacity $C$, to a bath with temperature $T_0$. A constant background power $P_0$ incident on the bolometer heats it to a temperature $T_1$,

$$T_1 = \frac{P_0}{G} \quad (2.1)$$

A variable input power

$$P_s(t) = \frac{dQ}{dT} = C \frac{dT_1}{dt} \quad (2.2)$$
incident on the bolometer leads to a differential equation for $T_1$. The total power $P_T$ is then

$$P_T = P_0 + P_s(t) = GT_1 + C \frac{dT_1}{dt}. \tag{2.3}$$

The behaviour of this equation is shown by setting $P_s(t)$ to a step function,

$$P_s(t) = \begin{cases} 0, & t < 0 \\ P_1, & t \geq 0 \end{cases}. \tag{2.4}$$

We then have, for $t \geq 0$,

$$\frac{dT_1}{dt} = \frac{P_0 + P_1}{C} - \frac{G}{C} T_1, \tag{2.5}$$

with initial condition $T_1(0) = P_0 / G$. This has solution

$$T_1(t) = \frac{P_0}{G} + \frac{P_1}{G} \left( 1 - e^{-t/(C/G)} \right). \tag{2.6}$$

From inspection, we see that the thermal time constant $\tau_T = C / G$. $T_1(t)$ approaches $(P_0 + P_1) / G$ for times long compared to $\tau_T$, and we see that the temperature of the bolometer is a direct measure of the input power.

To follow this theoretical argument further, we must account for electrothermal feedback (Mather 1982). In order to measure the temperature $T_1$, we run a current through the thermistor, which is in direct thermal contact with the absorber. The thermistor has a resistance $R(T)$ which depends strongly on temperature. Power is dissipated in the thermistor, heating the system. The load resistors (see Fig. 2.2) are chosen with resistance much greater than $R(T)$ so that, for a given bias voltage, the current through the system $I$ is, to first order, independent of thermistor resistance. The electrical power dissipated in the system is thus $P_I = I^2 R(T)$.

The primary mode of conduction in doped semiconductors at low temperature is by “hopping”, where electrons tunnel from one atom to the next without being promoted to the conduction band. The resistance of such a system is well described by

$$R(T) = R_0 e^{(\Delta/T)\xi}, \tag{2.7}$$

where $\Delta \sim 4-10 \text{K} \gg T$ and $\xi \simeq 0.5$. The temperature coefficient of resistance $\alpha(T)$ is then

$$\alpha(T) \equiv \frac{1}{R} \frac{dR}{dT} = -\frac{1}{2} \left( \frac{\Delta}{T^2} \right)^{1/2}. \tag{2.8}$$

We now modify Eqn. 2.1 by associating the background power $P_0$ with the electrical power $P_I$ and including the first-order correction to the feedback:

$$P_I + \frac{dP_I}{dT_1} T_1 = GT_1. \tag{2.9}$$
Figure 2.2: Schematic of the bolometer readout circuit. $V_i$ is the applied voltage and $V_o$ is the measured output. $R_B$ is the bolometer (thermistor) and $R_L$ is the load resistor. $C$ is the capacitance in the leads between the resistor and the amplifier, which has gain $G_{\text{amp}}$.

From Eqn. 2.8

$$\frac{dP_1}{dT} = I_1^2 \frac{dR}{dT} = \alpha(T) I_1^2 R(T), = \alpha(T) P_1$$

and Eqn. 2.3 becomes

$$P_T(t) = (G - \alpha(T) P_1) T_1 + C \frac{dT_1}{dt}$$

from which, by comparison with Eqns. 2.3 and 2.6 we see that the electrical time constant $\tau_e$ is

$$\tau_e = \frac{C}{G - \alpha(T) P_1}.$$  

Since $\alpha(T)$ is negative and $P_1$ is positive, $\tau_e$ is smaller than $\tau_T$, and hence the electrical response of the system is faster than the thermal response, ignoring the feedback.

### 2.1.3 Responsivity

The bolometer readout circuit is shown schematically in Fig. 2.2. An AC bias voltage

$$v_i(t) = V_i \sin(\omega t),$$

where $\omega$ is angular frequency of the AC signal and $V_i$ is the amplitude of the input signal, is applied across the system. The output voltage

$$v_o(t) = V_o(t) \sin(\omega t + \phi)$$
is read by a lock-in amplifier which isolates the signal from the broad-band noise. \( V_o \) varies in time due to the changing power incident on the bolometer and \( \omega \) is chosen to be much faster than the scale of changes in \( V_o \). The load resistance \( R_L \) is chosen to be much larger than the bolometer resistance \( R_B \) so that the current \( i(t) \) across the bolometer is nominally independent of input bolometer resistance and thus power.

The responsivity of a bolometer is defined as the change in voltage across the thermistor due to a change in input power,

\[
S = \frac{dV}{dP},
\]

where \( V \) is the voltage across the bolometer. When \( R_L \gg R_B \), the current \( I \) is nearly independent of the input power, so \( dV = I \, dR \). From the definition of \( \alpha(T) \) (Eqn. 2.15),

\[
dV = \alpha(T) \, R \, dT = \alpha(T) V \, dT.
\]

From Eqn. 2.11 for large time scales we have

\[
dT = \frac{dP}{G - \alpha(T) P}.
\]

Eqns. 2.15, 2.17 then imply that

\[
S = \frac{\alpha(T)}{G - \alpha(T) P},
\]

for \( R_L \gg R_R \). Given the thermal conductivity \( G \), temperature dependence of thermistor resistance, and loading on the bolometer, one could then calculate the responsivity. In practice, however, \( R(T) \) and the loading are not well known, and this theoretical formulation is of little practical use.

We can reformulate \( S \) in terms of measurable quantities, however; we measure a load curve for the detector by varying in the bias voltage \( V_i \) and measuring the thermistor resistance, given in Eqn. 2.25. At each point \((V_o, I)\) on the load curve, we can calculate the slope of the curve,

\[
Z = \frac{dV}{dT}.
\]

We can rewrite \( Z \) in terms of \( d(\log V) \) and \( d(\log I) \),

\[
Z = R \frac{d(\log V)}{d(\log I)} = R \frac{H + 1}{H - 1},
\]

where

\[
H = \frac{d(\log P)}{d(\log R)} = \frac{G}{\alpha(T) P}.
\]
We can then invert Eqn. 2.20 giving $H$ in terms of $Z$ and rewrite Eqn. 2.18,

\[ S = \frac{Z - R}{2V}. \]  

(2.22)

We now discuss how to infer $R$ from $V$ when the lead capacitance $C$ is not zero and the assumption $R_L \gg R_R$ cannot be made.

### 2.1.4 Read-out

A complication in the system is that the capacitance $C$ in the leads is significant. We show here how to deduce the bolometer resistance $R_B$ given the known quantities $V_i$, $R_L$, $\omega$, $C$ and the readout gain $G_{\text{amp}}$, and the measured output $V_o$.

If we define $R_N$ to be the impedance of the dotted region in Fig. 2.2,

\[ R_N = \frac{R_B \cdot j/\omega C}{R_B + j/\omega C} = \frac{R_B \omega C}{1 + (R_B \omega C)^2} \cdot \left( \frac{1}{\omega C} + jR_B \right), \]  

(2.23)

where $j$ is the imaginary unit $\sqrt{-1}$, we can write $V_o$ as a function of the bolometer voltage and circuit parameters:

\[ \frac{V_o}{V_iG} = \frac{R_N}{R_L + R_N} = \frac{(R_L/R_B + 1) + jR_L\omega C}{(R_L/R_B + 1)^2 + (R_L\omega C)^2}. \]  

(2.24)

Defining $V_o^* = \text{Re}(V_o)$ and taking the real part of Eqn. 2.24 we can solve for the bolometer resistance in terms of measured properties:

\[ R_B = \frac{R_L V_o^*}{2V_iG \left( 1 \pm \sqrt{1 - \left( \frac{2R_L \omega C V_o^*}{V_iG} \right)^2} \right) - 1}. \]  

(2.25)

The amplified bolometer signals $V_o$ are read by the data acquisition system, described in Sec. 2.6.1. This equation is used to infer $R_B$ when calculating load curves, discussed further in Sec. 3.3.1.

### 2.1.5 Noise Properties

The noise in a bolometer signal consists of electrical Johnson noise, thermal phonon noise and photon noise. The bolometers are designed such that photon noise dominates at high frequencies. The noise equivalent power (NEP) due to Johnson noise in the bolometer is

\[ \text{NEP}_J = \left( \frac{4kT}{P} \right)^{1/2} \frac{G}{|\alpha(T)|}. \]  

(2.26)
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Figure 2.3: The noise power spectrum for a single bolometer, measured in flight (solid curve). The spectrum is dominated by $1/f$ noise at low frequencies, with knee $f_0 \sim 0.1$ Hz. The part of the signal common to all bolometers (dotted curve) is subtracted, resulting in an uncorrelated noise power spectrum (dashed curve) with $f_0 \sim 0.02$ Hz. The excess power at $\sim 0.04$ Hz is due to the scanning frequency of the telescope and is common-mode to all detectors.

The excess power is due to the scanning frequency of the telescope and is common-mode to all detectors, and is thus proportional to $GT^2$ for $\alpha(T) \propto T^{-3/2}$ (Eqn. 2.8). Thermal phonon noise is described by

$$\text{NEP}_T = (4kT^2 G)^{1/2},$$

which is proportional to $\sqrt{GT^2}$, and thus acoustic noise dominates the electric Johnson noise for low $GT^2$. $G$ is then chosen to maximize signal-to-noise, given by $G \sim Q/T$ for a background load $Q$ (Mather 1984). The detectors are optimized for loads of 55, 40 and 30 pW at 250, 350 and 500 μm, respectively. The resulting NEP is $3 \times 10^{-17}$ W Hz$^{-1/2}$.

At high frequencies, the bolometers are described by white noise, but the noise increases strongly at low frequencies, a phenomenon known as $1/f$ noise (Mather 1982). The noise power spectrum for one bolometer, measured from in-flight data during observations of a blank field in the 2005 flight, is shown in Fig. 2.3. The knee frequency $f_0$ of this spectrum, where the power in the $1/f$ component is equal to the power in the white component, is $\sim 0.1$ Hz. We find, however, that a large fraction of the bolometer signals is common-mode between all bolometers and can be subtracted. The $1/f$ noise in the resulting uncorrelated signals is much reduced,
with $f_0 \sim 0.02$ Hz (dashed curve in figure).

### 2.2 Cryogenics

A long hold-time cryostat provides the cooling necessary for proper operation of the detectors. Liquid nitrogen (LN), liquid helium (LHe), a pumped $^4$He pot and a self-contained recycling $^3$He fridge keep the cold optics at $\sim 1$ K and the detector assemblies at $\sim 300$ mK. The JFETs, the first-stage amplifiers, are also located in the cryostat, and are kept at $T \sim 100$ K (see Fig. 2.4). The components are mechanically connected with thin-wall stainless steel and G10 fibreglass. The space around the components is evacuated to $< 10^{-6}$ atm for thermal isolation.

The LN tank holds $\sim 43$ L and sits in the outer part of the upper half of the cryostat. The LHe tank holds $\sim 32$ L and sits inside of the LN tank, also in the upper half of the cryostat. The bottom of the LHe tank is thermally connected to an optics bench, or “cold plate”, to which the $^3$He fridge and optics box assemblies are attached. The cold plate, $^3$He fridge and optics box are surrounded by a series of thermally isolating shields, each wrapped with multi-layer aluminized mylar insulation. The innermost shield is thermally connected to the cold plate and sits at $\sim 4$ K. The next, the *vapour-cooled* shield, is cooled by the vapour boil-off from the LHe tank and sits at $20–45$ K. The final shield is thermally connected to the LN tank and therefore sits at $\sim 77$ K. The optical beam passes through IR blocking filters attached to each shield (described in Sec. 2.3.3). The outermost layer, the cryostat shell, sits at ambient temperature, $\sim 0$ °C during flight.

A pumped $^4$He pot cools the optics box, including the cold optics and filters, to $\sim 1$ K. It is also necessary for cycling the $^3$He fridge. The pot is simply a chamber into which helium from the LHe tank is bled through a capillary and which is pumped to a low pressure. In the lab a roughing pump is used, but at float the pot is open to the atmosphere which is at $\ll 0.02$ atm.

The $^3$He fridge consists of a sealed chamber in two parts connected by thin-wall stainless steel, to which the pumped pot is thermally connected. Liquid $^3$He collects in the lower chamber. The other chamber is filled with charcoal which absorbs the evaporated gas and serves as a pump which reduces the pressure of the liquid and reduces the boiling temperature to $\sim 300$ mK. A cold plate is attached to the bottom of the liquid chamber, which serves as a heat sink for the detector assemblies. A copper rod attached at one end to the $^3$He fridge cold plate passes
Figure 2.4: Cut-away view of the cryostat. The upper half of the cryostat contains the liquid cryogen tanks, with the nitrogen tank surrounding the helium tank, the two separated by the vapour-cooled shield (VCS). The optics box containing re-imaging optics (M3, M4 and M5) and detectors is found in the lower half. Light enters the cryostat through the window on the left. The $^3$He fridge sits in front of the optics box and has been suppressed from this image.
through a hole in the optics box and connects to the three bolometer assemblies to provide cooling. The fridge cycle ends when all of the liquid has evaporated, typically after 3–4 days. The recycling process involves turning on a heater connected to the upper chamber. When the charcoal warms (to \( \sim 20 \) K), the gaseous \(^3\)He is released into the system, whereupon it reaches the cooled stainless wall, re-condenses and collects again in the lower chamber. After all of the \(^3\)He has been evaporated from the charcoal, the heater is turned off and the cycle is repeated. The recycling process takes \( \sim 1.5 \) hours.

A window cut in the side of the cryostat allows the optical beam to enter the cryostat. A 0.002 inch sheet of high density polypropylene, transparent to submillimetre radiation and strong enough to withstand the 1 atm pressure difference, covers the window.

Commandable valves capped with pressure regulators are installed on the cryogen tank fill lines to maintain \( \sim 1 \) atm pressure above the liquid cryogens. The pumped pot outlet is also capped with a commandable valve in order to maintain vacuum on the pot while on the launch pad and during ascent.

Low thermal-conductance electrical wires run from the detectors to the first-stage JFET amplifiers located in a cavity in the centre of the cryostat. The amplifiers have weak thermal connections to the vapour-cooled shield and the cryostat shell and sit at 120–140 K.

Two heat switches\(^1\) connect the optics box to the cold plate and enable the cooling of the optics box, a relatively large thermal load, during initial cool down from ambient temperatures. The heat switches remain off in normal operation.

The cryostat is designed to keep the detectors and optics cold for \( \sim 13 \) days. The load on the LN, including radiation, thermal conduction through wires and mechanical conduction through G10, is \( \sim 7.2 \) W. The load on the LHe is \( \sim 86 \) mW, from wires, mechanical, pumped pot and average load due to fridge cycle.

---

\(^1\)A heat switch is a cryogenic device that connects two cryogenically-cooled surfaces. It consists of a thin-wall stainless steel cavity filled with a cryogen. In its rest state, the cryogen, cooled by the lower surface, is in its liquid state and the cavity is a near vacuum and is thus a thermal insulator, isolating the upper surface from the lower. A voltage can be applied to a heater which boils the cryogen and the gas conducts heat from one surface to the other.
Figure 2.5: Diagram of the Cassegrain warm optics (left) and cold re-imaging optics (right).
Incoming light reflects off of the primary and is redirected by the secondary through a hole in the primary into the cryostat. The re-imaging optics direct the light to the beam splitters and to the detector arrays.

2.3 Optics

The BLAST optics consist of a Cassegrain telescope and a set of re-imaging optics. The telescope consists of a primary and secondary which are at ambient temperatures. Two re-imaging mirrors and a Lyot stop are cooled to ~1 K. A set of blocking, beam splitting and band-pass filters sit at various points along the beam. The BLAST optics are described further in Olmi (2002).

2.3.1 Warm Optics

The telescope consists of a 2-m primary and a 50-cm correcting secondary, producing an f/5 focus 18 cm behind the surface of the primary. A spherical 2-m carbon fibre with 2.5 \( \mu \)m surface RMS designed and built by Composite Optics Incorporated (COI) is on loan to BLAST from NASA. A carbon fibre primary backing structure and secondary support structure, as well as the aluminum secondary mirror, were provided by COI (see Fig. 2.5).

Additionally, a second primary mirror, aspherical and with diameter 1.8 m, machined from a single piece of aluminum, was constructed for use in the 2003 BLAST test flight in order to save the carbon fibre mirror for a later long-duration science flight. A corresponding 40-cm secondary mirror and support structure were also built.
2.3.2 Cold Optics

The cold optics consist of two re-imaging ellipsoidal mirrors which serve to correct aberrations in the main telescope and provide flat focal planes. A Lyot stop (or aperture), an image of the secondary, is located between the two re-imaging mirrors and provides side-lobe rejection. All three mirrors are machined from aluminum and are located in the optics box inside the cryostat. See Fig. 2.6 for geometry.

Each of the beams (after passing through or reflecting off of the beam splitters) is directed towards the detector array by a flat reflecting mirror.

A calibration lamp (cal-lamp) is mounted behind the Lyot stop (see Sec. 2.3.2) for the purposes of monitoring variation in bolometer responsivity throughout the flight. The commandable lamp is turned on for \( \sim 150 \text{ ms} \) every 10–15 min in normal operation.

2.3.3 Filters

A series of filters placed in the optical beam, within the cryostat, provide IR blocking, colour-splitting and bandpass definition (Ade et al. 2006).
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Figure 2.7: The BLAST bandpasses, measured using an infrared Fourier transform interferometer. For each bandpass, the curves measured from several detectors have been averaged together, the curves have been corrected for the spectrum of the source, and noise in the wings has been suppressed.

The first filters that the beam passes after it enters the cryostat are a stack of IR blocking filters, designed to reduce the thermal loading on the coldest parts of the cryostat. These filters are placed on each of the radiation shields (vapour-cooled, liquid nitrogen and liquid $^4$He) and on the front of the optics box.

Next, the beam passes through the two low-pass dichroic filters, which split the beam and direct the outputs towards the 3 detector assemblies. Finally, bandpass filters, placed on the front surface of each horn array, define the frequencies seen by the detectors. The transmission functions at each wavelength are measured by observing a cold source (absorber in liquid nitrogen) with an infrared Fourier transform spectrometer (Fig. 2.7).
2.4 Gondola Frame

The BLAST gondola frame provides a pointed platform for the telescope and the attachment point to the balloon flight train. The frame design is driven by the stringent pointing requirements of the science case. The design of a structure rigid enough to meet the pointing requirements and still light enough to fly on a balloon is a difficult task, and the mechanical engineering company AMEC Dynamic Structures, located in Port Coquitlam, B.C., was contracted to design and build the frame.

A target weight of $\sim 2000$ kg was set for the entire scientific payload (i.e. total payload excluding balloon-control electronics and ballast). The maximum width and height of the gondola are constrained by the geometry of the buildings in the field where the gondola is assembled. The launch procedure requires that, in the launch orientation, no part of the gondola intersects a plane $20^\circ$ from vertical passing through the launch vehicle/gondola attachment point. A final constraint on geometry is that all components of the gondola must fit into standard sea shipping containers. Additionally, the gondola must survive the vibrations and shocks suffered during launch and there must be no structural failure during any of the possible load cases under parachute shock, loads which can be as high as $10$ g along the vertical. Finally, the mechanical tolerances must account for temperatures which vary by as much as $\sim 100$ °C during the course of the flight while maintaining alignment and motion.

The frame consists of three components: an outer frame, suspended from the balloon flight train by four steel cables and a pivot motor; an inner frame which is attached to the outer frame at two points along a horizontal axis; and a set of sun shields that attach to the outer frame and keep the telescope shaded from solar radiation (see Fig. 2.8). The elevation range of the inner frame, which holds the 2-m primary mirror and the $\sim 200$ kg cryostat, is 25–60°. The entire gondola can rotate to any azimuthal angle.

As a consequence of the pointing requirements, the feedback rate of the control system is $\sim 10$ Hz. To accommodate this, the gondola is designed to be rigid, with a minimum resonance frequency at 14.4 Hz, corresponding to rotation of the inner frame about the elevation axis with the elevation motor fixed, due to bending and torsional stiffness of the inner frame structural components. All mechanical tolerances are set to minimize backlash. All bearings are low friction and low stiction with low-temperature grease. The center-of-mass of the system is on
Figure 2.8: The BLAST gondola frame, seen head on and from the left (port) side. The gondola is made up of 3 components: (i) a base frame suspended by cables from the balloon attachment point, which also serves as an azimuthal motor; (ii) an inner frame which supports the telescope, cryostat, readout electronics, and pointing sensors, free to move in elevation relative to the base frame; and (iii) a set of shields which enclose the telescope from the Sun and reflected light from the horizon.
the rotational axis so that translations of the gondola (by wind or from the balloon) do not generate torques which re-orient the telescope.

2.4.1 Outer Frame

The outer frame is the base of the telescope gondola. It is connected to the pivot motor, which is connected directly to the balloon flight train, by four 3/4 inch steel cables. The outer frame supports the inner frame, the sun shields, reaction wheel, attitude electronics, flight computers, batteries, support electronics and other smaller sub-systems. It is constructed primarily of aluminum I-beam and pipe and weighs 700 lb bare.

The outer frame provides pointing control in the azimuth (horizontal) direction. Fine control is provided by spinning the reaction wheel and coarse control by a motor in the pivot. The balloon is subject to torques through differential wind speeds, which causes it to rotate one way or the other. The gondola counters this torque by spinning up the reaction wheel. On long time scales, the pivot torques against the balloon to reduce the reaction wheel rotation speed and avoid very high speeds.

Two “pyramids” support the inner frame through a torque motor and free bearing. Since the torque to move the inner frame is applied on only one side of the gondola, the pyramid supporting the free bearing does not need to be as stiff as the motor side. The motor-side (port) pyramid is constructed of 6 inch aluminum I-beam. To save weight, the bearing-side (starboard) pyramid is constructed of lighter-weight 3.5 inch (outer diameter) aluminum pipe.

2.4.2 Inner Frame

The inner frame provides the pointed platform and supports the telescope, cryostat and detectors, detector electronics and various pointing sensors, including the integrating star cameras. It is attached to the outer frame at two points along a horizontal axis through a torque motor and a free bearing. It is constructed of aluminum square beam and weighs ~300 lb bare.

The inner frame provides pointing control in elevation through the torque motor. It is free to move from roughly 25–60° in elevation. Digitized detector signals and housekeeping signal cables are passed through to the outer frame electronics and flight computers through the free bearing.
2.4.3 Sun Shields

Sun shields are required because the telescope observes during the day. They surround the back and sides of the instrument, preventing solar radiation (direct and scattered) from entering the beam and thermally protecting the telescope and all of its sub-components. The shields consist of an aluminum frame to which foam sheets covered with aluminized mylar are attached with tape and string.

The shields are designed to block direct sunlight from striking any part of the instrument when it is pointed directly away from the Sun and the Sun elevation is no larger than 30°. This limit was chosen with the Antarctic night in mind; the Sun rises to higher elevations during northern hemisphere flights, and in this case, the secondary mirror suffers direct sunlight at certain telescope elevations. The shields also protect the telescope from radiation scattered from the ground, which is assumed to be at < −5°. The Sun can reach elevations greater than 30°, however, depending on the latitude and time of the flight, and pointing the telescope away from directly anti-Sun allows for direct illumination of the secondary and illumination of the primary through a single scattering. The front of the gondola is not completely sealed with shields to allow heat to escape to the sky.

2.5 Pointing System

Accurate pointing is one of the most difficult aspects of the BLAST gondola design. The 2-m primary mirror corresponds to a diffraction limit of ~30″ at 250μm. The required pointing accuracy is easily achievable from the ground where the telescope is mounted to a stable platform (the Earth), but a balloon-borne telescope is subject to torques from the balloon and wind. To achieve the desired in-flight pointing level of 30″ RMS and 5″ RMS for post-flight reconstruction, responsive motors and accurate pointing sensors are needed.

2.5.1 Active Pointing

The telescope is controlled primarily by three torque motors: the elevation, reaction wheel and pivot motors. The motors are all Kollmorgen Direct Drive DC Torque Motors using rare-earth permanent magnets. The housings are custom designed and incorporate the bearings. All torque motors are driven by 50 Amp PWM servo amplifiers. In addition to the three primary
motors, there is an inner frame fluid balancing system used for long-term balancing, and an active roll damper.

2.5.1.1 Elevation

The elevation motor housing and bearing provide the connection points between the inner and outer frames. These are shown on the left side of Fig. 2.8. The torque motor is 181 mm wide and 101 mm tall and provides 76 N·m of torque. It is housed along with a spherical roller bearing, 80 mm wide (outer diameter) and 23 mm tall. The spherical bearing lets the alignment of the motor be defined by the attachment point at the other end, the free bearing, which allows for some amount of mechanical imprecision, a certainty in an object as big as the BLAST frame.

A shaft passes through the centre of the motor and connects through a flexible coupling to a shaft encoder, mounted to the outside of the motor housing. This encoder is used to measure the relative angle between the inner and outer frames.

2.5.1.2 Reaction Wheel

A reaction wheel, a large disk with a high moment of inertia, provides control of azimuthal pointing. To achieve a large moment of inertia with as little weight as possible, we attempted to construct a large, stiff, light-weight structure with heavy “pucks” at its perimeter. An obvious light, strong material is aluminum honeycomb, a very thin aluminum hexagonal-celled structure sandwiched between thin aluminum sheet metal. Approximately 50 kg in brass pucks are attached to the perimeter.

The initial design was made by AMEC. It consists of a circular disk with 59 inch diameter made from 1 inch thick aluminum honeycomb (see Fig. 2.9). Cut-outs were made for lightweighting, producing a wheel-like structure with spokes. 48 brass pucks, cylinders with 3 inch diameter and approximately 1 inch high, each weighing ~1 kg, are screwed down to threaded inserts in the rim of the wheel.

The wheel was balanced using a rig consisting of a piece of 8 inch aluminum channel, a mount for the reaction wheel and a cross-piece for stabilization (Fig. 2.10). The mount holds the reaction wheel on a 1 inch steel rod supported by 2 bearings. An 18 inch arm attaches on one side to the shaft and on the other supports a 3/4 inch shaft which supports the reaction wheel. A small DC motor and a rubber band drive the wheel and a shaft encoder reads the
Figure 2.9: Dimensions of the original BLAST reaction wheel, given in inches. The final design uses 3 inch honeycomb, eliminates the cut-outs and embeds the weights inside the honeycomb.
angular position. The wheel is driven to a constant speed and an imbalance in the wheel will cause it to oscillate. A capacitive transducer (not shown) is attached to the arm and measures the arm’s motion. Custom Labview software reads the shaft encoder and capacitive transducer. The angular position of the wheel corresponding to the phase of the arm’s displacement tell us where to add weight. A small weight (~35 grams) is attached and the measurement is repeated, iterating until the wheel is as balanced as possible. The accuracy of the measurement is limited by noise in the capacitive transducer, ~0.1 V. This corresponds to a change in weight at the perimeter of the wheel of ~2 g.

After installing the balanced wheel on the gondola, we determined that the structure was far from stiff enough to support the weight with enough rigidity. This hadn’t been obvious during balancing since the balancing rig is far from rigid. It was only once the wheel was mounted on the gondola frame when we were able to test the active pointing system that this became clear. We modified the design to use 3 inch aluminum honeycomb, eliminated the cutouts and embedded the brass pucks inside the honeycomb. The new design proved to be sufficiently sturdy.

The reaction wheel torque motor is the same model as the elevation motor, but the housing is significantly different. It uses two parallel angular contact ball bearings, 72 mm wide and 17 mm high, which fix the alignment of the rotor compared to the outer housing. A shaft, attached to the reaction wheel hub with a flexible coupling, passes through the centre of the motor to a shaft encoder and connects to an optical shaft encoder, mounted to the bottom of the motor housing. The hub of the reaction wheel motor provides alignment of the wheel, ensuring that the vertical axis of the wheel is aligned with the gondola frame.

Since the reaction wheel sits on the floor of the base frame, directly bellow the inner frame, a plate made of 1 inch aluminum honeycomb is mounted a few inches above the wheel to protect it from dropped items and the weight of team members’ inadvertent footsteps during assembly of the instrument. This cover also proved a useful mounting spot for the flight batteries.

2.5.1.3 Pivot

The gondola is subject to external torque from wind shear on both the gondola and the balloon. Torque motors cannot run at high speeds, and the reaction wheel motor will eventually saturate trying to overcome the external torques. We therefore place an additional torque motor in
Figure 2.10: Detailed drawing of the rig for balancing the reaction wheel. The wheel is held off-axis by a 1 inch shaft, supported by ball bearings. The wheel is spun by a small DC motor connected with an O-ring. Threaded rods allow for tilting the system.
the pivot, at the interface between the gondola and the balloon. This motor reduces excess angular momentum built up by the external torques and keeps the reaction wheel speed within reasonable limits. In-flight performance of this system is described in Sec. 3.6.2.

The pivot is the only attachment point between the gondola and the flight train, and is thus a single point of failure. It must therefore be reliably strong, and be able to sustain the 10-g shock on parachute opening during descent. The BLAST pivot is made of hardened steel and consists of a central shaft, an outer casing, a torque motor and several ball bearings. The outer casing has four ears (machined from the same block of metal, not welded) which attach to the suspension cables which in turn attach to the four corners of the base frame, and the shaft is bolted to a universal joint above. The outer diameter of the shaft is larger than the hole in the casing, so even if the inner workings of the pivot fail, the shaft cannot pass through the casing, separating the gondola from the flight train. The pivot and universal joint are shown in the left of Fig. 2.11.

The original design for the pivot, shown in the right of Fig. 2.11, called for two motors, the main torque motor (hatched region, lower) and a smaller race motor (hatched region, upper). The race motor rotor is connected to the outer casing through a bearing (top) and to the shaft through a larger bearing (bottom). The main torque motor rotor is attached directly to the shaft. The stators of both motors are attached to the casing. The purpose of the race motor is to eliminate static friction associated with starting the main motor from stand-still; the race motor runs at all times at a nominal speed, greater than the largest possible main-motor speed, spinning both bearings. The main motor then never has to start the bearings spinning. In practice, however, we found that the friction in the system was large, even with the race motor system, and we therefore replaced the race motor with a block of aluminum, greatly simplifying the system. We also replaced the spherical bearings (to allow for misalignment) with regular axial ball bearings, greatly reducing the friction and improving the overall performance.

Also shown in Fig. 2.11 are the shaft encoder (rectangle at bottom of assembly), a flexible coupling (above the encoder) and a 18-circuit slip ring (above the coupling), which allows signals to be passed from the gondola to the flight train. The Columbia Scientific Balloon Facility (CSBF) uses these wires to send signals from the CSBF electronics to the flight train.

The race motor is 208 mm wide (outer diameter) and 27.7 mm tall, providing 12 N·m of torque. The main torque motor is 178 mm wide and 50.6 mm tall and provides 27.1 N·m of
Figure 2.11: The BLAST pivot with universal joint on the left, and in cross-section on the right. The pivot consists of a central shaft which connects to the balloon flight train and an outer housing which suspends the gondola. The shaft is captured by the housing. This is the original design which has since been modified by eliminating the race motor.

2.5.1.4 Reaction Wheel Alignment Test

The rotation axis of the reaction wheel should point at the pivot so that reaction wheel torque couples purely to azimuthal acceleration. While assembling the telescope before the 2005 flight, we tested the alignment of the wheel by measuring the variation in height of the bottom surface across one full rotation. The measured vertical deflection is shown in the top panel of Fig. 2.12. A sine wave is fit to the data and we find a peak-to-peak amplitude of 0.046 inch. This corresponds to a misalignment of $\sim 3'$.

To investigate whether this misalignment is significant, we compared the power spectrum of the gyroscope signals with the reaction wheel running with the reaction wheel stopped to
Figure 2.12: Alignment of the reaction wheel, as measured during the assembly of the instrument before the 2005 flight. Vertical deflection of the bottom of the wheel as a function of wheel rotation is shown (top), along with a sine wave fit to the data. The curve has a peak-to-peak amplitude of 0.046 inch. The fit residuals (bottom) show a sine wave with a period of half a rotation, showing flexure in the wheel.
see whether the motion was visible in the gyroscopes (Fig. 2.13). The test was performed with
the gondola hanging from the ceiling, with the reaction wheel spinning at 0.8 Hz, and with the
elevation frame locked at 18.5°. The reaction wheel is driven by a servo which aims to keep
the azimuth of the telescope stable. We thus do not see the signal of the reaction wheel in
the yaw gyro, since the wheel is driven to minimize motion in azimuth. The pitch servo was
turned off for this test, and we thus do not see damping in the elevation gyro. We note small
peaks at the reaction wheel rotation frequency in the pitch and roll gyroscopes that disappear
when the reaction wheel is turned off. It was decided that this amount of signal was too small
to worry about re-aligning the wheel, as the pointing system seemed to work well despite the
misalignment.

2.5.1.5 Fluid Pump System

In the course of a 10-day flight ~40 L of liquid nitrogen are boiled off. Since the cryostat is
~50 cm away from the inner frame rotation axis, this change in mass results in a fairly large
torque. To compensate, a pumped-fluid balancing system is installed on the inner frame. It
consists of two 10 L tanks mounted on the inner frame. The tanks are placed as far away from
the rotation axis as possible, while keeping the centre-of-mass of the tanks as near to the axis as
possible. The tanks are connected by a line that runs through a check-valve and a bi-directional
motor, and a fluid called Dynalene\textsuperscript{2} is transferred from one tank to the other so that the frame
is as balanced as possible. The balance of the inner frame, indicated by the current drawn by
the elevation motor, is monitored. If the balance is off one way or the other over a long time
period, fluid is pumped in the appropriate direction to correct the balance. Fluid transfer is
also required after a large slew in elevation, since the inner frame becomes unbalanced by the
redistribution of cryogens inside the cryostat.

2.5.1.6 Active Roll Damper

A small reaction wheel is used to actively damp gondola oscillations in roll, about the forward
direction. The wheel is made of aluminum, approximately 16 inch diameter, and weighs a few
pounds, most of which is at the perimeter. Given the wheel’s small moment of inertia, it is
not expected to damp oscillations in roll quickly, but is effective on long time scales which are

\textsuperscript{2}Chosen for its high density and low freezing point.
Figure 2.13: Power spectra of gyroscope signals with the reaction wheel running at 0.8 Hz (top row) and not running (bottom row).

The reaction wheel speed is marked with a dotted line in all plots. A small excess at 0.8 Hz is visible in the pitch (elevation) and roll gyroscopes, but not in the yaw gyro as the reaction wheel speed is controlled off of the yaw gyro.
otherwise unconstrained.

2.5.2 Pointing Sensors

BLAST uses many complementary and redundant pointing sensors. The primary pointing sensors are the rate gyroscopes and optical star cameras. Coarse sensors, in particular the differential GPS, are used as input to the fine sensors.

2.5.2.1 Gyroscopes

BLAST uses two sets of redundant optical rate gyroscopes. Each set consists of three orthogonal single-axis optical gyroscopes. The models used are ECORE 2000 analogue output and the DSP 3000 digital output gyroscopes from KVH Industries. The analogue and digital gyroscopes have noise levels of $5 \times 10^{-2}$ and $4'' \text{s}^{-1/2}$, respectively.

The two sets of gyroscopes are each mounted orthogonally inside of a rectangular box, which is in turn mounted on the inner frame. The axes correspond to yaw (azimuth), pitch (elevation) and roll (cross-elevation) of the inner frame. Mechanical misalignment is inevitable, however, such that the gyroscopes signals will not be pure measurements of the nominal rotation that they are intended to measure. This misalignment is measured after the boxes have been installed in their final positions on the inner frame. We rotate the gyroscopes in the three directions separately, first rotating only the inner frame, then rotating only the outer frame with the inner frame pointed at 0 and 90°. The coefficients of the corrective transformation matrix are measured from the resulting gyroscope signals during these measurements.

It was found during testing that the gyroscopes are sensitive to magnetic fields via the Faraday effect. For this reason, the gyroscope boxes were shielded with mu-metal, a material with a high value of $\mu$, the magnetic permeability, to shield the gyroscopes from the changing magnetic fields experienced by rotation of the gondola in the Earth’s magnetic field.

2.5.2.2 Integrating Star Cameras

Gyroscopes measure velocity and not position. To get a position from the gyroscopes the solutions must be integrated, which gives an error proportional to the square root of time. BLAST uses two integrating optical star cameras to correct the gyroscope drift on scales of $\sim 1\text{ min}$. The main purpose of having two cameras is for redundancy, in case one fails. For
historical reasons we refer to the cameras as the ISC and OSC. The cameras are identical except for the sensor: the ISC’s sensor has a much larger pixel well depth, but also a larger read noise. See Rex et al. (2006) for details.

Each camera is a self-contained pressurized unit consisting of the detector (CCD), lens, control computer and focus aperture and motors. A 4-foot baffle is attached to the front of the vessel to reduce the amount of light scattered into the detector. The ISC (OSC) uses a $1312 \times 1024$ ($1360 \times 1036$) CCD with 6.8 (6.45) $\mu$m pixels. Each camera uses a 200-mm f/2 Nikon lens, giving a $2^\circ \times 2.5^\circ$ field of view and a 7" pixel scale. A red filter is used to increase the contrast between stars and sky. Focus control is needed to compensate for expansions and contractions due to temperature changes, and aperture control is needed for ground tests.

The cameras work autonomously and communicate with the BLAST flight computers via ethernet. Power and a trigger signal come from the ACS. Custom software detects stellar objects in the image and compares their relative positions to the positions of stars in a catalogue to construct the orientation of the camera. Ecliptic coordinates of the optic axis and roll (rotation of image on the sky), as well as diagnostics such as number of stellar objects found in the image, are returned to the flight computers.

A proper pointing solution requires at least 3 stars. Integration times of up to 300 ms are required to obtain 3 stars in the field of view in bright daytime float conditions. Typical mapping slew speeds will spread the stars across many pixels in 300 ms, so the exposures are timed to occur during scan turn-arounds when the scanning velocity is very low. Short exposures ($\sim 60$ ms) are taken during slews; these exposures will usually detect only one or two stars (if any), but can be useful for updating the post-flight pointing solution during long slews.

2.5.2.3 Encoders

Each of the torque motor assemblies is equipped with a shaft encoder for use in the control servo loop. The elevation motor uses a Gurley Precision Instruments model A25S absolute encoder with 17-bit resolution, giving a resolution of $\sim 10''$ per count. Absolute position is unimportant for the pivot and reaction wheel, so relative encoders are used. The pivot encoder, Dynapar Series HS35, has 11-bit resolution giving $\sim 10'$ per count position resolution and the reaction wheel encoder, Dynapar Series M21, has 1000 counts per revolution giving $\sim 20'$ per count.
2.5.2.4 Differential GPS

BLAST uses a Differential Global Positioning System (DGPS), model ADU5 from Magellan Navigation. The system consists of four receivers that are mounted to the top of the Sun shield assembly. The DGPS provides latitude, longitude, altitude and orientation. The receivers are placed as far from each other as possible, in approximately a $4\text{ m} \times 2\text{ m}$ rectangular array, to provide higher accuracy.

2.5.2.5 Sun Sensor

A sun sensor consisting of two orthogonal linear $1 \times 2048$ CCDs is mounted to the back of the gondola. It reports the azimuth of the gondola relative to the Sun to an accuracy of $\sim 3'$.

2.5.2.6 Magnetometer

A three-axis flux-gate magnetometer, Applied Physics Model 534, is used as a rough absolute pointing sensor. It measures the azimuth of the gondola with respect to the Earth’s magnetic field and is accurate to a few degrees. Of course, the direction of the Earth’s magnetic fields are not well known near the (magnetic) poles, so we do not expect high accuracy from the magnetometer, but it provides a useful backup should the DGPS and sun sensor fail.

2.5.2.7 Clinometers

Several Model 900 biaxial clinometers from Applied Geomechanics are mounted on the gondola in order to measure tilt of the pivot, outer and inner frames. For the most part, they are ignored, since they measure local acceleration and not absolute angles. An accelerometer mounted at the center of gravity of a simple pendulum in free motion will measure no change in angle. A $\pm 45^\circ$ clinometer mounted on the inner frame is sometimes used to measure elevation in the lab.

2.6 Warm Electronics

The BLAST electronics, excluding the cold-stage amplifiers inside the cryostat, consist of three main sub-systems: (i) the data acquisition units, consisting of 2 boxes called the preamp and the data acquisition system (DAS), which process the bolometer and cryostat housekeeping signals; (ii) the attitude control system (ACS), a unit which drives motors and reads pointing sensors;
and (iii) the flight computers, which control the telescope and write the data to disk. The three systems communicate over serial RS485, in a proprietary architecture known as the BLASTbus. The BLASTbus transmits 32-bit requests and responses running at 4 MHz, providing a data bandwidth of 1 Mbit s\(^{-1}\). The electronics system is shown schematically in Fig. 2.14.

The DAS and ACS use analogue-to-digital converter (ADC) cards developed by the BLAST team. Each card consists of a digital signal processor (DSP), a watchdog circuit, 25 ADCs, 24 digital I/O lines and 4 pulse width modulation (PWM) outputs. The ADCs provide 24-bit sampling at 10 kHz. Each card monitors the BLASTbus, accepting commands and providing data upon request.

### 2.6.1 Data Acquisition System

The preamp and DAS units transform the analogue bolometer and housekeeping signals to the 100 Hz digitized time-streams that are written to disk. The preamp crate consists of 12 preamp cards, each operating on 24 data channels. The DAS crate consists of 12 DAC cards, each of which filters and digitizes 24 bolometer and diagnostic signals. It also houses the bias card (discussed below) and the power module that provides DC power to all of the preamp and DAS electronics.

The positive and negative bolometer signals output by the JFETs are fed to the preamp crate, where the pairs of signals are differenced, amplified and filtered. The filter is a 100 Hz bandpass centred on 208 Hz. The filtered signals are then passed on to the DAS where they are digitized to 24 bits at 10 kHz. The signals are then fed to the digital lock-in amplifier. A numerical sine wave is constructed at the bias frequency and with user-specified phase. The rectified signals are then low-pass filtered in the time domain, with the first null at 50.08 Hz, to eliminate aliasing in the final data product. The data are then decimated to 100.16 Hz and assembled into 5 Hz frames.

The DAS also houses the bias card, which generates the bolometer bias signals. A 200.32 Hz square wave is generated by dividing down the main 32 MHz clock. The amplitude is set by the user, selectable from 0–300 mV, and then filtered into a sine wave. Spare ADC channels on the DAS are used to read inner-frame thermometry, described in Sec. 2.8.3.

The preamp and DAS crates are mounted on the inner frame next to the cryostat. Care is taken to isolate the electronics in both crates from radio-frequency (RF) noise generated by
Figure 2.14: Schematic of the warm electronics. The line style indicates the communication protocol. Solid lines indicate analog signals. The Attitude Control System (ACS), Data Acquisition System (DAS), preamp and flight computers are described in Sec. 2.6.
the various motors on the gondola. RF-tight fittings are used on the face plates of each card in the preamp crate, RF filters are placed in-line on the preamp output connectors, and all of the seams in both crates are sealed with conductive aluminum tape.

2.6.2 Attitude Control System

The ACS drives the telescope pointing system and reads pointing sensors. It consists of 3 ADC cards and a power module, which provides DC power for all of the gondola subsystems, excluding the DAS and preamp. The ACS relays commands from the flight computers to various subsystems, including the pointing motors, lock motor, fluid balance system and cooling pumps. The ACS also reads several sensors, including the gyroscopes, magnetometer, inclinometers and temperature sensors. It also runs the temperature control loops for the gyroscope boxes. Finally, it sends the exposure request pulses to the star cameras.

2.6.3 Flight Computers

BLAST is controlled by the master control program (MCP), a custom-built multi-threaded program written in C. MCP creates the in-flight pointing solution and controls gondola motion, archives data, and manages telemetry, commanding and other miscellaneous functions. MCP operates on Intel Celeron computers running Slackware Linux. BLAST uses two redundant computers to reduce the likelihood of in-flight downtime. The computers run independently of each other, each running their own instance of MCP, but only one of them is in charge at a given time. A watchdog circuit monitors that the computer currently in charge is running properly and as soon as it is not, switches control to the other computer and reboots the inactive machine.

The computers communicate with the DAS and ACS over the BLASTbus, sending commands and requesting data, which it writes to disk. Additionally, they communicate with the star cameras and sun sensor over ethernet and with the GPS, SIP computers and lock motor using standard serial connections.

The computers, along with the hard drives, ethernet switch and watchdog circuit are housed in a pressurized vessel, as conventional hard drives require near-atmospheric air pressure for

\[\text{Since BLAST runs above 99% of the atmosphere, the computers are subject to a high flux of cosmic rays which disable the computers upon impact.}\]
proper operation. The pressurized vessel additionally provides temperature control for the computers via self-heating.

The pressurized vessel containing the flight computers is located on the outer frame of the gondola. It is the single most important item for recovery.

### 2.7 Power System

BLAST’s nominal power requirement is 540 W. An additional 160 W is required by the data transmitters during the initial line-of-sight portion of the flight (see Sec. 2.8.4). This power is provided by arrays of solar cells. When at full illumination, perpendicular to the Sun, the arrays provide a peak power of \( \sim 1250 \) W. When the telescope is pointed \( \geq 60^\circ \) away from anti-Sun, the power provided by the arrays can be less than that required. Rechargeable batteries are available when the arrays are not able to provide full power. In practice, the telescope does not spend large amounts of time observing at \( \geq 60^\circ \) from anti-Sun, and the batteries are rarely used, except for on the launch pad and during ascent.

The BLAST power system is divided into two completely distinct subsystems, nominally the inner frame and outer frame systems. The ACS and DAS thus do not share electrical grounds, which eliminates the possibility of line noise on the outer-frame system, due to the pointing system motors, feeding into the bolometer signals.

#### 2.7.1 Solar Cells

The BLAST solar arrays, designed and fabricated by MEER Instruments\(^4\), consist of 10 panels of 96 cells each. The system is overrated for the BLAST power requirements, providing significant redundancy. The power requirements of the inner- and outer-frame subsystems are not equal, with the outer-frame subsystem requiring more power to run the pointing motors, thus 6 of the 10 panels provide power for the outer-frame subsystem and the remaining 4 provide the inner-frame power.

The cells of each panel are connected in series, with strings of 4–8 cells connected in parallel with a set of diodes, so that should one cell die, the entire panel does not short out. Given constant illumination, the solar cells run at constant current up to a maximum output power

\(^4\)http://www.meer.com
where the current falls to zero. The BLAST power modules run at \( \sim 32 \text{ V} \), well below \( P_{\text{max}} \) for nominal illumination. Each panel is able to lose one string of 8 cells without loss of output power. Additionally, the redundancy is such that an entire panel can be lost and the system will still be able to provide the necessary power.

The BLAST solar arrays are mounted on either side of the gondola, along the back sun shield supports. The cells are bonded to a flexible plastic mesh, supported by a thin-walled aluminum tube frame. The backs of the cells are mostly unobscured, and are painted white so that they can efficiently cool by radiating to the sky.

### 2.7.2 Batteries

BLAST uses nickel metal hydride rechargeable batteries developed for use in hybrid vehicles. The units each provide 85 A-hr at 12 V and weigh 17 kg. Each of the 2 power systems uses 2 batteries connected in parallel to produce the required 24 V. This type of battery is subject to a positive feedback cycle as their charge reaches the top-off voltage: the charge efficiency decreases at high voltages which increases the temperature of the battery, which further decreases the top-off voltage. To avoid this catastrophic failure, the temperatures of the batteries are monitored by MCP which can set the top-off in the charge controller.

### 2.8 Miscellaneous

#### 2.8.1 Lock Motor

An inner-frame locking mechanism, which fixes the inner frame to a specific elevation relative to the outer frame, was installed. This prevents loss of control of the inner frame due to the large shocks experienced on launch, descent, and in the lab when the inner frame is significantly unbalanced. The locking mechanism should be retractable so that the frame is free to move once stable conditions have been reached.

Ideally, one would place the lock motor far from the elevation axis to reduce stress on the mechanism. In order to allow for several locking positions, however, we position the lock \( \sim 8 \text{ inches} \) from the elevation axis, on the bottom side of the free bearing side pyramid on the outer frame. The locking mechanism is a linear actuator, capable of 250 lb of force, which

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5Cobasys Model 9500
drives a 3/4 inch hardened steel shaft into a restraint attached to the inner frame. A linear ball-bearing block supports the shaft and takes the axial load. The restraining mechanism is a 1 inch-thick aluminum plate attached securely to the inner frame, with several holes, allowing for restraint of the inner frame at several elevations from 5°-90°. The hole at 45° is larger than the rest and is the position used for launch — should the lock motor fail to retract, there would be some small motion of the inner frame allowed. This is additionally useful for fine-tuning of the balance of the inner frame before launch.

2.8.2 Cooling System

In the high-altitude balloon environment, there is not enough air for efficient cooling of warm subsystems by convection. The primary cooling mechanism is radiation, and care must be taken to ensure that no sensitive component runs too hot or too cold.

One particular danger is that bare aluminum exposed to the Sun, in the absence of cooling by convection, will heat to in excess of 130°C. All aluminum which may be exposed to the Sun is therefore painted white. Care is taken so that the aluminized mylar which covers the sun shields is mylar side out for sun-facing surfaces. On the inside of the gondola where there is no danger of exposure, the aluminized side is exposed so that the shields can absorb thermal radiation from warm gondola components. Gaps in the forward-facing sun shields allow for radiative cooling to the sky.

The ACS, with its power unit and 3 ADC cards runs warm, but due to its location on the gondola can easily radiate to the sky. The ADC cards are shielded from each other by aluminum sheets which conduct heat to the exterior of the box.

The receiver and DAS run very warm and are shielded from the sky by the primary mirror. They must be actively cooled. A circulating liquid cooling system, filled with Dynalene, the same substance used by the liquid balancing system (Sec. 2.5.1), transfers heat from the receiver and DAS to the inner frame, which radiates to the sun shields. As with the ACS, aluminum sheets shield the ADC cards from each other and conduct heat to the outside of the box. A coil of pipe from the liquid cooling system is soldered to a copper plate on the outside of each box. Similar heat-transfer plates are mounted in-line in the cooling system to either side of the inner frame. The cooling system pump and tubing, excluding the heat transfer plates, are isolated from the inner frame with rubber grommets to minimize the transfer of mechanical vibrations.
to the inner frame.

The flight computers self-heat within the pressure vessel. The vessel is painted white and radiates to the sky.

The star camera computers self-heat, like the flight computers, but the lenses still run cold. Thermistors and heaters are mounted to the lenses — in each star camera, the computer monitors the temperature of the lens and runs the heater when necessary. The star camera vessels are not painted, as they would cool too much if that was done.

The gyroscopes must be kept within a certain temperature range, but due to their location behind the primary mirror on the inner frame, tend to run cold. Heaters are installed inside the gyroscope boxes and their environment is thermally controlled by the flight computers.

The primary mirror is thermally isolated from the warm components sitting directly behind it, as its temperature translates into thermal loading of the detectors. Several layers of mylar super-insulation ensure that it is not subject to heating by the warm electronics.

2.8.3 Thermometry

The temperatures of various gondola components are monitored throughout the flight, in some cases for thermal control and in some cases for in- and post-flight diagnostics. All of the spare ADC channels, 10 on the ACS and 9 on the DAS, are used to read AD590 temperature transducers. For the 2005 flight, the outer-frame thermometers were attached to: the sunshield frame (×4); lock motor; free elevation bearing; ACS; face of the flight computer vessel; Sun sensor; and back of the outer frame heat exchange plate. The inner frame thermometers were attached to: one of the star camera baffles; tip of the telescope baffle; back of the primary mirror; shell of the cryostat; pumped pot valve on the cryostat; face of the lower fluid balance tank; and inner frame (×3).

2.8.4 Telemetry

The BLAST gondola is equipped with radio-frequency transmitters for data transfer during the flight. These transmitters require line-of-sight (LOS) to the base station and are thus used only during the first part of flight, before the gondola passes over the horizon. From the launch site in Kiruna, Sweden, this happens approximately 12 hrs after launch. The data transmitter provides a 1 Mbit s\(^{-1}\) transfer rate, the full data rate of BLAST. Additionally, two
high-frequency video transmitters provide the ground station with the video output from each star camera computer, providing useful star camera diagnostics not otherwise available. These transmitters are provided and run by CSBF.

After the gondola passes over the horizon, when the LOS transmitters are no longer operable, communications with the ground stations are over satellite link. This provides both up- and down-links, but at a much lower rate than the LOS transmitters. The down-link bandwidth available over the satellite connection is $\sim 6\,\text{kbit s}^{-1}$, and thus only a small fraction of the data is available for diagnostic purposes. A few detector channels, with high data compression, along with all of the low-rate diagnostics (temperatures, currents, etc.) are transmitted over the satellite link.
Chapter 3

The Experiments

This chapter describes the BLAST flights, including the 1-day North American test flight in 2003, the first long-duration polar science flight from Sweden in 2005 and the second long-duration flight from Antarctica in 2006. The pre-flight planning, in-flight commanding and diagnostics, and flight performance are detailed.

3.1 Scanning Strategies

BLAST is a mapping experiment and must therefore observe the sky in a manner that will allow high-fidelity maps to be produced. Since the bolometer loading, and thus signal baselines, vary significantly at large time scales (Section 2.1.5), it is important that each point in the sky is visited several times by each bolometer. Measurements of a patch of sky are therefore most useful if the telescope rasters the sky several times.

It is of utmost importance that the thermal conditions for the bolometers be kept as stable as possible. This leads us to scan the sky in azimuth, since scanning in elevation causes several different effects on the bolometer environment and loading. When the elevation of the telescope is changed, the cryostat is tilted. The sloshing liquid cryogens can cause drastic drifts in the bolometer baselines. Additionally, varying elevation changes the airmass, the integral of air density along the line of sight, seen by the detectors. Airmass is lowest at the zenith and highest at the horizon. Increased airmass means increased loading on the detectors. It is not practical to completely eliminate motions in elevation, but it is desirable to limit it to slow drifts rather than fast scans.

In order to track the bolometer baseline drifts, it is desirable to re-observe a particular position on the sky, although not necessarily by the same detector, within the time scale corre-
sponding to the $1/f_{\text{knee}}$. For a given scan size $\theta_{\text{scan}}$, this gives a minimum scan rate

$$\omega_{\text{scan}} = \theta_{\text{scan}} f_{\text{knee}}.$$

(3.1)

However, the telescope should not scan so quickly that the sky signal becomes significantly smeared due to the finite bolometer response function, characterized by the time constant $\tau \sim 2\text{ ms}$ (see Sec. 2.1.2). This smearing can be described as a convolution of the sky, sampled at the appropriate rate, with the detector response function, $\exp(-t/\tau)$. The smearing is thus a function of scan speed. Scanning a point source at a rate $\omega_{\text{az}}$ with a telescope beam characterized by its full-width at half-maximum (FWHM), the resulting signal $s(t)$ is a Gaussian in time with width $t_{\text{FWHM}} = \text{FWHM}/\omega_{\text{az}}$. If we adopt a maximum acceptable smearing of $t_{\text{FWHM}}/\tau = 5$, then we find a maximum scan rate

$$\omega_{\text{az}} \approx \frac{\text{FWHM}}{5\tau}.$$  

(3.2)

The 250 $\mu$m beam size is the smallest, equal to $30''$, and thus, with $\tau = 2\text{ ms}$, specifies that $\omega_{\text{az}}$ should be no larger than $\sim 3000'' \text{s}^{-1}$. Using this value and $f_{\text{knee}} = 0.02 \text{ Hz}$ (Sec. 2.1.5), we see from Eqn. 3.1 that we can make scan over and back on angular scales of $\sim 20^\circ$. Larger scans can be made by mapping more quickly and accepting larger amounts of smearing, or by accepting correlated $1/f$ noise in the maps.

Additionally, the elevation drift speed $\omega_{el}$ should not be so large that, upon returning to a point on the sky one scan later, that elevation has changed by more than approximately half the height of the array. Since the arrays are $\sim 6.5'$ high, this corresponds to a maximum elevation drift speed of

$$\omega_{el} \approx \frac{200'}{\theta_{\text{scan}}/\omega_{\text{az}}}.$$  

(3.3)

Using the values for $\theta_{\text{scan}}$ and $\omega_{\text{az}}$ given above, this gives a maximum elevation drift speed of $\sim 60'' \text{s}^{-1}$.

A final requirement is that, if possible, each field is mapped with multiple scan directions on the sky. This is known as cross-linking and greatly improves the ability to produce high-fidelity maps in the map-making stage. Since we have eliminated the possibility of scanning the telescope in elevation, we must scan the field with varying sky rotations. Ideally, the field will be scanned once as the field is rising (before it transits, the earlier the better) and once as it sets (after it transits, the later the better).
Figure 3.1: Scan modes used by BLAST. From left to right, the cap, box and quad. [This figure appears in Pascale et al. (2007).]

The master control program (MCP) provides 3 useful scan modes: the cap, box and quad modes, illustrated in Fig. 3.1. All 3 of these specify boundaries which are filled by rastered scans. A minimum scan length is used, as the turn-arounds at either end of the scan takes $\sim 2$ s for the deceleration and re-acceleration. The shapes are thus slightly distorted at the top and bottom, most visible in small maps. In all modes, the user specifies the azimuth scan speed and the elevation drift speed. “Cap” is a circle with user-specified centre sky location ($\alpha, \delta$) and radius. “Box” is a rectangle in az/el coordinates with user-specified centre ($\alpha, \delta$) and side lengths. “Quad” is an arbitrary quadrilateral, with ($\alpha, \delta$) for all four corners specified by the user. At the beginning of each scan, the flight code calculates the target points at the end of each raster based on the requested scan velocities. When the telescope reaches the end of a raster and obtains a new star camera solution, any drift in desired pointing is corrected for by recalculating the elevation drift speed in order to reach the next target point.

3.2 Flight Planning

As described in Section 2.8.4, the ground station has full commanding capabilities while the gondola is within line-of-site, usually the first 12–24 hours of the flight. After this point, however, only intermittent communications are possible. For this reason, BLAST must be fully autonomous. Before launch, a full set of observation commands must therefore be provided. Scheduling such a large set of observations is a time-consuming and complex operation, so an automated flight scheduler was developed specifically for the BLAST flights. The scheduler
creates a list of commands to perform during the flight and the local sidereal time (LST) at which to perform them. Each entry specifies the MCP command to perform, the LST day and time to start the command, and any parameters associated with the command. If the command is a mapping command, the map continues until the LST of the next command is reached.

The scheduler takes as inputs: (i) a list of targets, complete with desired scan parameters, the desired map sensitivity and relative importance; (ii) a list of calibration targets that should be regularly observed; and (iii) launch time, location, expected flight length and path. Built into the scheduler are: (a) the knowledge of observational constraints, such as the telescope’s elevation limits, the locations of the Sun, Moon and planets which are to be avoided; (b) estimates of how often the $^3\text{He}$ fridge will need to be cycled and for how long; and (c) a model of the power system to ensure that the observations are not scheduled such that the batteries could become fully drained.

The observations are scheduled in LST, since it is the time needed, along with the observational latitude, in order to transform sky coordinates ($\alpha, \delta$) to telescope coordinates (az, el). During the flight, the current Universal Time (UT), longitude and latitude are reported to the flight computers by the GPS, from which LST is calculated. Specifying observation times in LST is complicated, though, by the fact that the telescope moves in longitude. In a northern-hemisphere summer flight, the balloon travels to the west, increasing the length of an LST hour. This effect is not large since the gondola travels slowly in comparison to the Earth’s rotation speed. Nevertheless, an estimate of the gondola’s speed of travel in longitude is used to calculate the anticipated LST duration corresponding to a given observation time.

An additional complication is that the gondola also travels in latitude. Visibility constraints, namely the rising and setting times of targets, depend on the observational latitude, which therefore must be estimated ahead of time. In order to optimize the observation schedule for whatever latitude the gondola happens to be at during a long duration flight, separate schedule files are generated for different latitude ranges. It is expected, based on previous balloon flights, that the gondola will stay within a $\sim 15^\circ$ band centred on the launch latitude. Schedule files are made for three $6^\circ$ ranges covering this band, with an overlap of $1^\circ$. The current latitude of the gondola, as determined by the GPS, is used to determine which of the schedules to use, and should the gondola happen to drift into a different latitude band, MCP switches to the appropriate schedule. The $1^\circ$ overlap adds “hysteresis” to the system, eliminating the possibility
of rapidly switching between 2 schedules, should the gondola travel along one of the dividing lines.

Another assumption made by the flight scheduler is the sensitivity of the instrument; the user specifies the depth to which they wish to observe a field and the observing time required to reach this depth then depends on the sensitivity. The detector sensitivity is not well known until measured in flight, so schedules are created for two likely scenarios: the nominal expected sensitivity; and, conservatively, sensitivities 2 times worse, due to either poor telescope focus or excess loading. The operator chooses which of the two cases to use based on sensitivity measurements made early in the flight. For each of the 2 possibilities, the full set of latitude ranges is calculated, resulting in 6 schedule files.

The flight scheduling program is written in IDL. The user provides a high-level description of the fields to be observed, including mapping parameters, desired integration time and relative importance. The program outputs the set of schedule files, in a format readable by MCP, along with several sets of diagnostic plots. These include: (i) a flight overview, showing the total amount of time spent on each source, and when in the flight the observations occur; (ii) azimuth and elevation of the telescope in time for the duration of the flight; and (iii) Sun angle and estimated battery charge for the entire flight. Should the user not feel comfortable with the schedule, the inputs can be adjusted and the scheduler re-run. The program takes approximately 15 minutes to create schedules for a 5-day flight at a time resolution of 1 min.

3.3 In-Flight Diagnostics

Throughout the flight, the telescope is monitored by a team of ground-station operators. Two software packages are available for quick-look analysis of the data: a graphical control panel, written by the BLAST team, which displays diagnostics such as pointing information, bolometer voltages, temperatures, star camera diagnostics, motor currents and battery voltages; and kst\[1\], an open-source time-stream plotter, which provides fast data access and easy manipulation of data visualization. Additionally, specialized routines written in IDL and C are used for analysis of specific in-flight tests (e.g. Sec. 3.3.1).

Commands are sent to the gondola by the ground station crew using a graphical user

\[1\]http://kst.kde.org\] originated by C. B. Netterfield, a member of the BLAST team.
interface written specifically for BLAST. All of the MCP commands, both high-level commands such as \textit{cap} or \textit{box} as well as low-level commands such as setting the cooling-pump speed, are available to the user through the interface. A program driven by the interface then relays the commands to the CSBF transmitter and are then received by MCP, running on the gondola flight computers.

The operator can override the flight schedule at any time by simply sending a new observation command, but, at times, particularly during the line-of-sight portion of the flight, the user can disable the pre-loaded schedule. Since the line-of-sight period is the only time during the flight when the full data set is transmitted to the ground, so it is important that the crucial diagnostics, such as bias and phase settings, sensitivity measurements, and measurement of the telescope/star camera alignment, be performed during this period. Additionally, since there is no guarantee that the data will be recoverable, it is important that a useful set of observations be scheduled during this time.

3.3.1 Bolometer Biases and Phases

The proper bias and lock-in phase settings must be determined in flight, as they depend on the in-flight loading conditions. The bias card has a \textit{ramp} setting which cycles the bias voltage from 0 to 300 mV periodically. This allows for the measurement of the sensitivity, as described in Sec. 2.1.3. An IDL routine was written to quickly process the load-curve data and determine the optimal bias voltages. A graphical interface allows the user to easily select which detectors to process, select the time ranges over which to calculate the curve, and visualize the resulting responsivity curves. The bias voltage corresponding to maximum responsivity is calculated for each detector. A single bias voltage must be applied to all bolometers of a given wavelength, however, so the median bias voltage corresponding to maximum responsivity is reported for each wavelength. The load curves and responsivities for the 350 \(\mu\text{m}\) bolometers taken during the 2005 flight are shown in Fig. 3.2. The distribution of bias voltages corresponding to peak responsivity of each bolometer is shown in Fig. 3.3. The responsivity curves are calculated from the raw bias and bolometer voltages as described in Secs. 2.1.3 and 2.1.4.

Since there is significant capacitance in the bolometer readout circuit (Sec. 2.1.4), a phase shift in the output signal is introduced. In order to get the maximum signal from the lock-in amplifier, this phase shift must be corrected. This is done by manually adjusting the phase of
Figure 3.2: The 350 μm load curves (top) and responsivities (bottom) taken at the beginning of the 2005 flight. The peak responsivity of each detector is marked with a filled circle. A histogram of the bias voltages corresponding to the maximum responsivities is shown in Fig. 3.3.
Figure 3.3: The distribution of bias voltages corresponding to maximum responsivity of each of the 350 µm detectors, as measured at the beginning of the 2005 flight. The dashed line indicates the mean of the distribution, the dash-dot line the median, and the dotted lines the standard deviation about the mean. In retrospect, a more useful diagnostic would have been to plot the distribution of sensitivities for several bias values.
the lock-in reference signal, while pulsing calibration lamp, and looking for maximum response. The lock-in phase is specified card-by-card, with 24 detector channels per card.

3.3.2 Pointing

Pointing measurements are performed regularly throughout the flight. This involves scanning a known bright point source. The map need only be deep enough to ensure that the source is well-detected and that the map is well-sampled.

The initial pointing measurement at the beginning of the flight provides the measurement of difference in the bore-sight pointings of the telescope and star cameras. A naive map (described in Sec. 4.4) can be quickly made and the position of the source read off from the map. The difference in the inferred location of the source from the known location, as determined from the literature, is the offset between the telescope and the star cameras. This value is used to correct the pointing of the star cameras, and all requested sky positions are adjusted by this offset.

Pointing measurements throughout the flight allow for the measurement of changes in this offset. This offset can change as a function of elevation due to flexing of the inner frame, although the amplitude of this effect is not known. Such variation is explored in post-flight analysis.

3.3.3 Flux Calibration

Flux calibrations are also performed throughout the flight. Flux calibration sources can be the same as those used for pointing calibration, with the additional requirement that the object’s submillimetre flux is well known, or in the absence of that, well-known infrared and millimetre fluxes which can be interpolated into the submillimetre.

The first calibration measurement at the beginning of the flight is used to estimate the detector sensitivity. We calculate the expected submillimetre flux density $S_b$ at each wavelength $b$, based on published measurements of the source. We scan the source multiple times, and, for each bolometer, take the largest peak signal $V$ (in Volts) as the measurement. The white (high-frequency) part of the noise-power spectrum is used to calculate the bolometer noise $n$, measured in V Hz$^{-1/2}$. The bolometer sensitivity is then $nS_b/V$, measured in Jy s$^{1/2}$. 
3.4 Test Flight (2003)

The BLAST science goals require multi-day observations, and hence a long-duration balloon (LDB) flight. Before this can be done, however, CSBF requires a successful North-American test flight. The primary purpose of this flight is to ensure proper operation of the instrument. Such a flight also provides an opportunity for scientific operations, however, so a secondary goal is to accomplish useful scientific observations.

In order to accomplish this secondary goal, it is desired that the flight be as long as possible. The primary limitation on the length of a North American flight is how far the gondola drifts. This clearly depends on stratospheric wind speeds, and thus we would like to fly when the wind speeds are lowest. This occurs during a period known as *turnaround*, occurring in the Spring and Fall, when the stratospheric winds change direction from west-to-east to east-to-west, and vice versa.

The BLAST test flight occurred in September 2003, launching from the CSBF balloon base in Fort Sumner, New Mexico, near the eastern border of the state. The launch occurred at 9:15 AM on September 28 and the flight was terminated at 11:57 AM the following day in the northwest corner of the state, 300 km away from the launch site.

The instrument was recovered with minimal damage. The gondola landed upright and suffered only a bent leg and minor damage to the GPS supports.

3.4.1 Flight Hardware

Due to delays in the bolometer array manufacturing process, only the 500 μm array was available for the test flight. Additionally, the array provided had been fabricated with thermal conductivity designed for observation from space with SPIRE, and thus would be of reduced sensitivity for the BLAST loading conditions. It was decided that, due to the missing wavelength coverage and the reduced sensitivity of the 500 μm detectors, the team would not risk damaging the carbon fibre mirror during the test flight. A 1.85-m parabolic aluminum primary was therefore fabricated for use during the flight.

Additional differences in the instrument compared to what is described in Chapter 2 include: (i) only one star camera was used, as concern about redundancy in this component did not come until later; (ii) no solar panels were used, as the flight was short enough to run on batteries;
and (iii) lithium batteries were used instead of rechargeable batteries, since there was no power source to charge them with. Additionally, the inner frame telescope baffle shown in Fig. 2.8 was replaced with a smaller styrofoam baffle to save total weight.

3.4.2 Performance

The instrument mostly performed well during the test flight, except that the star camera failed. We were thus unable to make maps of the sky. Instead, we concentrated on very bright point sources (e.g. Mars and Saturn) with which maps could be made based on the instantaneous detection in the bolometer signals. We were in fact unable to make maps, but the bolometer sensitivity was estimated by looking at the peak signal in observations of Mars and measuring the noise level in signal-free regions of the scan. We measured a peak signal of $s = 6.5 \times 10^{-3} \, \text{V}$ and a noise level of $n = 7 \times 10^{-6} \, \text{V}$ in one 100 Hz sample. The area-integrated flux density of Mars, a point source for BLAST, was 25 MJy at 500 $\mu$m, leading to a 500 $\mu$m NEFD of $\sim 2.5 \, \text{Jy} \, \text{s}^{1/2}$. This is an order of magnitude lower than the design value, but lower sensitivity was expected due to non-optimal bolometer parameters and a loss of sensitivity due to mis-focus of the submillimetre telescope.

3.5 First Science Flight (2005)

BLAST performed its first science flight in June 2005 from the Esrange rocket facility in Kiruna, Sweden, located north of the Arctic Circle. BLAST launched from Esrange at 1:10 UTC on June 12$^{th}$ and landed on Victoria Island in northern Canada at 6:15 UTC on June 16$^{th}$, 5 days later. The instrument was recovered in reasonably good condition, but with significantly more damage than during the test flight. The primary and secondary mirrors and the front sunshield panels were all damaged beyond repair.

3.5.1 Flight Hardware

Improvements to the instrument included: (i) a second star camera; (ii) full detector arrays, with correct thermal characteristics; and (iii) the addition of solar panels and rechargeable batteries. Additionally, extra care was taken in the preflight-measurement of the telescope focus and alignment.
Figure 3.4: A baseline-subtracted 250 μm bolometer time-stream taken during a scan of Saturn in BLAST’s first science flight, showing complex structure in the beam shape. The frames along the horizontal axis are at a rate of 5 Hz, thus the plot shows ~2.5 s of data. The expected beam shape is a Gaussian, with width approximately equal to the width of the central peak. Note that the signal is negative, since the bolometer resistance (and thus voltage) decreases with increasing input power.

3.5.2 Performance

The instrument mostly performed well, except that the beam-shape was significantly larger than designed. We first noticed this during scans of Saturn which showed complex structure (see Fig. 3.4). We concluded that the telescope was severely out of focus or that the primary mirror was damaged. The pointing system worked well, with the exception of the DGPS and the Sun sensor. The DGPS returned position and time information, but not orientation. It is thought that the mounting points for the antenna were not sufficiently rigid. The star cameras and gyroscopes functioned properly throughout the flight, however, so the failures did not affect the flight observations in any way. Ambient temperatures experienced during the flight are shown in Fig. 3.5.
Figure 3.5: Temperatures of various components (top panel) and gondola altitude (bottom panel) during the 2005 science flight. Curves are: (a) the pressurized vessel containing the flight computers; (b) inner frame; (c) lock motor; and (d) back of the primary mirror. We see rapid cooling during ascent, then a strong diurnal variation of 10–20°C. Diurnal variations in altitude, on the order of 2 km, are also visible.
3.5.3 Observations

Because of the large beams, the instrument’s resolution and point source sensitivity were significantly reduced. We decided, based on the measurements of Saturn, that the sensitivity was too low for useful extragalactic studies, and therefore concentrated on Galactic targets for the majority of the flight. This level of failure had not been anticipated, and the flight schedule could not accommodate such a change, so the observation commands had to be sent by hand from the ground station throughout the flight.

We chose to observe: (i) several point sources, both Galactic and extragalactic, some of which were targeted for calibration purposes; (ii) large maps of several fields in the Galactic plane and other specific Galactic features, including star forming regions and cold molecular clouds; (iii) a supernova remnant; and (iv) Solar System objects, including Saturn and the asteroid Pallas. The observations are summarized in Table 3.1.

Analysis of bright compact sources, specifically Pallas, NGC 4565, Mrk 231, Arp 220, K3-50, W 75N, IRAS 20126+4104, CRL 2688, LDN 1014, IRAS 22134+5834 and IRAS 23011+6126, is presented in Truch et al. (2007). The nearby edge-on Sb spiral galaxy NGC 4565 is discussed in further detail in Chapter 6 of this document. Reduction of the 4 square degree Galactic plane survey towards the constellation Vulpecula is described in Chapter 5 and in Chapin et al. (2007). Analysis of the other Galactic plane fields is underway and will be published at a later date. A paper discussing observations of the supernova remnant Cas A is in preparation.
Table 3.1: Summary of BLAST 2005 Observations

<table>
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<tr>
<th>Field</th>
<th>R.A.</th>
<th>Dec.</th>
<th>Size $^a$</th>
<th>Time $^b$</th>
<th>Mode</th>
<th>Object</th>
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<td>+25:59</td>
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<td>0.8</td>
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<td>Edge-on spiral galaxy</td>
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<td>Mrk 213</td>
<td>12:56</td>
<td>+56:52</td>
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<td>0.5</td>
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<td>AGN/ULIRG</td>
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<tr>
<td>IVCG86</td>
<td>14:41</td>
<td>+49:05</td>
<td>1.50</td>
<td>6.3</td>
<td>cap</td>
<td>Intermediate velocity cloud</td>
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<tr>
<td>Arp 220$^c$</td>
<td>15:35</td>
<td>+23:30</td>
<td>0.12</td>
<td>6.1</td>
<td>box</td>
<td>ULIRG</td>
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<td>ELAIS N1</td>
<td>16:15</td>
<td>+55:52</td>
<td>0.25</td>
<td>8.8</td>
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<td>Extragalactic blank field</td>
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<td>0.25</td>
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<td>High velocity cloud</td>
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<td>16:45</td>
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<td>1.50</td>
<td>3.9</td>
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<td>19:16</td>
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<td>6.1</td>
<td>quad</td>
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<td>6.5</td>
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<td>7.9</td>
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<td>Planetary nebula</td>
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<td>1.0</td>
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<td>1.3</td>
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<td>Bright high-mass protostar</td>
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<td>+47:34</td>
<td>0.40</td>
<td>1.5</td>
<td>quad</td>
<td>Low-mass star forming region</td>
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<td>+58:49</td>
<td>1.00</td>
<td>2.8</td>
<td>box</td>
<td>Medium/high-mass protostar</td>
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Continued on next page
### Table 3.1 – continued from previous page

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<td>(dd:mm)</td>
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<td>(hrs)</td>
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<td>0.1</td>
<td>box</td>
<td>Planet</td>
</tr>
</tbody>
</table>

\(a\) Requested map size is listed. Actual observed area will typically be larger.

\(b\) Requested observation time is listed. Actual integration time may be shorter.

\(c\) Flux calibration target
3.6 Second Science Flight (2006)

After recovering the BLAST instrument from Northern Canada after the 2005 flight, the BLAST team set about repairing the instrument for another flight, so that the extragalactic science goals could be accomplished. Since the carbon fibre primary was destroyed, the aluminum primary used in the 2003 test flight, which had a surface finish suitable for only the 500 μm array, was re-surfaced for use at 250 μm. All other repairs and upgrades were minor. The second BLAST science flight occurred in December 2006 from McMurdo Station in Antarctica. The telescope was launched at 1:54 UTC December 21st and the flight was terminated at 1:05 UTC January 2nd, after approximately 270 hours of observing time. The instrument landed 750 km southwest of McMurdo, but a failure in the parachute release mechanism caused it to be dragged for ~200 km. The gondola finally came to rest in a crevasse field with only the bare base frame still attached to the parachute. The rest of the instrument, including the sun shields, inner frame, cryostat, telescope and data vessel, were scattered along the drag path. The data vessel, cryostat and telescope were eventually recovered.

3.6.1 Flight Hardware

The 2 m aluminum primary mirror built for the test flight was resurfaced for the 2006 science flight. Due to the large diurnal temperature variations seen in the 2005 flight (Fig. 3.5), it was decided that the telescope should be focusable by command. The distance between the primary and secondary should be correct to within 1 wavelength; it was calculated that thermal contraction of the mirrors could cause the separation between the primary and secondary to change as much as 50 μm°C⁻¹, and thus change as much as 500 μm throughout the day. A motorized system for in-flight refocusing of the secondary was implemented. The secondary is coupled to its mounting structure via stainless steel leaf springs and its position is controlled by 3 stepper motor actuators, allowing for both translation along the optic axis and tip/tilt control.

A system for pre-flight focusing of the optics was also implemented. A satisfactory method of focusing the telescope and aligning the full optical system was never achieved in previous flights, due to the high emissivity of the atmosphere and the large distance of the far field limit (~10 km). It was realized that the telescope could be focused at a finite nearby distance
by increasing the primary-secondary distance. Spacing blocks were manufactured to offset the secondary by the correct amounts in order to focus the telescope at 50, 100 and 150 m. A cold source with a chopping shutter was mounted to a commandable stage with two degrees of freedom and placed at the near focus. The source was rastered across the image plane and beam maps were created. The focus of the telescope was adjusted by controlling the secondary and it was verified that correct focus could be achieved.

Additional changes to the instrument from the 2005 flight include upgraded sun shields, a new sun sensor and a replacement lock motor. Extensions to the upper part of the sun shields were designed and built, providing shading for the secondary mirror at azimuths other than directly anti-sun. A new sun sensor consisting of an array of photo-diodes replaced the orthogonal CCD sun sensor. The new sensor consumes significantly less power and allowed us to do away with the outer-frame cooling system.

3.6.2 Performance

The instrument performed well during the 12-day flight. The optics were in focus, providing the nominal diffraction-limited beam sizes. In-flight measurements of the sensitivity at each wavelength show NEFDs close to that expected. We measured 240, 239 at 157 mJy s$^{1/2}$ and 250, 350 and 500 µm, marginally worse at 250 and 350 µm and slightly better at 500 µm than nominal (see Table 1.2). The pointing system worked well, except for two isolated incidents: an anomaly with the gyroscopes and an automatic shutdown of the pivot motor controller due to overheating. The gyroscopes suffered several failures throughout the flight. The failures seemed to be short-lived, however, and we were always able to switch to the other set whenever a failure happened. About 2 hours were lost due to these failures. In certain telescope orientations and sun elevations, the new sun shields reflected sunlight onto the back surface of the pivot motor controller which was not painted. The temperature controller exceeded 62.5°C twice during the flight, at which point it automatically shut down and the system was not able to maintain pointing. Very little time was lost due to these shutdowns. The DGPS once again did not return gondola orientation, but the pointing system worked fine without it.

We have analysed a short section of pointing data during the 2006 flight to demonstrate the effectiveness of the pointing servo. In Fig. 3.6 we show a 12 minute section of data taken during a scan. The third panel shows the signal measured by the azimuthal gyroscope. The
Figure 3.6: A short section of pointing data from the 2006 flight, taken during a scan. The bottom plot shows the difference between the measured and requested azimuthal velocity. A sudden event in the pivot is not seen in the azimuthal gyroscope, demonstrating the effectiveness of the pointing servo in overcoming the high-stiction pivot motor. [This figure appears in Pascale et al. (2007).]
scan consists of constant velocity motion with constant acceleration at the turn-arounds. The requested azimuthal velocity is overplotted and is buried by the noisy signal. The difference between the measured and requested azimuthal velocity is shown in the fourth panel. Small spikes are visible in the difference at the corners in the scan speed. The second panel shows the speed of the reaction wheel, which drives the telescope in azimuth. We see that the period-averaged reaction wheel speed increases towards the 6 min mark due to an external torque on the balloon/gondola system. The top plot shows pivot position as measured by the pivot encoder. The requested pivot motor torque is proportional to reaction wheel speed, and at the 6 min mark the requested torque exceeds stiction in the pivot. The fact that this sudden event is not seen in the gyroscopes is proof of the excellent performance of the pointing servo. During the entire 2006 flight, except during failures of the pivot motor controller, the reaction wheel stayed between 0 and 0.3 revolutions per second.

3.6.3 Observations

The 2006 flight lasted $\sim 270$ hr, 94% of which was spent performing scientifically useful observations. Extragalactic science is the primary driver for BLAST, so observations during the flight were concentrated on extragalactic fields. Extragalactic observations comprised 69% of the flight, with 18% on Galactic fields, 7% on calibration and the remaining 6% wasted due to instrumental problems and operator error. The observations are summarized in Table 3.2.

A total of 160 hr was spent on extragalactic blank-field surveys. Two regions of the sky were imaged: the Chandra Deep Field South (CDFS) and the Akari South Ecliptic Pole (SEP). The CDFS has been observed extensively, including very deep observations in the IR, X-ray, optical and radio. A region of 0.6 square degrees was mapped to $5\, \text{mJy beam}^{-1}$, the expected confusion limit at 250 $\mu$m, and a large region of 9 square degrees was mapped to $15\, \text{mJy beam}^{-1}$. A similarly-sized field at the SEP was also observed. This region coincides with multi-band IR measurements by the Akari space telescope. In addition to the blank-field surveys, several biased surveys towards known high-density regions and several nearby resolved galaxies were observed.

The majority of the time allocated to Galactic observations were spent observing the Vela Molecular Cloud in the Galactic plane. This region includes a number of cold molecular clouds with the potential to form stars as well as sites of recent star formation containing massive
proto-stars and H II regions. A 40 square degree region was observed for 40 hours, reaching a depth of 70 mJy beam$^{-1}$, and a larger 150 square degree region encompassing the deep field was mapped to a depth of 210 mJy beam$^{-1}$. BLAST does not constrain large-scale structure in the wide map due to the length of the scans, but it is hoped that FIRAS data will be able to restore the low spatial frequencies that are unconstrained by BLAST. Small maps were also performed in the Gum Nebula, a region with known star formation, of the supernova remnant Pup A, and η Carina, a site of massive star formation.

Analysis of the 2006 data is underway, with the intention of publishing the first extragalactic maps in the fall of 2007.
Table 3.2: Summary of BLAST 2006 Observations

<table>
<thead>
<tr>
<th>Field</th>
<th>R.A.</th>
<th>Dec.</th>
<th>Size</th>
<th>Time</th>
<th>Mode</th>
<th>Object</th>
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<td>(dd:mm)</td>
<td>(deg²)</td>
<td>(hr)</td>
<td></td>
<td></td>
</tr>
<tr>
<td>NGC 1097</td>
<td>2:47</td>
<td>-30:13</td>
<td>0.07</td>
<td>0.9</td>
<td>cap</td>
<td>Resolved nearby galaxy</td>
</tr>
<tr>
<td>NGC 1291</td>
<td>3:18</td>
<td>-41:05</td>
<td>0.11</td>
<td>0.7</td>
<td>cap</td>
<td>Resolved nearby galaxy</td>
</tr>
<tr>
<td>Abel 3112</td>
<td>3:18</td>
<td>-44:13</td>
<td>0.12</td>
<td>4.2</td>
<td>cap</td>
<td>Biased survey, $z = 0.075$ cluster</td>
</tr>
<tr>
<td>CDFS W</td>
<td>3:32</td>
<td>-28:08</td>
<td>9.30</td>
<td>52.1</td>
<td>quad</td>
<td>Extra-galactic blank-field (wide)</td>
</tr>
<tr>
<td>CDFS D</td>
<td>3:33</td>
<td>-27:47</td>
<td>0.59</td>
<td>37.8</td>
<td>cap</td>
<td>Extra-galactic blank-field (deep)</td>
</tr>
<tr>
<td>NGC 1365</td>
<td>3:34</td>
<td>-36:07</td>
<td>0.05</td>
<td>0.6</td>
<td>cap</td>
<td>Resolved nearby galaxy</td>
</tr>
<tr>
<td>NGC 1512</td>
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<td>0.6</td>
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<td>Resolved nearby galaxy</td>
</tr>
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<td>1.0</td>
<td>cap</td>
<td>Resolved nearby galaxy</td>
</tr>
<tr>
<td>Akari SEP</td>
<td>4:44</td>
<td>-53:31</td>
<td>7.60</td>
<td>68.7</td>
<td>quad</td>
<td>Extra-galactic blank-field</td>
</tr>
<tr>
<td>NGC 1808</td>
<td>5:08</td>
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<td>0.05</td>
<td>0.6</td>
<td>cap</td>
<td>Resolved nearby galaxy</td>
</tr>
<tr>
<td>PKS 0529</td>
<td>5:31</td>
<td>-54:54</td>
<td>0.12</td>
<td>3.4</td>
<td>cap</td>
<td>Biased survey, $z = 2.575$ radio galaxy (deep)</td>
</tr>
<tr>
<td>PKS 0529 W</td>
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<td>-54:54</td>
<td>0.26</td>
<td>0.5</td>
<td>cap</td>
<td>Biased survey, $z = 2.575$ radio galaxy (wide)</td>
</tr>
<tr>
<td>Bullet</td>
<td>6:59</td>
<td>-55:58</td>
<td>0.12</td>
<td>2.4</td>
<td>cap</td>
<td>Biased survey, $z = 0.296$ cluster</td>
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<tr>
<td>VYCMa</td>
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<td>-25:47</td>
<td>0.09</td>
<td>6.2</td>
<td>box</td>
<td>Calibrator</td>
</tr>
<tr>
<td>GumGlob</td>
<td>8:07</td>
<td>-35:58</td>
<td>0.25</td>
<td>0.3</td>
<td>quad</td>
<td>Cometary Globules near Gum Nebula</td>
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<tr>
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<td>-36:10</td>
<td>0.09</td>
<td>0.3</td>
<td>box</td>
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<td>-42:09</td>
<td>0.09</td>
<td>0.1</td>
<td>box</td>
<td>Calibrator</td>
</tr>
<tr>
<td>Pup A</td>
<td>8:24</td>
<td>-43:01</td>
<td>3.15</td>
<td>1.2</td>
<td>cap</td>
<td>Supernova Remnant</td>
</tr>
<tr>
<td>IRAS 08247-4223</td>
<td>8:26</td>
<td>-42:34</td>
<td>0.09</td>
<td>0.1</td>
<td>box</td>
<td>Calibrator</td>
</tr>
<tr>
<td>IRAS 08307-4303</td>
<td>8:33</td>
<td>-43:15</td>
<td>0.09</td>
<td>0.1</td>
<td>box</td>
<td>Calibrator</td>
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Continued on next page
Table 3.2 – continued from previous page

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<td>(dd:mm)</td>
<td>(deg(^2))</td>
<td>(hrs)</td>
<td></td>
<td></td>
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<td>−40:40</td>
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<td>0.3</td>
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</tr>
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<td>−42:56</td>
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<td>6.9</td>
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<td>Calibrator</td>
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<td>0.1</td>
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<td>−44:32</td>
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<td>1.1</td>
<td>box</td>
<td>Calibrator</td>
</tr>
<tr>
<td>Vela D</td>
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<td>−45:04</td>
<td>42.80</td>
<td>40.4</td>
<td>quad</td>
<td>Survey of several star-forming regions (deep)</td>
</tr>
<tr>
<td>IRAS 09002-4732</td>
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<td>−47:46</td>
<td>0.09</td>
<td>0.1</td>
<td>box</td>
<td>Calibrator</td>
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<tr>
<td>IRAS 09014-4736</td>
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<td>−47:50</td>
<td>0.09</td>
<td>0.1</td>
<td>box</td>
<td>Calibrator</td>
</tr>
<tr>
<td>IRAS 09018-4816</td>
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<td>−48:30</td>
<td>0.09</td>
<td>0.7</td>
<td>box</td>
<td>Calibrator</td>
</tr>
<tr>
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<td>−47:58</td>
<td>0.09</td>
<td>0.1</td>
<td>box</td>
<td>Calibrator</td>
</tr>
<tr>
<td>Vela W</td>
<td>9:18</td>
<td>−47:42</td>
<td>154.00</td>
<td>10.6</td>
<td>cap</td>
<td>Survey of several star-forming regions (wide)</td>
</tr>
<tr>
<td>G284.35−0.42</td>
<td>10:24</td>
<td>−57:55</td>
<td>0.09</td>
<td>0.6</td>
<td>box</td>
<td>Calibrator</td>
</tr>
<tr>
<td>Eta Carina</td>
<td>10:45</td>
<td>−59:41</td>
<td>3.15</td>
<td>2.0</td>
<td>cap</td>
<td>Star-forming Region</td>
</tr>
<tr>
<td>PKS 1138</td>
<td>11:41</td>
<td>−26:31</td>
<td>0.12</td>
<td>1.1</td>
<td>cap</td>
<td>Biased survey (z = 2.156) radio galaxy</td>
</tr>
</tbody>
</table>

\(^a\)Requested map size is listed. Actual observed area will be larger.

\(^b\)Requested observation time is listed. Actual integration time may be shorter.
Chapter 4

Data Processing

In this chapter we describe the process of creating a calibrated map at any one of the 3 frequencies, given the raw data produced by the telescope. A specialized map-maker, described in Sec. 4.4, has been developed which addresses the particular peculiarities of BLAST-style instruments, namely large maps made from time-streams with strong common-mode signals between detectors. Before the map-maker can be applied, however, the bolometer data must be cleaned and the pointing solution, the telescope’s position on the sky as a function of time, must be calculated. Much of this chapter is described in greater detail in Truch et al. (2007, hereafter TRU07) and Patanchon et al. (2007, hereafter PAT07).

As with the design and assembly of the instrument, the data processing and analysis is a team effort. The entire data analysis procedure is described here and in the next chapters, but we point out that the work by the author was primarily focused on the pointing solution (Sec. 4.3), preliminary analysis of the Galactic plane field (Secs. 5.2 5.6.2), including image deconvolution and source extraction, and the deconvolution (Sec. 6.1) and profile-fitting (Sec. 6.2) of the edge-on spiral galaxy NGC 4565.

4.1 Data Cleaning

Before the raw bolometer data can be transformed into maps, several non-astronomical signals must be removed. Spikes in the time-streams due to cosmic rays and other anomalies must be detected and masked out. Additionally, the data are deconvolved to remove the effect of a low-pass filter applied during data collection.
4.1.1 Spike Removal and Filter Deconvolution

Several types of glitches are removed from the time-streams. We distinguish between single-sample spikes caused by the data acquisition system and broader features resulting from non-astronomical signals, such as calibration lamp pulses and cosmic ray hits. The order of data cleaning is to first remove the single-sample spikes, deconvolve the time-streams to remove the signature of the anti-aliasing filter (Sec. 2.6.1), and then flag and remove the broader time-scale glitches.

Single-sample spikes, due to glitches in the data acquisition system downstream from the filter stage, are first removed in order to avoid ringing due to the deconvolution. Single-point excursions of $> 8 \sigma$, where $\sigma$ is an estimate of the time-stream RMS, are flagged and replaced with the mean value of the samples on either side of the spike. We replace $\sim 2000$ single-point spikes in each hour of each detector data stream.

The data are deconvolved to remove the signature of the low-pass filter (see Sec. 2.6.1). This filter has a cut-off frequency of $\sim 35$ Hz and is designed to avoid high-frequency noise aliasing. The deconvolution is performed in Fourier space.

The calibration lamp (cal-lamp) pulses, used to track responsivity throughout the flight (see Sec. 4.2.1), are removed from the time streams. The lamp is pulsed for a duration of $\sim 150$ ms once every 15 min during the flight. There is no need to automatically detect the cal-pulses in the data, as their location in the data streams are indicated in a flag field. The time intervals corrupted by the cal-lamp pulses are flagged and filled with approximately constrained noise. These regions are $\sim 150$ ms in length, much smaller than the scale corresponding to the correlated-signal $1/f$ knee of the noise power spectrum ($\sim 10$ s, Sec. 2.1.5). White noise equal to the high frequency level in the power spectrum is thus a reasonable model, and we add to that a line fit to 20 samples (200 ms) on either side of the gap.

Cosmic rays are sometimes absorbed by a bolometer. The energy of the ray heats the bolometer and the event is seen in the time-stream. The signal falls off exponentially with time scale equal to the detector time constant, $\tau \sim 2$ ms (in fact, cosmic ray hits are used to measure the detector time constant). A 10-\(\sigma\) event thus falls of to the noise level in $\sim 2.3 \tau$, shorter than the sample spacing, 5 ms. Events are detected by searching for brief excursions from the noise level in each detector signal. To avoid over-detecting spikes on top of true bright signals, this operation is performed on signal-subtracted maps. This is an iterative process, with a new naive
Figure 4.1: The distribution of cosmic ray hits in a single 250 μm bolometer throughout the flight. 3 events with amplitude 10–15 mV fall outside of the range of the plot.

map (see Sec. 4.4) made at each stage. The pointing matrix \( P \) is used to create the signal-only time stream from the map \( m \). We start with a single-sample spike removed, deconvolved and cal-lamp pulse removed time-stream \( t \), and with an estimate of the sky being \( m_0 \) (all zeros). At each iteration, we perform the following steps:

1. calculate an anticipated signal-only time-stream \( s_i = P m_{i-1} \)
2. calculate residual time-stream \( r_i = t - s_i \)
3. detect spikes in \( r_i \)
4. form cleaned time-stream \( c_i \) from \( t \) by removing 500 ms interval at each spike
5. make map \( m_i \) from high-pass filtered \( c_i \) using naive map-maker

This is repeated until \( c \) has reasonably converged, usually after a few iterations. We detect \( \sim 100 \) cosmic ray hits per hour in a single 250 μm detector. The distribution of hits as a function of peak signal is shown in Fig. 4.1. In all, about 2% of the data are removed due to single-point glitches, cal-lamp pulses and cosmic ray spikes.
4.2 Calibration

The deconvolved and despiked bolometer signals are calibrated in 3 steps. First, the time-varying responsivity in each bolometer is corrected using measurements of the calibration lamp throughout the flight. Then, the global responsivity of each bolometer is compared to a reference bolometer (at the same wavelength) by measuring the response to an astronomical point source. Finally, the overall absolute calibration is calculated by fitting the model spectral energy distribution of a well-studied astronomical object to the BLAST measurement. These last two steps could be performed at the same time, but they are kept separate because the latter is model-dependent, while the former is not.

4.2.1 Time-varying Calibration

As described in Secs. 2.2 and 4.1.1, a calibration lamp located in the cryostat is used to track bolometer responsivity throughout the flight. The lamp is pulsed periodically, providing a repeatable step in load on the detectors; any change in the response of the detectors to the cal-lamp is a direct measure of change in the bolometer responsivity. The responses of six 250 μm detectors to a cal-pulse event are shown in Fig. 4.2.

The amplitude of each pulse in each bolometer time-stream is measured by subtracting a baseline from either side of the pulse and fitting a template profile to the pulse. The amplitudes of the pulses are interpolated in time and inverted to provide a multiplicative time-varying calibration for each bolometer time-stream. Fluctuations in responsivity of ~8% RMS are found over the flight, dominated by diurnal variations due to atmospheric loading on the detectors.

4.2.2 Flat-fielding

The relative calibration from one detector to another, known to optical astronomers as flat-fielding, is performed using measurements of flux calibrators. The point source CRL 2688 was observed using an exceptionally slow elevation scan speed so that fully-sampled maps could be made from single-bolometer data. The naive map-maker, described in Sec. 4.4 is used to make maps for individual detectors. The flux of the source in each map, measured in Volts integrated over the point spread function (PSF), is calculated using PSF photometry. This involves fitting the best estimate of the PSF to the map. We fit for the single parameter $S_\nu$, the flux density
Figure 4.2: The responses of 6 different 250 μm detectors to a single cal-pulse event. These events occur every 15 minutes throughout the flight and are used to track variation in bolometer responsivity.
of the point source, which is given by weighted least squares fitting:

\[ S_\nu = \frac{\sum_{x,y} [M(x,y)P(x,y)/\sigma^2(x,y)]}{\sum_{x,y} [P^2(x,y)/\sigma^2(x,y)]}, \]  

(4.1)

where \((x,y)\) are the 2 spatial directions, \(M(x,y)\) is the map of the calibrator, \(P(x,y)\) is the PSF and \(\sigma^2(x,y)\) is the noise variance in pixel \((x,y)\). We estimate \(P\) by averaging together maps of all point sources measured throughout the flight by all bolometers at each wavelength. The difference in the inferred flux density \(S_\nu\), measured for each detector, gives the bolometer-bolometer calibration coefficient. At each wavelength, we chose a reference bolometer and adjust the time-streams of all others by the appropriate scalar value (multiply by \(S_{ref}/S_i\)).

### 4.2.3 Absolute Calibration

The absolute calibration uses the same integrated flux densities measured in the previous section, after correction for the flat-fielding. The measured band flux \(\hat{S}_B\), in unknown units, is related to the true band flux density \(S_B\) by a single calibration coefficient \(\alpha\). The true \(S_B\) is the integral of the source spectral energy distribution (SED), \(S_\nu\), over the normalized BLAST band transmission function \(T_\nu\) (see Sec. 2.3.3).

\[ S_B = \int T_\nu S_\nu d\nu. \]  

(4.2)

Given the SED, \(S_B\) and thus \(\alpha\) can be determined. Detailed SEDs for submillimetre-bright astronomical objects are not well known, however, and a model must be assumed. The model is constrained by measurements of the source by other telescopes at infrared to millimetre wavelengths.

Since the calibration is only a single number, one only need use measurements of a single source. We use Arp 220, a bright Ultra-luminous Infrared Galaxy (ULIRG) at redshift \(z = 0.018\) and with a luminosity of \(\sim 2 \times 10^{12} L_\odot\). It is located well out of the Galactic plane at a Galactic latitude of \(+53^\circ\), so the field is relatively free of foreground contamination. We fit the commonly-used modified blackbody model,

\[ S_\nu = A\nu^\beta B_\nu(T), \]  

(4.3)

with \(B_\nu(T)\) the Planck function, to measurements in the FIR and submillimetre. Ancillary data from UKT14 at 350–1100 \(\mu m\) (Rigopoulou et al. 1996), SCUBA at 450 (Dunne and Eales 2001) and 850 \(\mu m\) (Lisenfeld et al. 2000; Dunne and Eales 2001; Seiffert et al. 2007), SHARC-II\(^1\) at

\[ \text{http://www.submm.caltech.edu/~sharc/analysis/calibration.htm} \]
350 μm, ISOPHOT at 60–200 μm (Klaas et al. 2001) and 12–200 μm (Spinoglio et al. 2002), and IRAS at 60 and 100 μm (Sanders et al. 2003) are used in the fit. Errors in the measurements are taken from the published values, typically 10%. The ISOPHOT data are considered to be 100% correlated and other measurements reported by a single group are assumed to be 10% correlated.

The fit is performed by $\chi^2$ minimization to all ancillary data, accounting for measurement and correlated errors (see Fig. 4.3). Errors in the parameters and in the model are calculated by Monte Carlo simulations of noise realizations described by the measurement and the correlated errors, and refitting the SED to each mock data set.

BLAST measurements of Arp 220 are determined from the maps by PSF photometry, as described above, providing an uncalibrated measurement $V$. The best-fit SED is integrated over the band transmission functions (Eqn. 4.2) and the calibration coefficients $\alpha = V/S_B$ are calculated and applied to all subsequent measurements. Errors in $\alpha$, termed calibration errors, are calculated using the Monte Carlo simulations. The calculation of the $\alpha$ is repeated for each mock SED and the variance of the $\alpha$ over the full set determines the error. Correlations between different bands are also measured, producing a full covariance matrix for calibration errors that can be used in later analysis. We find calibration errors of 12, 10 and 8% at 250, 350 and 500 μm. The Pearson correlation matrix showing the correlation of calibration errors between bands is listed in Table 4.1.

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<td>350 μm</td>
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<td>0.96</td>
<td></td>
</tr>
<tr>
<td>500 μm</td>
<td></td>
<td>1</td>
<td></td>
</tr>
</tbody>
</table>

### 4.3 Pointing Solution

The pointing solution is a critical part of the data pre-processing. An in-flight pointing solution is calculated by the Master Control Program for purposes of pointing control, but is improved upon in post-flight analysis. The inputs to the pointing solution calculation are the gyroscope
Figure 4.3: The spectral energy distribution fit of Arp 220, used as the absolute calibration source for BLAST. The BLAST data are marked with diamonds and have been scaled to the SED model fit to previous measurements of Arp 220 in the submillimetre and infrared, which are marked with triangles. The error enveloped is computed using Monte Carlo simulations. The resulting $1\sigma$ uncertainties in the calibration coefficients are 12, 10 and 8% at 250, 350 and 500 $\mu$m respectively. [This figure appears in Truch et al. (2007).]
signals and star camera solutions.

4.3.1 Gyroscopes

As described in Sec. 2.5.2, BLAST incorporates two sets of three orthogonal single-axis optical gyroscopes. Within each set, the three gyroscopes measure angular velocities about axes nominally aligned with elevation, roll and yaw of the inner frame. The individual gyroscope signals are orthogonalized using the measurements made pre-flight (Sec. 2.5.2). The angular velocities measured by the gyroscopes are then rotated into the star camera reference frames, as they are not necessarily aligned with the line-of-sight direction defined by the inner frame. This rotation is measured post-flight by integrating the gyroscope solution through large telescope slews, using the star camera solution before and after as a measure of the true camera pointing.

4.3.2 Star Cameras

The BLAST star cameras report boresight pointing in flight when a unique solution is found, which requires at least 3 detected stars. This is improved upon in post-flight analysis in an iterative manner. On the first pass, the star camera plate scale and rotation are fit along with boresight pointing for all multi-star frames. The measurements of plate scale and rotation are noisy, so they are smoothed in time.

After this first pass, the pointing solution is calculated, as described in the next section. The pointing solution is then used to further refine the star camera solutions, allowing for the use of one- and two-star images. The star camera plate scale and rotation are fixed to the interpolated values of filtered solutions determined in the first iteration.

4.3.3 Integration

The gyroscope angular velocity and star camera absolute references are combined using a Kalman filter, an algorithm for iteratively calculating the least-squares solution of a quantity (or set of quantities) given a time-ordered set of noisy measurements. In our case, the inputs (measurements) are the gyroscope signals and star camera solutions and the outputs are the bore-site pointings. The Kalman filter is described in detail in Kamen and Su (1999, hereafter KS99). Throughout this section, we use the notation of KS99.
We use the formulation of the quaternion, described below, instead of cumbersome rotation matrices. We first describe the linear Kalman filter and then describe its non-linear extension, since the conversion from the measured angular velocities to the quaternions is not linear. Finally, we describe Kalman smoothing and then discuss the results of applying the method to the BLAST pointing data.

4.3.3.1 Quaternions

Quaternions are commonly used in systems where rapid and frequent calculations of rotations in 3 dimensions are needed, such as computer graphics and attitude control. There are several advantages for the use of quaternions as opposed to rotation matrices in such systems: (i) storage — one need store only 4 numbers instead of 9; (ii) computation — many fewer calculations are needed to perform a given rotation; (iii) numerical stability — it is much simpler to re-normalize a numerically distorted quaternion than it is to re-orthogonalize a distorted rotation matrix; and (iv) simplicity — rotations about arbitrary axes are easily encoded in the quaternion formulation.

A quaternion is a four-element extension of a complex number, defined as

\[ q = (a, b, c, d) = (a, \vec{v}) = a \cdot 1 + b \cdot i + c \cdot j + d \cdot k \]  

(4.4)

where 1, i, j and k are the orthogonal basis quaternions. Addition of two quaternions is performed element-by-element, while multiplication is distributive, following the relations

\[ i^2 = j^2 = k^2 = ijk = -1. \]  

(4.5)

From these relations, it is easily shown that the multiplication of two quaternions \( q = (a, \vec{v}) \) and \( r = (b, \vec{w}) \) is given by

\[ qr = (a, \vec{v})(b, \vec{w}) = (ab - \vec{v} \cdot \vec{w}, a\vec{w} + b\vec{v} + \vec{v} \times \vec{w}), \]  

(4.6)

where ‘×’ is the usual 3-dimensional cross-product. The inverse of a quaternion \( q^{-1} \) is defined such that \( q \cdot q^{-1} = 1 \), therefore

\[ q^{-1} = q^* / |q|^2 = (a, -\vec{v}) / |q|^2, \]  

(4.7)

where \( q^* = (a, -\vec{v}) \) is the conjugation operator. A normalized quaternion \( \hat{q} \) has inverse

\[ \hat{q}^{-1} = \hat{q}^* = (a, -\vec{v}). \]  

(4.8)
The rotation of a vector $\vec{v}$ by an angle $\alpha$ about an axis $\hat{n}$ can be encoded into quaternion form by writing
\[
(0, \vec{v'}) = Q (0, \vec{v}) Q^*,
\] (4.9)
where
\[
Q = q_{\hat{n}}(\alpha) = \left( \cos \left( \frac{\alpha}{2} \right), \hat{n} \sin \left( \frac{\alpha}{2} \right) \right).
\] (4.10)
Note that since $\hat{n}$ is a unit vector, $q_{\hat{n}}(\alpha)$ is a unit quaternion. Multiple rotations can be stacked:
\[
Q = q_3 q_2 q_1 = q_{\hat{n}_3}(\alpha_3) q_{\hat{n}_2}(\alpha_2) q_{\hat{n}_1}(\alpha_1),
\] (4.11)
and we note that $Q^* = q_1^* q_2^* q_3^*$.

The quaternion rotation formulation is used for calculation of the BLAST pointing solution.

### 4.3.3.2 The Kalman Filter

The Kalman filter provides an optimal solution for estimating the state of a system given time-ordered noisy measurements of signals describing the state. We present the filter in its general form before applying it to the BLAST pointing solution.

The state/measurement model assumes that the signal $s(n)$ is linear transformation of the state $x(n)$,
\[
s(n) = C x(n),
\] (4.12)
where $C$ is the measurement matrix and $n$ is the (discrete) time index. Additionally, the state $x$ is propagated forward in time by the relation
\[
x(n + 1) = \Phi x(n) + \Gamma w(n),
\] (4.13)
where $\Phi$ is the time-update matrix, $w(n)$ is the process noise at time $n$ and $\Gamma$ is the noise propagation matrix. A noisy measurement of the signal is made:
\[
z(n) = s(n) + v(n),
\] (4.14)
where $v(n)$ is the measurement noise. Additionally, the error covariance matrix $P(n)$, defined as
\[
P(n) = \mathbb{E}[\hat{x}(n) \hat{x}^T(n)]
\] (4.15)
is calculated, where \(T\) is the transpose operator and

\[ \tilde{x}(n) = x(n) - \hat{x}(n) \]  

(4.16)

is the error in the estimate \(\hat{x}(n)\) of the true state \(x(n)\) and \(E[y]\) is the expectation value of \(y\).

The Kalman filter updates the estimate of the state of the system in two steps: the time update and the measurement update. At the time update step, also called the \(a \ priori\) estimate, the state is updated using the time-update matrix \(\Phi\). At the measurement update or \(a \ posteriori\) estimate, the measurement \(z\) is used.

Given the \(a \ priori\) solutions \(\hat{x}^-(n)\) and \(P^-(n)\) at time \(n\), the measurement \(z(n)\) at time \(n\) is acquired and the \(a \ posteriori\) quantities are calculated:

\[
K(n) = P^-(n)C^T[CP^-(n)C^T + R]^{-1}; \quad (4.17)
\]

\[
\hat{x}(n) = \hat{x}^-(n) + K(n)[z(n) - C\hat{x}^-(n)]; \quad (4.18)
\]

\[
P(n) = P^-(n) - K(n)CP^-(n). \quad (4.19)
\]

\(K(n)\) is known as the Kalman Gain and \(R\) is the measurement noise covariance matrix describing \(v\) (Eqn. 4.14). The state is then updated to the next time step through

\[
\hat{x}^-(n+1) = \Phi \hat{x}(n) \quad (4.20)
\]

\[
P^-(n+1) = \Phi P(n)\Phi^T + \Gamma Q \Gamma^T, \quad (4.21)
\]

where \(Q\) is the process noise covariance matrix describing \(w\) (Eqn. 4.13).

The Kalman filter is ideally initialized at time \(n = 0\) with

\[
\hat{x}^-(0) = E[x(0)] \quad (4.22)
\]

\[
P^-(0) = E \left[ (x(0) - \hat{x}^-(0)) (x(0) - \hat{x}^-(0))^T \right]. \quad (4.23)
\]

Lacking this information, a guess for \(\hat{x}^-(0)\) and the assumption that \(P^-(0) = \lambda I\) for some \(\lambda > 0\) usually prove sufficient.

### 4.3.3.3 The Non-linear Kalman Filter

When the state update function \(\Phi\) and/or signal transformation function \(C\) are not linear, the Kalman procedure must be modified. For BLAST, only the time-update matrix \(\Phi\) is non-linear, so we here ignore cases where \(C\) is also non-linear.
Eqn. 4.13 of the state/measurement model is rewritten as

\[ x(n + 1) = \phi(x(n)) + \Gamma w(n), \]  

where \( \Phi \) is replaced by \( \phi \), a non-linear function of \( x(n) \). It is assumed that \( \phi \) is reasonably smooth and is expanded about \( \hat{x}(n) \):

\[ \phi(x(n)) = \phi(\hat{x}(n)) + J_\phi(\hat{x}(n))[x(n) - \hat{x}(n)] + \ldots, \]

where \( J_\phi(x) \) is the Jacobian of \( \phi \) evaluated at \( x \):

\[ (J_\phi(x))_{ij} = \frac{\partial \phi_i}{\partial x_j}. \]

Since \( C \) has not changed, the measurement update (Eqns. 4.17-4.19) remains the same, but the time update is modified as follows:

\[ \begin{align*}
P^-(n + 1) &= J_\phi(\hat{x}(n))P(n)J_\phi^T(\hat{x}(n)) + \Gamma Q \Gamma^T; \\
\hat{x}^-(n + 1) &= \phi(\hat{x}(n)).
\end{align*} \]

### 4.3.3.4 Kalman Filter Smoothing

The in-flight pointing solution is calculated causally — a star camera measurement is made and the pointing solution is updated. In the post-flight reconstruction, however, we have the full data set available and can run the filter forwards or backwards. For a data set of length \( N \), the forward solution at time \( n \) depends only on data from time 0 to \( n \), while a backward solution at the same time \( n \) depends only on data from time \( N \) back to \( n + 1 \). These two solutions are thus independent and they can be combined to produced a further refinement of the estimate of the state \( x(n) \). The optimal solution to the smoothing for the linear model is given in KS99.

### 4.3.3.5 Application to BLAST

We calculate the BLAST pointing solution based on one set of gyroscopes and the set of pointings from one of the star cameras. We choose the digital gyroscopes and the ISC star camera (see Sec. 2.5.2) because of better noise properties and performance. In principal, solutions based on different combinations of the two sets of gyroscopes and star cameras could also be smoothed together, but that step is not undertaken here.
We identify the state $x$ of the system with the 3 gyroscope biases $b_i$, the zero-points in the gyroscope signals which vary in time, and the 4 components of the quaternion $q_i$, describing the current orientation of the telescope:

$$x(n) = \begin{pmatrix} \vec{b}(n) \\ q(n) \end{pmatrix}.$$ (4.29)

$x(n)$ is thus a 7-component vector. There is no direct measure of the gyroscope biases $\vec{b}$; knowing that they vary, however, we allow them to vary as parameters in the fit.

The signal $s(n)$ and measurement $z(n)$ is the pointing quaternion $q$ calculated from the yaw (right ascension, RA), pitch (declination, Dec.) and roll reported by the star camera. If we set up our coordinate system such that $\hat{z}$ is vertically up, $\hat{y}$ points to RA = 0 and $\hat{x}$ is chosen so that we have a right-handed system, an arbitrary pointing $q$ can be described by the 3 rotations $q_z(y)$, $q_x(p)$ and $q_y(r)$, where $y$, $p$ and $r$ are yaw, pitch and roll, respectively:

$$q = q_y(r)q_x(p)q_z(y).$$ (4.30)

$C$, the matrix that gives the signal $s$ from the current state $x$ (Eqn. 4.12) is then a $7 \times 4$ matrix given by

$$C = \begin{pmatrix} 0_{34} & I_4 \end{pmatrix}.$$ (4.31)

Here $0_{34}$ is the $3 \times 4$ zero matrix and $I_4$ is the $4 \times 4$ identity matrix.

From Eqn. 4.28 it is clear that the time-update function $\phi(x(n))$ depends on the angular velocities measured by the gyroscope. Given the state $q(n)$ and an angular velocity $\vec{\omega}(n)$ at time $n$, the state $q(n+1)$ at time $n+1$, one time step $\Delta t$ later, can be written

$$q(n+1) = \delta q(n) q(n),$$ (4.32)

where

$$\delta q(n) = \left( \cos \left( \frac{|\vec{\omega}| \Delta t}{2} \right), \frac{\vec{\omega}}{|\vec{\omega}|} \sin \left( \frac{|\vec{\omega}| \Delta t}{2} \right) \right) \simeq \left( 1, \frac{\vec{\omega} \Delta t}{2} \right).$$ (4.33)

Here we have used Eqn. 4.10 and the last equality follows from the first-order Taylor expansions of the trigonometric functions — the higher-order terms are small since $\Delta t$ is small. Writing

$$q(n+1) = q(n) + \dot{q}(n) \Delta t,$$ (4.34)

we see by comparison with Eqns. 4.32 and 4.33 that

$$\dot{q}(n) = \begin{pmatrix} 0, \vec{\omega} \\ \frac{\vec{\omega}}{2} \end{pmatrix} q(n).$$ (4.35)
We identify \( \tilde{\omega} \) with the gyroscope signals \( \tilde{\Omega} \), but also account for the varying biases \( \tilde{b} \). We therefore write

\[
\tilde{\omega} = \tilde{\Omega} + \tilde{b}.
\]  

(4.36)

Finally, given no other information about \( \tilde{b} \), we set

\[
\tilde{b}(n + 1) = \tilde{b}(n).
\]  

(4.37)

\( \phi(\hat{x}(n)) \) is thus written

\[
\phi(\hat{x}(n)) = \phi \left( \begin{pmatrix} \tilde{b}(n) \\ q(n) \end{pmatrix} \right) = \begin{pmatrix} \tilde{b}(n + 1) \\ q(n + 1) \end{pmatrix}
\]

\[
= \begin{pmatrix} \tilde{b}(n) \\ \left( 1, \frac{\Delta t}{2} \left( \tilde{\Omega}(n) + \tilde{b}(n) \right) \right) q(n) \end{pmatrix}.
\]  

(4.38)

The \( J_\phi(\hat{x}(n)) \) are calculated numerically at each time-step. The components of the input \( \hat{x}(n) \) are varied by a small amount and the effects on \( \phi \) are measured using the familiar Newton’s quotient:

\[
\frac{\partial \phi_i}{\partial x_j} \approx \left( \frac{\phi(x + h e^j) - \phi(x)}{h} \right)_i,
\]  

(4.39)

where \( e^j \) is the unit vector along direction \( j \),

\[
e^j_k = \begin{cases} 
0, & k \neq j \\
1, & k = j
\end{cases},
\]  

(4.40)

and \( h \) is chosen to be small compared to \( x_j \).

The noise propagation matrix \( \Gamma \) transforms the gyroscope measurement noise into process noise, the noise on the components of \( x \). Given the bias noise vector \( \tilde{n} \) and the angular velocity noise \( \tilde{N} \), both 3-component vectors, we form

\[
\tilde{w} = \begin{pmatrix} \tilde{n} \\ \tilde{N} \end{pmatrix},
\]  

(4.41)

and then

\[
\Gamma = \frac{\Delta t}{2} \begin{pmatrix}
I_3 & 0_{33} \\
-q_1 -q_2 -q_3 & q_0 q_3 -q_2 \\
-q_3 q_0 q_1 & q_2 -q_1 q_0
\end{pmatrix},
\]  

(4.42)
where $I_3$ and $0_{ij}$ are as defined above and the lower-right hand quadrant produces the quaternion multiplication $(0, \vec{N}) q$, with $q_i$ the elements of $q$, when operated on $\vec{N}$. The noise covariance matrices $Q$ and $R$, describing $\vec{w}$ and $\vec{v}$, respectively, are estimated from the expected noise properties of the gyroscopes and star cameras.

The gyroscopes are sampled at 100 Hz. The filter is updated at every time step, but the star camera measurements $z$ are not available at every sample. In such cases, the measurement update steps (Eqns. 4.17-4.19) are skipped and $\hat{x}^-$ and $P^-$ are simply copied onto $\hat{x}$ and $P$.

An additional complication is that occasionally the star cameras make mis-identifications and return incorrect solutions. Since at all times we have an estimate of the current pointing solution, a bad star camera solution can be rejected if it is sufficiently far away from the calculated solution. We define an error envelope based on the expected gyroscope error, which increases in time as $\sqrt{t}$. We set the scale of the error envelope to the nominal value of $\sigma_{\text{gyro}} = 4'' s^{-1/2}$. To quantify the concept of $1/f$ noise is the gyroscopes, the envelope is expanded exponentially after a delay. The error envelope $\epsilon(t)$ is defined as:

$$\epsilon(t) = \sigma_{\text{gyro}} \cdot \sqrt{t} \cdot f(t), \quad (4.43)$$

where $t$ is the time since the last star camera (SC) solution and $f(t)$ is the exponential envelope,

$$f(t) = \begin{cases} 1, & t \leq t_0 \\ \exp \left( \frac{t-t_0}{\tau_{\text{err}}} \right), & t > t_0 \end{cases}. \quad (4.44)$$

The exponential parameters $t_0 = 5s$ and $\tau_{\text{err}} = 3s$ are chosen empirically. Every time a star camera solution becomes available, the angular distance $\theta_{\text{SC}}$ between the SC solution and the most recent integrated solution is compared to the error envelope and is rejected if $\theta_{\text{SC}} > 5\epsilon$. In a few cases, errors in timing between the star camera and the flight computers cause incorrect SC solutions to be reported that are not rejected by this filter. In such cases, the bad solutions are flagged by hand and are ignored. As an example, a small part of the pointing solution, complete with error envelopes and an incorrect SC solution, is shown in Fig. 4.4.

Finally, as discussed above, the filter can be run backwards in time, producing a completely independent measure of the pointing solution. The two solutions can then be smoothed together to increase the accuracy of the estimate. We do not use the formalism of KS99, which only applies to the linear system, and instead use an approximation to the smoothing function — a weighted sum based on the error envelope $\epsilon$. If $q_f$, $q_b$ and $q_s$ represent the forward, backward
Figure 4.4: The forward pointing solution. The top and middle plots show declination and right ascension, respectively, of the telescope as determined by the Kalman-filtered pointing solution. The solid black are the integrated solution, the circles the pointings reported by the star cameras, and the dotted lines are 5 times the error envelopes. The bottom plot shows the number of stars found in the star camera image used to calculate the given star camera solution. A star camera timing error is evident for the solution near frame 904290 — this solution is rejected as it falls outside of the error envelope.
and smoothed solutions, and $\epsilon_f$, $\epsilon_b$ and $\epsilon_s$ their respective error envelopes, then the smoothed solution is given as

\[
q_s = \frac{\epsilon_f^2 q_f + \epsilon_b^2 q_b}{\epsilon_f^2 + \epsilon_b^2},
\]

\[
\epsilon_s = \sqrt{\frac{\epsilon_f^2 \epsilon_b^2}{\epsilon_f^2 + \epsilon_b^2}}.
\]

The smoothed solution $q_s$ is transformed to yaw, pitch and roll, which are written to disk. A small section of data showing how the smoothed solution interpolates the forward and backward solutions is displayed in Fig. 4.5.

In practice, the Kalman filter is started at the time for which the first star camera solution is available. The gyroscope biases $\vec{b}$ are initialized to 0 and the $q$ are initialized to the SC solution.

### 4.3.3.6 Results

The non-linear Kalman filter pointing solution was implemented in C++ in order to take advantage of the quaternion and matrix classes in Matpack. The processing time for a 1-hr chunk of data on a 3.2 GHz Pentium 4 desktop computer is approximately 20 s.

### 4.4 Map Making

In its most rudimentary form, map-making is simply a projection of samples into the associated map pixels (e.g. Wright et al. 1996; Prunet et al. 2001). If the data samples are completely independent and we assume, for the moment, that the noise is the same for all samples, then we can formulate the map-making problem in the language of linear algebra (we denote matrices and vectors with upper and lower case letters, respectively):

\[
Pm = d.
\]

Here, $d$ is the time-ordered data vector of length $n_d$, $m$ is the map vector, containing the value of each pixel in the map, of length $n_p$, and $P$ is the $n_p \times n_d$ pointing matrix, which projects the data onto the map. The data vector $d$ includes measurements from many bolometers, so is of length $n_d = n_t n_b$, where $n_t$ and $n_b$ are the number of time samples and number of bolometers.
Figure 4.5: A small segment of pointing data. Circles represent star camera solutions. The solid lines are the integrated solutions and the dashed lines are 10 times the error envelope (exaggerated for visibility). The thin lines represent the forward (lower) and backward (upper) solutions, while the thick line represents the smoothed solution.
respectively. The data samples for each detector are ordered in time, then stacked detector-by-detector. The pointing matrix $P$ can, in principle, be a complicated matrix which takes into account the shape of the telescope’s point spread function. If the pixel size is chosen to be small enough, however, typically 1/5\textsuperscript{th} of the beam width, it is a reasonable assumption to project the measurement $d_i$ entirely into a single pixel $m_j$. The rows of $P$ are then all 0 except for a 1 in the row corresponding to the correct pixel, as determined by the pointing solution.

Eqn. 4.47 has solution

$$m = (P^T P)^{-1} P^T d,$$

(4.48)

where $^T$ is the matrix transpose operator and $^{-1}$ denotes matrix inverse. We note that the $j^{th}$ element of the $n_p$-element vector $P^T d$ is the sum of the data points projected onto pixel $j$, and $P^T P$ is an $n_p \times n_p$ matrix containing the number of observations of each pixel on the diagonal. The $j^{th}$ element of the right side of Eqn. 4.48 is thus simply the average of the data points projected onto pixel $j$.

We can extend the model of Eqn. 4.47 to include errors associated with the data. We form the $n_d \times n_d$ matrix

$$N = \begin{pmatrix} \sigma_1^2 & \sigma_{12}^2 & \cdots & \sigma_{1n}^2 \\ \sigma_{12}^2 & \sigma_2^2 & \cdots & \sigma_{2n}^2 \\ \vdots & \vdots & \ddots & \vdots \\ \sigma_{1n}^2 & \sigma_{2n}^2 & \cdots & \sigma_n^2 \end{pmatrix},$$

(4.49)

where $\sigma_i$ is the noise variance associated with the measurement $d_i$ and $\sigma_{ij}$ is the noise covariance associated with the measurements $d_i$ and $d_j$. Eqn. 4.48 then becomes

$$m = (P^T N^{-1} P)^{-1} P^T N^{-1} d.$$  

(4.50)

This formulation, for pixel $m_j$, is completely equivalent to the weighted least-squares solution

$$m_j = \frac{\sum_k d_k w_k}{\sum_k w_k},$$

(4.51)

where the sum is over all data points projected onto the $j^{th}$ pixel and the weights $w_k \equiv \sigma^{-2}$. Under the assumption of uncorrelated data, the $\sigma_{ij} = 0$ where $i \neq j$ and $N$ is diagonal, and is thus trivially inverted. Since $P$ has only one element per row, $P^T N^{-1} P$ is also diagonal and Eqn. 4.51, which we call the \textit{naive} map-maker, is easily calculated. Given correlated data, the naive map-maker can still produce scientifically useful maps, provided that the input time
streams are first strongly high-pass filtered. This removes signals at spatial scales on the order of the scan speed times the cut-off frequency. When using the naive map-maker, we filter the data with a high-pass 8th order Butterworth filter with cut-off frequency $f_0 = 0.05\,\text{Hz}$. In order to prevent ringing at the start and end of the time streams when the Fourier transforms are applied, the first and last 2000 samples of input data are apodized; these samples are not projected into the map.

To retain the large spatial scales, $N$ can be further extended to account for correlations between data points, both correlations in time (due to the $1/f$ noise exhibited by bolometer signals) and between bolometers (due to e.g. sky noise, loading due to optics, etc.) by including the appropriate off-diagonal elements. Assuming perfect knowledge of the elements of $N$, Eqn. 4.50 gives the optimal solution to the map-making problem. In practice, however, it is not directly computable, due to limitations in computational power and memory. Matrix inversion is computationally expensive, with the number of operations needed to calculate the inverse of an $n \times n$ matrix being proportional to $n^3$. The matrices $N$ and $P^T N^{-1} P$ are both very large for BLAST maps, and simplifying assumptions and approximations must be made. See PAT07 for a detailed discussion of the map-maker used for the BLAST maps.

Even given the full covariance matrix $N$, the largest time scale bolometer drifts are poorly constrained by the map-maker. A 5th order polynomial is fit to the entire time stream and removed. The data are then high-pass filtered, as with the naive map-maker, but with a cut-off frequency of $f_0 = 5 \times 10^{-4}\,\text{Hz}$. The data are similarly apodized.

In principle, the map-maker could measure the noise covariances directly from the data, but care must be taken to accurately separate the noise from the signal. This is not always practical, and we instead use measurements of the extragalactic blank field ELAIS-N1 (specifically chosen for low foreground levels) as a measure of the bolometer noise power spectra. In most cases, we ignore bolometer-bolometer correlations, due to the increased complexity of the map-making process. For smaller maps (e.g. Cas A), these correlations can be taken into account.

The map-maker writes the resulting maps to file in FITS\textsuperscript{3} format. The outputs are: an intensity map; a hits map, giving the number of data samples projected into each pixel; and a corresponding variance map, giving an estimate of the noise variance in each pixel based on the diagonal elements of the pixel-pixel noise covariance matrix.

\textsuperscript{3}Flexible Image Transport System, a standard file format used in astronomy.
Simulations testing the effectiveness of the map-maker are described in [PAT07]. It is shown that the optimal map-maker retains power on large spatial scales while more naive algorithms do not, and that the transfer function of the map maker, defined as the ratio between the amplitudes of fluctuations in the output map relative to the input map, is usually within 3% of unity.
Chapter 5

Galactic Plane Survey

We now apply the data processing procedures described in Chapter 4 to observations made during the 2005 science flight. In this chapter we describe the analysis of a survey in the Galactic plane and in the next chapter present the analysis of a nearby edge-on spiral galaxy.

Observations of Galactic dust emission in the submillimetre are sensitive to the earliest stages of star formation (e.g. Shu et al. 1987). Both molecular cores lacking internal sources of radiation and proto-stellar stars obscured by large amounts of dust elude detection by mid-IR instruments such as IRAS and Spitzer, as the dust temperatures are too cold for significant emission at \( \sim 100 \mu m \). Of particular interest are high-mass proto-stellar objects (HMPOs), a class of dense clouds that may eventually form massive OB stars with luminosities in the range \( 10^2 - 10^5 \, L_\odot \). These objects are embedded in dense molecular clouds which they have not yet ionized and thus exhibit no H\( \text{\textsc{ii}} \) radio emission. They are seen only in cool thermal emission, to which BLAST is particularly sensitive. BLAST is also sensitive to a later stage of evolution known as ultra-compact (UC) H\( \text{\textsc{ii}} \) regions, in which hot stars have already formed in the core of a surrounding molecular cloud. These stars ionize the surrounding gas and are thus visible in the radio.

Observations at 850 \( \mu m \) with SCUBA (e.g. Johnstone et al. 2000) have detected large numbers of objects in Galactic molecular clouds, but with only 1 band, do not allow determination of temperature. Measurements at 450 \( \mu m \) would allow for temperature constraints, but are much less sensitive and thus mapping speeds are significantly longer in order to obtain reasonable signal-to-noise ratios. BLAST has the unique ability to conduct large-area unbiased searches for cold dust emission from pre- and proto-stellar objects, and with its 3-band measurements, is able to constrain the temperatures of the coldest objects with \( T \lesssim 20 \, K \).

As discussed in Sec. 1.4, the molecular cloud mass function, the number of objects found with a given mass range, can help to constrain star formation models. Given a sufficient number of
objects, a mass function for BLAST sources can be created, providing constraints for formation models.

Large amounts of time were available for Galactic studies during the 2005 flight, because of the reduced point-source sensitivity of the telescope, due to the partial failure of the warm optics, discussed in Sec. 3.5.2. Observations made during the 2005 flight are summarized in Table 3.1. Four large fields in the Galactic plane were mapped, as shown in Fig. 5.1. The observations are summarized here:

- Towards Aquila in the Galactic plane. The field mapped is a 2.5° × 1.6° rectangle (“quad”) aligned in Galactic coordinates, centred at ℓ = 45.7°, b = −0.2°. A total of 6.1 hours was spent observing this field. A 1-σ point-source sensitivity of 370 mJy was reached.

- Towards Sagitta in the Galactic plane. The field mapped is part of a 3° × 2° rectangle (“quad”) aligned in Galactic coordinates, centred at ℓ = 53.5°, b = 0.1°. The map was aborted half way through completing the scan to pursue another object. A little less than an hour was spent on this field, reaching a 1-σ point-source sensitivity of 215 mJy, but the coverage is very uneven and no cross-linking was achieved.

- Towards Vulpecula in the Galactic plane. The field mapped is a 1.5° × 1.5° rectangle (“quad”) aligned in Equatorial coordinates, centred at ℓ = 59.2°, b = 0.4°. A total of 6.5 hours was spent observing this field. A 1-σ point-source sensitivity of 320 mJy was reached.
reached.

- Towards Cygnus X in the Galactic plane. Two overlapping fields were observed at this location: a circle (“cap”) with 1° radius centred at \( \ell = 80.8^\circ, b = 0.8^\circ \) towards W 75N, a young protostar/compact H II region; and a \( 3.6^\circ \times 1.9^\circ \) rectangle (“quad”) roughly aligned in Galactic coordinates, centred at \( \ell = 80.6^\circ, b = 0.4^\circ \), encompassing W 75N and other star-forming regions. A total of 13.2 hours was spent observing this field. A 1-\( \sigma \) point-source sensitivity of 370 mJy was reached.

We have chosen one of the observations, the \( 1.5^\circ \times 1.5^\circ \) survey towards the constellation Vulpecula, to analyse in detail. The field contains the open cluster NGC 6823, a prominent H II region called Sh2-86 and a supernova remnant. The reason for choosing this field is that it is the smallest of the Galactic plane surveys, and is thus the easiest for which to produce optimal maps. Additional advantages include the facts that it contains many bright sources and was observed at several different times during the flight, at varying position angles and for a reasonably large total observing time.

This chapter details the analysis performed on the Vulpecula observations. We start by describing the observations in detail and discuss how the maps are made. We then describe attempts to spatially deconvolve the images to produce maps at a resolution nominally equal to the designed diffraction limit. Compact source detection and extraction, as well as simulations to quantify the accuracy of the reduction routines, are then described. Finally, we present scientific conclusions based on the BLAST compact source list by identifying the sources with observations of this field at other wavelengths.

## 5.1 Observations

The Vulpecula Galactic Plane field was observed 5 times throughout the flight for a total of 6.1 hours. Details of each observation are listed in Table 5.1. The sky coverage in each observation is shown in Fig. 5.2. We see that the field is not fully sampled in all observations. Observation 1 was aborted in order to perform a calibration map of CRL 2688, after which the Vulpecula observations were continued with Observation 2. The coordinates in the command for Observation 3 were entered incorrectly, resulting in the shape seen in the middle-left panel of
Table 5.1: Summary of Vulpecula Observations

<table>
<thead>
<tr>
<th>Observation Number</th>
<th>Start (hr)</th>
<th>Length (hr)</th>
<th>Number of Scans</th>
<th>Parallactic Angle</th>
</tr>
</thead>
<tbody>
<tr>
<td>1</td>
<td>22.9</td>
<td>0.4</td>
<td>0.5</td>
<td>$-14.6^\circ \rightarrow -12.8^\circ$</td>
</tr>
<tr>
<td>2</td>
<td>23.7</td>
<td>2.0</td>
<td>2.2</td>
<td>$-10.9^\circ \rightarrow -0.4^\circ$</td>
</tr>
<tr>
<td>3</td>
<td>29.7</td>
<td>0.9</td>
<td>0.5</td>
<td>$18.5^\circ \rightarrow 21.1^\circ$</td>
</tr>
<tr>
<td>4</td>
<td>30.5</td>
<td>0.4</td>
<td>0.4</td>
<td>$21.1^\circ \rightarrow 22.0^\circ$</td>
</tr>
<tr>
<td>5</td>
<td>46.9</td>
<td>2.8</td>
<td>2.8</td>
<td>$-21.1^\circ \rightarrow -12.0^\circ$</td>
</tr>
</tbody>
</table>

Fig. 5.2 The command was resent when this was noticed, resulting in Observation 4, but there was not enough time remaining to complete a full scan. Observation 5 completed successfully.

We note that the field was observed over a large range of parallactic angles, defined, for a given point on the sky, as the angle $\phi$ between the direction of azimuth and the direction of Right Ascension, and hence the direction of the scan lines of the observations. The parallactic angle depends on the observer latitude $\lambda$, time of observation LST, and position of the source, $(\alpha, \delta)$ via

$$\sin \phi = \frac{\sin H \cos \lambda}{\cos \delta},$$

where $H$ is the hour angle of the source, LST $- \alpha$. We see that the range of observed parallactic angles is $\Delta \phi \sim 40^\circ$, but that Observations 3 and 4, those at $\phi \approx +20^\circ$, do not cover that whole region. However, observations 2 and 5 together span $\phi \sim -20^\circ$ over most of the field, providing adequate cross-linking for constraining $1/f$ noise in the map-making process.

5.2 Telescope Beams

As discussed in Sec. 3.5.2, the telescope was significantly out of focus during the 2005 flight, producing strongly non-Gaussian asymmetric beams with power on scales much larger than the designed diffraction limit.

We measured the point spread function of the telescope during the flight using observations of point-like objects, using the full set of observations of the point source CRL 2688. Optimal maps are made using all bolometers at each waveband. The maps are created in telescope
Figure 5.2: Sky coverage for the 5 scans of Vulpecula together with the total cross-linked coverage in the bottom right panel. Note the substantially different scan directions due to sky rotation. Notice also that the range of rotations varies significantly across the combined map.
coordinates, centred on CRL but parallel to the telescope’s frame of reference, so that the asymmetric beam is always oriented in the same direction. The maps from each visit are averaged to produce a single measurement of the beam at each wavelength, shown in Fig. 5.3.

A telescope’s beamwidth is in most cases characterized by its full-width at half-maximum (FWHM), equivalent to $2\sqrt{2\ln 2}\sigma = 2.35\sigma$ for a Gaussian $g(x) \propto \exp(-x^2/(2\sigma^2))$. The 2005 BLAST beams are so complex, however, that they are not well described by their FWHM, nor is the FWHM even well-defined. We instead use the full-width at half-power (FWHP) as a diagnostic. We define FWHP as the diameter of the circle centred on the beam that encloses half of the beam’s power. For a Gaussian PSF, we can calculate the fractional power $P$ enclosed by a circle of arbitrary radius $r_0$:

$$P(r_0) = \int_0^{r_0} r \, dr \int_0^{2\pi} d\theta \, g(r, \theta),$$

(5.2)

where

$$g(r, \theta) = g(r) = \frac{1}{2\pi\sigma^2} e^{-r^2/(2\sigma^2)}$$

(5.3)

is a 2-dimensional circular Gaussian, normalized so that $\int_\Omega g(r) \, d\Omega = 1$, and thus $P(\infty) = 1$. Carrying out the integral in Eqn. 5.2

$$P(r_0) = 1 - e^{-r_0^2/(2\sigma^2)},$$

(5.4)
Table 5.2: Beam Widths

<table>
<thead>
<tr>
<th>Band</th>
<th>FHWM(^a)</th>
<th>FWHP</th>
<th>EWHP</th>
<th>(P_{cen})(^b)</th>
</tr>
</thead>
<tbody>
<tr>
<td>250 (\mu)m</td>
<td>40(')</td>
<td>3.4(')</td>
<td>3.1(')</td>
<td>4%</td>
</tr>
<tr>
<td>350 (\mu)m</td>
<td>58(')</td>
<td>3.5(')</td>
<td>3.2(')</td>
<td>6%</td>
</tr>
<tr>
<td>500 (\mu)m</td>
<td>75(')</td>
<td>3.5(')</td>
<td>3.2(')</td>
<td>14%</td>
</tr>
</tbody>
</table>

\(^a\)Nominal designed diffraction limit
\(^b\)Percentage of power in nominal beam

...and hence the FWHP for a Gaussian is \(2r_{1/2}\), where \(r_{1/2}\) is such that \(P(r_{1/2}) = \frac{1}{2}\). We find \(r_{1/2} = \sqrt{2 \ln 2} \sigma\) and see that the FWHP is equivalent to the FWHM for a Gaussian beam profile.

The BLAST beam FWHPs are calculated by numerically integrating the beam maps. The beams are normalized so that the total area under them in the map is 1.0, after first subtracting an estimated (constant) background by taking the average of pixels in an annulus surrounding the beam. The azimuthally-averaged radial profile is calculated by averaging the pixel values in 1-pixel wide annuli and multiplying by the radius of the annulus. The profile is integrated to produce the integrated beam power, and the curves interpolated to find the FWHPs. The results are shown in Fig. 5.4 and Table 5.2. The similarity of shape points to beam deformation rather than diffraction as the cause of the beam degradation.

An alternative approach to characterizing the beams is to ask how strongly the power is clustered. We devise the equivalent-width half-power (EWHP), defined as the diameter of the circle with area equal to \(N \Delta x^2\), where \(\Delta x\) is the pixel size and \(N\) is the number of pixels needed to make up half of the beam power. For a beam with a monotonically decreasing radial profile, this is equivalent to the FWHP, and thus the EWHP for a Gaussian is equal to the FWHM. The EWHP is easily calculated by sorting the map pixels in order of decreasing brightness, and, starting from the bright end, adding up pixels until half the power is reached. The EWHP is then \(2 \Delta x \sqrt{N/\pi}\). The resulting EWHPs are listed in Table 5.2 and are surprisingly not very different from the FWHP. This is due to the fact that, in all 3 beams, most of the power is found in the ring which is in fact fairly wide.

Finally, we can ask what fraction of the beam power is found within the nominal diffraction
Figure 5.4: Beam power profiles for the 2005 BLAST beams. The upper panel shows azimuthally-integrated power as a function of radius. The lower panel shows the integrated radial power. The thin grey band includes the full-width at half-power location of all three bands, coincidentally where the curves intersect.
limit. These numbers can be read off the integrated power profile, and we find that 4, 6 and 14% of the power falls within the diffraction-limited beam widths for the 250, 350 and 500 μm beams, respectively, compared to 50% for a true Gaussian beam.

Attempts were made to model the beam shapes seen in Fig. 5.3 in the optical modelling software ZEMAX\(^1\), used to design the optics. Various shifts and tilts of the primary and secondary mirrors were applied, but we were unable to reproduce the features seen in the 3 beams. We conclude that the beam degradation was due to some more serious defect, such as surface delamination of the primary. The 6-fold symmetry seen in the beams hints at such an effect, as the primary was constructed in 6 segments.

5.3 Alignment and Maps

Maps of the Vulpecula survey are made using the optimal map-maker described in Sec. 4.4. Before the data are fed into the map-maker, however, the 5 observations must be registered, since we expect variations in the star camera/telescope alignment throughout the flight, as discussed in Sec. 3.3.2. We assume that the star camera/telescope alignment does not change significantly within an observation.

Observations of pointing calibrators were performed throughout the flight, but the Vulpecula field contains several bright point sources which are suitable for self-registration and provide a more direct measure of the misalignment. We choose 7 bright point sources in the field that have been observed with SCUBA to perform the alignment. Publicly-available SCUBA observations, with pointing precision of \(\sim 3''\), are examined to determine the true coordinates of the 7 sources. For each source in each of the 5 observations, we make naive thumbnail maps with high-pass filtered data, as described in Sec. 4.4. We perform the alignment with the 250 μm data only, as this beam has the smallest central core, allowing for the best determination of the shifts. The relative alignments between bands are not expected to vary, as their relative aligned depends only on the cold optics which should remain fixed throughout the flight. The maps are constructed in telescope coordinates (az, el), since it is the natural frame in which to calculate the offsets. The thumbnails are centered on expected location of each source, as determined from the SCUBA maps, and are 12’ on a side and with 5’ pixels (see Fig. 5.5).

\(^1\)http://www.zemax.com
Table 5.3: Relative Misalignment Measurements in Vulpecula Observations

<table>
<thead>
<tr>
<th>Scan Differences</th>
<th>Sources</th>
<th>ave</th>
</tr>
</thead>
<tbody>
<tr>
<td></td>
<td>a</td>
<td>b</td>
</tr>
<tr>
<td>1–5</td>
<td>...</td>
<td>...</td>
</tr>
<tr>
<td>2–5</td>
<td>(0,5)</td>
<td>(1,5)^w</td>
</tr>
<tr>
<td>3–5</td>
<td>...</td>
<td>...</td>
</tr>
<tr>
<td>4–5</td>
<td>...</td>
<td>...</td>
</tr>
</tbody>
</table>

Note that not all sources are observed in all observations. Also, the fact that the beam shape is clearly visible and not significantly smeared in each thumbnail (where the source is observed) is assurance that our assumption of constant alignment within an observation is reasonable.

The relative misalignment between observations is measured by comparing observations 1–4 with observation 5. This is done for each source and the results are averaged. The misalignment is measured by cross-correlating the thumbnail images. The cross-correlation of 2 maps $M$ and $N$ gives the correlation map $C$:

$$C(i,j) = \sum_k \sum_l M(k,l)N(k-i,l-j),$$  \hspace{1cm} (5.5)

where $C$ is largest when $M$ and $N$ are properly aligned. In some cases, where the signal-to-noise ratio is low, this method does not provide a good measurement and a more simplistic approach is applied: the central pixels in each map are found by eye, and the misalignment is simply the difference of the two. The measured misalignments ($\Delta x, \Delta y$) are shown in Table 5.3 in units of pixels ($5''$). Sources for which the cross-correlation did not work and which were calculated by hand are denoted with “h”, and weak sources for which the measurement is questionable are marked with “w”. We see that the misalignments of sources within an observation are consistent with each other. We take an average of all measurements within an observation, excluding those marked with “w”. The average offsets are listed in the final column of the table.

The absolute offsets of the map compared to the sky are measured by finding the misalignment of the peaks of the 7 sources in observation 5. Sources b and f are ignored and we find an absolute shift of $(\Delta x, \Delta y) = (5.2, 0.6)$ in units of pixels.

Finally, the overall offset between bands is measured. Since the beam shapes are drastically
Chapter 5. Galactic Plane Survey

Figure 5.5: Thumbnail maps of 250 μm data, constructed in telescope coordinates, centred on the 7 alignment sources (labelled a–g) in each of the 5 observations (labelled 1–5). The maps are 12′ × 12′ with 5″ pixels (only the central 6.7′ × 6.7′ regions are shown). Note that not all sources are observed in all observations. The white targets denote the expected source centre, as measured from publicly-available SCUBA observations. Misalignment of the target with respect to the centre of the source are due to variations in the star camera/telescope boresight alignment. This misalignment is corrected by cross-correlating the thumbnail maps of observations 1–4 with observation 5, which serves as the reference.
different, the cross-correlation method used above is not applicable, and we use the by-hand method of finding the source centres. This method is not highly accurate, however, and the band maps are more accurately aligned later in the analysis process (Sec. 5.5.1).

The measured offsets, between observations, between bands and relative to the sky, are used to correct the pointing data. The offsets are input into the map-maker and the pointing fields read from file are shifted by the appropriate amounts.

Optimal maps at each wavelength are made using the method discussed in Sec. 4.4. The size of the maps are determined by the map-maker based on the pointing data in the specified frame ranges. The pixel size is limited by computer memory, as the $n_{\text{pix}} \times n_{\text{pix}}$ matrix $(P^T N^{-1} P)^{-1}$ (Eqn. 4.50) must be stored. We create maps with 18" pixels, which is near the limit of memory on readily available computers. We do not account for detector-detector cross-correlations in the map-making process, as the calculation time required for such a large map is prohibitive. The maps are shown in Fig. 5.6.

Many point-like objects are visible across the field, as well as diffuse emission throughout. The similarity of the maps is striking; this similarity, along with the fact that the brightness increases towards shorter wavelengths, indicates that the bulk of the emission in both the compact and diffuse sources has similar spectra in the BLAST bands.

The map intensities are calibrated using the results of the flux density calibration procedure described in Sec. 4.2.3.

5.4 Image Deconvolution

The raw Vulpecula BLAST maps contain significant bright structure, including many point-like objects, as can be seen by the characteristic beam shapes stamped across the maps (Fig. 5.6). However, due to the high source density, there is significant crowding across the map, making source extraction difficult. PSF fitting methods were attempted, but the source flux densities were not well determined and the extraction was far from satisfactory. We decided to attempt to deconvolve the images to recover the diffraction-limited resolution in order to resolve the significant confusion. The deconvolution process amplifies noise at high frequencies, but because of the long integration times and resulting high signal-to-noise ratios, the maps can be deconvolved to recover structure on scales near to the diffraction limit. We find that we can
Figure 5.6: Raw maps created using the optimal map-maker. Point-like and extended structures are visible across the map. The similarity of the maps and increase in brightness towards shorter wavelengths indicates that the bulk of the emission is thermal Rayleigh-Jeans across the BLAST bands. The complex beam structures are clearly visible in the brighter point sources.
successfully extract the sources from the deconvolved maps, despite the amplification of noise due to the deconvolution.

5.4.1 Deconvolution Methods

Deconvolution in astronomy has a long history (see the review by Starck et al. 2002, hereafter STA02). The convolution process can be stated, using the notation of STA02, as

\[ I(x, y) = (O \ast P)(x, y) + N(x, y), \]  

(5.6)

where \( I \) is the observed map, \( O \) is the true image, \( P \) is the instrument’s PSF, \( N \) is measurement noise and “\( \ast \)” is the convolution operator. Invoking the convolution theorem, this expression can be written in Fourier space as

\[ \hat{I}(u, v) = \hat{O}(u, v) \hat{P}(u, v) + \hat{N}(u, v), \]  

(5.7)

where \( \hat{X} \) is the Fourier transform of \( X \). A naive solution to the convolution problem is by simple division in Fourier space,

\[ \hat{\hat{O}}(u, v) = \frac{\hat{I}(u, v)}{\hat{P}(u, v)} = \hat{O}(u, v) + \frac{\hat{N}(u, v)}{\hat{P}(u, v)}. \]  

(5.8)

In practice, this direct inversion method injects structure into the map \( \hat{\hat{O}} \) due to the fact that we have a noisy measurement of \( P \). Many methods have been developed to solve Eqn. 5.6 iteratively, including least squares, maximum likelihood and wavelet-based methods. In particular, the well-developed Richardson-Lucy deconvolution algorithm (Richardson 1972; Lucy 1974) (which results from maximizing the likelihood assuming Poisson noise) has been used extensively, including application to images produced by the pre-refurbished Hubble Space Telescope (Hanisch and White 1994). The Richardson-Lucy algorithm in particular is not applicable to the BLAST, as the BLAST maps exhibit Gaussian noise. STA02 and Vio et al. (2005) discuss the Landweber algorithm, an iterative least squares solver, which accommodates Gaussian noise.

These algorithms typically require perfect knowledge of the PSF, which we do not have for the BLAST maps. In addition, the effective BLAST PSFs vary significantly across the maps due to angular asymmetries in the pattern, and the variation of the PSF for different BLAST bolometers. Over time the orientation of this pattern on the sky changes, such that the effective
shape of a point source depends on the amount of time spent observing at different parallactic angles and with each detector, along with instrumental noise variations.

Methods for simultaneously estimating the beam along with the deconvolved map exist (e.g. Thiebaut 2002), but typically the beam is assumed to be constant across the map. One could imagine using a varying PSF, but the deconvolution becomes significantly more complex, since convolution in Fourier space is no longer applicable. Also, the calculation of the effective PSF at each pixel is difficult, requiring detailed knowledge of pixel-pixel correlation matrix produced by the map-maker.

We conclude that without significantly more complex algorithms requiring large amounts of development time, strong artifacts due to convolution are inevitable, and hence we use the direct inversion method. It is not optimal, but produces useful deconvolved maps in reasonable amounts of time.

5.4.2 Application to BLAST

We use Eqn. 5.8 to calculate the deconvolution. It is crucial that our estimate of \( P \) is well behaved so as to not amplify noise in the solution.

Equation 5.8 can be re-written in real space as

\[
\hat{O} = I \ast \hat{P},
\]

where \( K \), the deconvolution kernel, is the inverse transform of \( \hat{P}^{-1} \). In order to suppress the amplification of noise at high frequencies we effectively re-convolve the map by a Gaussian \( G \) with width approximately equal to the designed diffraction limit of the telescope. This can be combined into the deconvolution formulation by writing

\[
\hat{K} = \hat{G}/\hat{P}. \tag{5.10}
\]

We use \( Gs \) with FWHMs of 40, 50 and 60" for the 250, 350 and 500 \( \mu \)m maps, respectively.\(^2\)

The PSFs are estimated by two methods. In the first case, a synthetic PSF is constructed based on measurements of point sources throughout the flight. This synthetic PSF attempts

\( ^2\)These values are smaller than the nominal diffraction limits at each wavelength, due to a misunderstanding in the design of the telescope — at the time when this analysis was performed, it was thought that the 2005 optical system had been designed with 30, 45 and 60" beams, but that was not the case.
to account for the range of parallactic angles over which the field was observed by averaging together appropriately-weighted beam rotations. The second method uses a measurement of a point source taken directly from the map. An isolated source with relatively low surrounding diffuse emission from the north-west (upper-right) corner of the map is used (see Figure 5.6). A polynomial is fit to the background and subtracted from the source to remove the diffuse emission. The deconvolutions are performed using both types of PSF. We find that the synthetic PSFs provide the best results at 350 and 500 µm, while the isolated source is better for the 250 µm map, based on the amplitude of the ripples seen in the resulting deconvolved maps.

In order to reduce edge effects caused by the Fourier transforms, both the maps $I$ and the PSFs $P$ are apodized with a cosine function and zero-padded prior to deconvolution. The map is apodized over a scale of 5' across a rectangle bordering the map and the PSF is apodized over 2' across a circle of radius 4'.

The power spectrum of $K$ for the 250 µm deconvolution is shown in Fig. 5.7. The radially-averaged profile is also given. The main features to notice are: the null at low frequency, caused by the large-scale structure in the PSF; two non-circular rings of high power at approximately 0.08 and 0.15 pix$^{-1}$; and high-frequency noise at $k > 0.3$ pix$^{-1}$, due to measurement noise in the PSF. Also note that the power falls off at large frequencies, due to the multiplication by $\tilde{G}$. We further reduce the small-scale noise by smoothing the kernel at high frequencies. At frequencies greater than 0.2 pix$^{-1}$, we set $K$ to a smoothed version of the radial profile.

The result of this deconvolution procedure is shown in Fig. 5.8. The improvement in resolution compared to the raw maps is clear. The point sources seen as stamps of the beam in the raw image are all bright, compact sources in the deconvolved map. Additionally, many fainter sources are seen in the deconvolved map that are only hinted at in the raw map, particularly in the lower-left quadrant. Also clearly visible in the deconvolved map are the ripples surrounding the brightest sources, due to the imperfect knowledge of the PSF. The ringing is worst around the sources in the lower-right corner. This corner of the map is observed with significantly different scan angles than the rest of the map (Fig. 5.3), producing significantly different effective beams.

Finally, the variance map produced by the map-maker is propagated through the deconvolution filter, providing a noise map for use in source-finding and fitting. Given the map $I$ with
Figure 5.7: The power spectrum 250 μm deconvolution filter $K$ before smoothing (top) and the radially-averaged spectrum (bottom). The complex structure at low frequencies is due to structure in the PSF and the spiky structure at high-frequencies is due to having a noisy measurement of the PSF. The filter is smoothed at high frequencies prior to deconvolution.
Figure 5.8: The 250 μm raw Vulpecula map (top) and deconvolved map (bottom). The improvement in resolution is clear, as is the ringing around bright objects caused by the deconvolution. The dashed box indicates the source-free region where the map variance is calculated (Sec. 5.5.1).
pixel variance $\xi_I^2$ and convolution kernel $K$, we write
\[ \xi_O^2 = \langle \hat{O}^2 \rangle = \langle (I * K)^2 \rangle. \] (5.11)

The $(i,j)^{th}$ element of $I * K$ can be written
\[ (I * K)(i,j) = \sum_{k,l} I_{k,l} K_{i-k,j-l}, \] (5.12)
and thus
\[ \xi_O^2(i,j) = \left\langle \left( \sum_{k,l} I_{k,l} K_{i-k,j-l} \right) \left( \sum_{m,n} I_{m,n} K_{i-m,j-n} \right) \right\rangle 
\]
\[ = \left\langle \sum_{k,l} I_{k,l}^2 K_{i-k,j-l}^2 \right\rangle + \left\langle \sum_{k,l} \sum_{k,l \neq m,n} I_{k,l} I_{m,n} K_{i-k,j-l} K_{i-m,j-n} \right\rangle 
\]
\[ = \sum_{k,l} \langle I_{k,l}^2 \rangle K_{i-k,j-l}^2 + \sum_{k,l} \sum_{m,n \neq k,l} \langle I_{k,l} I_{m,n} \rangle K_{i-k,j-l} K_{i-m,j-n}. \] (5.13)

We assume that the pixels in the map are uncorrelated, and thus the expectation value in the 2nd term of Eqn. 5.13 gives 0. The first term reduces to
\[ \xi_O^2(i,j) = \sum_{k,l} \xi_I^2(k,l) K_{i-k,j-l}^2, \] (5.14)
and we have
\[ \xi_O^2 = \xi_I^2 * K^2. \] (5.15)

The pixel variance map $\xi_O^2$ is calculated along with the deconvolved map $\hat{O}$.

We can now estimate the reduction in point-source sensitivity due to the deconvolution process. A point source of flux density $S_\nu$ which is observed by a telescope with a Gaussian PSF of width $\sigma_B$ results in a map $I$ with intensity at pixel $(i,j)$ of
\[ I_{i,j} = S_\nu G_{i,j} + n_{i,j}, \] (5.16)
where
\[ G_{i,j} = \left( \frac{1}{2\pi\sigma_B^2} \right) e^{-\left( i^2 + j^2 \right)/(2\sigma_B^2)} \] (5.17)
and $n_{i,j}$ is the measurement noise, characterized by the noise variance $\xi_n^2 \equiv \langle n_{i,j}^2 \rangle$. Assuming constant noise, not entirely accurate but good enough for the estimate, the least-squares estimate of the source flux density is
\[ \bar{S} = \frac{\sum_{i,j} I_{i,j} G_{i,j}^*}{\sum_{i,j} G_{i,j}^2}. \] (5.18)
Table 5.4: Point Source Sensitivity

<table>
<thead>
<tr>
<th>Band (µm)</th>
<th>Raw Maps</th>
<th>Deconvolved Maps</th>
</tr>
</thead>
<tbody>
<tr>
<td>250</td>
<td>88.2</td>
<td>17.0</td>
</tr>
<tr>
<td>350</td>
<td>87.8</td>
<td>24.7</td>
</tr>
<tr>
<td>500</td>
<td>90.4</td>
<td>31.9</td>
</tr>
</tbody>
</table>

The point source sensitivity \( \xi_s \) is then

\[
\xi_s^2 = \langle \xi^2 \rangle = \frac{\sum_{i,j} \left( \tilde{I}_{i,j}^2 \right) G_{i,j}^2}{\left( \sum_{i,j} G_{i,j}^2 \right)^2} = \frac{\xi_p^2}{\sum_{i,j} G_{i,j}^2}. \tag{5.19}
\]

The denominator is simply the 2-dimensional integral of \( G^2 \), equal to \( 1/(4\pi\sigma_b^4) \), and we find

\[
\xi_s = 2\sqrt{\pi}\sigma_b\xi_p. \tag{5.20}
\]

We present the average \( \xi_p \) in the raw and deconvolved variance maps and the resulting point source sensitivities in Table 5.4. We see that even though the noise per pixel increases by a factor of 10–20 due to the deconvolution, the point source sensitivity decreases by only a factor of 3–4, due to the decreased beam sizes. This is a small price to pay given the huge increase in resolution.

### 5.5 Compact Source Extraction

We now describe the process used to extract flux densities from the deconvolved maps. We first used the Mexican Hat Wavelet (MHW) technique (e.g. Barnard et al. 2004, described below) for locating point source objects obscured by large amounts of diffuse emission, but found that the sources were not well described by a single width across the map. We determined that this was due to the varying PSF, and so instead we perform least-square fits of Gaussians to objects found using the MHW.
5.5.1 Source Identification

The MHW has been used to detect point sources in the presence of large-scale foregrounds. It is a circularly symmetric function with radial profile

$$\psi(r) = \left[ 2 - \left( \frac{r}{\sigma_0} \right)^2 \right] e^{-r^2/(2\sigma_0^2)}, \quad (5.21)$$

which effectively subtracts a local background, much like aperture photometry. The MHW with $\sigma_0$ matched to the PSF is therefore simply just a “compensated” PSF with compact support, which returns zero for a uniform background. The 2-dimensional MHW is

$$\Psi(x, y) = \psi \left( \sqrt{x^2 + y^2} \right). \quad (5.22)$$

A map containing a point source with flux density $S$ smeared by a Gaussian beam with width $\sigma_b$ convolved with the MHW produces a map with value

$$\frac{2S}{\left(1 + \sigma_b^2/\sigma_0^2\right)^2} \quad (5.23)$$

at the position of the source. Setting $\sigma_0 = \sigma_b$, the value of the pixel in the convolved map at the location of the source is $S/2$. We convolve the deconvolved maps with $\psi(r)$, with $\sigma_0$ chosen to correspond with FWHMs of 40, 50 and 60$''$ for the 250, 350 and 500 $\mu$m maps, the widths of $G$ in the deconvolution.

It should be noted that the MHW convolution technique is sensitive to compact sources that are not in fact points, that is, with widths larger than the telescope FHWM. The sensitivity decreases with increasing source width, however, as can be seen by increasing $\sigma_b$ in Eqn. 5.23.

Peaks in the maps are identified by finding all local maxima, defined as pixels with values larger than the 8 surrounding neighbors, above the threshold $\Gamma = N\xi$, where $\xi$ is the average noise per pixel in the maps. $\xi$ is measured by taking the variance of pixel values in a relatively source-free region of the map (dashed box$^3$ in Fig. 5.8). We find $\xi = 140$ mJy at 250 $\mu$m. We also calculate the pixel variance using Eqn. 5.15 setting $K$ equal to $\Psi$. We find results consistent within 50%, but use the former estimate, since it is a direct measurement. The level $N$ is chosen such that, given uniform, uncorrelated Gaussian noise, we expect only about 1 false detection in the map. With $N_p$ map pixels, we wish to find $N$ such that

$$\frac{1}{\sqrt{2\pi}} \int_{-\infty}^{\sqrt{N}} e^{-t^2/2} \, dt = 1 - \frac{1}{N_p} \quad (5.24)$$

$^3$The box is rotated because the analysis was performed in Galactic coordinates.
This integral is the normal cumulative distribution function. We find its inverse by rational approximation\(^4\) (i.e. using a ratio of 2 polynomials) and, with \(N_p \sim 10^5\), find \(N \approx 4.3\).

The deconvolved maps are not accurately described by uniform noise, however. In particular, the noise ripples produced by the deconvolution process lead to many false detections found near bright sources. No robust method to automatically reject these false peaks was found, since the ripples are not uniform across the image. Therefore we reject by hand any peak that appears to be associated purely with the noise ripples. This process is applied to all 3 maps, producing lists of 59, 78 and 50 sources (with significant overlap), in the 250, 350 and 500 \(\mu\)m maps. A complete list of 87 sources is made by taking the union of the 3 lists.

The 3 maps are then co-aligned given the list of point sources. The 350 and 500 \(\mu\)m maps are aligned with the 250 \(\mu\)m map by performing a cross-correlation (Eqn. 5.5) of the central pixels of each source. This alignment procedure is more robust than the one described in Sec. 5.3, as the beams in the deconvolved maps are all centrally peaked, and significantly more compact than in the raw maps.

\subsection*{5.5.2 Source Fitting}

We fit circularly-symmetric Gaussian profiles to each of the sources found in the previous section. As well as fitting for integrated flux density at each wavelength, we allow the width of the Gaussians to vary to accommodate the varying PSF. It is not thought that the varying profile widths needed to fit the sources is an indication that the sources are extended, as they mostly appear point-like in higher-resolution observations (e.g. with MSX, see Sec. 5.6.2.3). We also fit for source position, providing sub-pixel precision on source location. We fit the model to all 3 maps simultaneously, since each source should be characterized by a single width. The model for intensity at pixel \((i,j)\) at wavelength \(b\) is then

\[
\tilde{I}_b(i,j) = \sum_s \frac{S_{b,s}}{2\pi \sigma_{b,s}^2} \exp \left( -\frac{(i-x_s)^2 + (j-y_s)^2}{2\sigma_{b,s}^2} \right),
\]

where \((x_s,y_s)\) is the position of source \(s\), \(\sigma_{b,s}\) is the width of source \(s\) at band \(b\), and the sum is over all \(N_s\) sources. The \(\sigma_{b,s}\) depend on a single source width \(\sigma_s\), given as the width of the convolution of 2 Gaussians with width \(\sigma_b\) (beam) and \(\sigma_s\) (source):

\[
\sigma_{b,s} = \sqrt{\sigma_b^2 + \sigma_s^2}.
\]

\(^4\)Using the algorithm presented at \url{http://home.online.no/~pjacklam/notes/invnorm/}
We solve for the best-fit parameters $S_{b,s}$, $x_s$, $y_s$, and $\sigma_s$ by least-squares. We form $\chi^2$ as the sum of the squares of the residuals in the 3 maps,

$$\chi^2 = \sum_b \chi_b^2$$

$$\chi_b^2 = \frac{\sum_{i,j} \left( I_b(i,j) - \tilde{I}_b(i,j) \right)^2 w_b(i,j)}{\sum_{i,j} w_b(i,j)},$$

where $I_b$ is the data, $\tilde{I}_b$ the model and $w_b$ the pixel weight, equal to the inverse of the variance map $\xi^2$, all at band $b$. To better constrain these non-linear fits, we extend $\chi^2$ to include constraints on source position $(x_s, y_s)$ and width $\sigma_s$. We add the terms $\chi_{\text{pos}}^2$,

$$\chi_{\text{pos}}^2 = \sum_s \frac{(x_s - \hat{x}_s)^2 + (y_s - \hat{y}_s)^2}{\alpha_{\text{pos}}^2},$$

which adds a penalty for source positions significantly different that the initial guesses $(\hat{x}_s, \hat{y}_s)$, and $\chi_{\text{sig}}^2$,

$$\chi_{\text{sig}}^2 = \sum_s \left( \frac{\sigma_s}{\alpha_{\text{sig}}} \right)^4,$$

which constrains the source widths to small values. This gives the final form

$$\chi_T^2 = \sum_b \chi_b^2 + \chi_{\text{pos}}^2 + \chi_{\text{sig}}^2.$$

We treat $\alpha_{\text{pos}}$ and $\alpha_{\text{sig}}$ as tuned parameters, set to 0.08 and 5 pix, respectively.

In principle, we could solve for all $N_s \times 6$ parameters simultaneously. This is unnecessary, however, as the parameters for sources sufficiently far apart are independent, that is, $\tilde{I}_b(i,j)$ depends only on sources near pixel $(i,j)$. We group the nearby sources and fit each set independently. The groups are defined by drawing a circle centred on each source, and sources with overlapping circles are grouped together. The size of each circle depends on the brightness of the source, and is meant to enclose all of the map pixels for which the source contributes significantly to $\tilde{I}_b$. If the signature of each source was purely Gaussian, we could simply use the radius at which the flux density falls to some fraction of the map noise. Because of the deconvolution ripples, however, we need a function that falls off much more slowly. We choose the function

$$R(S) = \begin{cases} \sqrt{S/(n \xi)} \sigma_b, & S > S_0 \\ R_0, & S \leq S_0 \end{cases}$$

(5.32)
where $S$ is an estimate of the source flux density, $\xi$ is the map noise, $\sigma_b$ is the nominal beam width, $n$ and $R_0$ are tuned parameters, and $S_0$ is chosen such that $\sqrt{S_0/(n \xi)} \sigma_b = R_0$. We choose $n = 1.7$ and $R_0 = 3$ pix, and produce 78 regions, each containing 1–3 sources.

We perform the fit to each group of sources over a region of the map large enough that the local background can be estimated. A square of size $N_{\text{reg}} \times N_{\text{reg}}$ is centred on each source in the group, and the region is defined as the union of the overlapping squares. We use $N_{\text{reg}} = 40$ pix. Sources not in the group but falling within the region are masked out by excluding pixels within $R(S)$ of the source. A 4th order 2-dimensional polynomial is fit to the map region with all sources masked out, whether they are in the group or not. The background fit is subtracted from the map. $\chi^2_T$ is then minimized, where the sums in Eqs. 5.25, 5.29 and 5.30 are over all sources $s$ in the group, and the sums in Eqn. 5.28 are over all pixels $(i, j)$ in the region but not within $R(S)$ of sources not in the group. We use the routine AMOEBA (Press et al. 1998), based on the multidimensional downhill simplex method (implemented in IDL), to minimize $\chi^2_T$. The parameters are initialized to our best estimates of their true values: $S_{b,s}$, $x_s$ and $y_s$ from the MHW convolution, and $\sigma_s = 0$. A fit for one example region is shown in Fig. 5.9.

We see structure in the residuals, but on a scale much smaller than the flux density in the sources. The remaining flux density $S_r$ in the residual maps $r$, calculated as

$$S_r = \sqrt{\sum_{i,j} r(i,j)^2}, \quad (5.33)$$

where the sum is over the pixels in the map, is 32, 11 and 7.4 Jy compared to 480, 175 and 78 Jy within the two sources, in the 250, 350 and 500 $\mu$m maps, respectively. We calculate the reduced $\chi^2$ in the fit, defined as

$$\chi^2_{\text{red.}} = \sum_b \frac{1}{N_b} \sum_{i,j} \left( \frac{r_b(i,j)}{\xi_b(i,j)} \right)^2, \quad (5.34)$$

where we sum over bands $b$, $N_b$ is the number of pixels in the band $b$ map, and $\xi_b(i,j)$ is the noise in pixel $(i,j)$ at band $b$. The result is $\chi^2_{\text{red.}} = 2.7$, which is significantly different than 1.0, but not terribly surprising, since the deconvolved maps are not well described by independent Gaussian noise.

Finally, we perform a second cut on the source list, as the fits do not converge for some of the faintest sources as well as a few of the sources located near the edge of the map or within
Figure 5.9: The results of the source-fitting algorithm for a sample region. The left column shows the deconvolved maps, the middle column the best-fit model to all 3 bands, and the right column the residuals of the fit. A background polynomial is removed from the map before the fit is performed and is not included in the residuals. Note that the map contrast levels have been clipped to show structure at small scales. The peak intensities are 4100, 700 and 350 MJy sr$^{-1}$ in the 250, 350 and 500 $\mu$m raw maps.
the deconvolution ripples of brighter sources. We calculate a signal-to-noise ratio across all maps for each source,

\[ r_s = \sqrt{\sum_b \left( \frac{\hat{S}_{b,s}}{\xi_b} \right)^2}, \]

(5.35)

where \( \hat{S}_b \) is the estimate of the flux density for source \( s \) at band \( b \) from the MHW convolution and \( \xi_b \) is the noise at band \( b \). We reject all sources with \( r_s < 5.75 \) and one source just above this limit, but located near the ripples around a brighter source, producing a final list of 60 sources. The value 5.75 is chosen as it represents the limit above which (nearly) all of the fits are reliable. The source locations are shown on the 350\,\mu m map in Fig. 5.10.

5.5.3 Simulations

We characterize the source identification and flux density fitting routines described above by a series of simulations. Ideally, one would like to “observe” a simulated sky with the BLAST beams using the real pointing solutions and create a map using the optimal map-maker. The computations of each of these steps are prohibitively expensive, however, as the first requires a convolution at every time step and the second has already been discussed. We instead used a much simpler simulation, that of inserting a source (convolved with the PSF) into the raw maps, running the deconvolution, source identification and flux density extraction routines, and comparing the output positions and flux densities with the input. We run this simulation many times with the input source inserted into a different random location in the map. We can thereby test for: (i) flux density bias, whether the output flux density tends to be systematically different from the input flux density, and obtain uncertainties in flux density at the same time; (ii) positional accuracy; and (iii) completeness, the fraction of the input sources that are found by the identification routine, as a function of input flux density. The simulations are run over a range of input source flux densities spanning the range of flux densities for the detected sources.

The source extraction routine is modified slightly from that described above in order to avoid the slow procedure of having to manually reject unreliable peaks in residual ripples around bright sources. We start with the full list of peaks identified by the MHW convolution. The list is examined one item at a time, starting with the brightest. The threshold \( \Gamma \), previously set to \( N\xi \), is modified in the neighborhood of the bright source. We start with \( \Gamma_0 = N\xi \) and
Figure 5.10: Positions of the 60 BLAST-detected sources overlaid on the 350 μm deconvolved map. The different symbols indicate the distances we determine to each source, either associated with the molecular cloud surrounding the open cluster NGC 6823 at 2.3 kpc (circles), the Perseus arm at 8.5 kpc (triangles), or at 14 kpc in the outer galaxy (square). See Sec. 5.6.4. [This figure appears in Chapin et al. (2007).]
update it as we step through the list,

$$\Gamma_{n+1}(i,j) = \begin{cases} f(i,j;S), & f(i,j;S) > \Gamma_n(i,j) \\ \Gamma_n(i,j), & f(i,j;S) \leq \Gamma_n(i,j) \end{cases}. \quad (5.36)$$

Here, $f(i,j;S)$ is a function which depends on the source flux density $S$, as determined by the MHW convolution. Any source falling below $\Gamma_{n+1}$ is rejected from the list, and the next-brightest source is found. This is repeated until there are no remaining unexamined sources.

We choose a Gaussian as the functional form,

$$f(i,j;S) = \eta S \exp \left( -\frac{(i-x_0)^2 + (j-y_0)^2}{2\sigma^2} \right), \quad (5.37)$$

where $(x_0,y_0)$ is the source location and $\eta$ and $\sigma$ are tuned parameters. We find $\eta = 0.22$ and $\sigma = 9.3$ pix provide the best results. This modified routine is able to reproduce the initial by-hand source list to an accuracy of $\sim 10\%$.

### 5.5.3.1 Flux Bias

We expect the flux densities extracted by the routine described above to be biased for two reasons: (i) the deconvolution transfers power from the source into the ripples, biasing the extracted flux low; and (ii) allowing the source width $\sigma_s$ to vary biases the reconstructed flux high. The first effect is of the order 5%, and the second as large as $\sim 20\%$.

In order to extract the flux density bias and errors, 500 iterations are performed at each of several input fluxes. At each iteration, a 250 $\mu$m flux is chosen and the 350 and 500 $\mu$m fluxes are assigned based on the median source colours found in the field, i.e. $S_{350} = S_{250}/2.40$ and $S_{500} = S_{250}/5.14$. The simulated point source is convolved with each beam and inserted into a randomly-determined location in the maps. The maps are deconvolved, as described above, and the MHW source identification routine with the modified threshold $\Gamma$ is applied. A simulated source is considered detected if: (i) the new list contains an additional entry compared to the original list; and (ii) the new list contains no new entries, but the location of the input source coincides with a real source and the new source is brighter than the original.

If the source is detected, the flux extraction routine of Sec. 5.5.2 is applied. The sources are divided into groups as before, but only the region containing the simulated source is fit. The best-fit parameters are recorded and the next iteration performed. If the source is not detected, the iteration is not counted against the 500 and we try again with another input location.
Figure 5.11: The distribution of flux density biases at 250\,\mu m resulting from 500 iterations of $S_{250} = 100\,\textrm{Jy}$ sources inserted into the raw maps. The dashed lines show the 68\% confidence limits and the dotted line is the midpoint between these limits. A Gaussian (dash-dotted line) centred on the midpoint and with width equal to the 68\% limits describes the histogram well.

The resulting output flux density distributions are roughly Gaussian with a slight positive tail. The distribution for one of the sets of simulations is shown in Fig. 5.11. The Bayesian 68\% confidence limits are found and the distributions are fit by the Gaussian that passes through these points. We find that the measured flux density is biased high by a factor of about 1.2 at 250\,\mu m, no bias at 350\,\mu m, and a factor of about 1.1 at 500\,\mu m, nearly independent of input flux density (Figure 5.12). We model these results as a constant bias in flux and with a 2-component error envelope. The bias increases at low flux, but only by a small fraction of the error envelope, so we are justified in treating it as constant. The bias is calculated as the mean of the biases at each input flux, weighted by the errors, giving biases of 1.19, 1.02 and 1.08 at 250, 350 and 500\,\mu m, respectively. The error envelopes are fit by the function

$$\sigma_{\text{bias}} = \sqrt{\left(\frac{A}{S_{\text{in}}}ight)^2 + B^2}, \quad (5.38)$$

where $A$ and $B$ are free parameters. The first term represents error due to noise in the map, while the second is due to the fitting process. The models are shown as dashed lines in the figure. We use these bias measurements to correct the extracted fluxes and we use the error
Figure 5.12: Results of the flux bias simulations. Thick error bars are the 1-σ widths of the distributions. The thin error bars are the 1-σ widths scaled by a factor of 5 for visibility at the bright end. We see that the reconstructed flux is biased high at 250 and 500 μm, but that the bias is mostly independent of flux. A small increase at low flux is found, but this is well within the error envelope. The error envelopes (dashed lines) are fit with a 2-component phenomenological model.
model as the uncertainty of the measurement. The resulting calibrated fluxes and statistical errors for each source are listed in Table 5.5.
### Table 5.5: Fluxes and Uncertainties for BLAST-detected Sources

<table>
<thead>
<tr>
<th>BLAST ID</th>
<th>Source name</th>
<th>$F_{250}$ (Jy)</th>
<th>$\sigma_{250}$ (Jy)</th>
<th>$F_{350}$ (Jy)</th>
<th>$\sigma_{350}$ (Jy)</th>
<th>$F_{500}$ (Jy)</th>
<th>$\sigma_{500}$ (Jy)</th>
<th>$V_{lsr}$ (km s$^{-1}$)</th>
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<tbody>
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<td>V01</td>
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<td>4.2</td>
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Table 5.5 – continued from previous page

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$^a$These sources are located on ripples in the deconvolved map and the colors are considered unreliable.

$^b$V07 is believed to lie in the outer galaxy.

$^c$These sources are associated with a molecular cloud in the Perseus arm.

$^d$Also has a comparable component at $-15$ km s$^{-1}$.

$^e$Also has a comparable component at $23$ km s$^{-1}$. 
Figure 5.13: Positional uncertainty distribution for 500 simulations. Simulated sources have fluxes of 200 Jy. An azimuthally-integrated 2-d Gaussian profile is fit to the data and overplotted.

5.5.3.2 Positional Uncertainty

We also investigate the positional accuracy of the source extraction routine. The distance between the input and fit source positions $\Delta r$ are calculated for each of the simulations described above. We fit the resulting distributions with a 2-d circular Gaussian, integrated over the azimuthal angle $\theta$,

$$
\psi(r) = N_{\text{sim}} \frac{r}{\sigma_{\text{pos}}^2} \exp \left( -\frac{r^2}{2\sigma_{\text{pos}}^2} \right),
$$

(5.39)

where $N_{\text{sim}}$ is the number of simulations run and $\sigma_{\text{pos}}$ is the parameter describing the width of the distribution. See Fig. 5.13 for an example distribution. The positional errors for all input fluxes are shown in Fig. 5.14. We find that the positional errors vary from $\sim 3''$ at the bright end to $\sim 30''$ at the faint end.

5.5.3.3 Completeness

We explore the completeness of the BLAST-detected compact sources using the same simulations described above, except that the flux extraction fits do not need to be carried out. At each of the input fluxes, 1000 simulations are performed and the number of times that the input is detected is counted. The results are shown in Figure 5.15.
Figure 5.14: Positional uncertainty for each of the input fluxes. The values plotted are the $\sigma_{\text{pos}}$ estimates from Eqn. 5.39.

Figure 5.15: Source extraction completeness as a function of source flux. The completeness is nearly 1 at high flux densities and drops off at low fluxes, as one might expect. [This figure appears in Chapin et al. (2007).]
5.6 Results and Discussion

We now attempt to constrain the intrinsic properties of the BLAST-detected sources. We first look at what can be said with BLAST data alone, then attempt to cross-identify the sources with measurements at infrared wavelengths. The spectral energy distributions (SEDs) are then calculated, distances to the sources estimated, and luminosities and masses determined.

5.6.1 Colours

We use a colour-colour plot as a diagnostic of the emission properties of sources that BLAST detected in the field. We model the submillimetre emission of the sources with the modified blackbody function

$$S_\nu = A \left( \frac{\nu}{\nu_0} \right)^\beta B_\nu(T),$$

where $S_\nu$ is the flux density at frequency $\nu$, $A$ is proportional to flux, $\beta$ and $T$ are intrinsic properties of the source, and $B_\nu(T)$ is the Planck function giving the intensity at frequency $\nu$ for a blackbody of temperature $T$. The emissivity index $\beta$ is typically found to fall in the range 1–2. The colour

$$\frac{S_{\nu_1}}{S_{\nu_2}} = \left( \frac{\nu_1}{\nu_2} \right)^\beta \frac{B_{\nu_1}(T)}{B_{\nu_2}(T)}$$

then allows for the discrimination of sources based purely on the intrinsic properties $\beta$ and $T$.

Before the colours of the sources can be compared, however, we must correct the flux densities for the fact that the BLAST bands have large spectral widths. Because the bands are broad, the conversion from the band-average flux density $S_B$ to the true flux density $S_{\nu_1}$ at frequency $\nu_1$ (nominally the centre of the band) depends on the SED of the source. Given the model SED $\tilde{S}_\nu$ (described in Sec. 5.6.3), we calculate the model band-averaged flux density

$$\tilde{S}_B = \int \tilde{S}_\nu T_\nu d\nu,$$

where $T_\nu$ is the area-normalized band transmission function. The true measured flux density is then

$$S_{\nu_1} = S_B \frac{\tilde{S}_{\nu_1}}{\tilde{S}_B}.$$  

We show $S_{250}/S_{350}$ against $S_{250}/S_{500}$ for each source in Fig. 5.16. Errors on the colours are calculated based on the 1-σ statistical errors determined by the simulations (Sec. 5.5.3.1). 3 sources are not shown in the figure as they fall outside of the plot, due to faint 350 and
Figure 5.16: A colour-colour plot for the BLAST-detected sources. Sources with anomalous colours due to the varying deconvolution ripples are not included. Error bars are based on the 1-σ statistical errors determined by the simulations, and are shaded so that the more significant detections are darker. 3 sources with large colours, due to faint 350 and 500 μm flux densities, lie outside of the plot range. Also plotted are modified blackbody models with: constant $\beta$ equal to 0, 1, 2, and 3, increasing towards the bottom, and temperatures ranging from 5 to 200 K (solid lines); and a model with constant temperature of 20 K and with $\beta$ ranging from −1 to 3 (dashed line). [This figure appears in Chapin et al. (2007).]
500 µm flux densities. Eleven sources, nine of which would lie in the plot, are located within the deconvolution ripples and thus have suspect colours. These sources are omitted from the plot and are denoted in Table 5.5; they are treated separately in the analysis that follows.

We overplot model SEDs based on the modified blackbody function (Eqn. 5.40). Curves with constant $\beta$, equal to 0, 1, 2 and 3, and with $T$ ranging from 5 to 200 K are shown as solid lines. A curve with a constant temperature $T = 20$ K and $\beta$ ranging from −1 to 3 is shown as a dashed line. We see that sources lie in significantly distinct regions of the parameter space, but that $\beta$ and $T$ are nearly indistinguishable. We note, however, that $\beta$ ranging from 1–2 fits the majority of the sample. The BLAST data alone cannot accurately constrain the SED corresponding to thermal emission in many of the sources from this survey. To proceed further, measurements of the BLAST sources at other wavelengths are needed.

### 5.6.2 Cross-identification

We now examine published mid-infrared (MIR) observations of the Vulpecula field in order to better constrain the SEDs of the BLAST-detected sources.

#### 5.6.2.1 IRAS

We use the all-sky MIR observations by the Infra-red Astronomical Satellite (IRAS) at 12, 25, 60 and 100 µm. We search the IRAS Point Source Catalog version 2.0 (PSC) (Helou and Walker 1988) for sources coinciding with BLAST detections; specifically, we look for all sources in the PSC within a certain radius of the BLAST sources. This radius is the quadrature sum of the BLAST uncertainties, as determined by the positional uncertainty simulations (Sec. 5.5.3.2), with the semi-major axes of the PSC error ellipses. We conservatively set the BLAST error at the bright end to $13''$ (equivalent to FWHM = 30''). We find identifications in the PSC for 23 of the 60 BLAST sources.

For the remaining sources, we extract flux densities directly from the IRAS maps. We use the IRAS Galaxy Atlas (IGA) (Cao et al. 1997) at 60 and 100 µm and the Mid-Infrared Galaxy Atlas (MIGA) (Kerton and Martin 2000) at 12 and 25 µm, produced by the resolution-enhancing algorithm HIRES (Aumann et al. 1990). Flux densities are measured by performing aperture photometry in the maps at the BLAST source locations. In some cases, a clear detection is made. In cases where there is no apparent source or in crowded fields, an upper
limit is quoted. A total of 28 of the 60 BLAST sources are detected at 60 or 100 μm by IRAS, either from the PSC or by direct measurement in the IGA.

5.6.2.2 MIPS

In cases where IRAS 60 or 100 μm detections are not available, we search the 70 μm MIPSGAL (Spitzer Space Telescope) (Carey et al. 2005) images. Aperture photometry is performed in the map at the BLAST source locations, as with the IRAS maps. MIR detections for an additional 10 BLAST sources are made from the MIPSGAL maps.

5.6.2.3 MSX

We additionally search the Midcourse Space Explorer (MSX) catalogs at 8, 12, 14 and 21 μm, compiled from maps with higher angular resolution than that of BLAST. We use the MSX Point Source Catalog version 2.3 (Egan et al. 2003), searching for counterparts in the same manner as with the IRAS PSC. We find identifications for 40 of the 60 BLAST sources. In some cases, 2 or more MSX counterparts are found — here, we take the sum of the flux densities as an upper limit.

5.6.3 Spectral Energy Distributions

Given the BLAST flux densities along with MIR measurements, we attempt to constrain the SEDs of all 60 sources. We fit the modified blackbody function (Eqn. 5.40) to data at \(\lambda > 100\) μm and interpolate data at \(\lambda < 100\) μm in log-log space. We calculate the fit by minimization of \(\chi^2\), using a full covariance matrix as weights. For the BLAST data, we use the covariance matrix calculated in Sec. 4.2.3 describing the errors due to calibration, and add the statistical uncertainties of Sec. 5.5.3.1 along the diagonal. For the MIR data, we use the measurement errors given in the catalogues, or as determined from the noise in the maps. We include upper limits in the fit using “survival analysis” (see CHA07). Since \(\beta\) is not well-constrained, it is fixed to 1.5 in all fits.\(^5\) For the 11 sources whose colours are not well-determined by BLAST, \(T\) is fixed to 20 K. Example fits for two sources are shown in Fig. 5.17. The distribution of inferred source temperatures, excluding the 11 sources with unreliable colours, is shown in Fig. 5.18.\(^6\)

\(^5\)Values of 1–2 are seen in the literature (e.g. Calzetti et al. 2000). Adopting \(\beta = 2\) would decrease our inferred temperatures by \(\sim 5\) K and increase cloud masses (Sec. 5.6.5) by a factor of \(\sim 2\).
Figure 5.17: Example SED fits for two of the BLAST sources. The fit to the brightest source (V30) is shown on the left and to the coldest source (V11, $T \simeq 10$ K) on the right. The error envelopes are for the modified blackbody component only, which is fit to data with $\lambda \geq 100$ $\mu$m. [These figures appear in Chapin et al. (2007).]

Figure 5.18: The distribution of dust temperatures for the 60 BLAST-detected sources, excluding 11 with unreliable colours. If we choose $\beta = 2$ instead of 1.5, the inferred temperatures decrease by $\sim 5$ K. [This figure appears in Chapin et al. (2007).]
Errors in the parameters and 1-σ error envelopes (see figure) are calculated by Monte Carlo simulations. Mock data are generated from realizations of Gaussian noise, including both correlated and uncorrelated errors described by the covariance matrix, scattered about the measurement. The fits are performed at each iteration, and the 68% confidence envelope is found. Best-fit temperatures and 1-σ errors are listed in Table 5.6.

Finally, we calculate an estimate of the bolometric flux $S$ of each source by integrating the SED $S_\nu$ from 2–5000 μm.
### Table 5.6: Source Parameters From SED Fits

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<th>$L_{\text{FIR}}$ ($L_\odot$)</th>
<th>Lum. Fraction$^a$ (%)</th>
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<td>84</td>
<td>26.5 ± 8.8</td>
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<td>233 ± 45</td>
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<td>107 ± 11</td>
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<td>145 ± 43</td>
<td>...</td>
<td>1.5 ± 0.7</td>
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<td>0.4 ± 0.4</td>
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<td>...</td>
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<td>103 ± 21</td>
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Continued on next page
### Table 5.6 – continued from previous page

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<th>BLAST ID</th>
<th>Temperature (K)</th>
<th>$M_c$ (M$_\odot$)</th>
<th>$L_{\text{FIR}}$ (L$_\odot$)</th>
<th>Lum. Fraction</th>
<th>$L_{\text{FIR}}/M_c$ (L$<em>\odot$ M$</em>\odot^{-1}$)</th>
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<td>3400 ± 390</td>
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<tr>
<td>V60</td>
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<td>65 ± 28</td>
<td>42 ± 21</td>
<td>34</td>
<td>0.6 ± 0.5</td>
</tr>
</tbody>
</table>

$^a$The percentage the bolometric luminosity ($L_{\text{FIR}} + L_{\text{IR}}$) made up of the FIR luminosity.

$^b$V07 is believed to lie in the outer galaxy.

$^c$These sources are associated with a molecular cloud in the Perseus arm.
Figure 5.19: Sample $^{13}$CO spectra (Brunt and Heyer 2007) for two BLAST sources. V09 (solid line) is assigned a velocity of 34.4 km s$^{-1}$. V20 (dashed line) is assigned a velocity of 0.8 km s$^{-1}$.

5.6.4 Distances

In order to determine luminosities and masses from the inferred bolometric fluxes, we must first find the distance to each source. We examine $^{13}$CO($1\rightarrow0$) spectra of the Vulpecula region (Brunt and Heyer 2007), with spatial and velocity resolutions of 46'' and 1 km s$^{-1}$, respectively. If the line-of-sight velocity of a source can be determined, its distance can be inferred by referring the velocity to a Galactic rotation curve, assuming that the rotation models accurately describe the disk structure and that peculiar velocities are negligible.

The $^{13}$CO spectrum at the location of each BLAST source generally shows a single compact peak. See Fig. 5.19. In such cases, we identify the centre of the peak as the source’s velocity. For sources with multiple $^{13}$CO peaks, it might be possible to correlate the peaks with the multiple MSX sources (Sec. 5.6.2.3), but this was not attempted. The velocities determined for each source are listed in Table 5.5. Our velocity measurements agree well with previous measurements in CS (Bronfman et al. 1996; Beuther et al. 2002) and NH$_3$ (Molinari et al. 1996; Zinchenko et al. 1997; Sridharan et al. 2002), on average within 1 km s$^{-1}$. In cases where multiple peaks are found, other measurements of H\textsc{i} absorption spectra from the VLA Galactic Plane Survey (VGPS) (Stil et al. 2006) together with C$^{18}$O spectra (Brunt and Heyer 2007)
are used to further constrain the ambiguities.

The distribution of velocities for the 60 BLAST sources is shown in Fig. 5.20. We conclude that 49 of the 60 sources lie in a molecular cloud complex, within which the open cluster NGC 6823 has already formed. The radio recombination line velocity and width of the H\(\text{II}\) region Sh2-86, clearly associated with the open cluster, are 29.4 km s\(^{-1}\) and 24.2 km s\(^{-1}\), respectively. This curve is shown in the left panel of the figure and clearly encompass the grouping of BLAST sources. We adopt the distance to the cluster of \(d = 2.3\) kpc (Massey et al. 1995) for these sources. A second grouping of 10 sources at \(-5\) km s\(^{-1}\) are associated with the Perseus arm, determined to be at a distance of 8.5 kpc using the Galactic rotation curve of Brand and Blitz (1993). A final source with \(v = -54\) km s\(^{-1}\) is inferred, again using the rotation curve, to be at \(d = 14\) kpc, in the outer Galaxy. Sources associated with the Perseus arm and the outer Galaxy are indicated in Table 5.5.

5.6.5 Luminosities and Masses

Given the SED fits of Sec. 5.6.3 and the distances determined in the previous section, we can now calculate the mass and luminosity of each of the BLAST-detected sources. The bolometric thermal luminosity is simply calculated from the integrated flux, \(L = 4\pi d^2 S\). Mass is determined from \(A\), the amplitude parameter in Eqn. 5.40,

\[
A = \frac{M_c \kappa_0}{R d^2} \tag{5.44}
\]

where \(M_c\) is the total clump mass, \(\kappa_0\) is the dust mass absorption coefficient evaluated at \(\nu_0\), \(R\) is the gas-to-dust mass ratio and \(d\) is the distance to the source. We use \(\kappa_0 = 10\) cm\(^2\) g\(^{-1}\) at \(\nu_0 = c/250\) \(\mu\)m and assume \(R = 100\) (Hildebrand 1983). The derived luminosities and masses are listed in Table 5.6. Luminosity vs. mass is plotted for the 49 objects associated with NGC 6823 in Fig. 5.21. The 10 objects associated with the Perseus arm are shown in Fig. 5.22. Errors on luminosity and mass are determined using the same Monte Carlo simulations used to calculate the errors on parameters in the SED. Loci of constant temperature and 250 \(\mu\)m flux density are overplotted, assuming modified blackbody emission with \(\beta = 1.5\) and the dust mass parameters as described above. Note that the 11 sources with unreliable colours (10 in NGC 6823 and 1 in the Perseus arm) have temperatures fixed to 20 K and thus lie along the 20 K isotherms.
Figure 5.20: The distribution of velocities for all 60 BLAST sources (left panel, histogram), excluding one source at $-54\text{ km s}^{-1}$. The dashed line shows the integrated spectrum across the whole $^{13}$CO map, including more diffuse gas. The scale of the spectrum is given on the right vertical axis, in antenna temperature. The solid curve is a Gaussian profile with centre and width corresponding to radio recombination line measurements of the H II region Sh2-86 (Lockman 1989), which encompasses the open cluster NGC 6823. The vertical scale of this curve is arbitrary. The distribution of the central cluster of sources is consistent with the emission profile from Sh2-86. The remaining sources are believed to lie in the Perseus arm. At right, the inferred distances corresponding to the source velocities as given by the Galactic rotation model of Brand and Blitz (1993). The dark band contains the NGC 6823 sources and the light band the Perseus arm sources. The dotted line at $-54\text{ km s}^{-1}$ indicates an object which we place at 14 kpc, based on this model. [These figures appear in Chapin et al. (2007).]
Figure 5.21: Luminosity vs. mass for the 49 objects associated with NGC 6823. The numbers correspond to the BLAST IDs listed in Tables 5.5 and 5.6. Uncertainties in luminosity and mass are determined by Monte Carlo simulations. Loci of constant temperature and constant 250 μm flux density are overplotted. Sources without MSX and IRAS identifications are labelled with circles and squares, respectively. Objects identified as ultra-compact H II regions are marked with triangles. 10 objects with unreliable BLAST colours have been assigned a temperature of 20 K. [This figure appears in Chapin et al. (2007).]
Figure 5.22: Luminosity vs. mass for the 10 objects associated with the Perseus arm. Annotations are as in Fig. 5.21. One object (V19) with unreliable BLAST colours has been assigned a temperature of 20 K. [This figure appears in Chapin et al. (2007).]
We see that the BLAST-detected sources in NGC 6823 cover a broad range in luminosity and mass. The potentially most interesting objects detectable by BLAST are high-mass protostellar objects (HMPOs), massive heavily obscured regions in the very early stages of star formation. These objects are seen through their cold dust emission and thus are undetectable in the MIR. Objects found in the middle-lower region of Fig. 5.21 without MSX or IRAS detections, are potential HMPOs. Another class of object found by BLAST is ultra-compact (UC) H\textsc{ii} regions, marked with triangles in the plots. Such objects are in a much later stage of star formation and are thus hotter and show significant MIR and radio flux. These objects are identified by searching for counterparts in radio data, using the NRAO VLA Sky Survey (Condon et al. 1998) at 1.4 GHz, together with the second and third MIT-Green Bank 5 GHz surveys (Langston et al. 1990; Griffith et al. 1990) and with the Arecibo Telescope (Watson et al. 2003). We find 2 UC H\textsc{ii} objects (sources V02 and V05) in NGC 6823 and 1 (source V23) in the Perseus arm. A further two objects in NGC 6823 V18 and V30, are identified as Hyper Compact (HC) H\textsc{ii} objects, a less-extreme case of the UC H\textsc{ii} (Keto 2007). From Figs. 5.21 and 5.22, it can be seen that these correspond to most, but not all, of the brightest and hottest sources in the field.

5.6.6 Mass Function

We present the molecular cloud mass function for the sources measured by BLAST. The mass function for the 49 NGC 6823 sources is plotted in Fig. 5.23. The bins are logarithmically spaced in mass. Poisson error bars are indicated. At the high-mass end, our results are consistent with measurements of molecular clouds in CO emission by Kramer et al. (1998), who found a power law \( \frac{dN}{dM} \propto M^{-\alpha} \) with \( \alpha = 1.7 \) over a similar mass range. This relation is scaled and overplotted in Fig. 5.23. We see that the BLAST data are consistent with the Kramer et al. (1998) data at high masses, but show a turn-over at \( M = 200–400 \, M_\odot \). It is not entirely clear whether this turn-over is physical or whether it is due to flux density completeness (Sec. 5.5.3.3). In order to convert flux density completeness to mass completeness, we must assume a temperature for the sources that were missed. The mass completeness is shown in Fig. 5.23 for a variety of temperatures. We see that if the missed sources are mostly 10–15 K, then the turn-over is easily explained by source completeness. Beltrán et al. (2006) (and references therein) find evidence for a steeper power law (\( \alpha \sim 2.1 \)) at high mass and a shallower slope (\( \alpha \sim 1.5 \)) at lower masses. We note that Beltrán et al. (2006) assume much higher dust
Figure 5.23: Mass function of the 49 sources associated with NGC 6823. A power-law found by Kramer et al. (1998) found for molecular clouds in a similar mass range is shown as a dashed line. The BLAST data are consistent with the Kramer et al. power law at high masses, but show a turn-over at 200–500 $M_\odot$. This turn-over is possibly physical, but is likely due to source detection completeness. Estimates of the completeness assuming a range of source temperatures are plotted as dotted lines. [This figure appears in Chapin et al. (2007).]
temperatures than found by the BLAST observations.

5.7 Discussion

Observations of star forming regions with BLAST are unique in their ability to directly constrain temperature. Previous observations with SCUBA either assume a canonical dust temperature (e.g. Johnstone et al. 2000) or assign temperatures based on gas emission lines (e.g. Moore et al. 2007), where available. Since the shape of the mass function depends on temperature of each object, accurate constraints of the mass function require accurate measurements of clump temperatures. The measurements presented here are an important step in the direction of precise mass function determination, which will lead to a better understanding of star formation processes. 20 of the 60 objects found by BLAST are cooler than 20 K; many of these are not seen by previous observations at shorter wavelengths. BLAST’s ability to detect these cold molecular clouds, and in particular the radio-quiet high-mass proto-stellar objects, will help us to understand the role of these objects in massive star formation (e.g. Beltrán et al. 2006; Thompson et al. 2006).

Follow-up observations at other wavelengths will help to further classify the types of objects detected by BLAST. High-resolution spectral observations of molecular lines will provide more accurate temperature and mass estimates, as well as determine chemical composition, structure and dynamics of the cores, all of which will help to constrain star formation models.

We have demonstrated BLAST’s ability to detect the coolest star-forming objects in Galactic fields, despite significant degradation of the telescope optics. The optics were fixed for the second science flight and, as discussed in Sec. 3.6.3, several Galactic fields were observed. In particular, a 40 deg$^2$ region in the Galactic plane was mapped to a depth similar to the that reached in the Vulpecula region, but with the nominal telescope resolution. At 250 µm, we will be $\sim$ 5 times more sensitive to point sources. This is increase in point source sensitivity, along with the increased map area, will allow us to produce catalogues of 10–100 times as many sources, providing many more HMPO candidates and vastly improving statistics in the mass function. Analysis of these data is underway.
Chapter 6

NGC 4565

As well as studying star formation in our own Galaxy, star formation occurring in local galaxies can also be probed with BLAST. Interstellar dust found in the disks of quiescent spiral galaxies is intimately related to star formation, and understanding the amount and distribution of dust will further our understanding of star formation in quiescent systems.

During the 2005 flight, BLAST observed the nearby edge-on Sb spiral galaxy NGC 4565, located at $\alpha = 12:36:21$, $\delta = +25:59:13$. NGC 4565 is well-studied at a variety of wavelengths, including optical, near-IR, mid-IR, far-IR, millimetre and radio. The galaxy is about $10' \times 2'$ in the optical, and thus should be well resolved by BLAST along the major axis. It is at an inclination angle of $87^\circ$ (de Vaucouleurs et al. 1991) and has a warp in the disk at a radius of about $7.5'$ (Rupen 1991). The warp is visible on both sides of the disk, but is more dramatic in the North. We adopt a distance of 10 Mpc, determined by Simard and Pritchet (1994) using the surface brightness fluctuation method, although others find a distance 50% larger, using the Tully-Fisher relation (Wu et al. 2002).

The surface brightness of spiral disks is often fit with a profile that falls off exponentially along the major axis and with a sech$^2z$ profile along the minor axis. BLAST does not resolve the minor axis, so we consider only the major axis profile,

$$\rho(x) = \rho_0 e^{-|x|/\alpha}, \quad (6.1)$$

where $x$ is a spatial variable aligned along the major axis and $\alpha$ is the disk scale length. Based on optical measurements, van der Kruit and Searle (1981) find a scale length of $\alpha = 113''$, while Wu et al. (2002) quote a thick disk scale length of $115''$ and a thick disk scale length of $157''$. In the near-IR, Rice et al. (1996) find scale lengths of 141, 121 and 119'' at J, H and K bands. Measurements in the far-IR with the Kuiper Airborne Observatory at 100, 160 and 200$\mu$m Engargiola and Harper (1992) lead to a scale length of 100''.
Neininger et al. (1996) mapped NGC 4565 continuum emission at 1.2 mm and line emission from the $^{12}$CO $J = 1-0$ and $2-1$ transitions using IRAM. They detect a molecular ring, a concentration of molecular material in a ring around the centre of the galaxy, in CO, peaking at $1-1.5'$. The continuum emission shows similar structure with a superimposed weaker component extending to high radius. Observations in the mid-IR by IRAS (Rice 1993) find no evidence of structure along the galaxy profile, but point out that the telescope scanned perpendicularly to the major axis and is thus not very sensitive to structure along it.

The 0.12 deg$^2$ field was mapped for a total of 50 min near the end of the flight. Minimal cross-linking was achieved. The maps created using the optimal map-maker (Sec. 4.4) are shown in the left-hand column of Fig. 6.1. The structure of the galaxy is significantly smeared by the large BLAST beams.

### 6.1 Deconvolution

As with the maps of the Vulpecula region, we attempt to deconvolve the NGC 4565 maps to recover structure at the nominal diffraction-limited beam size. However, the signal-to-noise (S/N) ratios in these maps is low. The pixel-to-pixel noise is 1.1, 1.0 and 0.6 MJy sr$^{-1}$ at 250, 350 and 500 μm, while the peak signal at each band is 17.1, 10.2 and 6.7 MJy sr$^{-1}$, giving peak S/N of 15, 10 and 10.

The deconvolution process described in Sec. 5.4.2 is applied to the raw NGC 4565 maps. We use the synthetic PSF (see Sec. 5.4.2) for all three bands. The resulting maps are shown in Fig. 6.1. A noticeable improvement in resolution is seen, with most of the structure along the minor axis in the raw maps disappearing, and hence being due to the extended beam shapes. The noise levels in the deconvolved maps are increased by a factor of $\sim 10$, while the peak signals increase by a factor of $\sim 5$, so the S/N in the deconvolved maps is about half that in the raw maps.

### 6.2 Radial Profile

We calculate radial disk profiles at each of the BLAST bands in order to determine the scale length of dust in the galaxy. The position angle of the major axis is 135°, measured counterclockwise from North (Rupen 1991). We rotate the deconvolved maps about the nominal centre...
Figure 6.1: Raw maps of NGC 4565 at 250, 350 and 500 μm (left column), created using the optimal map-maker described in Sec. 4.4. The raw maps are deconvolved to the nominal diffraction-limited resolution (right column) using the process described in Sec. 5.4.2. Improvement in resolution is clear, with the galaxy clearly resolved along the major axis, although the signal-to-noise ratio is lower in the resulting maps by a factor of ~2. Pixels here are 20'' on a side.
of the galaxy to bring the major axis horizontal. We average along the minor axis, over a box of size 60–80″ centred on the middle of the galaxy. The zero-flux level of the map, which is not zero in map units since the map-maker sets the mean of the map to zero, is estimated in a blank 20 pixel × 20 pixel region of the map. The profiles are shown in Fig. 6.2. We note a central dip in the 250 μm profile, possibly corresponding to the molecular ring seen in CO, as the scales are about the same. The dip is hinted at in the other wavelengths, but the low S/N and lower resolution conceal it.

We fit an exponential model (Eqn. 6.1) to each of the profiles (thin solid curves in Fig. 6.2). We find scale lengths of 190, 206 and 228″ at 250, 350 and 500 μm, respectively. We note, however, that a pure exponential is not a good model at the centre, due to the dip seen in the 250 μm data and hinted at in 350 and 500 μm. We therefore fit the profile to the wings of the data, ignoring the central 3′ region. These fits are shown in the figure with a thin dashed line. The scale lengths corresponding to these fits are 118, 156 and 142″ at 250, 350 and 500 μm, respectively. The scale length measured at 250 μm, the highest S/N map, agrees well with that seen at other wavelengths.

### 6.3 Spatially Integrated Flux Densities

We calculate spatially-integrated flux densities at each of the BLAST wavelengths by performing photometry on the raw maps using an 18′ × 8′ aperture. We use the raw maps instead of the deconvolved maps due to the higher S/N. We find integrated flux densities of 48±7.5, 27.2±3.8 and 14.2±1.9 Jy, at 250, 350 and 500 μm, respectively.

A spectral energy distribution, assumed to be described by a modified blackbody (Eqn. 5.40), is fit to the BLAST and IRAS 100 μm data (Alton et al. 2004) (see Fig. 6.3). The best-fit parameters are $T = 16.8 \pm 1.2$ K and $\beta = 2.0 \pm 0.2$. It is noted, however, that although these errors are quite small, there is a strong correlation between the two parameters, with a Pearson correlation coefficient of $\rho = -0.89$. If we assert the prior that $\beta = 1.5$, as we did in Chapter 5, we find $T = 19.5 \pm 0.5$ K, which agrees well with the cold dust temperature of 20 K found by both Engargiola and Harper (1992) and Alton et al. (2004). We also note that in SED fits to other BLAST objects (see Truch et al. 2007), we often find that the IRAM 1.2 mm data points are low compared to the BLAST data points. This is an additional source of uncertainty not
Figure 6.2: The major axis of profiles NGC 4565 at each BLAST band, extracted from the deconvolved maps. Exponential profiles are fit to the data (thin solid and dashed curves). The solid curve is a fit to all data, while the thin curve is a fit ignoring the central 3'. The data points marked with filled circles are used in the latter fit.
Figure 6.3: Modified blackbody fit to the BLAST and IRAS 100 μm data. The 60 μm point is not included in the fit, as it does not appear to be well-described by the single-temperature model. The 1.2 mm IRAM data point is included (Alton et al. 2004). The best-fit parameters are $T = 16.8$ K and $\beta = 2.0$. 
reflected in the quoted errors.

For the above-listed best-fit parameters, the corresponding integrated far-IR flux is \( S_{\text{FIR}} = (1.2 \pm 0.1) \times 10^{-12} \text{W m}^{-2} \). Assuming a distance of 10 Mpc to the galaxy, we find a bolometric luminosity of \( L_{\text{FIR}} = (4.3 \pm 0.4) \times 10^9 \text{L}_\odot \). With \( \kappa = 10 \text{cm}^2 \text{g}^{-1} \), this gives a dust mass of \((7.8 \pm 1.9) \times 10^7 \text{M}_\odot\). Again, with the caveat given above, these values agree reasonably well with Engargiola and Harper (1992), who find \( L_{\text{FIR}} = 9.8 \times 10^9 \text{L}_\odot \) and \( M_{\text{dust}} = 3.4 \times 10^7 \text{M}_\odot \), the latter corrected for differing values of the assumed value of \( \kappa \). These values are similar to those found for other quiescent spirals (e.g. Chini et al. 1995).

6.4 Discussion

Engargiola and Harper (1992) proposed a two-temperature component model to the disk profile, a warm \((T = 30 \text{K})\) spiral arm feature superposed on a cooler \((T = 18 \text{K})\) exponential disk. The model is shown visually in Fig. 6.4, both viewed face on and at an inclination of 85°. The warm component dominates at 100 micron, while the cooler exponential component dominates at 160 and 200 \(\mu\text{m}\). Engargiola and Harper (1992) find that star formation is strongly correlated with the warm dust component. The BLAST data appear to support this model, as the two exponential fits to each profile (Fig. 6.2) appear to differ more significantly at short wavelengths, which are more sensitive to the warm dust. The BLAST measurements at 250–500 \(\mu\text{m}\) can further help to constrain this model. We plan to perform this analysis in the near future.

We have made the first high-resolution observations of a nearby spiral galaxy at 250–500 \(\mu\text{m}\). We fit the major-axis profiles and find disk scale lengths consistent with those found at a variety of other wavelengths. However, the low resolution of our raw maps and poor signal-to-noise ratios in the deconvolved maps hinder the probing of substructure in the profiles. BLAST observed 6 nearby resolved galaxies during the 2006 flight, and with the nominal beam sensitivities, will provide detailed measurements of the distribution of dust in these objects.
Figure 6.4: The two-component dust model of Engargiola and Harper (1992) at 100, 160 and 200 μm. The model is a superposition of a warm spiral arm component and a cool exponential profile. The solid contours show the model face on while the dashed contours show the model as viewed from an inclination angle of 85°. [This figure appears in Engargiola and Harper (1992).]
A 2-m balloon-borne submillimetre telescope observing simultaneously at 250, 350 and 500 μm was built and has completed two science flights. The instrument performed well during the first flight from Sweden in 2005, except that the telescope was not properly focused, leading to a significantly non-Gaussian point spread function and a resulting factor of ~7 loss of resolution and point source sensitivity at 250 μm. We concentrated the observing time during the flight on Galactic studies, due to this reduced performance. The analysis of two of these observations has been presented.

A 4 square degree field toward the constellation Vulpecula in the Galactic plane was observed for 6 hr in the 2005 flight. The high signal-to-noise ratio in the map allows us to deconvolve the maps to higher resolution. We extract 60 compact sources from the map, displaying a large range of intrinsic properties. We identify the BLAST-detected sources with catalogues produced from mid- and far-IR observations by IRAS, MIPS and MSX, constraining dust temperature in the spectral energy distribution (SED) fit for the warmer sources. From the SED fits, we derive FIR-integrated flux and luminosity-to-mass ratios. $^{13}$CO spectra at the location of each source are examined. Source velocities are compared to the velocity profile of the H II region Sh2-86 and to a model Galactic rotation curve. Based on this comparison, we associate 49 of the sources with Sh2-86 at a distance of 2.3 kpc, 10 with the Perseus arm at 8.5 kpc and 1 in the outer Galaxy at 14 kpc. Using these distances, we calculate luminosities and masses for each of the sources. Based on these properties and whether or not the objects are identified in the IR and radio, we find source properties consistent with a range of evolutionary stages, from cool high-mass, low-luminosity clumps with no evidence of star formation, up to high-mass proto-stellar objects and ultra-compact H II regions.

In addition to the data presented here, we have observed the Vulpecula field at 1.1 mm with Bolocam on the CSO. The addition of this longer-wavelength data will help to better constrain...
the SED fits, in particular breaking the degeneracy between temperature $T$ and emissivity index $\beta$. This analysis, along with a study of the diffuse structure in the map, will be performed in the near future. Similar analysis on the other Galactic Plane fields is underway.

We have analysed observations of the nearby edge-on spiral galaxy NGC 4565. The object is resolved along the major axis by BLAST. We deconvolve the images, as with the Galactic plane field, to recover structure at the nominal diffraction limit of the telescope. The major axis profile measured at 250 $\mu$m shows the molecular ring at radius 1.5$'$ seen in $^{12}$CO and continuum 1.2 mm emission. We fit exponential light curves to the profiles and find scale lengths similar to those found in optical, near- and far-IR. We also fit for integrated FIR flux and, assuming a distance of 10 Mpc, find total luminosity and dust mass comparable to those found using FIR data at 100–200 $\mu$m. We anticipate exciting results from high-resolution measurements of nearby galaxies in the 2006 BLAST flight.

BLAST undertook its second science flight from Antarctica in December, 2006. The carbon fibre primary mirror used in the 2005 flight was replaced with an aluminum mirror and an active focusing mechanism was installed. Preliminary measurements point to near-nominal sensitivities and beamsizes. A total of 270 hours observing time was accomplished, with 69% of the time spent on extragalactic sources and 18% on Galactic fields. Nearly all of the extragalactic time was spent on 3 blank field surveys in 2 locations, the CDFS and Akari SEP. CDFS was observed in 2 modes: a deep 0.6 square degree map to the confusion limit and a large shallow 9 square degree map. The large/shallow strategy was used for the Akari SEP observation. Most of the Galactic time was spent making large maps of the Vela molecular cloud in the Galactic plane. The large maps (40 and 150 square degrees), 10 and 40 times larger than the Galactic Plane map presented in this thesis and with the nominal BLAST resolution, will provide an unprecedented data set. Preliminary analysis of the 2006 data is underway.

A number of new submillimetre and millimetre instruments will begin operating in the next few years. The SPIRE instrument about the Herschel Space Observatory, using the same detectors as BLAST but with a 3-m mirror and observing from space, is scheduled to launch in 2008. SPIRE will continue where BLAST has left off, performing the same types of observations but with more available observing time, faster mapping speeds due to the lack of noisy sky emission and higher angular resolution. SCUBA-2 is a new instrument under development for use at the JCMT (Holland et al. 2006). It features ~10 000 transition edge sensor (TES)
superconducting bolometers at each of 450 and 850 μm. The field of view is 7′ × 7′, 16 times that of SCUBA, and features a mapping speed 100 times. The instrument should be operational in 2008. The Large Millimeter Telescope is a 50-m telescope, situated at Volcán Sierra Negra in Mexico, designed to observe at 0.85–4 mm in continuum and heterodyne modes. First light is scheduled for 2008. The Atacama Large Millimeter Array (ALMA) presents a tremendous leap forward in submillimetre/millimetre astronomy, featuring at least 50 12-m telescopes and interferometric baselines of 150 m to 18 km (Wootten 2003). The array will be located on the Chajnantor plain of the Chilean Andes in the District of San Pedro de Atacama in Chile, at an altitude of 5000 m. The instrument will observe at wavelengths from 300 μm to 9.6 mm and will have spatial resolution as fine as 0.005″. Full operation is expected by 2012.

It is an exciting time in submillimetre astronomy, particularly as this waveband develops as an effective window for cosmological studies. BLAST provides a huge advance in sensitivity and mapping speed, compared to existing submillimetre telescopes, and gives us a preview of the type of data that will be available in the coming years.
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