Progress in Globular Cluster Research: Insights from NGC 6397 and Messier 4

by

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A THESIS SUBMITTED IN PARTIAL FULFILMENT OF THE REQUIREMENTS FOR THE DEGREE OF

Doctor of Philosophy

 $_{\mathrm{in}}$

The Faculty of Graduate Studies

(Astronomy)

The University Of British Columbia

(Vancouver)

October, 2008

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Abstract

Globular clusters are extreme stellar populations. They have the highest stellar density, and host both the oldest and most metal-poor stellar populations in the Galaxy. Their densities make them excellent testbeds for stellar dynamics, while the properties of their stars allows us to test our understanding of old and metal-poor stellar evolution. This thesis is comprised of three projects studying the two nearest globular clusters, NGC 6397 and Messier 4.

By examining high-quality HST photometry of NGC 6397, we have constrained the binary fraction in both the central regions, and beyond the halflight radius. We find a binary fraction of ≈ 0.05 in the core and ≈ 0.015 in the outskirts. In the context of recent N-body simulations by Hurley et al. [54], we interpret the observed binary fraction in the outer field as the primordial binary fraction. This value is lower than typically assumed, and has implications for cluster dynamics and N-body modeling.

We report the discovery that young white dwarfs are dynamically hotter than their progenitors. Using the same photometry as mentioned above, and archival *HST* photometry of Messier 4, we have found that young white dwarfs have an extended radial distribution, and therefore a higher velocity dispersion, compared with older white dwarfs and their progenitors. This implies the existence of a "natal kick". Implications for cluster dynamics and stellar evolution are discussed.

Finally, we present the spectra of 23 white dwarfs in Messier 4 obtained with the *Keck/LRIS* and *Gemini/GMOS* spectrographs. We find that all white dwarfs are of type DA. Assuming the same DA/DB ratio as is observed in the field, the chance of finding no DBs in our sample due to statistical fluctuations is 6×10^{-3} . This suggests DB formation is suppressed in the cluster environment. Furthermore, we constrain the mass of these white dwarfs by fitting models to the spectral lines. Our best estimate of the masses of the white dwarfs currently forming in Messier 4 is $0.51 \pm 0.02 M_{\odot}$. This extends the empirical constraint on the initial-final mass relation over the entire range of initial masses that could have formed white dwarfs in a Hubble time.

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List of Abbreviations

ACS: Advanced Camera for Surveys

AGB: Asymptotic Giant Branch

AO: Adaptive Optics

CADC: Canadian Astronomical Data Center

CCD: Charge Coupled Device

CFHT: Canada France Hawaii Telescope

CMD: Colour-Magnitude Diagram

COSTAR: Corrective Optics Space Telescope Axial Replacement

DR1: Data Release 1

ELT: Extremely Large Telescope

GC: Globular Cluster

GMOS: Gemini Multi-Object Spectrograph

GMT: Giant Magellan Telescope

HB: Horizontal Branch

HST: Hubble Space Telescope

IFMR: Initial-Final Mass Relation

IMF: Initial Mass Function

KS: Kolmogorov-Smirnov test

LRIS: Low-Resolution Imaging Spectrograph

MDF: Mask Definition File

- **MEF:** Multi-Extension Fits
- ML: Maximum Likelihood
- MOS: Multi-Object Spectrograph
- **MS:** Main Sequence
- MSRL: Main-Sequence Ridge Line
- **MSTO:** Main-Sequence Turn Off
- **OSCS:** Open Star Cluster Survey
- **PMS:** Pre-Main Sequence
- $\ensuremath{\mathbf{PSF}}\xspace$: Point Spread Function
- ${\bf RGB}{:}\ {\rm Red}\ {\rm Giant}\ {\rm Branch}$
- ${\bf RS:}\ {\bf Rank}\ {\bf Sum}\ {\rm test}$
- **SDSS:** Sloan Digital Sky Survey
- ${\bf SGB:}$ Sub-Giant Branch
- **SM4:** Servicing Mission 4
- **TMT:** Thirty-Meter Telescope
- **WFOS:** Wide-Field Optical Spectrograph
- WFPC2: Wide-Field and Planetary Camera 2

Acknowledgements

There were many people who helped me throughout my time performing this research, who I would like to thank.

First of all, I would like to express my gratitude to my supervisor, Harvey Richer. You have given me many things over my six years at UBC, but perhaps most importantly, you have passed on you enthusiasm for astronomy to me. You have given me guidance and ideas for projects. You have provided me with a great deal of high quality data, without which, these projects would not have been possible. Finally, when you were not able to answer a particular question for me, you were always able to put me in touch with someone who could. Thanks for everything.

I would like to thank all the members of my committee. In particular, Jasper Wall helped me immeasurably with all matters of a statistical nature. Thanks for letting me bounce ideas off you, and for your guidance on the kicks and binary fraction projects in particular. I owe a big thanks to both Jeremy Heyl and Brett Gladman. Thanks for all the time you two spent discussing the kicks project with me.

I would like to thank Jason Kalirai for all his help on all my projects. You always took the time to read my papers thoroughly, and give helpful comments. Thank you in particular for all your help with the white dwarf spectra. This project obviously would not have been possible without all your input.

I would like to thank Michael Rich and David Reitzel for admitting me to the Keck M4 spectra research project. I am grateful to have had access to such a world-class facility, and recognize that you two put in a lot of work to the project before I came on board.

I would like to thank all my other collaborators. In particular, Ivan King was very thorough while giving feedback with respect to the kicks project. Discussions with Jarrod Hurley regarding the binary fraction were very helpful.

Finally, thanks to Jason Rowe and Matthew Hasselfield for putting up with all my silly Linux questions over the years.

Dedication

This thesis is dedicated to Céline Cressman. Thanks for making that last seven years so much fun.

Chapter 1

Introduction

The gross effects of the laws governing our Universe are manifest in the everyday world. We feel the pull of gravity and see the light from the Sun on a daily basis. Yet, to investigate the more subtle effects of the same physical laws that shape the world we are familiar with, a more extreme environment is typically required. Many modern physics experiments are performed in extreme environments, such as near-perfect vacuums, or at temperatures approaching absolute zero or nearing those present during the Big Bang.

Compared to the typical distance between stars in the Galactic disk (10^{16} meters) , stars have very small radii (generally ~ $10^9 \text{ meters})$. Gravitational interactions between stars in the disk are very weak, and collisions are essentially non-existent. However, in a dense stellar environment, interactions that are weak, or very rare, in the disk become stronger and far more common. The densest stellar environment in our Galaxy is found in cores of the large, roughly-spherical conglomerations of stars known as globular clusters. A typical Galactic globular cluster will be composed of hundreds of thousands of stars. The bulk of these stars are concentrated within the half-light radius of the cluster—a distance that is typically less than ten parsecs. The stellar density in the cores of globular clusters without collapsed cores is increased over that of the galactic disk by a factor of 10^{2} - 10^{4} . In globular clusters with collapsed cores, the factor can be greater than 10^5 [47]. This extreme density makes many phenomena, such as stellar relaxation, mass segregation, and stellar systems formed by collisions, observable. Thus, globular clusters are the ideal location to study the process, and effects of, stellar interactions and dynamics.

1.1 Dynamics

The most obvious dynamical effect in a globular cluster is relaxation. Dynamical relaxation is the effect which describes a stellar population's tendency to distribute kinetic energy roughly evenly between its members [12]. Energy is exchanged between stars via gravitational interactions. In the absence of primordial mass segregation, a cluster is born with the property that the velocity of a particular star is uncorrelated with its mass [102]. More massive stars will therefore tend to have greater kinetic energies. As the relaxation process develops, more massive stars tend to give up their energy to less massive ones. This process leads massive stars to have lower velocities than lower-mass stars. A star's velocity determines the average radius of its orbit in the cluster potential. More massive stars will therefore segregate to the core of the cluster, while low-mass stars will be preferentially removed through evaporation [96]. This differentiation according to mass is called mass segregation, and the timescale on which this occurs is called a relaxation time, t_{rel} . According to Binney and Tremaine [12], the number of times a star crosses a stellar system before encountering enough stars to relax is given by

$$n_{\rm relax} = \frac{N}{8\ln\Lambda},\tag{1.1}$$

where N is the number of stars in the system, and $\ln \Lambda$ is the Coloumb logarithm. The Coloumb logarithm is expressed as

$$\ln \Lambda = \ln \left(\frac{R_{\text{system}}}{b_{\min}} \right), \qquad (1.2)$$

where R_{system} is the radius of the stellar system, and b_{\min} is the impact parameter that will deflect a star by 90°. The relaxation time is then simply the crossing time, t_{cross} ,

$$t_{\rm cross} = \frac{R_{\rm system}}{v},\tag{1.3}$$

times n_{relax} .

Another relevant timescale for stellar populations is the evaporation timescale, t_{evp} [96]. Every stellar population exists in a tidal potential. Characterizing the tidal force in terms of the escape velocity of a given population, these tides may be strong (as in the case of an open cluster orbiting near the Galactic center [98]) or weak (as in the case of the Magellanic Clouds orbiting the within the influence of the Milky Way [105]). The distribution of velocities in a stellar population is well described by a Maxwellian distribution [96]. This means that a small number of stars will have velocities many times the mean velocity of the population. These stars will typically become unbound, and lost, from the population, and hence the population will slowly be depleted of its members. Of course, as stars are removed from the population, the gravitational potential of the population weakens, lowering the escape velocity. The timescale on which a population of stars dissipates in referred to as the evaporation timescale.

The final timescales with which one can describe a stellar population are the evolutionary timescale (the timescale on which a star will evolve) and the Hubble time (the age of the Universe). Like globular clusters, open clusters also exhibit relaxation [71]. In fact, due to the lower number of stars in open clusters, they typically have relaxation times far shorter than those of globular clusters. However, open clusters also dissolve relatively quickly [62]. The most massive pressure supported stellar populations are elliptical galaxies. While elliptical galaxies have extremely long evaporation timescales, their relaxation times are far longer than a Hubble time. Globular clusters are the only stellar population which has the convenient property that $t_{rel} \ll t_{evl} \sim t_{Hub} < t_{evp}$, though indeed a fraction of the Galaxy's initial population has evaporated. This means that while globular clusters formed early in the history of the Galaxy, and have long since relaxed and evolved, they have not yet dissolved, and are therefore observable.

The rate for a particular encounter in a stellar system only depends on the number density of stars, n, the cross-section for encounter, σ , and the velocity of the stars, v. Combining these quantities, the encounter rate is given by

$$t_{\rm col} = \frac{1}{n\sigma v}.\tag{1.4}$$

As mentioned above, close stellar encounters, especially stellar collisions, are very rare in the disk of the Galaxy. Plugging in numbers typical for the disk of the Galaxy, one finds a collision should occur for a given star only once every 10^{19} yrs, or 10^9 Hubble times. Considering there are $\sim 10^{11}$ stars in the Galaxy, only hundreds of collisions are likely to have taken place in the Milky Way over the age of the Universe.

A close stellar encounter is defined as one that changes a star's velocity by a factor of order unity, and will alter the star's trajectory by roughly 90°. For this to occur, the gravitational potential energy between the two interacting stars at the point of their closest approach must be roughly equal to their relative kinetic energies. Therefore, assuming the two stars are of similar mass, this implies the impact parameter necessary for a gravitational deflection of this magnitude, b_{90} , is

$$b_{90} = \frac{Gm}{v^2}.$$
 (1.5)

Again, using numbers for the disk, and normalizing by the radius of the Sun, we find $b_{90}/R_{\odot} = 400$. Substituting this impact parameter as the crosssection in Equation 1.4, we find that close stellar encounters are 1.6×10^5 times more common that actual collisions. This number implies that 10^7 stars have had close encounters since the formation of the Galaxy. Considering the number of stars in the Milky Way, this is represents only one star in 10^4 , and close encounters are not important for the average star. However, with the increased density in globular clusters, close encounters are relatively common. In particular, three-body interactions, in which a preexisting binary becomes temporarily bound with another star, can create stellar systems that are very rare in the disk. Products of these stellar encounters, such as cataclysmic variables, low-mass X-ray binaries, blue stragglers, and millisecond pulsars are all far more common in globular clusters than they are in the disk [101].

1.2 Stellar evolution

Perhaps the single most commonly used diagram in optical observational stellar astronomy is the colour-magnitude diagram (CMD). This is because it is simultaneously both relatively easy to construct, and very informative. The CMD is the observational analog of the famous Hertzsprung-Russel diagram (HRD). While the location of a star on the vertical axis of the HRD depends only on the star's luminosity, a star's vertical location on a CMD depends on both its luminosity and its distance. If one wishes to compare a theoretical HRD with an observational CMD in a meaningful way, one must make an adjustment for the distance of each star. Because all the stars in a star cluster are at essentially the same distance, this correction for a cluster is very simple, i.e., the distance modulus for every star in the cluster can be accurately represented by a single value. They are therefore the ideal populations to test our theoretical understanding of stellar evolution. The loci of stars in both the CMD and HRD are functions of the age of a stellar population and its abundance of both iron peak elements ([Fe/H]) and so-called alpha elements ($[\alpha/Fe]$), such as Ne, Mg, Si, S, Ar, Ca, and Ti. Other than a dynamical laboratory and factory for stellar exotica, globular clusters serve another great purpose—as templates with which to test our understanding of old and metal-poor stellar populations. Shown in Figure 1.1 is a family of isochrones showing the variation as a function of age. Figure 1.2 shows the variation of a family of isochrones as a function of metallicity.

A star will pass through many stages throughout its life. There is a qualitative difference between high-mass stars $(M > 8M_{\odot})$, that will undergo a supernova explosion and end their lives as neutron stars or black holes, and low- and intermediate-mass stars $(M < 8M_{\odot})$ that will end their



Figure 1.1: A CMD constructed of isochrones of various ages. The metallicity for these isochrones is [Fe/H] = -2.0, with $[\alpha/Fe] = 0$. The ages vary from 2 Gyr (blue, left) to 15 Gyr (red, right) in 1 Gyr increments. The isochrones are by Dotter et al. [26].



Figure 1.2: A CMD constructed of isochrones of various metallicities. The age for these isochrones is 12 Gyr, with $[\alpha/Fe] = 0$. The metallicities vary from [Fe/H] = -2.5 (blue, left) to [Fe/H] = +0.5 (red, right) in increments of 0.5 dex. The isochrones are by Dotter et al. [26].

lives as white dwarfs [100]. Due to the functional form of the initial mass function (IMF), far more low-mass stars are created than high-mass stars. For example, consider a population of stars of the same size as a typical globular cluster (10⁵ stars). Assuming the population has an IMF consistent with the IMF of Salpeter [91] (i.e., $dN/dM \propto M^{-\alpha}$ where $\alpha = -2.3$), with a minimum mass equal to the 0.08 M_{\odot} hydrogen-burning limit, only ~ 300 stars will have an initial mass over 8 M_{\odot} . High-mass stars will also have much shorter lives than low-mass stars. The main-sequence lifetime of a star can be estimated from the following formula

$$\tau_{\rm star} = 10 \,\rm Gyr \, \left(\frac{M_{\rm star}}{M_{\odot}}\right)^{-3.5}.$$
 (1.6)

From this equation we can see that an 8 M_{\odot} mass star will live only 7 Myrs. So, while high-mass stars are certainly dramatic, and lead to many exciting phenomena, there are very few of them, and in a old stellar population, they have all long since evolved. The following discussion will therefore focus only on low- and intermediate-mass stars. The various evolutionary stages a star will pass through are shown in Figure 1.3.

A star cluster will form when a giant molecular cloud becomes unstable according to Jeans criterion [57]. Physically, this means that the time it takes for a sound wave to cross the cloud becomes longer than the free-fall collapse time scale. The sound crossing time can be written as

$$t_{\rm sound} = \sqrt{\frac{mR^2}{\gamma kT}},\tag{1.7}$$

where m is the mass of the gas particle, R is the cloud radius, γ is the adiabatic index, k is the Boltzmann constant, and T is the temperature of the gas. The free-fall time can be written as

$$t_{\rm ff} = \sqrt{\frac{1}{G\rho}},\tag{1.8}$$

where G is the gravitational constant, and ρ is the density of the cloud. By setting these two equations equal to one another, and replacing the mass density, ρ , with number density, $n = \rho/m$, we can determine the Jeans length,

$$R_J = \sqrt{\frac{\gamma kT}{Gnm^2}}.$$
(1.9)

Length scales longer than this are unstable, and a shock will cause clouds larger than this to collapse. As clouds begin to collapse, density will typically



Figure 1.3: A CMD constructed from the evolutionary track of a star with [Fe/H] = -2.0 and $[\alpha/Fe] = 0$, and $M_{\rm ZAMS} = 0.8 M_{\odot}$. In order of evolutionary sequence, the pre-main sequence (PMS) is shown in red, the main sequence (MS) is shown in purple, the sub-giant branch (SGB) is shown in yellow, the red-giant branch (RGB) is shown in dark blue, the horizontal branch (HB) is shown in green, the asymptotic-giant branch (AGB) is shown in light blue, and finally, the cooling sequence for a 0.5 M_{\odot} white dwarf is shown in orange. The logarithm of the age along the white-dwarf-cooling sequence is indicated. The main-sequence stellar track is by Dotter et al. [26]. The white-dwarf-cooling sequence is by Fontaine et al. [38].

increase faster than the temperature, and the Jeans length will decrease. Smaller and smaller volumes of the cloud will become unstable, causing the cloud to fragment as it collapses. Each of the fragments may fragment further, and eventually, the clouds will have a spectrum of masses that mimic the IMF of the stars. When the clouds become optically thick radiation emitted from within the cloud will now be trapped, and the temperature will increase. When T^3 increases more quickly than ρ , the Jeans mass will increase, and fragmentation will stop [18].

These fragments will continue to collapse, with the temperature in the center of the fragments (now called protostars) steadily increasing [48]. Once the temperature and pressure in the center of the fragments is sufficient to fuse hydrogen, the energy released from the nuclear reactions will resist further collapse, and the star will stabilize.

A good review of stellar evolution is given in Lebreton [67]. Once the stars form, they will lie in a sequence on the CMD. Higher-mass stars will be bluer and brighter, and lower-mass stars will be redder and dimmer. Stars will remain on the main sequence as long as they have hydrogen to fuse in their cores. To first approximation, a star will remain stationary on the CMD for the duration of its main-sequence lifetime. In contrast, once a star has exhausted its store of hydrogen in the core it will undergo a series of evolutionary changes and move to wildly different locations on the CMD.

Throughout the main-sequence lifetime of a star, the product of hydrogen fusion, helium, accumulates in the core of the star. The concentration of helium increases in the core, and eventually hydrogen fusion ceases there. Hydrogen fusion continues in a shell surrounding the core, and the star will ascend the red giant branch (RGB). As a star ages on the RGB, the core temperature and luminosity will steadily increase. In intermediate-mass stars ($8 M_{\odot} > M > 2 M_{\odot}$), the helium core will ignite in a steady fashion. However, in low-mass stars, the helium core will become degenerate before it ignites. The degenerate matter is an excellent heat conductor, and hence the entire core will become nearly isothermal. When an degenerate isothermal core ignites, it does so almost instantaneously in a "helium core flash".

As helium begins to burn in the core, the core expands. This expansion causes the outer layers of the star to contract, and the luminosity of the star actually decreases as the star contracts and becomes hotter. The star will then be on the horizontal branch (HB) or red-clump (depending on mass and metallicity). After helium ignites in the core, the evolution is analogous to that on the main sequence, with helium taking the place of hydrogen, and carbon/oxygen taking the place of helium. A carbon-oxygen core will begin to accumulate on the HB, and eventually extinguish fusion in the core. The star will again brighten and redden as it now burns helium in a shell. The star will then ascend the asymptotic giant branch (AGB). A rough rule of thumb is that hydrogen shell burning has a duration of approximately 10% of hydrogen core burning, and helium burning has a duration of 10% of hydrogen shell burning. For low-mass stars, this is their final state. The pressure at the center of a star is directly related to its mass. Furthermore the nuclear reaction rate, and hence the temperature are also controlled by the mass of the star. Low-mass stars can simply not generate the temperatures and pressures needed to ignite carbon and oxygen. In a final dramatic act, a star will eject its entire envelope after it can no longer sustain helium shell burning and form a planetary nebula.

1.3 White dwarf formation

While a star loses mass throughout its lifetime, the bulk of the mass loss occurs on the RGB and AGB. Stars will eject the entirety of their envelope mass during these stages. After envelope ejection the remaining degenerate carbon-oxygen core is now referred to as a white dwarf. Though the precise relation is slightly uncertain (as will be explored in Chapter 4), a large range of main-sequence masses will lead to a relatively small range of whitedwarf masses [46]. The maximum degenerate helium-core mass that can be sustained before igniting is $\approx 0.46 M_{\odot}$ [14, and references therein]. This sets the dividing mass between CO-core and He-core white dwarfs. A core less massive than this cannot burn to CO, and therefore must be composed of He; conversely, once a core achieves this mass, it will burn completely to CO, and hence can not be composed of He. Though CO-core white dwarfs are the most common type observed, white dwarfs are observed to have a mass less than this minimum mass. Stars with masses low enough to form heliumcore white dwarfs have not had sufficient time in the age of the Universe to evolve in isolation. Therefore, these stars are thought to have evolved via a binary channel. The idea is the envelope of a RGB star is stripped by a companion, and its evolution is truncated. The premature cessation of hydrogen-shell burning limits the helium mass that can accumulate in the degenerate core, and hence the helium core never achieves the critical temperature and density needed to ignite. When the envelope is ejected, the remaining degenerate core becomes a He-core white dwarf. The upper mass limit of a white dwarf is set by the famous Chandrasekhar mass [16] (explored in greater detail in Section 1.4).

Observationally, the mass of a white dwarf can be estimated via several

methods. The most direct method of estimating a white dwarf's mass is through a binary system. This is the most direct way of estimating the mass of most astronomical objects within the solar neighbourhood. If the semi-major axis and period of an orbiting system is known, the mass of the system can be determined directly from Kepler's third law. If the companion is a luminous star, its mass may be estimated by determining its spectral type. The mass of the white dwarf is then the difference between the system mass and the luminous-star mass. In fact, this method was used to determine the mass of Sirius B before it was understood to be a white dwarf. Other methods include measuring the gravitational redshift or modeling the atmospheric absorption lines, however, both of these method require a massradius relation. In Chapter 4 we will constrain the masses of young white dwarfs in the globular cluster Messier 4.

1.4 Mass of a white dwarf

An order-of-magnitude estimate of both the mass-radius relation and the maximum mass of a white dwarf can be determined from first principles independently of any observations. An argument of this type was first presented by Landau [66]. First, for an object in equilibrium, its total kinetic energy, E_k , must be less than its gravitational binding energy, E_g . If this were not the case, the object would not be, by definition, bound, and hence would experience mass loss. Nor can the gravitational binding energy be far greater than the kinetic energy, or else the object would contract further. Thus, the only possibility is that the kinetic energy roughly equals the binding energy. So, we start with two basic equations. The gravitational energy may be written as:

$$E_g = \frac{GM^2}{R},\tag{1.10}$$

where G is the gravitational constant, M is the mass of the star, and R is the radius and the star. The kinetic energy of the star may be written as the average kinetic energy per particle multiplied by the number of particles, or:

$$E_k = \frac{1}{2}mv^2 \times N, \qquad (1.11)$$

where m is the mass of the average particle, v is the velocity of the average particle, and N is the number of particles.

In the case of a white dwarf, degenerate electrons provide the support, and hence in this case 'particle' refers to an electron. Using p = mv, Equation 1.11 becomes

$$E_k = \frac{Np^2}{2m}.\tag{1.12}$$

White dwarfs are degenerate objects, meaning that all quantum states in the star are full. The momentum of a given particle is therefore of the same order as the uncertainty of the momentum, and hence can be related to the average length scale that a particle will occupy via the Heisenberg uncertainty principle, $p = \hbar/\Delta x$. The number density of particles is related to Δx by $\Delta x = n^{-1/3} = (3N/4\pi R^3)^{-1/3}$. We can replace p^2 in Equation 1.12 with $\hbar^2 (N/R^3)^{2/3}$ yielding

$$E_k = \frac{N\hbar^2 n^{2/3}}{2m} = \frac{\hbar^2 N (N/R^3)^{2/3}}{2m} = \frac{\hbar^2 N^{5/3}}{2mR^2}.$$
 (1.13)

Finally, the number of electrons in a star can be estimated to within a factor of 2 as M/m_p , where m_p is the mass of a proton. With this substitution Equation 1.13 becomes

$$E_k = \frac{\hbar^2 M^{5/3}}{m_e m_p^{5/3} R^2}.$$
(1.14)

Relating the kinetic energy to the gravitational energy we have

$$\frac{\hbar^2 M^{5/3}}{m_e m_p^{5/3} R^2} = \frac{GM^2}{R}.$$
(1.15)

Rearranging we have the mass-radius relation,

$$R = \frac{\hbar^2}{Gm_e m_p^{5/3}} M^{-1/3} = 10^4 \, km \left(\frac{M}{M_\odot}\right)^{-1/3}.$$
 (1.16)

As the mass of a white dwarf increases, the gravitational force increases. This decreases the radius of the star, and the electrons are "squeezed" more intensely. Since this decreases the uncertainty in the position of the electron, to satisfy the Heisenberg uncertainty principle, the velocity of the electrons must increase. Eventually, the electrons attain velocities such that the classical expression for their energy, $E_k = p^2/2m$, is no longer valid, and we must instead use the relativistic formulation, $E_k = pc$, where c is the speed of light. Following the same line of argument as before, and equating gravitational and kinetic energy we have

$$\frac{\hbar c M^{4/3}}{R m_p^{4/3}} = \frac{G M^2}{R}.$$
(1.17)

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Rearranging, we find that

$$M = \left(\frac{\hbar c}{G m_p^{4/3}}\right)^{3/2} \simeq 1 \, M_{\odot}.$$
 (1.18)

The line of reasoning presented above makes rather crude assumptions, yet manages to derive the same functional form as solutions taking into account a more realistic equation of state. A more precise solution assuming a carbon-oxygen core due to Chandrasekhar [16] yields

$$M_{ch} = 1.4 M_{\odot} \text{ and } R = 9 \times 10^3 \, km \left(\frac{M}{M_{\odot}}\right)^{-1/3}.$$
 (1.19)

1.5 A brief history of globular clusters

The first globular cluster was discovered in 1665 by Abraham Ihle, a German amateur astronomer. This cluster, in the Sagittarius constellation, appeared to Ihle as simply a nebulous smudge. It was later listed as the 22nd object in the famous catalogue by the French astronomer Charles Messier. In fact, it was Messier who first resolved globular clusters, and serendipitously discovered they were composed of stars [109]. The first thirteen globular clusters discovered are listed in Table 1.1.

This thesis will focus on the two globular clusters nearest to the Sun, Messier 4 and NGC 6397. Despite the proximity of these two clusters, and their bright apparent magnitudes (5th and 7th apparently brightest clusters respectively in the Galaxy), both these clusters were not discovered until a century after the first globular cluster. This is most likely due to the fact that they are Southern clusters.

The first key discovery that was made using globular clusters was the famous result of Harlow Shapley. Shapley noted that the distribution of globular clusters is not even on the sky, but rather heavily concentrated toward the Galactic center. If, as indeed is the case, the Galaxy is cylindrically symmetric, this implies that the Sun is not close to the center of the Galaxy. A similar argument using stars had previously come to the opposite conclusion. The reason that stars appear to be roughly evenly distributed around the sun is because only the nearest stars are bright enough to be easily observed, and the density of stars in the Galaxy changes rather slowly. Globular clusters do not suffer from the same fate because they are intrinsically bright, and because their orbits are not confined to the disk like stars. Using globular clusters, Shapley made the first correct order of magnitude estimate of the size of the Galaxy.

Cluster	Discoverer	Year
Messier 22	Abraham Ihle	1665
ω Centauri	Edmond Halley	1677
Messier 5	Gottfried Kirch	1702
Messier 13	Edmond Halley	1714
Messier 71	Philippe Loys de Chéseaux	1745
Messier 4	Philippe Loys de Chéseaux	1746
Messier 15	Jean-Dominique Maraldi	1746
Messier 2	Jean-Dominique Maraldi	1746
NGC 104	Abbé Lacaille	1751
NGC 4833	Abbé Lacaille	1751
Messier 55	Abbé Lacaille	1751
Messier 69	Abbé Lacaille	1751
NGC 6397	Abbé Lacaille	1751

Table 1.1: The first 13 globular clusters discovered. This thesis will focus on Messier 4 and NGC 6397. Despite the proximity of these two clusters, and their bright apparent magnitudes (5th and 7th brightest clusters respectively in the Galaxy), both these clusters were not discovered until a century after the first globular cluster. This is most likely due to the fact that they are Southern clusters.
As observational instrumentation and techniques improved throughout the 20th century, our understanding of globular clusters has steadily improved. The development of large-aperture telescopes allowed the detection of progressively fainter stars. However, observations continued to be limited in the cores of globular clusters by the extreme crowding found there. For many years, there was nothing that could be done to improve this situation. Without adaptive optics, the performance of a telescope will be limited by the Earth's atmosphere. Natural seeing typically limits resolution to slightly less than 1 arc-second, the diffraction limit of a 20 cm telescope operating at red wavelengths. Uncorrected telescopes larger than this will not benefit from an improvement in resolution. With naked-eye observations or photographic plates, the flux from star with overlapping seeing disk is simply lost. Observations of the cores of globular clusters were limited by this. Hence, measurements of core radii in distant compact clusters were dominated by seeing errors.

In the 1980s there were several technological developments that greatly improved observations of globular clusters. The replacement of photographic plates with charged-coupled devices (CCD) allowed the development of the technique of point-spread-function (PSF) photometry. This technique determines the magnitude of a star by modeling the two-dimensional profile of a star. Crucially, this allows a bright star to be subtracted from the image, revealing faint stars that would otherwise be obscured. PSF photometry allowed the first detailed investigations of the inner regions of globular clusters.

The second critical development around this time was the launch of the Hubble Space telescope (HST) in 1990, and the installation of the Corrective Optics Space Telescope Axial Replacement (COSTAR) in 1993. In the absence of the Earth's atmosphere, the PSF of the HST gives roughly an order-of-magnitude improvement over ground-based telescopes. With the improved seeing, the cores of many globular clusters, while still crowded, could be resolved. In fact, two of the projects discussed in this thesis make use of HST observations, and would simply be impossible from the ground using present technology.

Finally, recent developments in adaptive optics (AO) have made highresolution observations of globular clusters possible from the ground. AO is still a developing field. Though AO observations are presently restricted to rather small fields of view and reddish wavelengths, future developments will likely make wide-field, high-resolution, optical observations possible from the ground. Perhaps the most important niche that will be filled by ground-based large-aperture AO-enabled telescopes is that of spectroscopy of crowded fields. The multiplexing abilities of fiber-fed and multi-slit spectrographs is a key component to many projects. These types of instruments are impractical in space. Hence, multiplexed spectroscopy of faint targets remains a serious challenge. The definitive answers to problems requiring faint multi-object spectra in crowded fields most likely await the development of the 30-meter-class AO-enabled telescope such as the Thirty Meter Telescope (TMT), European Large Telescope (ELT), or the Giant Magellan Telescope (GMT).

1.6 Thesis format

In this thesis, I will describe three individual projects related to Messier 4 and NGC 6397. All three of the projects were somewhat independent from one another, and therefore use different astrophysical techniques, such as imaging and spectroscopy, and ground- and space-based observations. As such, I have written three long chapters describing each study in turn, as opposed to individual chapters on the data reduction, calibration, results, etc. Each of these three chapters is self-contained in that they describe the data reduction methods and scientific results unique to that project.

The first project, presented in Chapter 2, reports on an attempt to place tight constraints on the binary fraction of NGC 6397. This is a spacebased imaging study involving the cluster main-sequence stars. As will be discussed in more detail in Chapter 2, a primordial binary fraction of only several percent has a dramatic effect on the dynamics of a globular cluster. Primordial binaries delay core collapse by a factor of ten over a cluster with no primordial binaries. Our estimate of the primordial binary fraction, 0.012 ± 0.004 , is lower than is typically assumed. Taken in isolation, this result would suggest that globular clusters should undergo core collapse earlier than is typically assumed.

Interestingly, in some Galactic globular clusters, modeling suggests that there are not enough binaries to have postponed core collapse until today, yet these clusters have not collapsed. Drukier et al. [28] found that M 71 is one of these clusters, and conclude that "models in which the heating rate is artificially enhanced are able to reproduce the observations, but, in the absence of an identified source for this extra heating, such models are not physically justified". The second project reports the discovery of a potential source for this dynamical heat. By examining the radial distributions of various populations of stars, we have found observational evidence for the existence of a white dwarf natal kick. This study, presented in Chapter 3, is a space-based imaging study, and examines both the white dwarfs and main-sequence stars of both NGC 6397 and Messier 4.

Chapter 4 will present new results from a spectroscopic study of white dwarfs in Messier 4. In this chapter we examine the spectral types of the white dwarfs, and attempt to constrain their masses. The spectral types of white dwarfs allows us to explore subtleties of low-mass stellar evolution, and highlight the difference between that evolution in the cluster and field environments. The masses of white dwarfs provide insight into stellar evolution, and provide a link to the mass function of the cluster beyond the main-sequence turn off.

Finally, in Chapter 5, I conclude the thesis and link these projects together within a broader astrophysical picture.

Chapter 2

The Binary Fraction of NGC 6397

2.1 Introduction

Globular clusters (GCs) are fascinating dynamical test-beds. A key parameter to understanding the evolution of GCs is the binary fraction. A common parameter to characterize a globular cluster is the half-mass relaxation time, i.e., the relaxation time as calculated from parameters measured at the halfmass radius. The presence of primordial binaries can delay the onset of core collapse from on the order of 20 half-mass relaxation times to over 150 half-mass relaxation times [49].

In a sense, the evolution of a globular cluster is analogous to something familiar to almost every astronomer—the evolution of a star. This is, of course, not a perfect analogy, but I will present it here. In the case of a star, a bound cloud of gas begins to collapse. The density in the core of the cloud increases until it becomes optically thick, at which point the energy generated from the collapse can no longer be instantaneously radiated away. The star will continue to contract until the density and temperature are such that the fusion of hydrogen atoms in the core becomes important. Hydrogen fusion is exothermic, and this energy goes into heating the surrounding atoms. The star contracts until the gas is heated sufficiently to set up a pressure gradient to resist further contraction. If we consider the collapse phase and contraction phase as one, and call it "collapse", something similar happens in a globular cluster.

Three-body interactions in the core of a globular cluster—typically the collision of a single star with a hard binary system, resulting in an even harder binary, and a single star with higher velocity—are the analog of hydrogen fusion in stars. While wide binaries with orbital speeds less than the velocity dispersion of the cluster will tend to be disrupted in three-body interactions, hard binaries will tend to have their orbits hardened. In a typical three-body interaction, a single star become temporarily bound to

a pre-existing binary system. The bound triple is a typically chaotic and unstable system, and usually results with the least massive member of the triple being ejected at high speed from the system. The kinetic energy of the ejected member is extracted from the gravitational potential energy of the binary system, and the semi-major axis of the orbit is reduced. The energy extracted from these systems is sufficient to support cluster cores from collapse [55]. Energy may be extracted from a binary system until the obit is so hard that the stars begin to raise tides upon each other. Stars that orbit very closely to one another will emit gravitational radiation, shrinking the semi-major axis of their orbits. The stars will typically then spiral into each other and form a single more massive star, known as a blue straggler. The stable binary-burning phase will last until the energy stored in primordial binaries has been exhausted. Core collapse then proceeds until the stellar densities increase to such an extent that binary formation via three-body interactions become important [49].

Additionally, the formation rates of stellar exotica, such as blue stragglers, cataclysmic variables, and low-mass X-ray binaries, are all strongly influenced by the binary fraction [6]. For instance, Mapelli et al. [70] showed that while in the core of globular clusters blue stragglers form primarily by direct collisions, in the outskirts binary coalescence is the dominant formation mechanism. However, Leigh et al. [68] recently found that in the cores of 57 globular clusters the frequency of blue stragglers did not scale with the collision rate, as would be expected if direct collisions were the dominant formation process. If binary coalescence is the dominant formation channel, we would expect the frequency of blue stragglers to scale tightly with the binary fraction. Despite the key role the binary fraction plays in our understanding of globular clusters, it remains a weakly observationally constrained parameter in general.

The first searches for binaries in GCs examined stellar spectra for radial velocity variations [42]. This technique relies on identifying individual binary systems, and hence, for a statistically significant sample, it is a time consuming method. Furthermore, searches of this type are only sensitive to certain orbital periods, inclinations, eccentricities, and mass ratios. Results for the whole cluster must be determined by extrapolating from a particular sample, making use of the uncertain distributions of the above mentioned parameters.

As large area CCDs and sophisticated photometric techniques became increasingly widespread, photometric accuracy allowed the search for binaries via their distinct colours and magnitudes [88]. A main sequence star with a main sequence secondary will be brighter and redder than a single star with a mass equal to that of the primary. The maximum deviation from the main sequence is for equal-mass systems, which will have the same colour as the primary, but will be brighter by ~ 0.75 magnitudes. By modeling the distribution of stars in colour-magnitude space as a function of the binary fraction, one can use the CMD as a statistical constraint on the binary fraction. Methods that make use of the CMD investigate the entire cluster in the field of view at once. This makes them intrinsically much less observationally intensive than methods that rely on studying individual systems one at a time, i.e., the detection of radial velocity variations or photometric variability.

There is a consensus that the present day binary fraction in globular clusters is clearly lower than in the disk of the Galaxy (for instance, compare the estimate for the disk of $\approx 50\%$ due to Duquennoy and Mayor [29] to the results for globular clusters compiled in Table 5.2). However, the primordial binary fraction is still a matter of debate. Recently published papers on simulations of GCs include primordial binary fractions ranging from 5% [54] to 100% [56]. While it is unclear precisely how the binary fraction will evolve in the core of GCs, N-body simulations of Hurley et al. [54] show that the binary fraction beyond the half-mass radius remains stable. Thus, in the context of these simulations, a measurement of the present day binary fraction beyond the half-mass radius will constrain the cluster's primordial binary fraction. Here, we present a precise determination of the binary fraction both within and outside the half-mass radius.

2.2 Observations

The observations reported in this chapter were taken as part of the Hubble Space Telescope program GO-10424 [85]. With the primary science goals of: (1) observation deeper than the theoretical termination of the main sequence in a globular cluster, (2) the first clear observation of the collision-induced absorption "blue hook" feature of the white dwarf cooling sequence, and (3) the determination of the white dwarf cooling age of the cluster, 126 orbits were dedicated to imaging one ACS field 3'.17 E and 3'.93 S of the cluster core. For a sense of scale, R_{hm} , the half-mass radius of the cluster, is 2'.33 [47]. The field was chosen for the wealth of archival data in that location. Archival WFPC2 images were taken in this field in 1994 and 1997, giving us a baseline of over a decade for the measurement of proper motions. The proper-motion cleaned CMD was critical for several projects. Unfortunately, the archival data only covered approximately 60% of the ACS field, effectively reducing

the size of the observed field. For the remainder of this chapter we will refer to the overlap region between the *archival* WFPC2 images and the *new* ACS image as the "outer field". The layout of the instruments used in this project is shown in Figure 2.1.

During this campaign, each orbit was divided into three exposures, with the first and last exposure of each orbit taken with the F814W broad-band filter, and the middle exposure (with the darkest sky value) taken with the F606W broad-band filter. In several orbits, short exposures were taken to measure bright stars that were saturated on the longer exposures. The total exposure time was 179.7 ks with the F814W filter and 93.4 ks with the F606W filter. The data were reduced using a novel technique designed to find and measure the faintest possible stars in a field where there are also very bright stars (see Anderson et al. [4], for full details). Sources are detectable to a magnitude of F814W \simeq 30, with the 50% completeness limit being F814W $\simeq 28$. The proper motion measurements are severely limited by the archival WFPC2 data. Proper motion measurements are only reliable to a depth of F814W $\simeq 26.3$. This is *not* an impediment for this project, as we are concerned only with main sequence stars, the vast majority of which are brighter than F814W = 25. Figure 2.2 shows the colour-magnitude diagram (CMD) and the proper motions for the outer field, and demonstrates the utility of proper motions to isolate the cluster. For the remainder of this chapter, we will restrict ourselves to studying a section of colour-magnitude space covering 16 < F814W < 24 and 0.6 < F606W - F814W < 2.4 in this field. The F814W = 24 limit was chosen because the number of mainsequence stars drops very rapidly beyond this. The F814W = 16 limit was chosen because the main sequence is vertical brighter than this, making binaries photometrically indistinguishable, and furthermore, few cluster stars are brighter than this magnitude. The colour limits were driven by the minimum and maximum colours of the main sequence in the magnitude range chosen above.

During the primary target observations, the roll angle of the telescope was controlled in order to simultaneously obtain WFPC2 images centered on the core of the cluster. For the WFPC2, each orbit was also divided into three exposures. Though not used for this particular project, one exposure with the F336W filter was obtained, in addition to one with the F814W filter and one with the F606W filter. Similar to the outer field, there were archival WFPC2 images in this field. Archival images obtained in 1996, 1999, and 2001 with both the F606W and F814W filters enabled the measurement of proper motions over $\sim 75\%$ of the field. The overlap between the *new* WFPC2 images and the *archival* WFPC2 images will henceforth be referred



Figure 2.1: The footprint of the instruments used in this project. The figure is centered on the core of NGC 6397. The circle denotes a radius of 5'.0. For a sense of scale, R_{hm} , the half-mass radius of the cluster, is 2'.33 [47]. The primary target for this project was the ACS field (shown as the bold skewed rectangle in the lower left of the figure). The secondary field acquired for this project is the bold WFPC2 outline (the chevron shape) in the center of the figure. All archival WFPC2 images are denoted with lighter outlines.

to as the "inner field".

The inner-field images were reduced with standard DAOPHOT/Allstar [97] techniques. The observations in the inner field had several limitations compared with the outer field. The stellar density in the inner field is significantly greater than that in the outer field, causing much more scattered light. Furthermore, the larger pixel size and lower sensitivity of the WFPC2 camera prevented the detection of sources fainter than F606W $\simeq 27$ and F814W \simeq 26. The degraded photometry of the measured stars resulted in a significantly broadened main sequence compared with the outer field. Finally, short exposures were not taken, and hence all stars brighter than $F814W \simeq 17$ are saturated. The CMDs and proper motions of the inner field are shown in Figure 2.3. Because of the inferior quality of the photometry and astrometry of the inner field compared with the outer field, we restrict our study in this field to a smaller area of colour-magnitude space: from 17 < F814W < 22 and 0.7 < F606W - F814W < 1.6. For the remainder of the chapter, figures will be shown for the outer field only, except when a difference between the fields is of particular interest.

2.2.1 Optical binaries and field contamination

In our proper-motion cleaned sample there are two sources that could contribute "false" binaries: optical binaries, and field stars that share the mean proper motion of the cluster. We now show that these sources of false signals will make only a minor contribution to the observed binary fraction.

To address field star contamination we examine the proper motion diagrams using the setup shown in Figure 2.4. Our criterion for classifying a star as a cluster member was that the magnitude of its proper motion relative to the mean motion of the cluster was less than a certain radius, $|\mu| < r_c$. The mean proper motion of the field was offset from the cluster by several pixels in the x and y directions, namely x_f and y_f respectively. In order to calculate the contamination from the field, the proper motions were re-centered on the field population. This change of coordinates modifies the cluster selection criterion for cluster stars. Cluster members now have the property that:

$$\sqrt{x_f^2 + y_f^2} - r_c < |\mu| < \sqrt{x_f^2 + y_f^2} + r_c.$$
(2.1)

Selecting only stars in this annulus of proper motion space, we calculate the angle, θ , of the proper motion vector for each star, with $\theta = 0$ directed to the mean cluster proper motion. Stars in this annulus should come from both



Figure 2.2: The top panels show the observed proper motions of the stars in the outer field, while the bottom panels show colour-magnitude diagrams for the same stars. The left panels show all the data. In the upper middle panel the cluster population is circled, and the lower middle panel shows the CMD of the circled population only. In the right panel, the field population is circled, and again in the lower right plot the CMD of the circled population in shown. 24



Figure 2.3: The CMDs and proper motion diagram of the inner field. The various panels are analogous to Figure 2.2.



Figure 2.4: The setup to calculate the contamination from field stars that share the cluster proper motion.

the cluster and the field populations. While the field population should be distributed uniformly with angle, the cluster population should be strongly concentrated on $\theta = 0$. The distribution of angles is modeled as a Gaussian, representing the cluster, and a constant, representing the field population. The number of contaminating stars is estimated as the product of the total number of field stars in the annulus, n_f , and the fraction of the annulus that is occupied by the cluster, f_c , i.e.,

$$f_c = \frac{A_c}{A_a} = \left\{ \left(\frac{1}{r_c}\sqrt{x_f^2 + y_f^2} + 1\right)^2 - \left(\frac{1}{r_c}\sqrt{x_f^2 + y_f^2} - 1\right)^2 \right\}^{-1}$$
(2.2)

The best-fit parameters from our two component model suggest that there are 39 field stars in the annulus for the outer field. For the outer field, $f_c = 0.068$, implying approximately 2.7 field stars should have the same proper motion as the cluster.

However, field stars are only likely to have an effect on the binary fraction if they lie close to the binary sequence, i.e., with in 0.75 magnitudes brighter than the single-star main sequence. As shown in figure 2.5, only ~ 10% of the field stars lie within the binary sequence in our colour-magnitude space. We therefore expect less than one field star to contaminate the binary sequence. There are ~ 2.1×10^3 proper motion-selected cluster stars in the outer field. Given this number of cluster stars, one contaminating star will have a negligible effect on the measured binary fraction. A similar analysis was performed on the inner field. We expect the inner field CMD to be contaminated by ~ 8 stars (3.6×10^3 total), but again, less than one star will lie on the binary sequence.

Optical binaries are stars that are aligned on the sky, but not physically associated. In order to get a sense of the level of contamination from optical binaries, we perform a rough calculation to assess the crowding of the field. The ACS camera consists of $2 \times 2048 \times 4096$ pixels for a total of 16.8×10^6 pixels. We detected 8600 stars over the entire field, which implies we have one star for every 1950 pixels, or equivalently, one star per 44 square pixels. Our detection method requires that a source must be the brightest source within 7.5 pixels, and thus each star will occupy roughly 177 pixels. This implies that our field could have had an order of magnitude more stars before being confusion limited. Due to the uncrowded field, we expect the number of optical binaries to be small. Because of this, we can treat the observed stellar number counts as a good approximation of the true underlying number counts. The preceding statement is *not* meant to be an estimate of the contamination due to chance alignment, but rather serves to justify our subsequent analysis.



Figure 2.5: The derived model of the density of the field population (blue) with the density of the cluster superimposed (red). Note that only rare field star contaminants will have photometry that could be confused for a cluster binary.

We next proceed to quantify the filling fraction as a function of magnitude, f(m). This function quantifies what fraction of the CCD is occupied by stars of a given magnitude. For this study, it does not matter if two field stars overlap; these systems will be removed by their proper motions. However, two cluster stars that overlap, or a cluster and field star that overlap in the current epoch data will not be removed by proper motion measurement. We therefore calculate the filling fraction for cluster stars only. Although we require a star to be the brightest within 7.5 pixels in order to be detected, another star will interfere with the measurement of the magnitude of a star only if it lies within the FWHM of its PSF. This is due to the design of both Allstar, used to reduce the inner field, and the original software used to reduce the outer field. Allstar fits only the pixel within the radius defined by the fitrad parameter, which is generally set to the FWHM of the PSF. Although the reduction package used for the outer field fits the PSF to a greater radius, the number of pixels beyond the FWHM of the PSF mean that a contaminating star beyond this radius can only affect a small number of pixels, and therefore will not affect the measured magnitude. A star is assumed to occupy πr^2 pixels, with r = 2.

Only if two stars are of similar magnitude will their superposition affect the measured magnitude of the brighter star. From the width of our main sequence, we estimate that an increase in flux less than ~ 10% would be lost in the general photometric scatter. Accordingly, for each magnitude, m, we calculate the number of stars with magnitudes between m and m_1 , where $m_1 = m + 2.5$. We call this quantity n(m). The number of contaminating stars as a function of magnitude, c(m), is simply the product of the two previously introduced quantities, c(m) = f(m)n(m). The total number of contaminating stars is then the integral of c(m) over the appropriate range of magnitude. The total contamination from optical binaries in the outer field is expected to be three stars. This contamination corresponds to an overestimation of the binary fraction of ~ 0.15%. Due to increased crowding, the optical binary contamination in the inner field is expected to be more serious. We predict roughly 25 optical binaries in the inner field, corresponding to an increased binary fraction of ~ 0.8%.

2.3 Colour-magnitude space distributions

2.3.1 The single-star sequence

With a small number of simple assumptions one can create a model of the distribution of single stars in a CMD. One only needs a determination of

the main sequence ridge line (MSRL), a determination of masses along the MSRL, a model for the photometric error as a function of magnitude, and a model of the mass function. Save for the masses along the MSRL, these can all be determined from the observed stars.

The MSRL is determined by binning the CMD finely in magnitude, and determining a sigma-clipped mean colour for each bin. In order to derive a relation between stellar mass and location on the MSRL, we make use of stellar evolutionary models. The models we used, referred to as isochrones, model the luminosities and colours of a population of stars of a given age, metallicity, and alpha enhancement as a function of mass. We used isochrones by Dotter et al. [26]. Shown in Figure 2.6 is the difference in colour between the theoretical isochrones and the MSRL.

For magnitudes brighter than $F814W \simeq 19$ the discrepancy between the MSRL and the isochrone is less than 0.02 magnitudes. However, the match between isochrone and MSRL degrades for stars fainter than this. This is most likely due to missing stellar atmosphere opacities at low temperatures. In order to correct the discrepancy, we assumed that the isochrone described an accurate relation between stellar mass and F814W. The mismatch between the isochrone and the MSRL only becomes important for masses less than $\sim 0.1 M_{\odot}$. We only need to use an isochrone in order to assign particular locations on the MSRL a particular mass. Due to the small difference in mass between this point and the hydrogen burning limit, the mass assigned to a particular location on the MSRL would not have been affected if it was assumed that the error was due entirely to the F814W magnitudes, and the F606W magnitudes were correct. The F606W values of the isochrone were adjusted to accord with the MSRL. The maximum deviation of the isochrone from the MSRL in the region we are examining is ~ 0.3 mags. The match of these isochrones to our data is as good as with any other publicly available set.

The observed width of the main sequence is a combination of both intrinsic effects, such as differential reddening and minor differences in metallicity, and observational effects, such as noise, scattered light, and PSF variations. One can further subdivide the observational sources of error into those expected to be correlated and uncorrelated in the two observed bands. Sources of correlated errors, such as defocus, scattered light, and confusion, will have a smaller effect on the measured colour of a star, and hence the broadening of the main sequence, than sources of uncorrelated errors, such as photon noise. For the purposes of our model, the source of the broadening is not important. We expect the photometric error (i.e., noise) to dominate the width of the main sequence of NGC 6397, and therefore we will refer to



Figure 2.6: The difference between the colours of the empirically derived MSRL and the theoretical isochrone. The difference is very small brighter than F814W=20, and reaches a maximum deviation of ~ 0.3 mags at F814W=24. Note that sole purpose of the isochrones is to assign masses to the MSRL. The deviation only becomes significant for $M \leq 0.1 M_{\odot}$, after which there are few stars, and not much mass before the hydrogen-burning limit, and hence it is not a concern.

everything that broadens the main sequence as photometric error.

High-mass-ratio binaries will only have a small deviation from the MSRL, and will likely be lost in the photometric scatter of the single-star sequence. They will therefore also increase the width of the main sequence. However, all main sequence-main sequence binaries will lie to the red side of the cluster main sequence, and thus scatter to the blue side should provide a reliable measure of the photometric error. The only sources of contamination here will be main sequence-white dwarf binaries and field stars that happen to share the mean proper motion of the cluster. Both of these are expected to have a negligible influence on the derived error. In order to model photometric error, we subtract the colour of the MSRL from the measured colour for each star. Then, for those scattering to the blue of the MSRL, we make a histogram of the number of stars and difference from the MSRL. The distribution was noted to have faint, but extended tails. Without in-depth analysis, the best we can do is to model the error as a superposition of Gaussians. Our error model is *not* intended to represent anything physical, but is rather an empirical fit. It was found that if we determined the best-fit parameters for a superposition of two Gaussians, the tails were still not extended enough to accurately represent the observed scatter. A superposition of three Gaussians was found to model the observed blue-side scatter well. The blue-side scatter for both fields is shown in Figure 2.7.

Although we need the error as a function of magnitude, the best-fit parameters were determined for the entire sample. This is because if we divide the sample into magnitude bins, the widths and heights of the bestfit Gaussians vary wildly, whereas we expect the underlying sources of error to vary smoothly and nearly monotonically with magnitude. In order to give our error model some sensitivity to magnitude, the widths of the Gaussians were fixed, and their relative heights were allowed to vary as a function of magnitude.

While all clusters may be born with a "universal" initial mass function, dynamical evolution will alter the mass function and make it a function of position in the cluster. Furthermore, observational complications, such as incompleteness, make the determination of the "true" mass function difficult. As with the photometric error, for the purposes of this project we do not need to disentangle physical effects from observational effects; we are solely concerned with the *observed* mass function. By interpolating between points on the isochrone, we may determine the mass of each star. As shown in Figure 2.8 and 2.9, both the fields were satisfactorily fit by a log-normal



Figure 2.7: The blue-side scatter of (F606W-F814W)-MSRL, i.e., a "straightened" CMD, showing our observed photometric error. The fit of a model consisting of three superimposed Gaussians is shown with dotted and dashed lines for the WFPC2 and ACS fields respectively.

function:

$$N(M) = N_0 \exp\left[\left(\frac{-\log_{10}(M/M_0)}{\sqrt{2}\sigma}\right)^2\right].$$
(2.3)

The best fit values for the outer (inner) field were: $N_0 = 132.6$ (56.56), $M_0/M_{\odot} = 0.22$ (0.55), and $\sigma = 0.30$ (0.50) yielding a reduced $\chi^2 = 1.42$ (1.25).

With the determination of the mass functions, photometric errors, and isochrones, we are now in a position to determine the probability distribution of the single star sequence in colour-magnitude space, $F_s(c,m)$. We pixelize the colour-magnitude plane finely enough that the minimum width of the main sequence is sampled by two pixels. This amounts to 450×450 pixels for the colour-magnitude space in the outer field, and 200×200 pixels in the inner field. Over-sampling the colour-magnitude plane does not inhibit the determination of the binary fraction, but adds computational time. Under sampling the colour-magnitude plane will degrade the sensitivity of the experiments, but will not skew the results in a particular direction [88].

The isochrone is interpolated so the mass points are evenly spaced. We found we needed to have a mass interval of $10^{-4}M_{\odot}$ in order to ensure points would fall in continuous pixels at the low-mass end of the main sequence. In order to calculate F_s , the colour and magnitude are calculated for each mass, and the value of the pixel corresponding to that colour and magnitude is increased by the amount proportional to the value of the mass function for that mass. For example, from Figure 2.8, one can see that the value of the colour-magnitude pixel corresponding to a 0.2 M_{\odot} star would be increased by 130 units, while the pixel corresponding to a 0.7 M_{\odot} would be increased by 35 units. After the single star sequence has been added to the colour-magnitude plane, it is convolved with the model for the photometric error. Finally, the resultant array, F_s , is normalized so that integrating over the entire colour-magnitude space examined yields a probability of unity. Figures 2.10 and shows F_s for the outer field.

2.3.2 The binary sequence

An unresolved binary system will photometrically lie off the main sequence in the colour-magnitude plane. The deviation from the MSRL is a function of the mass ratio, $q = M_2/M_1$. The deviation from the main sequence in magnitude increases monotonically, with the maximum deviation of ~ 0.75 magnitudes for q = 1. The deviation in colour is always to the red, with a maximum deviation for $q \simeq 0.8$ and no deviation for q = 0 and q = 1. Shown in Figure 2.11 is the CMD of the outer field with the MSRL overlaid,



Figure 2.8: The mass function of the outer field, with the log-normal fit, $N(M) = 132.6 \exp\left[\left(\frac{-\log_{10}(M/0.22M_{\odot})}{0.3\sqrt{2}}\right)^2\right]$, is shown with a line. The reduced χ^2 for the fit was 1.42.



Figure 2.9: The mass function of the inner field, with the log-normal fit, $N(M) = 56.56 \exp\left[\left(\frac{-\log_{10}(M/0.55M_{\odot})}{0.5\sqrt{2}}\right)^2\right]$, is shown with a line. The reduced χ^2 for the fit was 1.25.



Figure 2.10: The CMD showing the probability distribution of the single star sequence. The outlined area shows the "footprint" of the binary sequence, i.e., the area in which the probability of finding a star from the binary distribution is greater than 2% of the peak value of the binary distribution.

and the deviation from the MSRL of binaries with various mass ratios. Mass ratios greater than 0.4 are clearly distinguishable from the main sequence for most of the colour-magnitude space that we are exploring.

The simulation of the distribution of binary sequence in colour-magnitude space, $F_b(c, m)$, requires only two additional assumption over the simulation of the single-star sequence: the distribution of q's, referred to here as the q function, as in analogy to the mass function, and the binary fraction as a function of mass of the primary. In the field, it has been observed that high-mass stars are more likely to be found in binaries than low-mass stars. For this project we only have access to a relatively small range of primary masses, from 0.1 M_{\odot} to 0.8 M_{\odot} . We will therefore assume that there is no mass dependence over the range we are examining. Both dynamics [40, and references therein] and observations [37, 77] suggest that the q function should favor mass ratios close to unity over extreme mass ratio objects. Pinsonneault and Stanek [77] suggest a population with 55% of the q's drawn from a flat distribution, with the remainder drawn from a "twin" population (q>0.95). However, Pinsonneault and Stanek [77] are exploring a very different regime in terms of mass ratio and primary masses than we are in NGC 6397; for all the systems included in their study q > 0.45 and $M_1 > 7$ M_{\odot} . Fisher et al. [37] constrain the q function with $M_1 \gtrsim 1 M_{\odot}$. While they do not provide a functional form for the q function, they also find the distribution to be strongly peaked for q's close to unity. The simulation in Hurley et al. [54] was well fit over the range 0.18 < q < 1 by a power law, $N = N_0 q^{\alpha}$, with $\alpha = 1.7$. We are not suggesting that a power law has a physical basis; however, we require an analytical form in order to model the data.

We modeled the q function with a power law with three values of the exponent, $\alpha = -3$, 0, 5/3, and 3, i.e., the value that fits the simulations, a flat distribution, and two extreme bracketing values. The method for determining the probability distribution of binaries in colour-magnitude space, F_b , is very similar to that for F_s . Each star in the isochrone is chosen sequentially as the primary mass. Then, for all q's that imply a secondary mass greater than the hydrogen-burning limit, the luminosities of the primary and secondaries are combined, and the magnitude and colour of the system is calculated. The value in the pixel corresponding to that colour and magnitude is then increased according to the value determined from both the mass function (with respect to the primary) and the q function. Similar to the single star sequence, after the probabilities for all binaries have been calculated, the composite probability is convolved with a model of the photometric error and normalized to unity.



Figure 2.11: The CMD of the outer field. Overlaid is the MSRL, as well as the expected ridge lines for binary star sequences of several mass ratios.

 F_b is shown in Figures 2.12, 2.13, 2.14, 2.15 for q values of -3, 0, 5/3, and +3 respectively. Note, the binary sequence is well separated from the main sequence for low masses, but overlaps for high masses, particularly for negative values of q, i.e., equal mass binaries are suppressed. This is due to both the steeper MSRL, and the possibility of more extreme mass-ratio binaries with a secondary still above the hydrogen-burning limit at higher masses. The areas of high probability are coincident with areas in which binary sequences of varying mass ratios lie close to one another as shown in Figure 2.11. In fact, as is shown in Figure 2.11, the binary sequence for q = 0.2 is *never* distinguishable from the photometric scatter of the singlestar sequence. Our data are simply not sensitive to binaries with mass ratios more extreme than approximately 0.3.

2.4 Results

During this exercise, F_s and F_b need to be calculated only once for each value of α . Then, for any given binary fraction, f, the probability of observing a star can be written as:

$$F_t(c,m|f) = (1-f)F_s(c,m) + fF_b(c,m).$$
(2.4)

Our model of the observed colour-magnitude diagram is therefore a family of two parameters: the binary fraction, f, explicitly through the dependence on Equation 2.4; and α , implicitly through its influence on F_b .

For a given α , the best fitting f is estimated by the maximum likelihood (ML) method. The probability of finding n stars in a particular colourmagnitude range is simply given by the Poisson distribution. For this study, the explicit form is:

$$P(c,m \mid n) = N_{stars} F_t(c,m \mid f)^n e^{-F_t(c,m \mid f)} / n!.$$
(2.5)

The likelihood of the data given the model is then:

$$L(f) = \prod_{i=0}^{N} P(c, m \mid n_i).$$
(2.6)

The absolute value of L(f) is meaningless; however, as N becomes large, the value of f that maximizes L(f) will tend to the true value of f, namely f_0 . As shown in Romani and Weinberg [88], for one degree of freedom, the "1- σ " confidence in f will be the value of f such that $\ln f$ drops from its peak value by 0.5. The likelihoods for the various models for the outer field



Figure 2.12: The CMD showing the probability distribution of the binary sequence. A value of $\alpha = -3$ was assumed for this binary sequence. The outlined area shows the footprint of the single star sequence.



Figure 2.13: The CMD showing the probability distribution of the binary sequence. A value of $\alpha = 0$ was assumed for this binary sequence. The outlined area shows the footprint of the single star sequence.



Figure 2.14: The CMD showing the probability distribution of the binary sequence. A value of $\alpha = 5/3$ was assumed for this binary sequence. The outlined area shows the footprint of the single star sequence.



Figure 2.15: The CMD showing the probability distribution of the binary sequence. A value of $\alpha = +3$ was assumed for this binary sequence. The outlined area shows the footprint of the single star sequence.

and inner field are shown in Figures 2.16 and 2.17 respectively. The results are relatively insensitive to the choice of α —the difference between the most likely values is ~ 0.001 and ~ 0.01 for the inner and outer fields respectively. The derived values are tabulated in Table 2.1. Because of the observed α in the models of Hurley et al. [54], we prefer the $\alpha = 5/3$ value. The results for $\alpha = -3$ are not included in the previously mentioned figures and table, however, are shown in Figures 2.18, 2.19, and 2.20.

	outer	inner
α	f	f
0	0.021 ± 0.005	0.052 ± 0.010
5/3	0.012 ± 0.004	0.051 ± 0.010
3	0.008 ± 0.003	0.051 ± 0.010

Table 2.1: The best fitting values of the binary fraction for both the inner and outer fields.

In order to investigate the sensitivity of the result to the chosen value α in the outer field, we attempt to detemine the best fitting value of α and f simultaneously. If we allow all values of q, we find that the best-fitting value of α is ≈ -3 , and the 1σ -value of f is 0.04 ± 0.01 for the outer field. The probabilities for this series of models is shown in Figure 2.18. If a value of $\alpha = -3$ is used, much of the binary sequence lies directly over the single-star sequence as is shown in Figure 2.12, accounting for the higher allowed value of f. However, negative values of α are not likely to occur in a relaxed cluster. This is because, though the primordial binary frequency has been preserved, many of the binary systems have undergone three-body interactions over the lifetime of the cluster. In a three-body interaction, typically the lightest member of the system is eventually ejected, driving the typical q values close to unity.

Our data can not distinguish between single stars and binary systems with mass ratios more extreme that ≈ 0.2 . Where therefore simulate the binary sequence requiring the mass ratios to be greater than 0.2 and 0.5. These results are shown in Figures 2.19 and 2.20. These results are almost completely independent of α . This can be understood by noting that binary sequences of many values of q lie very close to one another in Figure 2.11. In order to recreate the number of stars in the region of the binary sequence we need a modest binary fraction, but the stars may be of almost any mass ratio. Note that even if we assume no constraint on α and simulate all values of q, the value of f is always less than 0.1, but with very reasonable



Figure 2.16: The likelihoods of the models for the outer fields.



Figure 2.17: The likelihoods of the models for the inner fields.



Figure 2.18: The probabilities obtained from fitting α and f simultaneously. All values of q are allowed in this case. If a value of $\alpha = -3$ is used, much of the binary sequence lies directly over the single-star sequence as is shown in Figure 2.12, accounting for the higher allowed value of f. The counters show the 1, 2, and 3 σ levels.

constraints on q or α , the derived value of f is 0.01 ± 0.005 .

2.5 Conclusions

This chapter presents a precise determination of the binary fraction of NGC 6397. The binary fraction in the outer field (ranging from 1.3–2.8 R_{hm}) is 0.012 ± 0.004 , and 0.051 ± 0.010 in the inner field (< 1 R_{hm}). The simulations of Hurley et al. [54] suggest the present day binary fraction beyond the half-mass radius is very close to the primordial binary fraction. Due to the competing creation and destruction processes that operate in the core of the cluster, the binary fraction in the central regions of the cluster is enhanced over the primordial value. In this context, our results suggest the primordial binary fractions for this project will be discussed in the concluding chapter of this thesis.



Figure 2.19: The probabilities obtained from fitting α and f simultaneously. Only values of q > 0.2 are allowed in this case. The result is relatively insensitive to the value of α . The counters show the 1, 2, and 3 σ levels.


Figure 2.20: The probabilities obtained from fitting α and f simultaneously. Only values of q > 0.5 are allowed in this case. The result is relatively insensitive to the value of α . The counters show the 1, 2, and 3 σ levels.

Chapter 3

White Dwarf Natal Kicks

3.1 Introduction

The large proper motions of some pulsars indicate that they have space velocities of hundreds of kilometers per second, which were presumably imparted to them when they became neutron stars (for recent reviews see Podsiadlowski et al. [78], Romani [87]). The question naturally arises, could something of this sort happen to white dwarfs? They are, of course, not observed to have comparably high velocities, and their births do not involve anything as energetic as a supernova.

Assuming that a kick does occur during the creation of a white dwarf (or indeed in some earlier stage of low- or intermediate-mass stellar evolution), what is its expected signature, and where would we expect to find it? In various stellar populations, white dwarf natal kicks would have different observational signatures. The dominant population of nearby white dwarfs are in the field. Though the Galactic disk is dynamically cold, it has a non-negligible component of random motions. When observing kicks from pulsars, these random motions are not a problem because the magnitude of the kicks are so high. The magnitude of the white dwarf kicks are expected to be smaller than random motions of disk stars, making it almost impossible to isolate this effect. Natal kicks would therefore be invisible in the dominant local population of white dwarfs. If natal kicks do indeed exist, this explains how they could remain undetected for so long.

Nearby open clusters have moderately sized populations of white dwarfs. However, the escape velocity from an open cluster is typically just several kilometers per second. The hypothesized kicks could easily have a magnitude of this order. Therefore, in an open cluster, the existence of a natal kick will manifest itself as a reduction in the number of observed white dwarfs. This approach has several limitations. Fundamentally, due to their low escape velocities, open clusters can only provide lower limits on the effect. Furthermore, it is far simpler to confirm the presence of a particular star than to convincingly argue that a star is truly missing—a "missing" star may simply be missed due to incompleteness, or could be locked in a binary system. With these limitations in mind, it is interesting to note that Weidemann [104] and Williams [106] have already found a hint that open clusters are deficient in white dwarfs. To investigate this question further, Fellhauer et al. [33] performed N-body simulations, and found that a kick with a magnitude of approximately twice that of the cluster velocity dispersion would deplete a loosely bound open cluster of almost all its white dwarfs.

The hypothesis that white dwarfs begin their lives with a small velocity excess leads to scenarios that are particularly conducive to testing in globular clusters for the following reasons:

- Large escape velocities Globular clusters have relatively large escape velocities, and hence will retain most of their white dwarfs.
- Large numbers of stars If the kicks to be of the same magnitude as the velocity dispersion, and randomly oriented with respect to a star's orbital elements, they will be very difficult to definitevely dectect for an individual star. We are therefore forced to make a statistical measurement. These clusters have hundreds of white dwarfs, so that tests can be made in a statistically significant way.
- **Relaxed population** Furthermore, globular clusters are relaxed stellar populations, implying that all stars will have approximately the same kinetic energy, i.e., $m_1v_1^2 = m_2v_2^2$. Because a star's velocity is related to the average distance it orbits in a gravitational potential, globular clusters have the extremely useful property that their stars are sorted by mass and velocity dispersion. Because of this, single-epoch photometry can reveal the relative velocity dispersions of various populations.
- Short crossing time and long relaxation time The progenitors of the newest white dwarfs in a typical globular cluster were recently the most massive main-sequence stars in the cluster, and therefore would have had the most concentrated radial profile. In a typical Galactic globular cluster, these stars had a mass of ~ 0.8 M_{\odot} , but have become white dwarfs with a mass of ~ 0.5 M_{\odot} . Assuming a quiescent birth, they begin their lives with the spatial distribution that is appropriate for their progenitors. Over the course of a relaxation time—of the order of a few times 10^8 years—they will acquire the more extended spatial distribution that goes with their new, lower mass. However, if white dwarfs are given natal kicks, they then start their lives with a velocity dispersion that is too large for their original spatial distribution. Over

the course of a crossing time—only a million years or so—they acquire a more extended spatial distribution, perhaps even more extended than they will have after relaxation has made their spatial distribution correspond to their new masses.

- White dwarfs exist in distinct dynamical states In a typical globular cluster, stars well down the white dwarf cooling sequence have been white dwarfs for dozens of relaxation times. These stars are almost certainly relaxed. On the other hand, the youngest white dwarfs have changed their masses very recently, and are almost certainly not relaxed.
- White dwarfs over a range of ages have similar mass Finally, globular clusters are old. As populations of stars age, they change ever more slowly. The rate of change of turn-off mass as a function of time is extremely high in a young population, but after several Gyr, it is low. In this chapter we will study white dwarfs with a range of cooling times. This means that they came from progenitors of different masses. If one was studying a young stellar population, the difference in mass of the progenitors would be significant, and would create a range of white dwarf masses. Because globular clusters are old stellar populations, the turn-off mass changes very slowly. The oldest white dwarfs examined in the study had cooling ages of 3.5 Gyr. The mass difference between the progenitors of these stars and the progenitors of the youngest white dwarfs be less than a tenth of a solar mass. Though the white dwarf initial-final mass relation is uncertain [34, 59], the mass difference between the white dwarfs must certainly be much less than the mass difference between their progenitors. Current constraints on the IFMR near the low-mass end of the relation [61] suggest the difference in mass between white dwarfs in a globular cluster with an cooling age of 3.5 Gys and one born today should be approximately 0.01 M_{\odot} . This allows us to treat all white dwarfs as essentially equal mass objects, greatly simplifying the analysis.

The question then is, do young white dwarfs in a globular cluster have a spatial distribution that is more extended than that of their progenitors? Until recently, the problem with addressing this question has been that white dwarfs in a globular cluster—particularly the old, fainter ones—are at a magnitude where they cannot be easily observed. Today, however, deep imaging with HST is able to reach faint enough to image the entire white dwarf cooling sequence in very nearby globular clusters [46, 85].

	NGC 6397	Messier 4
Distance Modulus (true)	12.03 ± 0.06 [46]	11.18 ± 0.02 [82]
A_{F814W}	$\dots 0.33 \ [46, 93]$	$\dots \dots $
[Fe/H]	-2.02 ± 0.07 [65]	-1.20 ± 0.15 [27]
$R_{half mass}$	140''[47]	$\dots 219''[47]$
$t_{half mass}$	$\dots 0.29 \text{ Gyr} [47]$	$\dots 0.66 \text{ Gyr} [47]$
Observed field radius	179 - 391''	

Chapter 3. White Dwarf Natal Kicks

Table 3.1: Cluster properties: References: Hansen et al. [46]; Sirianni et al. [93]; Kraft and Ivans [65]; Richer et al. [82]; Drake et al. [27]

In this chapter we examine the radial distributions of the white dwarfs in the two nearest globular cluster, Messier 4 and NGC 6397. We will show here that in these two clusters the radial distributions of their young white dwarfs are significantly more extended than that of their progenitors and old white dwarfs.

3.2 Observations

In NGC 6397 we used HST's ACS Wide Field Channel to image a single field centered 5' SE of the cluster center for 126 orbits (for details see Anderson et al. [4], Richer et al. [85]). As discussed in Chapter 2, this project was not the primary science driver for these data. The primary science driver for these data were to obtain images to such a depth that the end of the white-dwarf cooling sequence would be observable. To achieve this goal, observed area was sacrificed for depth, and though more multiple ACS fields would have been valuable for this aspect of the project, only a single field was observed. The exposure time was divided between the F814W and F606W filters. The field ranges from 179'' to 391'' from the cluster center, corresponding to 1.3–2.8 half-mass radii [47]. Proper motions, for the elimination of field stars, were measured on archival WFPC2 images. We were able to extend our magnitude limit considerably by blind measurement of the WFPC2 images at every point where the ACS images told us that there was a star, but the shorter exposure time of the WFPC2 images was still what set the limit of our completeness. Even so, artificial-star tests showed that at the magnitude that corresponds to a white dwarf age of 5 Gyr our completeness was still 92%.

In M4 we used positions, magnitudes, and proper motions from previous



Figure 3.1: The CMD of NGC 6397. The white dwarf ages, according to the Wood/Bergeron white dwarf cooling models, and the main-sequence masses, as determined from models of Baraffe, are indicated. The magnitude range of the old (lower) and young (upper) white dwarfs are shown with arrows.



Figure 3.2: The CMD of Messier 4. The white dwarf ages, according to the Wood/Bergeron white dwarf cooling models, and the main-sequence masses, as determined from models of Baraffe, are indicated. The magnitude range of the old (lower) and young (upper) white dwarfs are shown with arrows.

work [83, 84]. Our data cover two fields, at radii 29–220" and 270–430", respectively. Although the large radial range was helpful, the increased crowding in the inner field (compounded by the fact that these data were obtained solely with WFPC2) made completeness a strong function of radial position for magnitudes fainter than F814W \simeq 25.3, limiting us to white dwarfs younger than 1.8 Gyr. Table 5.3 summarizes the major properties of NGC 6397 and Messier 4.

The essence of this Chapter is to compare the radial distributions of white dwarfs of various ages with those of main-sequence stars of various masses. Neither mass nor age is directly observable; in what follows we will need models of both main-sequence stars and cooling white dwarfs. To assign a mass to each main-sequence star, we use main-sequence models of Baraffe et al. [7], with the metalicity appropriate for NGC 6397 and Messier 4. To assign an age to each white dwarf, we use a set of cooling models that were developed using interiors from Wood [110] and atmospheres from Bergeron et al. [10]. Our cooling models are for 0.5 M_{\odot} stars with pure carbon cores and thick hydrogen-rich atmospheres. Shown in Figures 3.1 and 3.2 are the colour-magnitude diagram (CMD) of NGC 6397 and Messier 4 with the main-sequence masses and white dwarf cooling ages superimposed.

3.3 Schematic research plan

The methods used in this chapter were developed specifically for this research, and hence may be unfamiliar to the reader. We will therefore first describe our plan without reference to the data, in order to put the discussion including data in context.

In this chapter we hope to test if white dwarfs receive a kick some time during their formation. We will be investigating this by way of comparing the radial distribution of white dwarfs younger than a relaxation time with the radial distribution of their main sequence progenitors. Furthermore, we use the white dwarfs older than a relaxation time as a "control" population by comparing them to the same-mass main sequence stars, and to the young white dwarfs themselves.

This will take place in two stages. The first step will be to determine how well we can distinguish the aforementioned "young" and "old" populations. To do this, we test if the radial distributions of the two white dwarfs populations are consistent. In fact, doing this, we make the zeroth-order assumption that we can detect mass segregation over the range of radii that we are investigating. A priori, we can generalize the results of this test into three categories: 1) old white dwarfs are more centrally concentrated than young white dwarfs, 2) the radial distributions of the old and young white dwarfs are statistically indistinguishable, or 3) old white dwarfs are less centrally concentrated than young white dwarfs.

- Young white dwarfs are dynamically cold Finding that young white dwarfs are more centrally concentrated, or dynamically colder, than old relaxed white dwarfs supports the hypothesis of a quiescent birth, or at least implies that the magnitude of a natal kick is very small. Remembering that absence of evidence is not evidence of absence, a result in this category makes the proof of the existence of a natal kick extremely difficult.
- Young and old white dwarfs are indistinguishable Finding that young and old white dwarfs have radial distributions which are indistinguishable tells us very little. The radial distributions could truly be similar, or we could simply have too few stars to detect a subtle difference between the means of the distributions. Even if we could determine that the distributions are physically similar, we still can not determine if this is due to a kick that coincidentally has the magnitude needed to heat the young population to its equilibrium value, or if relaxation happens faster than we assume for some other reason. Note, a significant result from this test is *not* necessary to prove the existence of a natal kick.
- Young white dwarfs are dynamically hot While this test yielding a significant result is not a *necessary* condition to prove the existence of a kick, it is *sufficient*. If we find the young white dwarfs are dynamically hotter than the old (i.e., their radial distribution is statistically less centrally concentrated), then we have proved the existence of a kick. This follows from two unassailable assumptions: main sequence stars lose mass as they become white dwarfs, and white dwarfs much older than a relaxation time will have relaxed. If these two assumptions are valid, it follows that old white dwarfs will be dynamically hotter than their main-sequence progenitors. If we observe young white dwarfs becoming colder as they age, rather than hotter, there is necessarily some mechanism that heats white dwarfs to greater than their equilibrium value at some post-main-sequence evolutionary stage!

While this test can prove (but not disprove) the existence of a kick, it is a rather blunt tool. We have at our disposal a yardstick with which to measure velocity dispersion. This is, of course, the radial distributions of populations of cluster main-sequence stars of varying mass. We will make use of this property of the globular clusters in the second set of tests. By comparing a population of white dwarfs with a sequence of populations of main sequence stars of varying masses, we can determine which matches the most closely. Unfortunately, due to the nature of statistical tests, we are limited to specifying which mass ranges of main sequence stars have radial distributions that are inconsistent with that of a given white dwarf population. However, if we show that all main sequence stars of mass less than m_1 and greater than m_2 have radial distributions inconsistent with a given white dwarf population, we have effectively constrained the velocity dispersion of that population to match that of main sequence stars with $m_1 < m < m_2$. Though this can not be rigorously justified, this seems more akin to a measurement rather than strict hypothesis testing, and hence, we will quote both 1σ and 2σ significance levels, rather than just 2σ .

Unfortunately, main sequence stars with masses equal to that of the white-dwarf progenitors, $m_{\rm prog}$, obviously are no longer present in the cluster. We are therefore limited to comparing the white-dwarf radial distributions with populations of main sequence stars that are hotter than the population we are actually interested in. However, if we find that young white dwarfs are hotter than a population of main sequence stars with a characteristic mass m_0 , where $m_0 < m_{\rm prog}$, we have effectively shown that young white dwarfs are hotter than their progenitors, and hence have proved the existence of a natal kick. In this test, the old white dwarfs serve as a control sample. We know the mass of these white dwarfs is $\sim 0.5 M_{\odot}$, and that they are relaxed. They should therefore have radial distributions that are consistent with $\sim 0.5 M_{\odot}$ main sequence stars. Though there are subtle effects due to sample size and completeness, the accuracy with which we can constrain the velocity dispersion of the old white dwarfs should be indicative of approximately how well we can constrain that of the young white dwarfs. In the following section we will do both tests, i.e., we will directly compare the radial distributions of the young and old white dwarfs to eachother, and compare both radial distributions to those of a series of main-sequence stars.

3.4 Radial distributions

To test whether two distributions differ significantly, we make use of two statistical tests: the familiar Kolmogorov-Smirnov (KS) test and the Wilcoxon rank-sum (RS) test. The KS test uses the maximum difference between the normalized cumulative distributions, while the RS test (equivalent to the Mann-Whitney U-test) compares rankings within the merged, sorted distribution. Each of the tests gives the probability that two random samplings from the same parent population will show differences as large as the observed ones. However, the RS test has several advantages over the KS test. Most importantly, it is designed to be applied as a one-tailed test, and is therefore more sensitive. Furthermore, in contrast to the KS test, the RS test is ostensibly sensitive only to differences between the means of two distributions, not their shapes. This makes results with small sample sizes more robust. The advantages of the KS test are the fact that people are generally familiar with it, and hence comfortable with it, and secondly, it is simpler to visualize. While we generally show the probabilities derived from the both tests, we use the results from the RS test exclusively to draw conclusions.

Mass segregation

Mass segregation is known to be present in NGC 6397 and Messier 4 [5, 63]. However, all subsequent tests we perform rely implicitly on our ability to detect mass segregation, so following the procedure outlined in Section 3.3, we first establish that we can detect mass segregation in our fields. We test the sensitivity of our data sets to mass segregation by comparing the radial distributions of main sequence stars of a variety of masses. Figures 3.3 and 3.4, show the cumulative radial distributions for these groups for NGC 6397 and Messier 4 respectively. The power with which the RS test can distinguish between them is indicated in Tables 3.2 and 3.3 for NGC 6397 and Messier 4 respectively.

Our data are clearly able to detect mass segregation in Messier 4. Note that the plateau in number counts in Messier 4 is caused by a gap in the two fields in which we are observing. We do not expect this to have an effect on our subsequent analysis. The picture is not as clear in NGC 6397. Due to the decreased range of radii we examine in this field the power of the RS test to distinguish between the radial distributions is decreased. However, even over the limited radial extent of our fields we can clearly detect segregation between well separated mass bins. As a sanity check, we test the effect of restricting the Messier 4 field to just the outer field. This should mimic the NGC 6397 observations, though because the WFPC2 field is smaller than the ACS field, and because the brightest stars are saturated (and therefore not measured) in the Messier 4 field will be slightly worse than with NGC

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			NGC	6397			
Bin	$0.2~M_{\odot}$	$0.3~M_{\odot}$	$0.4~M_{\odot}$	$0.5~M_{\odot}$	$0.6~M_{\odot}$	$0.7~M_{\odot}$	$0.8~M_{\odot}$
$0.1 M_{\odot}$	+1.66	+0.92	+0.93	-0.69	-1.42	-0.66	-0.99
$0.2~M_{\odot}$		-0.88	-0.63	-2.44	-3.21	-2.11	-2.28
$0.3~M_{\odot}$			+0.14	-1.59	-2.38	-1.56	-1.76
$0.4~M_{\odot}$				-1.68	-2.34	-1.46	-1.78
$0.5~M_{\odot}$					-0.68	-0.14	-0.42
$0.6~M_{\odot}$						+0.51	+0.12
$0.7~M_{\odot}$							+0.33

Table 3.2: The significance with with the RS test can distinguish between the main-sequence samples of various mass in NGC 6397. The results are given in sigmas. The sign of the result indicates if the test distribution is more or less extended than sample distribution. For instance, the first row of the first column indicates that the 0.1 M_{\odot} distribution is *more* concentrated than the 0.2 M_{\odot} distribution with a significance of 1.66 σ . Given how the table is set up, we would expect to see all results with a negative sign. The long dashes indicate the test is comparing the same mass bin. The ellipsis indicate the comparison for these mass bins is tabulated in an alternate column. Setting our confidence limit at 2 σ , all the significant results are shown in bold.

6397. Table 3.4 shows the results for the restricted field. By limiting the range or radii we examine, we drastically reduce the statistical significance of our result. The reduced statistical significance of the NGC 6397 results with respect to the Messier 4 results is almost certainly due to the restricted range of radii we examine in this field.

As a final sanity check, for both clusters, we sort the data by mass, and test the radial distributions of the odd-numbered indices against the evennumbered indices. We expect these tests to give non-significant results, and indeed, these tests give a 0.3 and 0.2 σ result for NGC 6397 and Messier 4 respectively.

Comparing "young" and "old" white dwarf populations

The white dwarfs that have had time to undergo dynamical relaxation should have a distribution similar to that of the $\sim 0.5 M_{\odot}$ main-sequence stars. In contrast, unrelaxed white dwarfs, assuming a quiescent birth, should have a more centrally concentrated distribution, like their more massive



Figure 3.3: A comparison of the radial distributions of various masses of main sequence stars in NGC 6397. The significance with which the distributions can be distinguished is listed in Table 3.2. The number of stars in each bin is: 0.08–0.15 M_{\odot} (272), 0.15–0.25 M_{\odot} (430), 0.25–0.35 M_{\odot} (371), 0.35–0.45 M_{\odot} (273), 0.45–0.55 M_{\odot} (237), 0.55–0.65 M_{\odot} (247), 0.65–0.75 M_{\odot} (180), 0.75–0.78 M_{\odot} (119).



Figure 3.4: A comparison of the radial distributions of various masses of main sequence stars in Messier 4. The significance with which the distributions can be distinguished is listed in Table 3.3. The number of stars in each bin is: 0.09–0.15 M_{\odot} (190), 0.15–0.25 M_{\odot} (423), 0.25–0.35 M_{\odot} (483), 0.35–0.45 M_{\odot} (498), 0.45–0.55 M_{\odot} (660), 0.55–0.60 M_{\odot} (315).

		Mess	sier 4		
Bin	$0.2~M_{\odot}$	$0.3~M_{\odot}$	$0.4~M_{\odot}$	$0.5~M_{\odot}$	$0.6~M_{\odot}$
$0.1 \ M_{\odot}$	-3.79	-4.47	-5.99	-6.63	-10.8
$0.2~M_{\odot}$		-0.94	-2.67	-3.35	-8.96
$0.3~M_{\odot}$			-1.93	-2.76	-8.81
$0.4~M_{\odot}$				-0.60	-6.83
$0.5~M_{\odot}$					-6.70

Table 3.3: The significance with with the RS test can distinguish between the main-sequence samples of various mass in Messier 4. The columns are the same as in Table 3.2.

	MICODICI 7		liciu oniy	
Bin	$0.2~M_{\odot}$	$0.3~M_{\odot}$	$0.4~M_{\odot}$	$0.5~M_{\odot}$
$0.1~M_{\odot}$	+0.54	-0.60	+0.87	+0.08
$0.2~M_{\odot}$		-1.36	+0.23	-0.61
$0.3~M_{\odot}$			+1.64	+1.00
$0.4~M_{\odot}$				-0.88

Messier 4—Outer field only

Table 3.4: The significance with with the RS test can distinguish between the main-sequence samples of various mass in the outer field Messier 4. The columns are the same as in Table 3.2.

progenitors. NGC 6397 and Messier 4 exist in the Galactic potential, and periodic changes in the lumpy potential could affect the relaxation process. In fact, the orbit of NGC 6397 has a disk-crossing time of ~ 100 Myr [60, 75], and one might naively expect disk shocking to play a role in the internal cluster dynamics. However, in relaxed clusters, such as NGC 6397, we expect these effects to be most important close to the tidal radius, and insignificant at the half-mass radius [41].

To study the distributions of the white dwarfs, we need a "young" and an "old" sample. When we refer to age here, we mean the time since the star became a white dwarf. We want all the stars of the young sample to be young enough that they have not experienced appreciable relaxation since they acquired the lower mass that goes with being a white dwarf. Therefore, we need to look at the relaxation time in our region—bearing in mind, however, that relaxation time is only an approximate time scale rather than an exact quantity. The Harris [47] catalog gives a relaxation time of 0.29 Gyr at a half-mass radius (2'3) for NGC 6397. Given that even the inner edge of our field is at a radius where the stellar density is lower than it is at 2'3, our stars should remain unrelaxed for ~ 1 Gyr. For Messier 4, the situation is slightly more complicated. The Harris [47] catalogue gives a relaxation time of 0.66 Gyr at the half-mass radius (3'7). Our sample has regions where relaxation is faster than this, and regions where it is slower.

For the choice of the old-sample boundaries there are two considerations. The lower age boundary should be high enough such that the stars have had enough time to have relaxed to the distribution that is appropriate for their white dwarf mass. In contrast, the upper age limit is driven by observational constraints rather than dynamical considerations. White dwarfs become dimmer as they age. The oldest white dwarf in our sample must be at a magnitude such that completeness is not a function of crowding (and therefore radial position). In NGC 6397, we take the old group to be from 1.4–3.5 Gyr (62 stars). In Messier 4, we are severely limited by position-dependant completeness at faint magnitudes, and are forced to only include stars with ages from 1.1–1.8 Gyr (35 stars). This will limit the significance with which we can distinguish between young and old white dwarfs in Messier 4.

For the choice of the young-sample boundaries there is a different set of considerations. The lower boundary should not be less than a crossing time. However, this timescale is of the order of Myrs—shorter than the youngest inferred white dwarf age in our sample—so this boundary can be set to zero Gyr. In principle it would be advantageous to set the upper boundary to select only the very youngest white dwarfs for this sample, and thereby

ensure that they would have had very little time to undergo interactions with other stars. In practice, it is necessary to set the upper boundary to an age comparable to a relaxation time in order to ensure a viable sample size, while maintaining a sample largely free from the effects of relaxation. In NGC 6397, we choose a young sample that goes from 0.0–0.8 Gyr (22 stars). In Messier 4, due to the large range in radii that we are examining, there is no single age that represents the relaxation time for the entire sample. We note that we are in no danger of inducing a false signal by choosing an upper boundary that includes some relaxed white dwarfs, but we will rather dilute any signal that is truly there. In Messier 4, we choose the young white dwarf sample to go from 0.0–0.9 Gyr (115 stars).

While our definitions of young and old were guided by physical principles and observational constraints, they remained somewhat arbitrary. In fact, there is some sensitivity to the upper boundary of the young group. For example, in NGC 6397, if that boundary is placed anywhere from 0.4 to 0.8 Gyr the result remains strongly significant; however, increasing the boundary beyond this decreases the significance of the result. Figures 3.5 and 3.6 compare the radial distributions of the young samples with the old ones in both clusters.

Contrary to what one would expect supposing a quiescent birth scenario, the younger white dwarfs tend to be farther from the cluster center than old white dwarfs. It is unclear which cluster we expect to get a stronger signal from a priori. We have more area, and a greater range of radii to examine in Messier 4, which would suggest we should see a stronger signal there; however, the old sample in Messier 4 has fewer members than in NGC 6397, and this should weaken the effect. The observation that the young white dwarfs have a more extended radial distribution is significant with a probability of 6×10^{-3} in NGC 6397 (2.8 σ in a normal distribution). This is an acceptably robust statistical result and hence, according to Section 3.3, we can conclude the existence of a natal kick in this cluster. In Messier 4, the result of this test is only significant with a probability of 9×10^{-2} (1.7 σ in a normal distribution). This is not a robust result in itself. As mentioned in Section 3.3, finding young white dwarfs are hotter than old white dwarfs is a sufficient, but not necessary, condition to proving the existence of a kick. Furthermore, we note that while Messier 4 does not provide compelling evidence for the existence of a kick in isolation, it does provide corroborating evidence when taken in the context of our results in NGC 6397.



Figure 3.5: A comparison of the radial distributions of the young and old white dwarfs in NGC 6397. The upper panel shows the RS test and the lower panel the KS test. The vertical lines in the upper panel indicate the radial position of each star, with the vertical bars denoting the locations of the old (upper) and young (lower) white dwarfs. Surprisingly, the young white dwarf population is *less* centrally concentrated than the old population.



Figure 3.6: A comparison of the radial distributions of the young and old white dwarfs in Messier 4. The panels are the same as in Figure 3.5, with the vertical bars denoting the locations of the old (upper) and young (lower) white dwarfs. Surprisingly, the young white dwarf population is *less* centrally concentrated than the old population.

Comparing white-dwarf and main-sequence populations

Additionally, we would like to know the magnitude of the kick. In order to address this question, we attempt to constrain the white dwarf velocity dispersion. By comparing the radial distribution of the white dwarfs with those of a series of main sequence populations of varying masses we can constrain the velocity dispersion of the white dwarfs.

The results are shown in Figures 3.7 and 3.8. Setting the acceptance level at 0.05 (2σ in a normal distribution), we can *reject* the hypothesis that the young white dwarfs have a radial distribution consistent with that of main-sequence stars more massive than ~ 0.5 M_{\odot} in both NGC 6397 and Messier 4. Furthermore, we note that we cannot reject the hypotheses that these same stars have radial distributions consistent with that of lower mass main-sequence stars. Taken together, these two results are proof of the existence of a natal kick for both clusters. Lending confidence to our method, Figures 3.7 and 3.8 also show that the radial distributions of the old white dwarfs are not inconsistent with that of main-sequence stars of their supposed mass, but are inconsistent with those of low mass (M < 0.25 M_{\odot}) main-sequence stars. These results are summarized in Table 3.5.

We have made the assertion that young white dwarfs have a more extended radial distribution that high-mass main-sequence stars. However, this is only a subtle difference, and is difficult to detect, particularly over the limited range of radii we examine in NGC 6397. Obtaining a significant result is further hampered by small numbers. The fact that it is difficult to predict *a priori* the statistical significance of a population that is genuinely more extended clouds the interpretation of Figures 3.7 and 3.8. However, we have at our disposal populations with known radial distributions, i.e., the main sequence stars. In order to get a sense of the effect of the sample size on Figures 3.7 and 3.8, we create analogous figures by drawing a sample of the same size from the main sequence rather than white dwarfs. We drawn from the main sequence with masses centered around 0.2 M_{\odot} and 0.5 M_{\odot} for the young white dwarfs and old white dwarfs respectively. This is shown in Figures 3.9 and 3.10. We are certain of the masses of the stars in this sample (by design), but can not determine the main-sequence bin to which they match to any better accuracy than for the white dwarfs. This gives us confidence that the somewhat noisy shapes of the lines in Figures 3.7 and 3.8 are simply due to the low numbers of stars in these samples.

Finally, we would like to know exactly at what evolutionary point the kick occurs. Schematically, to associate the kick with a particular stage of stellar evolution, we simply need to compare the radial distributions of stars



Figure 3.7: The RS probability that the radial distribution of a given white dwarf population could be drawn from the same parent distribution as the radial distribution of main-sequence star of a given mass for NGC 6397. The dashed lines indicate the equivalent of one-tailed 1σ , 2σ , and 3σ results. Note: the mass value of the points are the mean values of the mass bins. Adjacent points making up the lines are not independent, and are composed of approximately 90% the same stars. Each main-sequence bin has the same number of stars, and therefore, does not have precisely the same range of masses. Each mass bin has a range of approximately $\pm 0.1 M_{\odot}$.



Figure 3.8: The RS probability that the radial distribution of a given white dwarf population could be drawn from the same parent distribution as the radial distribution of main-sequence star of a given mass for Messier 4. The dashed lines indicate the equivalent of one-tailed 1σ , 2σ , and 3σ results. Note: the mass value of the points are the mean values of the mass bins. Adjacent points making up the lines are not independent, and are composed of approximately 90% the same stars. Each main-sequence bin has the same number of stars, and therefore, does not have precisely the same range of masses. Each mass bin has a range of approximately $\pm 0.05 \ M_{\odot}$.



Figure 3.9: The RS probability that the radial distribution of a mainsequence star population of the same size as a given white dwarf population could be drawn from the same parent distribution as the radial distribution of main-sequence star of a given mass for NGC 6397. The 0.2 M_{\odot} sample has 22 stars, and is supposed to mimic the young population. The 0.5 M_{\odot} sample has 62 stars, and is supposed to mimic the old population. The significance levels and main-sequence mass bins are the same as in Figure 3.7.



Figure 3.10: The RS probability that the radial distribution of a mainsequence star population of the same size as a given white dwarf population could be drawn from the same parent distribution as the radial distribution of main-sequence star of a given mass for Messier 4. The 0.2 M_{\odot} sample has 115 stars, and is supposed to mimic the young population. The 0.5 M_{\odot} sample has 35 stars, and is supposed to mimic the old population. The significance levels and main-sequence mass bins are the same as in Figure 3.8.

Chapter 3. White Dwarf Natal Kicks

	young		old	
	1σ	2σ	1σ	2σ
Messier 4	0.22–0.38 M_{\odot}	$< 0.52 \ M_{\odot}$	0.37–0.55 M_{\odot}	$> 0.21 \ M_{\odot}$
NGC 6397		$< 0.49~M_{\odot}$	$> 0.56~M_{\odot}$	$> 0.43 \ M_{\odot}$

Table 3.5: The young and old white dwarfs are constrained to have velocity dispersions equivalent to main-sequence stars in the mass range indicated (i.e., have radial distributions that are inconsistent with all main sequence stars out of the listed ranges). The σ indicated represents the level of significance assuming a normal distribution.

bracketing that evolutionary stage with the radial distribution of the mainsequence progenitors of the youngest white dwarfs. In evolutionary order, the mechanisms that may be responsible for a kick are: asymmetric winds on the red giant branch, the helium core flash, thermal pulses on the asymptotic giant branch, or mass-loss during the planetary nebula phase. In practice, constraining where the kick occurs is made more difficult by the brevity of post-main sequence evolutionary lifetimes. The short duration of these stages corresponds to low numbers of stars observed in them. Unfortunately, our dataset is not suitable to answer this question; we simply have too few post-main-sequence stars. Wide-field ground based observations could be used to get a handle on this problem.

3.5 Are white dwarfs born with a kick?

We have shown that the radial distribution of young white dwarfs in NGC 6397 and Messier 4 are more extended than would be expected given a quiescent birth. Other than a natal kick, is there an alternate explanation for this observation? One might imagine a mechanism related to the white dwarfs being in binary systems, either through binary disruption or the inhibition of white dwarf formation through common envelope evolution or mergers. However, the binary fraction in NGC 6397 is low [19, 23], and any explanation invoking binaries is implausible. Alternatively, the newly born white dwarfs could undergo very close encounters during their first passage through the inner regions of the cluster with their reduced masses, and have their velocities increased to greater-than-equilibrium values through stellar interactions alone. We have produced N-body simulations that accurately model both binary evolution and stellar encounters, but omit natal kicks; we

do not observe the young white dwarfs to have extended radial distributions in these simulations. A full description of these N-body models can be found in Hurley et al. [54]. NGC 6397 has a collapsed core, and one may consider if interactions with the central density cusp could somehow create this effect. However, the white-dwarf progenitors experienced the same potential as white dwarfs, so it is difficult to see how this could have an effect.

Assuming the existence of a natal kick, we can naively estimate its size in the following way. At the radial distances that we are considering, the measured velocity dispersion of giants in NGC 6397 is 3.3 km s⁻¹ ($M_{\text{giant}} \sim 0.8 M_{\odot}$)[80]. In accordance with equipartition of energy in a relaxed system, we expect the velocity dispersion of stars of mass M to be given by $\sigma = \sigma_{\text{giant}} \sqrt{M_{\text{giant}}/M}$. Thus, σ for 0.2 M_{\odot} stars should be about twice that of the giants. Because velocity dispersions add in quadrature, the magnitudes of impulsive kicks that would transform the radial distribution of a 0.8 M_{\odot} population to that of a 0.5 M_{\odot} or 0.2 M_{\odot} population are 3 km s⁻¹ and 6 km s⁻¹, respectively. A similar analysis yields essentially the same results for Messier 4. These values are comparable to the velocity dispersion of a typical globular cluster, but still well below their escape velocities.

The fact that we are examining a relatively small range of radii, and that we have a small number of stars in the white dwarf sample, prevents us from putting tight constraints on the velocity dispersion of the young white dwarfs. However, we note that the young white dwarfs have an extended radial distribution, implying an extra source of velocity dispersion. The most natural explanation for this observation is the existence of a natal kick. The recent developments, and directions for future research are discussed in the concluding chapter of this thesis.

Chapter 4

Spectra of White Dwarfs in Messier 4

4.1 Introduction

As we have seen in Chapter 3, much can be gleaned from the photometric study of white dwarfs, particularly those found in star clusters. However, as with main sequence stars, a detailed study of an individual white dwarf requires spectral information. Due to their intrinsically low luminosities, the spectra of white dwarfs, particularly those in globular clusters, have been poorly studied in comparison. In many ways, white dwarfs are much simpler objects than their main sequence progenitors. The intense gravitational field of a white dwarf typically stratifies the star into three distinct layers with minimal overlap between them. They are: 1) a 0.6 M_{\odot} degenerate carbonoxygen core, 2) a $10^{-2} M_{\odot}$ non-degenerate helium envelope, and 3) a nondegenerate $10^{-4} M_{\odot}$ hydrogen atmosphere. Due to gravitational settling, white-dwarf atmospheres are generally devoid of metals, and hence have very simple spectra, similar to that of A-type main-sequence stars. These are the well known DA white dwarfs, and comprise approximately 3/4 of all known white dwarfs. A relatively low-signal-to-noise-ratio, low-resolution spectrum is typically all that is required to determine the spectral class of a white dwarf. However, if one can obtain a high-signal-to-noise-ratio spectrum, one can additionally constrain the mass and luminosity of the star.

In this Chapter we use spectra of white dwarfs in the globular cluster Messier 4 in order to investigate two issues: the ratio of DA/DB stars in star clusters, and the spectroscopic mass of Population II white dwarfs.

4.1.1 Spectral types

Due to their degenerate natures, white dwarfs have extreme surface gravities, and therefore have very broad atmospheric lines. The two most common spectral types of white dwarfs are named after the main sequence spectral

Chapter 4. Spectra of White Dwarfs in Messier

Spectral Type	Primary Characteristics
DA	Only Balmer lines; No He I or metals present
DB	He I lines; No H or metals present
DC	Continuous spectra, no lines deeper than 5%
DO	He II strong, no He I or H present
DZ	Metal lines only, no H or He lines
DQ	Carbon features, either atomic or molecular

Secondary Characteristics

Р	Magnetic white dwarfs with detectable polarization
Н	Magnetic white dwarfs without detectable polarization
E	Emission lines are present
V	Optional symbol to denote variability
X	Peculiar or unclassifiable spectra
? or :	Uncertain assigned classification

Table 4.1: The white dwarf classification scheme developed by McCook and Sion [72]. The temperature class is defined as 50400 K divided by the effective temperature of the star.

type they most closely mimic, but are prepended with a 'D'. DAs therefore show prominent hydrogen absorption lines, like A stars, while DBs show prominent helium I lines. However, unlike main-sequence stars, white dwarf spectral types are not indicative of temperature. DAs are commonly found with temperatures ranging from 60000 K to 4000 K. Types other than DA or DB are relatively rare. The most commonly used white-dwarf classification scheme, that of McCook and Sion [72], is outlined in Table 4.1. The scheme consists of the letter D (indicating that the star is degenerate), a primary spectral class (indicating the dominant atomic species), and secondary characteristics (e.g., magnetization or polarization). Following the spectral class, a temperature index is often included. This temperature is defined as 50400 K divided by the effective temperature. Stars that are primarily in one class, but have weak features of another class have both spectral types in their names, with the primary type first. For instance, a star with strong H I features and weak He I features would be referred to as DAB. Spectra of each of these classes can be found in the first data release (DR1) of the Sloan Digital Sky Survey (SDSS) [64]. Examples of the primary spectral classes from SDSS-DR1 are shown in Figures 4.1, 4.2, and 4.3. Figure 4.1 shows

the temperature sequence of DA white dwarfs. Figure 4.2 shows a more restricted temperature sequence of DBs. Characteristic spectra of types DC, DO, DZ, and DQ are shown in Figure 4.3.

Perhaps the greatest difference between white-dwarf and main-sequence spectral classification is that, unlike main-sequence stars, the spectral types of white dwarfs can change multiple times throughout a star's life. Hansen and Liebert [43] outline the presumed multiple evolutionary sequences of white dwarfs with "thick" and "thin" hydrogen envelopes. A "thick" envelope has a mass of $10^{-4} M_{\odot}$. A "thin" envelope can have a hydrogen envelope as small as $10^{-10} M_{\odot}$. These stars become white dwarfs as DAs and DOs (showing He II absorption lines) respectively. As the stars cool, the DOs will transform into DBs. At these temperatures (150000 K–30000 K), DAs far outnumber DBs. In fact, from 45000 > T > 30000 there is the "DB gap" [69]. This temperature range was, until recently, completely devoid of DB stars. With DR4, SDSS had accumulated 10^4 white-dwarf spectra. Within the DR4 sample, approximately 10 DB white dwarfs were found in this temperature range [31]. While this temperature range is no longer strictly a "gap", the DA/DB ratio is approximately 2.5 greater at 30000 K than it is at 20000 K. This implies that an atmospheric transformation takes place in approximately 10% of DAs as they cool through this range[31].

Below 30000 K, a helium convection zone is established. Convective velocities can become high enough to overshoot, and mix the helium layer with the hydrogen atmosphere, converting a type DA to a type DB. This occurs in $\sim 25\%$ of white dwarfs in this temperature range in the field. As the stars cool toward the ZZ Ceti-instability strip, the ratio of DA/DB stars approaches unity[43].

For temperatures lower than this, the situation becomes even more complicated and uncertain [11]. At these low temperatures the variety of spectral types increases, but in the non-DA gap (6000K > T > 5000K) non-DA white dwarfs have yet to be observed. In cool white dwarfs (< 5000 K) with hydrogen-dominated atmospheres, or slightly hotter white dwarfs with helium-rich atmospheres, neither hydrogen nor helium lines are excited, leading to an almost featureless spectrum in these DC stars . Even more rare are white dwarfs showing metal lines. These are the DQs, for those showing carbon features, and the DZs, for other atomic species. Finally, (DH) DP white dwarfs show evidence of having (non)-polarized magnetic fields.

The spectrum of a star on the main sequence is typically very complicated, and by comparing with model spectra, one can determine the star's mass and metallicity. The information is all but erased in the process that transforms the star to a white dwarf. When studying white dwarfs in the



Figure 4.1: Spectra of DA1 to DA7 white dwarfs. These spectra were obtained from SDSS DR1.



Figure 4.2: Spectra of DB2 to DB6 white dwarfs. These spectra were obtained from SDSS DR1.



Figure 4.3: Spectra of a DC, DO, DQ, and DZ white dwarf. These spectra were obtained from SDSS DR1.

field, the information is most likely unrecoverable. In contrast, in a cluster, one can determine the progenitor mass of the white by examining isochrones. Furthermore, the metallicity for a given cluster white dwarf can be constrained by studying the abundances in the remaining main sequence stars. However, it appears that the use of cluster white dwarfs as field-white-dwarf analogs may not be justified. In the field, in the temperature appropriate for our white dwarf candidates, the ratio of DA/DB stars is about 3.5:1 [64]. The situation in open clusters is radically different.

Prior to the year 2000, cluster white dwarf spectra had been obtained in a piecemeal fashion—one or two stars for each cluster, typically by different authors. The Canada-France-Hawaii Telescope (CFHT) Open Star Cluster Survey (OSCS) changed this. The idea behind this survey was to obtain high-quality wide-field images of many open clusters. The richness of the open clusters could then be identified, and rich clusters (with well-populated white dwarf cooling sequences) could be followed up spectroscopically. Previously, when spectral types were obtained for only several white dwarfs at a time, the absence of a particular spectral type was neither particularly surprising nor interesting.

In 2005, Kalirai et al. [58] reported 21 spectral identifications in NGC 2099. If the same DA/DB ratio observed in the field held for this cluster, several DB white dwarfs would have been found. Contrary to expectations, none were found. From Poisson statistics, finding a DB/DA ratio of 0/21 would occur approximately 2% of the time simply due to statistical fluctuations. This finding prompted Kalirai et al. [58] to examine all white dwarfs that had been spectroscopically identified in young open clusters at that point, all were of type DA. This had a vanishingly small chance of occurring due to a statistical fluctuation under the hypothesis that the same DA/DB ratio held as in the field.

An obvious explanation for this observation was not apparent. The clusters hosting these white dwarfs span a range of metallicity. Invoking metallicity as a causal agent for this effect therefore seems unjustified. One early explanation for this effect was the re-accretion of residual intra-cluster gas in open clusters. This explanation did not hold up on closer examination. The escape velocity in open clusters is typically very low ($v_{\rm esc} \sim 1 \, km \, s^{-1}$), and gas ejected by stellar winds ($v_{\rm wind} \sim 10 \, km \, s^{-1}$) would be expected to quickly escape. It was found that the accretion rate of intra-cluster gas in open clusters should be no greater than that of the ISM in the disk. Finally, Kalirai et al. [58] postulated an explanation based on the mass of the white dwarfs; because NGC 2099 was a young cluster, the young white dwarfs had high masses (~ 0.8 M_{\odot}) compared to the average disk population (~ 0.6 M_{\odot}). The idea was that He convection would be inhibited in high-mass white dwarfs, and therefore even white dwarfs with thin hydrogen atmospheres would remain type DA in this temperature range. Globular clusters are the oldest stellar population, and should be creating among the lowestmass white dwarfs in the Galaxy. It is of interest to determine whether the same paucity of DB white dwarfs exists in globular clusters. Early indications are that it does indeed hold. In 2004, Moehler et al. [76] obtained the spectra of 12 white dwarfs in NGC 6752 and another 8 in NGC 6397. Taken in isolation, the chance of finding no DB white dwarfs given the number observed in NGC 6752 and NGC 6397 is 0.07 and 0.17 respectively. We will consolidate, and add to, these numbers in this thesis.

4.1.2 Initial-final mass relation

Only the most massive stars will generate a type II supernovae. All other stars will end their lives as white dwarfs. Stars evolving to become white dwarfs will shed their outer layers, and decrease in mass. Interestingly, a large range of initial progenitor masses yield a relatively small range of final white-dwarf masses. The relation between a star's initial ZAMS mass and its eventual white dwarf mass is known as the initial-final mass relation. This is an important relation in many areas of astrophysics, and yet, until recently was quite poorly constrained. White dwarfs discovered with the CFHT Open cluster survey, and followed up with spectroscopy from Gemini/GMOS and Keck/LRIS have recently tightened up the constraints on this relationship [59, 61].

Kalirai et al. [61] derived an empirical relation for the initial-final mass relation with the following form:

$$M_{\text{final}} = (0.109 \pm 0.007) M_{\text{initial}} + 0.394 \pm 0.025 M_{\odot}.$$
 (4.1)

The initial-mass of the progenitors of our stars is essentially the same as the turn-off mass on the cluster. According to the isochrones of Dotter et al. [26], the turn-off mass of M4 is 0.78 M_{\odot} . Equation 4.1 predicts a white-dwarf mass of 0.48 \pm 0.02 M_{\odot} .

This empirical relation is in line with theoretical expectations. From stellar evolutionary models, one may estimate the white dwarf mass for a particular initial mass by examining the mass of the carbon-oxygen core of an AGB star of the same initial mass. A lower limit may be determined by assuming that none of the helium shell that is burning while the star is on the AGB is incorporated into the white dwarf. In this case, the models of Dotter et al. [26] predict a mass of 0.43–0.46 M_{\odot} . This range depends on how much mass-loss is assumed on the RGB (i.e. where the star ends up on the HB). A bluer HB star has a lower carbon-oxygen core mass by the end of the AGB. Messier 4 has some stars on either side of the instability strip so this entire range of white-dwarf carbon-oxygen core masses could be present. An upper limit can be determined by assuming that all the mass in the helium shell is incorporated into the white dwarf. In this case, the models of Dotter et al. [26] predict a mass of 0.51–0.53 M_{\odot} , with the lower mass again corresponding to lower initial HB mass. The functional form of the theoretical relationship is not linear, but considering the uncertainties that exist with the theory of thermal pulses on the AGB, and in the absence of high-quality data, this is sufficient for our purposes.

Due to their extreme ages, and their subsequent low main sequence turn off (MSTO) masses, globular clusters provide a window on the extreme lowmass end of the relation. The initial-final mass relation is of interest to the following areas: dating star clusters with white-dwarf cooling models, galactic chemical-enrichment models, and Type II supernovae rates.

Cluster dating

The white-dwarf cooling sequence, when investigated in detail, contains a great deal of information regarding a cluster. Most closely linked to our research group is the comparison of the white-dwarf cooling sequence with model cooling sequences of varying ages in order to constrain the age of a cluster [44–46]. In a star cluster, the white-dwarf cooling sequence is much fainter than the stars usually used to constrain the age—main sequence and post-main sequence stars. The white-dwarf-cooling sequence would naively seem to be an inferior method. However, using stellar isochrones to constrain absolute ages is notoriously difficult for old stellar populations[86]. The fundamental limitation is that cluster isochrones evolve very slowly beyond an age of several Gyr. Compounding this difficulty is uncertainty of cluster distance and metallicity. At a given distance and metallicity, an older cluster will have a fainter and redder MSTO. This effect can be mimicked by increasing both distance and metallicity. For globular clusters, an uncertainty of merely a few tenths in the distance modulus or metallicity can render ages uncertain up to several Gyrs. Finally, evolutionary tracks from different theoretical groups differ, exacerbating this problem. If all the uncertainties associated with this method are taken into account, the age constraints are often so loose as to be uninteresting.

As mentioned above, the determination of the white-dwarf cooling age

requires a far more intensive observational campaign than for MSTO ages. Until recently, the depth required to use this technique was simply unobtainable for any globular cluster. The Hubble Space Telescope (HST) equipped with large format imagers, such as the Wide Field and Planetary Camera 2 (WFPC2) and the Advanced Camera for Surveys (ACS), have made it possible to image the entire main sequence and the majority of the white-dwarf cooling sequence for the nearest globular clusters. In 2001 Richer et al. [83] used 123 HST orbits with WFPC2 to image M4. Hansen et al. [44] found a white-dwarf cooling age of 12.7 ± 0.7 Gyr. In 2005, Richer et al. [85] used 126 HST orbits with the ACS to image the white-dwarf cooling sequence of NGC 6397. Hansen et al. [46] obtained an age of 11.5 ± 0.5 Gyrs. In 2008, Richer et al. were approved for 121 HST orbits to attempt to determine the cooling age for 47 Tucanae. Assuming the upcoming HST servicing mission is successful, this group will have determined the white-dwarf cooling age for the nearest globular clusters spanning a range of metallicities and dynamical states. MSTO dating methods are largely limited by theoretical uncertainties. While important theoretical uncertainties certainly exist, the barriers to using the white-dwarf cooling method remains largely limited by observational constraints. The determination of the white-dwarf cooling age of these clusters required imaging to an approximate V-band magnitude of ~ 30 . These magnitudes are simply unobtainable from the ground with the present generation of ground-based telescopes.

If photometry of sufficient depth exists, dating with the white-dwarf cooling sequence has the potential to be more precise than than any MSTO method. Just as importantly, the systematic uncertainties of this method of cluster dating are different from those associated with the MSTO methods. The mass of a white dwarf determines its radius, and therefore its luminosity, which directly controls the cooling rate. One must adopt a particular initialfinal mass relation when attempting to determine the white-dwarf cooling age of a cluster. A significant uncertainty that remains unresolved with the white-dwarf cooling age is the mass of the white dwarfs in the cooling sequence. The most recent attempts to date clusters via the white dwarf cooling sequence assume an extrapolation of the initial-final mass relation constrained at high masses holds for low masses. By measuring the masses of these white dwarfs directly, we hope to be able to reduce this uncertainty such that it becomes unimportant for the age constraint. Again, the 2004 study of Moehler et al. [76] has made a start in this direction. Based on the average of fits to 20 individual spectra, they concluded the mass of the young white dwarfs in NGC 6397 and NGC 6752 is $0.53 \pm 0.03 M_{\odot}$. Unfortunately, the spectra are of low signal-to-noise ratio, and the fits to the individual
spectra give wildly different results.

Finally, globular-cluster white dwarfs provide the only access to information about the mass function of population II stars more massive than $\sim 1 M_{\odot}$. If one can disentangle the initial-final mass relation and cooling models for white dwarfs of various masses, one can assign a main-sequence mass to each observed white dwarf, and thereby study the mass function of a cluster beyond its main-sequence turn off.

Chemical enrichment

As mentioned above, stars that end their lives as white dwarfs shed much of their hydrogen envelope during the final stages of their evolution. The mass that is returned back to the interstellar medium will be enriched with metals. The amount of enriched gas that is fed back to the ISM depends directly on the initial-final mass relation. Globular clusters host the oldest, and therefore lowest mass, hydrogen burning stars. Constraining the initialfinal mass relation in a globular cluster will therefore tell us how much mass is returned to the ISM for any stellar population that exists today. This is a key parameter for understanding the chemical enrichment of galaxies.

Supernovae rates

If the initial-final mass can be tightly constrained, it is possible to determine the progenitor mass that will yield a white dwarf of the Chandrasekhar mass. This is, by definition, the most massive white dwarf that can exist. This progenitor mass is therefore the dividing line between evolutionary branches leading to type II supernovae and white dwarfs. Knowing this limit precisely has implications for galaxy energetics and feedback [99]. This is an important unknown parameter in many active fields of research [95].

4.2 Observations

When attempting to obtain spectra of white dwarfs, a key stage of the project is the selection of candidates. The project will make use of preimaging from GMOS on the Gemini South Telescope, and WFPC2 on the Hubble Space Telescope. The object select from the pre-imaging will be followed up with the GMOS spectrograph on Gemini South and the LRIS spectrograph on Keck I. When selecting field white dwarfs from a magnitudelimited photometric survey, the typical strategy is simply to select blue objects. Samples selected this way will be heavily contaminated with blue main-sequence stars. White dwarfs are intrinsically faint, thus in a magnitudelimited survey, they will be found preferentially nearby. Therefore, if available, high proper motions can be an effective selection criterion [73]. If multi-colour photometry is available, candidates may be selected in colourcolour space (for examples see Eisenstein et al. [30], and references therein). Finally, if one is selecting white dwarfs in a cluster, the white-dwarf cooling sequence highlights possible white dwarfs.

After the targets have been selected, spectroscopy of white dwarfs is generally similar to that of other targets. Yet, this particular project provided us with several challenges. Most importantly, globular clusters are generally quite distant, and therefore the stars in them are faint. Our targets generally have F555W magnitudes of approximately 23. Obtaining high signal-to-noise ratio spectra at these magnitudes limits us to rather lowresolution spectroscopy. Adding to the difficulty of this project is the fact that globular clusters are extremely crowded environments. Performing a adequate sky-subtraction in these environments is a challenge.

4.2.1 Target selection

For this project, we used two sources of pre-imaging to select stars that lie on the white-dwarf cooling curve of Messier 4. Our most likely contaminants here are field white dwarfs at the approximate distance of Messier 4 and bright compact blue galaxies. The most secure targets came from the HST photometry published in Richer et al. [84]. Stars in these fields have their cluster membership confirmed by their proper motions. Although there is a small chance of contamination by field stars that happen to share the proper motion of Messier 4, or mis-measured proper motions, these identifications are quite certain. However, both the GMOS and the LRIS fields of view were much larger than the WFPC2 field. In order to maximize the number of spectra obtained, our targets were required to be spread roughly evenly over the GMOS and LRIS fields, and therefore some had to be outside the HST field of view. These stars could only be selected in colour-magnitude space, and not proper-motion space. The targets selected only from the GMOS imaging were therefore less secure. All of the instruments used have different fields of view. Some areas have been studied with all three instruments, while others have only been studied by one or two. The footprints for the various instruments are shown in Figure 4.4.



Figure 4.4: The footprints of the various instruments used in this thesis.

4.2.2 HST imaging

Our study of this cluster began with an HST imaging campaign. Using archival images it was possible to select only stars that share the proper motion of Messier 4. This leads to quite a clean CMD. This CMD is shown in Figure 4.5.

From this CMD, the white dwarf-cooling sequence can be clearly seen. The only point requiring caution when examining the HST photometry is the fact that the extremely sharp PSF and low sky background of HST makes faint stars easily visible. The excellent observing conditions present when imaging faint stars with HST will not be shared during their ground-based follow up. Naively, one would assume the best targets are simply the brightest white dwarfs. However, on the ground, crowding is a very important effect. Care must be taken to select candidates that are isolated on the WFPC2 images. Shown in Figure 4.6 are the right ascensions and declinations of all white dwarf candidates brighter than F555W=24.

4.2.3 Gemini target selection

As can be seen from Figure 4.4, due to the small area of WFPC2, only a fraction of the GMOS field can be proper motion cleaned. In order to effectively select white dwarf candidates in the entire GMOS field, GMOS pre-imaging was taken in both the g' and i' filters. The images were corrected for bias and flat fielding by the Gemini pipeline. The pre-processed images were then reduced with the standard DAOPHOT/Allstar reduction techniques [97]. Again, a white dwarf sequence can be clearly seen blue-ward of the main sequence. This photometry was not used for anything other than the target selection, and hence was not rigorously calibrated. The only calibration that was applied was to ensure than the tip of the GMOS white dwarf-cooling sequence has the same g' magnitude and g'-i' colour as the WFPC2 white-dwarf cooling sequence did in F555W and F555W-F814W. Hence, the quoted g' magnitudes are effectively equivalent to F555W magnitudes, while the quoted g'-i' colours are approximately equal to F555W-F814W magnitudes in the white dwarf region, but become progressively less accurate to redder colours. The CMD constructed from the GMOS imaging is presented in Figure 4.7. Like the WFPC2 photometry, a clear white dwarf cooling sequence can be seen starting around g'=22.5 and g'-i'=0.3.

Figure 4.8 shows the footprint of the white dwarf candidates selected from the GMOS photometry (red points). The WFPC2 candidates are shown in green. Note the paucity of GMOS points in the inner WFPC2



Figure 4.5: The colour magnitude diagram from the HST-WFPC2 photometry $% \mathcal{A}$



Figure 4.6: Astrometry of the WFPC2 white dwarf candidates. A bright white dwarf candidate was defined as any star brighter than F555W < 24 and with F555W-F814W < 1.3.



Figure 4.7: The CMD constructed from the GMOS photometry. There is a clear white dwarf cooling sequence extending from 22 < g' < 24 with an approximate colour of g'-i'=0.3. In contrast to the proper-motion cleaned CMD constructed from the WFPC2 photometry, this CMD is contaminated by a field population immediately blue-ward of the main sequence. Of more concern is the possible contamination by field white dwarfs and perhaps blue compact galaxies.

fields. Though there are many white dwarfs here, the crowding is such that they become very difficult to detect with ground-based telescopes, such as Gemini. The red GMOS points are those stars with magnitudes between g'=20 and g'=24 and colours less than g'-i'=0.9. The priority 1 objects were believed to be white dwarfs. The priority 2 and 3 objects were objects that were likely not white dwarfs, or were difficult to observe. They were included to fill the slit mask. The GMOS photometry, WFPC photometry, astrometry (showing the position relative to the various instrumental footprints), and a postage stamp image of a small section of the GMOS image centered on the target are shown for each star in Appendix 1. An example of this, for target LRIS09 is shown in Figure 4.11. Finally, all the relevant candidate selection measurements are tabulated in Table 4.2.

4.2.4 Keck target selection

Though LRIS on Keck is a formidable spectroscope, the imaging produced by this instrument is generally of surprisingly poor quality. We constructed a CMD from LRIS photometry, but it did not have a coherent white-dwarfcooling sequence. Furthermore, by the time we were selecting targets for the LRIS mask, the GMOS spectroscopy has already been performed. We therefore had spectroscopic confirmation of many white dwarfs. The selection for LRIS was therefore performed with the same inputs as that for the GMOS photometry. The locations of the LRIS targets shown in the GMOS colour-magnitude space are shown in Figure 4.9. The astrometry of the LRIS targets are shown in Figure 4.10.



Figure 4.8: The footprint of the white dwarf candidates selected from the GMOS photometry. The WFPC2 candidates are shown in green. Note the paucity of GMOS points in the inner WFPC2 fields. Though there are many white dwarfs here, the crowding is such that they become very difficult to detect with ground-based telescopes, such as Gemini. The red GMOS points are those stars with magnitudes between g'=20 and g'=24 and colours less than g'-i'=0.9. The priority 1 objects were believed to be white dwarfs. The priority 2 and 3 objects were object that were likely not white dwarfs, or were difficult to observe for some reason. They were included to fill the slit mask.



Figure 4.9: The colours and magnitudes of the targets selected for LRIS spectroscopy are shown in blue. The targets for GMOS are shown in red. Most of the targets are in common, though there are several targets that are studied in only one of the datasets.



Figure 4.10: The footprint of the LRIS white dwarf candidates selected from the GMOS photometry. The WFPC2 white dwarfs are shown in green, while GMOS white dwarfs are shown in red. Those objects for which spectra was obtained are shown in orange. Note the paucity of GMOS points in the inner WFPC2 fields. Though there are many white dwarfs here, the crowding is such that they become very difficult to detect with ground-based telescopes, such as Gemini. The red GMOS points are those stars with magnitudes between g'=20 and g'=24 and colours less than g'-i'=0.9.



Figure 4.11: The GMOS photometry (upper left), WFPC2 photometry (lower left), astrometry showing position relative to the instrumental footprints (upper right), and postage stamp section of the GMOS CCD centered on the target (lower right) for object LRIS-09 (GMOS-14).

F555W					25.209						22.597	22.280	26.964	23.477	24.311	24.987		23.525	23.158	23.492	23.488	22.943			22.765
F555W-	F'814W				0.925						0.240	2.028	1.314	0.341	2.535	0.847		0.276	0.300	0.063	0.311	0.352			0.339
°°,		23.253	22.765	21.250		22.607	22.624	22.566	22.084	22.709	22.415			23.412			22.639		23.136	23.542	23.307	22.944	22.394	23.298	22.538
g'-i'		0.290	1.183	0.277		0.247	0.251	0.255	0.723	0.594	0.238			0.202			0.241		0.294	0.227	0.386	0.290	1.070	0.183	0.458
Υ		4448.0820	4215.0210	4167.0000	4100.6360	4061.5210	3921.2340	3729.6640	3581.0910	3491.5250	3163.0410	3116.3560	2971.5050	2731.8940	2669.8440	2607.0240	2474.5990	2411.8440	2264.9050	2160.6570	2000.7850	1985.3230	1879.6510	1721.1880	1512.7830
Х		2792.5150	4280.1540	3554.6580	2358.3820	2528.5750	4355.6940	4188.6270	1622.3460	1979.2760	3493.4700	4993.1560	1559.5270	3459.2080	2066.8700	1587.8570	2588.1240	4127.4700	3068.6450	3046.9740	4105.5270	3422.0510	4179.1110	3045.0370	3737.0720
DEC (J2000)		-26.558850	-26.553958	-26.553072	-26.551100	-26.551060	-26.548010	-26.544160	-26.541440	-26.539590	-26.532780	-26.528100	-26.530000	-26.524056	-26.524000	-26.521400	-26.518950	-26.517503	-26.514660	-26.512550	-26.509192	-26.508958	-26.506740	-26.503661	-26.499360
${ m RA}~({ m J2000})$		245.959350	245.925637	245.942104	245.971000	245.965350	245.923950	245.927700	245.985900	245.977800	245.943450	245.912000	245.986700	245.944237	245.976400	245.989000	245.964000	245.929067	245.953050	245.953650	245.929558	245.945062	245.927850	245.953608	245.937900
GMOS #		GMOS-21				GMOS-20	GMOS-19	GMOS-18	GMOS-16	GMOS-13	GMOS-14						GMOS-10		GMOS-06	GMOS-08			GMOS-09		GMOS-04
LRIS $\#$		LRIS-00	LRIS-01	LRIS-02	LRIS-03	LRIS-04	LRIS-05	LRIS-06	LRIS-07	LRIS-08	LRIS-09	LRIS-10	LRIS-11	LRIS-12	LRIS-13	LRIS-14	LRIS-15	LRIS-16	LRIS-17	LRIS-18	LRIS-19	LRIS-20	LRIS-21	LRIS-22	LRIS-23

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F555W	22.795								23.620				
F555W- F814W	0.361								0.391				
°°,	22.638	23.170	24.529	23.475	22.743	22.863	22.855	22.777	22.944	23.355	22.657	21.483	
g′-i′	0.242	0.353	1.891	0.515	0.277	0.275	1.073	1.163	0.290	0.359	0.231	0.969	ction.
Υ	1469.2190	1250.9860	858.3190	842.7700	759.3000	491.9880	155.4430	856.8970	1985.2870	2399.3340	3652.2490	3239.9650	r target sele
Х	4390.8490	3392.3070	4666.6750	5315.9280	4548.6590	2368.1940	3850.8800	4550.3750	3422.2570	2619.4380	3871.3800	3191.9450	nents used fo
DEC (J2000)	-26.498406	-26.494110	-26.486017	-26.485622	-26.484030	-26.478870	-26.471894	-26.486000	-26.508960	-26.517420	-26.542630	-26.534360	: The measuren
${ m RA}~({ m J2000})$	245.923075	245.945700	245.916808	245.902079	245.919450	245.968950	245.935296	245.919450	245.945100	245.963250	245.934900	245.950350	Table 4.2
GMOS #		GMOS-03			GMOS-02	GMOS-01		GMOS-05	GMOS-07	GMOS-11	GMOS-15	GMOS-17	
LRIS #	LRIS-24	LRIS-25	LRIS-26	LRIS-27	LRIS-28	LRIS-29	LRIS-30						

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4.2.5 Spectral reduction

A spectrograph decomposes light from a source into its component wavelengths. Multi-object spectrographs (MOS) can generally act as both imagers, with two spatial dimensions, and spectrographs, with one spatial dimension and one spectral dimension. There are several steps that must be taken in order to transform a raw CCD output from a MOS into a form which can be analyzed. In very general terms the steps are as follows:

- **Overscan correction** When CCDs are read out, they are biased to a small non-zero voltage. This signature is the first thing that must be removed when processing a CCD image.
- **Spectral flat fielding** Each pixel in a CCD has a slightly different sensitivity. This effect must be corrected if one is to extract useful data from them. This pixel-to-pixel variation is a function of wavelength. In an imaging mode, each pixel is exposed to roughly the same wavelength of light, and the pixel-to-pixel sensitivity variations can therefore be characterized by exposing the entire CCD to a even light source. In a spectroscopic mode, each pixel will typically be exposed to a different wavelength of light. The aforementioned technique will therefore not work. A spectroscopic flat field is obtained by exposing the spectrograph to a bright source with a simple spectral shape, such as a halogen lamp. The low-order spectral shape is then divided out of the signal for each slit in a MOS individually, leaving the high-order pixel-to-pixel variation. This pixel-to-pixel correction is then applied to the CCD.
- Extraction and sky subtraction The process of transforming the spectrum from a two dimensional image to a one dimensional one is called extraction. Each slit on a MOS slitmask has a spectral and spatial dimension. In order to maximize the signal-to-noise ratio of the final spectrum, the slit must be as narrow as possible to eliminate scattered light, yet not so narrow that the star light is excluded. The spectral dimension of the slit should roughly match the seeing in which the observation was obtained. The slits will generally be several times larger in the spatial dimension. For our observations, the slits were generally 5".0 in the spatial dimension, and either 0".8 or 1".0 in the spectral dimension. The slit is extended in the spatial direction in order to correct for the emission of the sky. For each pixel in the spectral direction, the spectrum is extracted by summing the signal

over the pixels in the spatial direction for an extent of approximately the seeing. The sky value is calculated for each pixel in the spectral direction by determining the average of the pixels outside the region in which the signal was extracted from. The sky is subtracted from the signal from the source in the final extracted spectrum.

- Wavelength Calibration In order to get useful information from the spectrum, we must now calibrate the spectral direction to an absolute wavelength scale. This is accomplished by taking the spectrum of a lamp with known emission lines, and processing it in an identical manner to the science spectra. The mapping from pixel to wavelength is then calculated for each slit from the calibration spectrum, and applied to the science spectra.
- **Flux calibration** Finally, the response of the system as a function of wavelength is corrected. This is accomplished by observing a "flux standard". Flux standard stars are bright stars with very smooth and well-characterized spectra. They are often hot white dwarfs of type DB or DC. The measured signal as a function of wavelength is compared with the known values, and the correction from one to the other is calculated. This same correction is then applied to the science spectra.

While the general steps for reducing MOS spectra are similar regardless of the particular spectrograph, each instrument has its own subtleties. The steps used for both the GMOS and LRIS reductions will now be briefly summarized.

GMOS Reductions

GMOS is a low-resolution spectrograph with a $(5'.5 \times 5'.5)$ field of view. We used a 1".0 wide slit, and the B1200 grating (0.47 Å/pixel), which simultaneously covers 1500 Å, from 3700 Åto 5300 Å. The plate scale of the blue CCD is 0.47 arcsec/pixel.

In total, we obtained approximately 14.5 and 9 hours of science exposure with GMOS in 2005A and 2006B respectively. Because of the queue system we were able to require all exposures to be obtained in sub-arcsecond seeing. The GMOS observations obtained are listed in Table 4.3. Messier 4 is a southern target, and due to Gemini South's latitude, it was possible for all science images to be obtained with very low airmass. The slits were cut as to be aligned with the North/South axis, and the angle of the slit axes to

exposure	date	number of
times (s)		exposures
3600	06/06/2005	2
3600	06/07/2005	3
1687	06/08/2005	1
3600	06/09/2005	5
3600	08/04/2005	2
3600	08/08/2005	1
620	08/08/2005	1
3600	08/09/2005	1
3600	$08/19/2006^{\dagger}$	3
3600	$08/20/2006^{\dagger}$	1
3600	$08/21/2006^{\dagger}$	2
3600	$09/15/2006^{\dagger}$	1
3600	$09/16/2006^{\dagger}$	1
3600	$09/20/2006^{\dagger}$	1

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Table 4.3: GMOS observing times. The observations marked with † were obtained as part of a program for which the author was the principal investigator.

the zenith may vary. However, because of the low air mass at which the images were obtained the effects of differential atmospheric refraction are thought to be minimal.

The raw data frames were downloaded from the Canadian Astronomy Data Center (CADC) in multi-extension FITS (MEF) format. We reduced the data using the Gemini IRAF Package, version 1.4. GMOS is composed of three separate chips. The dispersion axis is perpendicular to the long axis of the chips, and therefore the dispersed spectra will cross the gaps between the chips, leaving gaps in the spectral coverage. The precise spectral coverage for a given star depends on its position in the detector, and is therefore slightly different for each star. Because the spectral range is different for each star, we are unable to choose a central wavelength such that no star will have a Balmer line that falls on a chip gap. To avoid a Balmer line falling on a gap and rendering the spectrum useless, we obtain the spectrum at two different grating offsets. This is equivalent to dithering a camera when obtaining imaging. The central wavelength was shifted from the default

exposure	date	number of
times (s)		exposures
2700	04/21/2007	2
900	04/21/2007	1
2700	04/22/2007	3
1800	04/22/2007	1
1800	$07/14/2007^{\dagger}$	2
2100	$07/14/2007^{\dagger}$	1
1800	$07/15/2007^{\dagger}$	4
1800	$04/12/2008^{\dagger}$	5

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Table 4.4: LRIS observing times. Observations obtained by the thesis author personally are indicated with † .

value of 4620 Å to 4720 Å for half of the exposures. These two sets of spectra are handled separately until the very last step of combining the data.

When reducing GMOS multi-slit data, the first IRAF task is to "prepare" the data for subsequent tasks using the GPREPARE task. Reducing multi-slit GMOS data using the Gemini IRAF package is partially automated. This means that the locations of the slits must be communicated to IRAF without user input. The positions of the slits are contained in the Mask Definition File (MDF). GPREPARE attaches the MDF to each science and calibration frame, and updates certain keywords in the headers of the frames, such as those for gain and read noise.

The next step is to create a master bias. This is accomplished using the GBIAS task. Several times per semester, Gemini creates calibration images. Instead of creating our own master bias, we used the one created by Gemini. The spectral flat images are then normalized using the GFLATTEN task. The master bias image and appropriate flats were applied to the science data using the GSREDUCE task. This task completed several functions, including subtracting the overscan, trimming the overscan region, subtracting the bias frame, cleaning the image for cosmic rays (analogously to the COSMICRAY task), applying the flat field correction, mosaicing the three chips together, interpolating the pixels within each chip gap, and cutting the MOS slits into separate spectra.

For the wavelength calibration, we obtained multiple spectroscopic frames from a CuAr lamp exposure. The automatic wavelength fitting routine, GSWAVELENGTH, was good enough to find a rough wavelength calibration, however, the calibration had to be verified interactively. The resultant residuals in the fit to the emission lines were all well behaved, typically at the 0.04 Å level (0.08 pixels). With this template, we then used the GSTRANSFORM task to apply the wavelength calibration to the science frames.

The sky subtraction was completed with the GSSKYSUB task. The central ~ 1 "0 of each slit was assumed to contain the stellar signal, and the sky was estimated from the remaining pixels. This typically gave us 6 pixels for the star, and 11 pixels both above and below the star for the sky estimation (the stars are centered in the slits). The sky was estimated using the usual IRAF task. A low-order polynomial (typically or degree 1 or 2) was fit to the pixel values outside the extraction region, and interpolated to determine the sky value at all pixels. There are only 37 spatial pixels due to our factor of two binning in this direction. A number of pixels at both the top and bottom of each slit were discarded.

Due to the faintness of our stars, the automatic routines for extracting the spectra were generally unsuccessful. We manually extracted each of the spectra into a 1D format with the GSEXTRACT task, using the "variance" weighting (i.e., pixels are weighted by the variance based on the data values and a Poisson/ccd model using the "gain" and "readnoise" parameters. This was a particularly challenging and uncertain aspect of the reductions. The sensitivity of GMOS-South decreases to wavelengths bluer than H_{ϵ} , and the already faint stars became almost invisible. Delineating the trace in these parts of the spectra was very uncertain. Unfortunately, in order to determine a reliable spectroscopic mass, H_{ϵ} is crucial, and H_8 is very useful. The lack of flux at these blue wavelengths ultimately limited the utility of the GMOS spectra. The resulting manually extracted spectra were combined in a weighted average based on the signal-to-noise of each spectrum.

LRIS Reductions

We furthermore obtained multi-object spectroscopy of white dwarfs in Messier 4 using LRIS on the Keck I telescope. LRIS is a dual-beam, low-resolution spectrograph with a $(5'.0 \times 7'.0)$ field of view. We operated with the D560 dichroic, which splits the light at roughly 5600 Å. The red side was not used in this study, and will not be discussed further. For the blue side, we used a 1''.0 wide slit, and the 600/4000 grism (0.63 Å/pixel), which simultaneously covers 2580 Å, from 3300 Åto 5880 Å. The plate scale of the blue CCD is 0.135 arcsec/pixel.

The target selection was performed using GMOS photometry, and used

the information from the preceding GMOS spectra. The final target selection and mask design was performed by David Reitzel and Jason Kalirai (at UCLA and UCSC respectively). The reduction of the LRIS data was performed by the author entirely within IRAF. In total, we were granted 7 half-nights with LRIS, and obtained 10.6 hours of exposure. However, the data were of highly variable quality. The observations were collected with seeing ranging from 0.8 to over 2.0. In an uncrowded field, the signal-tonoise ratio for equal exposure times will be higher for observation obtain in good seeing. In a crowded field this effect is exacerbated due to the scattered light from nearby bright stars that is incident upon the slit in poor seeing conditions. The final signal-to-noise ratios of the spectra are dominated by the signal-to-noise ratios of just several exposures obtained in the best seeing conditions. For two nights in April 2008, we obtained no data. On the third night of that observing run we obtained 2.5 hours of science exposure, but it was of limited value. At our resolution and central wavelength, only H_{α} , which is a rather poor mass indicator, landed on the red side. Hence, the red-side spectra were not reduced. Messier 4 is a Southern target and, from Mauna Kea, the zenith will always be to the North. The slits for the LRIS mask were cut to be only slightly off the North-South axis. We therefore think that differential atmospheric refraction will have at most a very small effect on the LRIS spectra.

The blue-side LRIS spectra were reduced as follows.

The IMRED > CCDRED > CCDPROC task was used to do the overscan correction, bias subtraction, and overscan trimming on all the images. Though LRIS is composed of two chips, the spectral axis is parallel to the long side of the chips, and hence, the spectra do not cross chip boundaries. The observations were therefore all taken at the same central wavelength.

The flat field normalization was done with the NOAO > TWODSPEC > APEXTRACT > APNORMALIZE task. The quartz lamp has very low signal at blue wavelengths. To improve the S/N ratio, we stacked several flats before determining the flat field correction. The flux at blue wavelengths is still low, but the correction is applied nonetheless. The trace, sky subtraction, and extraction were all performed with the APALL task, again using "variance" weighting. Bad pixels were interpolated over.

LRIS has better blue sensitivity compared to GMOS, and the trace at blue wavelengths was therefore far more certain. This field is extremely crowded. Our ability to obtain a reliable trace, and perform accurate sky subtraction is dependent on the particulars of each individual slit. This varies widely from slit to slit. Figures 4.12, 4.13, and 4.14 show the flux from the individual slits as a function of spatial pixel, integrated along the spectral dimension. Note that all slits have the same horizontal and vertical scales (54 pixels and 10⁵ ADU respectively), expect for LRIS-02, which has twice the vertical range. Note that some of the stars have very clean slits, and many pixel with which to calculate the sky values (e.g. LRIS-04), while others have noisy backgrounds or bright stars on the slits (e.g. LRIS-09). While it is still possible to extract the spectra from the noisy slits, it is not clear how it will affect the mass estimate from the spectral line fits. For example, excess flux from the bright companion could fill in some of the Balmer lines. The mass constraints from stars with noisy backgrounds are therefore suspect.

The wavelength calibration was calculated from spectra of three lamps containing Hg, Zn, and Cd. There are not all that many transitions at these wavelengths, and furthermore, at the resolution we used, some of the lines had odd shapes. The residuals were not as well distributed as with the GMOS data, and the final dispersion of the wavelength residuals of the line fits was typically ~ 0.1 Å (0.15 pixels). While there are not as many transitions as with the CuAr lamp used for the GMOS data, and the residuals had a higher dispersion, the wavelength solution is sufficiently precise for our purposes. The wavelength calibration was calculated with the NOAO > ONEDSPEC > IDENTIFY task, and was applied to the science images with the DISPCOR task.

The flux calibration for these stars was calculated from the flux standard HZ44. The calibration was calculated using the tasks STANDARD, and SENS-FUNC, and were applied to the science spectra using the task CALIBRATE. Finally, the individual spectra were combined using SCOMBINE, with the weights according to their signal-to-noise ratios.

4.3 Results

For both spectral identifications and the fitting of the Blamer lines, the continuum is not important. It has therefore been subtracted from both the GMOS and LRIS data in subsequent analysis. Shown in Figure 4.15 is the flux-calibrated spectrum of object LRIS-02 *before* the continuum subtraction.

Figure 4.16 shows the results for all the GMOS spectra that could be reliably extracted. GMOS objects 05, 12, and 17 have been omitted due to the difficulty in locating the star on the slit. These spectra clearly have quite low signal-to-noise ratios, and hence are not adequate to determine spectral masses. They are therefore only useful to identify the spectral type. Despite



Figure 4.12: The cross-sections of the slits of the stars LRIS-00, 02, 04, 05, and 06. The sections are displayed such that the white dwarf is in the center of the figure. The cross-sections were calculated from the first science exposure on April 22, 2007. This was one of the best nights in all our runs in terms of seeing.



Figure 4.13: The cross-sections of the slits of the stars LRIS-08, 09, 15, 16, and 19. The sections are displayed such that the white dwarf is in the center of the figure. The cross-sections were calculated from the first science exposure on April 22, 2007. This was one of the best nights in all our runs in terms of seeing.



Figure 4.14: The cross-sections of the slits of the stars LRIS-20, 22, 23, 24, and 29. The sections are displayed such that the white dwarf is in the center of the figure. The cross-sections were calculated from the first science exposure on April 22, 2007. This was one of the best nights in all our runs in terms of seeing.



Figure 4.15: The flux-calibrated spectrum of object LRIS-02

the low signal-to-noise ratio of these spectra, they are *all* clearly type DA. Of the 21 slits observed, 18 were white dwarfs of type DA, and the 3 objects not shown here proved to be too faint to reliably extract.

Though there was more measurable flux in the LRIS data, some of the spectra still suffered from low flux levels. The LRIS data were obtained after the GMOS data had been reduced. Therefore, the spectral type of all the objects in common between the two datasets were already known. Many of the faint stars were difficult to extract, and it was clear from preliminary reductions that the final signal-to-noise ratio for these stars would be too low to determine a mass. If the spectral types of these stars were already known, it was not worthwhile to fully reduce the data. The following objects had confirmed spectral types, and from preliminary reductions, were too dim, or had another star contaminating their slit making the extraction difficult: LRIS-7, 12, 17, 18, 21, 25, 26, and 28. Several objects that did not lie on the white dwarf cooling sequence were included on this slit mask simply to fill it. These objects were CV candidates, or white dwarf-main sequence star binary candidates. After preliminary reductions, none of these objects appeared to be interesting, and hence were not pursued further. These objects include: LRIS-1, 3, 10, 11, 13, 14, 27, and 30. The spectra of all the stars that were fully reduced are shown in Figure 4.17. Out of the 31 objects observed, 23 were strong white dwarf candidates. Eighteen of the objects were previously identified with the GMOS data. Ten of these objects had their spectral types confirmed, and an additional five were identified as type DA. Of the 23 strong candidates, all have now been confirmed as DAs.

It is unclear precisely how to characterize the signal-to-noise ratio of an observed spectrum in a straightforward manner. While many of our objects have strong signal in the vicinity of H_{β} and H_{γ} , the combination of atmospheric transmission and CCD sensitivity decreases the signal at bluer wavelengths, and almost all flux bluer than H_8 is lost. The widths of the Balmer lines in these stars are such that there is no continuum in which to measure the signal-to-noise ratio between the high-order lines. The signalto-noise ratio for the spectra, as measured between H_{β} and H_{γ} , are listed in Table 4.5. The values do not characterize the signal-to-noise ratio of the entire spectra, but are useful as a metric to compare the relative quality of the individual spectra. All stars listed in Table 4.5 are of sufficient quality to determine the spectral type. However, it is not clear how scattered light from nearby stars affects the spectroscopic determination of the mass. We determine spectroscopic masses for all the stars, but the masses determined for those stars without isolated slits should not be trusted.



Figure 4.16: The spectra for all the GMOS objects with reliable traces. The continuum has been subtracted. GMOS objects 05, 12, and 17 have been omitted due to the difficulty in locating the star on the slit. Due to low luminosities of the stars, the spectra are of too low signal-to-noise ratio to determine the mass. Despite the low signal-to-noise ratio of these spectra, they are *all* clearly type DA.

Object	SNR	Isolated	Composite	HST	
		on slit	candidate	selected	
LRIS 00	17.1	no	no	yes	
LRIS 02	25.4	yes	no	no	
LRIS 04	17.0	yes	yes	no	
LRIS 05	19.7	no	yes	no	
LRIS 06	19.0	yes	yes	no	
LRIS 08	13.1	yes	no	no	
LRIS 09	11.7	no	no	yes	
LRIS 15	13.2	yes	yes	no	
LRIS 16	4.1	no	no	yes	
LRIS 19	11.8	no	no	no	
LRIS 20	11.5	yes	no	yes	
LRIS 22	16.7	no	no	no	
LRIS 23	18.1	no	no	yes	
LRIS 24	15.9	yes	yes	yes	
LRIS 29	13.5	no	no	no	
Composite	27.3	n/a	n/a	n/a	

Table 4.5: The signal-to-noise ratio of the spectra determined between 4500 Å and 4750 Å. Because the signal-to-noise ratio varies with wavelength, the quoted value is not a characterization of the signal-to-noise ratio of the spectra, but can be used as a measure of the relative quality of the spectra.



Figure 4.17: The spectra for all the LRIS objects with reliable traces. The continuum for all objects has been subtracted. The objects not shown here were omitted due to the difficultly in locating the star on the slit, or if they had a confirmed spectral type and were too faint to obtain a mass. As with the GMOS data, these spectra are *all* clearly type DA.

4.3.1 Spectral types

In total, we have spectroscopically typed 23 white dwarfs. They are all of type DA. At this point it is of interest to examine if the same DA/DB ratio observed in the field holds for the cluster. We can use binomial statistics to determine the probability of observing no DB white dwarfs. The binomial distribution has the following form:

$$f(k;n,p) = \frac{n!}{k!(n-k)!} p^k (1-p)^{n-k}$$
(4.2)

where k is the number of observed "hits" (in this case, the observation of a DB white dwarf), p is the probability of any one event being a "hit", and n is the total number of events. When no hits are observed, Equation 4.2 takes a particularly simple form:

$$f(0;n,p) = (1-p)^n \tag{4.3}$$

Under the assumption that the DA/DB ratio, r, is the same in the cluster environment as it is in the field, the probability of observing a DB is:

$$p = \frac{1}{1+r}.\tag{4.4}$$

From SDSS-DR1, Eisenstein et al. [30] found $r \sim 3.5$. Putting the preceding equations together, we have an expression for the probability, f, of observing no DB white dwarfs:

$$f = \left(\frac{r}{1+r}\right)^n \tag{4.5}$$

Assuming all 23 white dwarfs are cluster members, the probability of observing no DB white dwarfs is 3×10^{-3} . There is almost certainly a small level of contamination by field white dwarfs. For instance GMOS-16 (LRIS-07) has distinctly narrower Balmer lines than the other stars in this sample. It is also brighter and redder than the other white dwarfs in this sample. These properties are consistent with a white dwarf that is both cooler and more nearby than the cluster white dwarfs. However, even if only 12 of the observed stars are genuine Messier 4 white dwarfs, this observation is still significant at the 2σ level. We should also take into account the white dwarfs identified by Moehler et al. [76]. If we add the 20 white dwarfs identified by them in two other cluster, the total number of globular cluster identifications rises to 43. The chance of observing no DBs in a sample this large is only 2×10^{-5} . This is not the first observation recognizing the DA/DB ratio is different in clusters and the field. In 2005, Kalirai et al. [58] reported finding 21 DA white dwarfs in the young open cluster NGC 2099. This prompted them to examine the spectral types of all ~ 65 previously identified white dwarfs in young open clusters. They found all were type DA. The chance of this occurring due to a random statistical fluctuation was vanishingly small. Several explanations for this were investigated, but the most promising explanation of this discrepancy focused on the high mass of these white dwarfs. The white dwarfs in NGC 2099 were approximately $0.8M_{\odot}$. In Messier 4, the mass of the main-sequence turn off mass is only $0.8M_{\odot}$. Clearly the stellar mass is not the variable controlling this effect.

Since 2005, a substantial number of white dwarfs in open clusters have had their spectral types determined. There are now a handful of non-DA white dwarfs identified. DBs have been found in NGC 6633 and NGC 6819 by Williams and Bolte [107] and Kalirai et al. [61] respectively. A DBA has been identified in the Hyades, and a DQ has been identified in NGC 2168 [108]. While it no longer appears to be impossible to form a non-DA white dwarf in the cluster environment, the formation mechanism is clearly strongly suppressed.

We have no convincing mechanism to suggest for this suppression. Kalirai et al. [58] explored the idea of re-accretion of intra-cluster hydrogen. The youngest white dwarfs in our sample have cooling ages of 10^7 years. A "thick" hydrogen atmosphere has a mass of $10^{-4} M_{\odot}$. The conversion of a "thick" atmosphere to be converted to a "thin" atmosphere requires an accretion rate of at least $10^{-11} M_{\odot}/yr$. But to serve as an explanation for the differing DA/DB ratios, the accretion rate in a cluster must be significantly greater than that in the field. This idea has difficulties. The winds from evolved stars typically have velocities of the order of $10^1 km/s$, while cluster escape velocities are only of the order of $10^0 km/s$. This means that the clusters will generally not retain any gas. Assuming a star is near the outskirts of the cluster, the flux of gas through any spherical shell will simply be equal to the rate of mass loss from the main-sequence stars. In the cluster, the wind velocity is greater than the velocity of the stars. However, this can be treated in an identical manner to field white dwarfs moving through the ISM in the field. The expression to calculate the accretion rate of gas is

$$\dot{M} = 4\pi\rho \frac{G^2 M^2}{V^3}.$$
(4.6)

Kalirai et al. [58] and concluded that it was not a possible mechanism. In light of these latest findings, and the absence of any other plausible mechanism, perhaps this scenario needs to be reexamined. Perhaps the model of the gas flowing smoothly away from the cluster is too simple. Clearly the wind from different stars will be traveling in different directions, at least in the inner parts of the cluster. Could these collisions cause shocks, and increase the accretion rate?

4.3.2 Spectroscopic masses

The precision of spectral-line fits depends directly on the signal-to-noise ratio of the spectrum. While we observe the spectrum between H_{β} to beyond H_8 , not all the lines are equally good at discriminating between changes in log(g) and T_{eff} . For stars in the temperature range which we are observing, an increasing temperature and gravity both decrease the strength of the loworder lines. It is only the combination of the higher- and lower-order lines which are sensitive mass indicators. The higher-order lines are suppressed much more strongly at high temperature than they are for high gravity (see Figure 3 in Bergeron et al. [9]). The best of our spectra have signal-tonoise ratios that are only marginally high enough for a reliable spectroscopic fit. Shown in Figure 4.18 are synthetic spectra of varying signal-to-noise ratios. Note that these simulations include only Poisson noise. They are, in effect, simulations of pre-atmospheric spectra. There is no sky or instrument signature in the simulations. Unlike the observed spectra, the signal-tonoise ratio of the synthetic spectra is even across the spectra. From visual comparison with Figure 4.18, most of our spectra seem to have signal-tonoise ratios in the H_{ϵ} region of approximately 5–10, with a signal-to-noise ratio of ~ 15 for our brightest objects.

To increase the quality of our spectra we explore co-adding individual spectra to create a single high-signal-to-noise-ratio spectrum. If indeed all the stars are genuine members of Messier 4, they should all have very similar masses. This is because we are observing stars exclusively near the tip of the cooling sequence, and hence have similar cooling times, and therefore similar progenitor masses. However, we need to proceed here with some care; as mentioned above, the profile of the Balmer lines change both with mass *and* temperature [9]. If we are going to combine spectra from more than one object, we need to ensure that they have both similar masses and temperatures. Five of our white dwarfs, listed in Table 4.6, have very similar GMOS photometry. The most significant magnitude discrepancy in the magnitudes of these stars is between LRIS 06 and LRIS 15 (0.073 mags). We can estimate the difference in temperature in the following way. A difference in magnitude of 0.07 is equivalent to a luminosity ratio of 1.07.



Figure 4.18: The precision with which white dwarf masses can be constrained using noisy spectra. Note, the critical region for constraining the mass of a star is the higher-order Balmer lines. Our spectra have a typical signal-tonoise ratio of 5–10 in this region. Note that these simulations include only Poisson noise. They are, in effect, simulations of pre-atmospheric spectra. There is no sky or instrument signature in the simulations. This figure was created by Jason Kalirai at UC Santa Cruz.

Chapter 4. Spectra of White Dwarfs in Messier 4

Object	g'-i'	\mathbf{g}'
LRIS 04	0.247	22.607
LRIS 05	0.251	22.624
LRIS 06	0.255	22.566
LRIS 15	0.241	22.639
LRIS 24	0.242	22.638

Table 4.6: The photometry of 5 white dwarfs with very similar photometry. These stars, excluding LRIS 05, will be referred to as the composite sample.

Assuming the stars are well represented by a blackbody, and have the same mass (and therefore radius), $T_1/T_2 = (L_1/L_2)^{\frac{1}{4}}$. This implies the stars in this sample all have the same temperature to within 2%. This is better than the accuracy of the fit that we will derive. All of these stars are isolated on their slits except for LRIS-05, and it is therefore excluded from the composite. Figure 4.19 shows the composite spectrum, and the 4 spectra that were co-added to create it. Note, that while these stars neither are the brightest photometrically, nor have the highest signal-to-noise ratios, they are among the best spectra (i.e., the composite candidates represent 5 of the 10 highest signal-to-noise ratio spectra). However, luminosity and signal-to-noise ratio were not the criterion upon which these stars were selected. Rather, these stars were the only ones with photometry similar enough that they could be assumed to have the same temperature.

Spectral fits were calculated by Jason Kalirai at UC Santa Cruz using the techniques described in Bergeron et al. [9]. The fitting technique attempts to directly constrain the effective temperature, T_{eff} , and the logarithm of the gravitational acceleration, $\log(g)$. The profiles of all Balmer lines are quite sensitive to T_{eff} variations, due to varying ionization levels, and $\log(g)$, due to linear Stark broadening that follows variations in the atmospheric pressure. The effects of varying T_{eff} and $\log(g)$ are different for the different Balmer lines. For instance, for $T_{\text{eff}} > 20000$, the equivalent widths of the low-order Balmer lines (H_{β} and H_{γ}) increase proportionally with $\log(g)$ due to the Stark effect, while the opposite effect is observed for lines of higher order due to the quenching of atomic levels (i.e., the fact that excited states merge with the continuum for very dense media). It is therefore essential to be able to fit high-order and low-order lines simultaneously.

The spectrum is first divided into sections containing one Balmer line each. Each section of the spectrum is then individually normalized using



Figure 4.19: The composite spectrum, and the spectra of the 4 objects that were co-added to create it.

two points on the continuum on either side of the absorption line. Thus, barring the existence of a "kink" or an unexpected change in the slope of the continuum in the middle of a Balmer line, the spectral-line fit is independent of the flux calibration. The fitting of the line shapes uses the nonlinear leastsquares method of Levenberg-Marquardt [79] based on the steepest descent method. The χ^2 statistic is calculated and minimized for combinations of $T_{\rm eff}$ and $\log(g)$, using normalized model line profiles of all absorption lines simultaneously. Masses and WD cooling ages ($t_{\rm cool}$) are found by using the updated evolutionary models of Fontaine et al. [38] for thick hydrogen layers ($M_H/M = 10^{-4}$) and helium layers of $M_{He}/M = 10^{-2}$. The core is assumed to have a 1:1 C/O ratio.

For illustrative purposes, the fit to the composite spectrum is shown in Figure 4.20. Fits for the rest of the stars are shown in Appendix II. The results for the stars that were analyzed are listed in Table 4.7.


Figure 4.20: The spectral line fit for the composite spectrum.

Age (yr)		$1.705 imes 10^7 \pm 6.249 imes 10^5$		$1.945 \times 10^7 \pm 1.358 \times 10^6$	$1.755 imes 10^7 \pm 1.758 imes 10^6$	$2.086 imes 10^7 \pm 2.224 imes 10^6$	$1.612 imes 10^7 \pm 9.053 imes 10^5$		$4.031 imes 10^7 \pm 5.954 imes 10^6$	$4.435 imes 10^7 \pm 1.237 imes 10^7$	$4.863 imes 10^7 \pm 8.113 imes 10^6$	$2.059 imes 10^7 \pm 3.167 imes 10^6$	$1.612 imes 10^7 \pm 9.053 imes 10^5$		$1.275 imes 10^8 \pm 1.078 imes 10^7$	$1.230 imes 10^7 \pm 7.017 imes 10^5$	$3.316 imes 10^7 \pm 3.700 imes 10^6$	-it is both an HST proper-motion gnitude is calculated assuming the sted here, LRIS-08, 16, 19, and 22,	
V_{the}		22.48	۵.	22.40	22.59	22.57	22.33	dwarfs	22.82	23.10	22.89	22.72	22.33		23.56	21.93	22.32	twice- und mag s not lis	
V_{obs}	lwarf	:	omposite	22.61	22.57	22.64	22.64	l white	23.25	22.42	22.94	22.54	22.64		21.25	22.62	22.86	is listed he V -be e object	, .
$M(M_{\odot})$	mposite white o	0.509 ± 0.017	used to create co	0.469 ± 0.030	0.560 ± 0.043	0.515 ± 0.044	0.480 ± 0.029	-motion selected	0.488 ± 0.039	0.751 ± 0.052	0.477 ± 0.045	0.605 ± 0.045	0.480 ± 0.029	Other stars	0.718 ± 0.025	0.449 ± 0.025	0.376 ± 0.039	Note, LRIS 24 site sample. T r Messier 4. Th	•
log(g)	Co	7.769 ± 0.038	stars 1	7.677 ± 0.078	7.868 ± 0.078	7.787 ± 0.093	7.697 ± 0.072	HST proper	7.748 ± 0.089	8.187 ± 0.088	7.733 ± 0.101	7.947 ± 0.074	7.697 ± 0.072		8.155 ± 0.040	7.588 ± 0.077	7.444 ± 0.107	to the spectra. er of the compc 1 appropriate fo	• • •
$T_{eff} (K)$		25140 ± 249		24331 ± 534	25567 ± 503	24317 ± 631	25668 ± 471		20637 ± 613	25078 ± 566	19716 ± 638	25300 ± 484	25668 ± 471		18847 ± 242	28196 ± 448	20900 ± 733	results of fits t d was a membe s and extinctior	
Ð		Comp.		LRIS 04	LRIS 06	LIRS 15	LRIS 24		LRIS 00	LRIS 09	LRIS 20	LRIS 23	LRIS 24		LRIS 02	LRIS 05	LRIS 29	4.7: The ed star, and ce modulus	

Table 4.7: The results of fits to the spectra. Note, LRIS 24 is listed twice—it is both an HST preselected star, and was a member of the composite sample. The V-band magnitude is calculated distance modulus and extinction appropriate for Messier 4. The objects not listed here, LRIS-08, 1 did not converge to a solution, or had too low a signal-to-noise ratio, and a fit was not attempted.

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For a given value of log(g) and T_{eff} , the mass and radius of the star can be determined. Given a distance modulus, the apparent magnitude of the star can then be predicted. One can then perform a sanity check by comparing the predicted magnitude with the observed magnitude. Figure 4.21 shows the difference between the predicted and observed magnitudes. The predicted magnitudes are generally more accurate for the isolated slits, lending confidence to these mass constraints. The difference between the predicted and observed magnitudes for the stars with isolated slits is generally less that 0.2, corresponding to an approximate luminosity difference of 20%. Assuming the determined temperature is independent of the determined log(g), a 20% error in flux corresponds to a 10% error in radius, and a 3% error in mass.

LRIS-02 was more than a magnitude brighter than all other stars in the white dwarf cooling sequence. If this star was a bona fide cluster member, we estimate that its cooling time is approximately one Myr (see Figure 1.3). Because of the very short lifetime of a star in this stage, we would be somewhat surprised to happen to have observed one. Furthermore, by using the theoretical absolute magnitude derived from the fit, and the distance modulus of the cluster, we can calculate the absolute magnitude of the star if it were in the cluster. It was found that LRIS-02 was two magnitudes brighter than predicted if it was indeed a cluster member. We therefore believe that this star is not a cluster member, and hence it will not be discussed further.

Our most reliable estimate for our final mass, $0.509 \pm 0.017 \ M_{\odot}$, comes from the fit to the composite spectrum. The stars that were combined to form the composite spectrum (LRIS–04, 06, 15, and 24) should have nearly identical masses. The mean mass derived from this sample is $0.506 \pm$ $0.020 \ M_{\odot}$. If we weight the spectra by the signal-to-noise ratio as measured between H_{β} and H_{γ} we obtain a mean mass of $0.507 \pm 0.014 \ M_{\odot}$. Note that this mass is very similar to the mass derived for the composite spectrum, and very close to the mass predicted from the initial-final mass relation (IFMR). The standard deviation of the masses of the stars that were combined to form the composite spectrum is $0.040 \ M_{\odot}$. This gives us another estimate of our error. Reassuringly, this error is similar to the error returned by the fitting routine. The (weighted) mean mass determined from all the spectra that have isolated slits is $(0.503 \pm 0.012 \ M_{\odot}) \ 0.500 \pm 0.017 \ M_{\odot}$.

Figure 4.18 shows that while the derived masses from low-signal-to-noise spectra have significant scatter, they do not show a significant bias. Figure 4.22 shows the masses derived in this thesis for both the individual spectra (in the inset) and for the composite mass (in the main panel). The other



Figure 4.21: The difference between the apparent magnitude predicted by the fitting routine, and the observed magnitude. Note, the theoretical predictions generally match more closely for the objects with isolated slits. The brightest object, LRIS-02, is the only object with an isolated slit with a strongly discrepant prediction. We therefore believe that this star is not a cluster member.

IFMR points come from Kalirai et al. [61] (and references therein). Note that the points for the other clusters are the mean masses of several white dwarfs within the cluster. The error bars represent the standard deviation of those masses. All clusters have white dwarfs with a range of cooling ages. For young clusters the difference in cooling age will translate into a nonnegligible difference in initial, and therefore final mass. The error bars for these points will therefore be larger than for a single star in that cluster. The least-squared fit for the IFMR including the mass determined for the composite spectrum,

$$M_{\text{final}} = (0.420 \pm 0.014) M_{\odot} + (0.101 \pm 0.006) M_{\text{initial}}, \qquad (4.7)$$

is shown in red and the IFMR from Kalirai et al. [61] is shown in black. The IFMR determined in this thesis is marginally consistent with that derived in Kalirai et al. [61]. Typically, semi-empirical initial-final mass relations are not simply linear. Most show a decrease in the slope toward low masses (for example Salaris et al. [90], Figure 3). While the data do not require a more complex empirical relation, the three lowest-mass points in Figure 4.22 do show a very mild hint of a shallow slope.

While our constraints on the initial-final mass relation from the individual spectra are weak, they are in line with the relation derived empirically at higher masses. The composite spectrum provides a solid constraint at low mass. This mass is consistent with the photometric estimate of Moehler et al. [76], $0.53\pm0.03 M_{\odot}$, and in line with what one would expect from mainsequence models. Because globular clusters are the oldest Galactic stellar populations, this effectively extends the initial-final mass relation over the entire mass range of stars that could have formed white dwarfs. The future prospects of this project will be discussed in the concluding chapter.



Figure 4.22: The IFMR for the stars observed in this study as well as those published for other clusters. The error bars for the M4 point are purely statistical, and have no systematic component. Assuming a linear relation, our IFMR is shown in red, while the relation derived by Kalirai et al. [61] is shown in black. The error bars of the relation derived by Kalirai et al. [61] are denoted with dotted lines. Though our mass values have a fair amount of scatter, those with isolated slits (diamonds) form a much tighter group. The values for these stars are very close to those predicted by the IFMR. The different initial masses shown in the inset are *not* meant to indicate a difference in mass. Offsets were added to the assumed initial mass, so the points would be spaced out enough to be visible.

Chapter 5

Conclusions and Future Work

We have presented three projects involving the two nearest globular clusters, and discovered a number of interesting results. First, in Chapter 2, we placed tight constraints on the binary fraction of NGC 6397, both in a field centered on the core, and in one beyond the half-light radius of the cluster. According to the recent N-body models of Hurley et al. [54], the primordial value should be preserved beyond the half-light radius, while interior to this, we should observe an enhancement in the binary fraction. Our observations dovetail neatly with this finding, suggesting a primordial binary fraction in the cluster of only $\sim 1\%$. In Chapter 3 we reported the first direct evidence that young white dwarfs are dynamically heated upon birth. Our preferred source of heat is an impulsive kick that occurs in some post-main-sequence evolutionary stage. Finally, in Chapter 4, we presented spectra of 23 white dwarfs in Messier 4. All these white dwarfs were of type DA, confirming that the DA/DB ratio is clearly different in the cluster environment than it is in the field. The spectra were of lower quality that we had hoped for, and consequently we were only able to put weak constraints on the white-dwarf mass in this cluster with the individual spectra. However, using a composite spectrum, we determined that mass of the white dwarfs currently forming in this cluster is $0.51 \pm 0.02 \ M_{\odot}$. This mass is consistent with an extension of the initial-final mass relation derived at higher masses. It is now interesting to look to the future, and speculate in which direction future research may bear fruit. The three different projects will be discussed in sequence.

5.1 The binary fraction of NGC 6397

The work presented in this thesis suggests that the primordial value of the binary fraction of NGC 6397 is close to 1%. Compiled in Table 5.2 are literature values for previous attempts to constrain the binary fraction in NGC 6397 and other Galactic globular clusters. The values determined in-

$R > R_{hm}$							
cluster	f	reference					
NGC 288	$0.0^{+0.1}_{-0.0}$	Bellazzini et al. [8]					
NGC 2808	$0.24\pm0.04^{\dagger}$	Ferraro et al. [35]					
NGC 3201	< 0.1	Cote et al. [21]					
NGC 6752	$0.02^{+0.16}_{-0.02}$	Rubenstein and Bailyn [89]					
NGC 6792	"low"	\dots Catelan et al. [15]					
M3	"low"	\dots Gunn and Griffin [42]					
	~ 0.04	Pryor et al. [81]					
	0.02 ± 0.01	\dots Zhao and Bailyn [112]					
M4	~ 0.02	\dots Richer et al. [84]					
M22	$0.03^{+0.16}_{-0.03}$	Cote et al. [22]					
M30	0.05 ± 0.02	\dots Alcaino et al. [1]					
M92	$0.00^{+0.03}_{-0.00}$	\dots Anderson [2]					
47 Tucanae	> 0.05	de Marchi and Paresce [25]					
	~ 0.02	\dots Anderson [2]					
ω Centauri	< 0.05	Elson et al. $[32]$					

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Table 5.1: Literature binary fraction constraints outside the half-mass radius. [†] This result was questioned by Walker [103].

terior to $R_{\rm hm}$ vary widely. This can be understood in the context of varying dynamical states. In sharp contrast to the values found in the core, the values determined beyond $R_{\rm hm}$, are generally quite low. Whether this implies that globular clusters formed with a universal low binary fraction requires a tight constraint in more clusters, however, it seems clear that to construct accurate dynamical models of many Galactic globular clusters a low primordial binary fraction must be used. The consequence of a low primordial binary fraction will generally be for globular cluster to have shorter first core collapse times.

In order to interpret our observations, the opportunity to compare with N-body models was extremely valuable. The limiting factor with these comparisons is 1) the limitation of the models themselves, and 2) the extreme computational cost of the calculation of the models. These two factors are inextricably entwined. Straight N-body models scale as N^2 . However, tight binary systems take far more computation than the average star.

cluster	f	reference				
NGC 288	0.15 ± 0.05	Bellazzini et al. [8]				
	> 0.06	Sollima et al. [94]				
NGC 362	0.21 ± 0.06	Fischer et al. [36]				
NGC 4590	> 0.09	Sollima et al. [94]				
NGC 5053	> 0.08	Sollima et al. [94]				
NGC 5466	> 0.08	Sollima et al. [94]				
NGC 5897	> 0.07	Sollima et al. [94]				
NGC 6101	> 0.09	Sollima et al. [94]				
NGC 6362	> 0.06	Sollima et al. [94]				
NGC 6397	< 0.07	\dots Cool and Bolton [19]				
NGC 6723	> 0.06	Sollima et al. [94]				
NGC 6752	0.27 ± 0.12	Rubenstein and Bailyn [89]				
NGC 6981	> 0.10	Sollima et al. [94]				
M3	0.14 ± 0.08	Zhao and Bailyn [112]				
M4	$0.23^{+0.34}_{-0.23}$	Cote and Fischer [20]				
M15	~ 0.07	Gebhardt et al. [39]				
M55	> 0.06	Sollima et al. [94]				
M71	$0.22^{+0.26}_{-0.12}$	\dots Yan and Mateo [111]				
Arp 2	> 0.08	Sollima et al. [94]				
Terzan 7	> 0.21	Sollima et al. [94]				
Palomar 12	> 0.18	Sollima et al. [94]				
Palomar 13	0.30 ± 0.04	Clark et al. [17]				

 $R < R_{hm}$

Table 5.2: Literature binary fraction constraints inside the half-mass radius.

Recent N-body models are only now getting to the point of creating realistic models of large open clusters with 10^5 stars , i.e., one N-body particle for each star in the cluster. Globular clusters require an order of magnitude more stars. The question is then in the context of the computational expense which direction to push to get to more realistic simulations. This result suggests that one should not sacrifice other physics in order to include large numbers of binaries—a very modest primordial binary fraction should be sufficient to realistically simulate a globular cluster.

5.2 White-dwarf-natal kicks

Since the observations discussed in this thesis were first reported, there have been several developments on the theoretical front.

5.2.1 Theoretical developments

Heyl has recently published four papers regarding the topic. The subject matter of the papers are summarized here:

In the first paper, Heyl [50] calculated how an impulsive kick affects the radial distribution of a population of stars. The problem is parameterized in terms of the mean velocity dispersion of the cluster, σ_c , the velocity dispersion of the white-dwarf progenitors, σ_{TO} , the depth of the gravitational potential, $\Psi(0)$, and the concentration of the cluster, c. The main conclusion was that kicks are slightly less effective at increasing the velocity dispersion than one would guess by simply adding the velocities in quadrature. This is primarily due to the fact that the kicked population has a different distributions of orbital elements than a non-kicked population. The orbits will typically be more eccentric, and will have less angular momentum than those with similar energy on undisturbed orbits. While the exact conditions of our clusters were not explored in this paper, it seems as if the kicks would increase the velocity dispersion by ~ 0.7 of what one would naively expect by adding the velocities in quadrature (as shown in Figure 5.1).

Heyl [51] explored how an impulsive kick affects binary and planetary systems in the second paper. The conclusion was that kicks raise the eccentricity of these systems, and can disrupt them.

The search for alternate observational signatures of kicks was performed in the third paper, Heyl [52]. The conclusion was that while the radial distribution is the most obvious effect, there are other subtle signatures. The eccentricity of the white dwarf orbits within the globular cluster is increased by the kick, and hence they tend to be on more radial orbits. This



Figure 5.1: A figure from Heyl [50], showing the final values of the bestfitting for the phase-mixed stellar distributions as a function of the initial value of σ_{TO}/σ_c , $\Psi(0)/\sigma_c^2$ and σ_k/σ_{TO} . From top to bottom the sets of curves are for $\sigma_{TO}/\sigma_c = 0.8$, 0.7, 0.6 and 0.5. Within each set the curves give $\Psi(0)/\sigma_c^2 = 4$, 6 and 8 in black, red and blue respectively. For the two smallest initial velocity dispersions the results for $\Psi(0)/\sigma_c^2 = 6$ (red) and 8 (blue) are indistinguishable. In Messier 4, the line-of-sight velocity is determined from giant stars, and therefore is almost equal to that of the turn-off stars, so, $\sigma_{TO}/\sigma_c = \sigma_{LOS}/\sigma_c = 3.9 \, km s^{-1}/6.4 \, km s^{-1} = 0.6$. In Messier 4, $\Psi(0)/\sigma_c^2=7.4$ so the most appropriate curve to look at is the bottom red line. Although, a kick of the magnitude we expect is not plotted on this graph, for the largest kick values, the final velocity dispersion is approximately 0.7 the value that one would naively calculate by adding the velocities in quadrature.

	NGC 6397	Messier 4
Velocity dispersion (σ_{LOS})	$\dots 3.5 \ kms^{-1} \ [47]$	$3.9 \ kms^{-1} \ [47]$
Velocity dispersion (σ_c)	$\dots 5.7 \ kms^{-1} \ [47]$	$6.4 \ kms^{-1} \ [47]$
<i>c</i>		$\dots 1.59[47]$
$\Psi(0)/\sigma_c^2$		7.40 ± 0.11 [74]

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Table 5.3: Cluster properties. σ_{LOS} refers to the line of site velocity, derived from spectra of giant stars. The mean velocity dispersion for the cluster, σ_c , is calculated assuming the mean stellar mass of the cluster is 0.3 M_{\odot} and the mass of the giants is 0.8 M_{\odot} . References: Harris [47]; McLaughlin and van der Marel [74]

has the effect of skewing the angle the proper motion vectors of these stars will have with respect to the mean proper motion of the cluster. This is a subtle effect, and most likely will be very difficult to detect with today's generation of instruments, but could be nice corroborating evidence in the future.

In the fourth paper, Heyl [53] focused on the effect the energetics of the kicked white dwarfs have on the dynamics of the cluster. In essence, the white-dwarf kicks convert the nuclear energy of the white-dwarf progenitor into kinetic energy. While the cluster has no access to the energy while it is in the form of nuclear potential energy, globular clusters efficiently redistribute kinetic energy to all stars. The kicks are therefore a heat source for the cluster. Over the lifetime of the cluster, kicks provide roughly one fifth as much energy to the cluster as binaries systems do. Amazingly, at early times in the cluster evolution, the energy from kicks can dominate over that from binary systems.

The implications from the fourth paper are truly exciting. Globular cluster orthodoxy suggests that the only energy source in a globular cluster are binary systems. A small primordial binary fraction can delay the onset of cluster core collapse from 20 half-mass relaxation times to 150. If there is indeed another energy source, core collapse may be delayed even further.

Finally, most stellar encounters take place in the core of the cluster. Therefore, the excess energy that the kicked white dwarfs have is likely to be deposited in the core, and thereby increase the velocity dispersion there. Searches for black holes in globular clusters that rely on stellar dynamics could possible get a false positive signal from this, but a full suite of N-body models is required to explore this.

5.2.2 Observational developments

In order to improve confidence in our result, one could perform observations in other clusters, or in a different area in the same cluster in order to increase the number of observed stars. In this vein, young white dwarfs with an extended radial distribution have recently been reported additionally in ω Centauri [13]. No paper has reported white dwarfs with a radial distribution consistent with that of their progenitors, though it is difficult to tell if this is significant. Whether or not this implies that this phenomenon is universal remains to be seen.

Though this is quite speculative, our preferred mechanism for the impulsive kick is currently an off-center helium core flash. This occurs at the tip of the red giant branch. This hypothesis is testable in several ways. The AGB or red clump stars in a globular cluster are stages of evolution that occur after the helium flash. Accordingly, one could compare the radial distributions of these populations to see if a difference can be measured. Using these populations has the advantage that they are roughly 10 magnitudes brighter than their white dwarf counterparts. This makes them observable from a single-orbit HST observation, or even from the ground. Furthermore, we are not restricted to the nearest globular clusters. Sarajedini et al. [92] recently observed 65 globular clusters with the ACS. A large fraction of these clusters have well populated post-main-sequence evolutionary branches, and could be used to determined at what stage the kick occurs. Care must be taken to account for field star contamination with observations like these. With the number of stars that existed in our observations, proper-motion cleaning was essential. Even though contamination is relatively small, our observed signal was weaker in the non-proper-motion cleaned sample, despite the fact that we increased the number of stars in our sample by 50%.

Having a large number of post-main sequence stars is essential when trying to constrain at what stage the kick occurs. The field of the ACS is generally too small to image an entire Galactic globular cluster with one pointing. However, large-format ground-based CCD imagers do not suffer from this problem—typically the bulk of a Galactic globular cluster may be imaged with one pointing from the ground. The problems with ground based observations are pushing observations to very deep magnitudes, and obtaining the extremely precise astrometry needed to perform proper motion cleaning. Anderson et al. [3] recently showed that proper-motion cleaning with only a several year baseline is becoming feasible from ground-based observations. If this proves to be feasible for a number of globular clusters, ground based observations may be a fruitful avenue with which to constrain the evolutionary stage in which the white dwarf natal kick occurs.

Finally, our research group has been granted an additional 12 orbits with HST-ACS to re-image the same field in NGC 6397. The observations for this proposal were originally planned for early 2007. Unfortunately with the failure of the ACS on January 27, 2007, the ACS was rendered inoperable. Servicing mission 4 (SM4) includes a plan to fix the ACS. Our group has been re-approved to obtain these data, and if indeed SM4 is successful, we should obtain the data sometime in 2009. The follow-up observations were originally planned to have a 2-year baseline. This would have been sufficient to proper-motion clean the entire field (rather than just the \sim 60% that was possible with the archival WFPC2 fields) to the truncation of the white dwarf cooling sequence. The failure of the ACS caused at least an additional 2 years of delay before the follow-up observations. With this increased baseline, it will be possible to resolve the internal propermotions of the stars. This will allow the direct measurement of the velocity dispersion of white dwarfs of various masses without referring to their radial distributions.

5.3 Spectra of white dwarfs in Messier 4

The first conclusion we can draw from the spectra we have obtained in Messier 4 is that they are all of type DA. With an additional 23 whitedwarf spectral-type identifications, we have roughly doubled the number of spectral identifications in globular clusters. It is now incontrovertible that the DA/DB ratio is different in the field than in the cluster environment. Unfortunately, there is no obvious mechanism for this. A large range of metallicities and masses has been explored, and these are not causal agents in this effect. The only other difference between the field and cluster environment that springs to mind is the mean density of the ISM. The idea that DBs are suppressed in the cluster environment due to the re-accretion of residual intra-cluster hydrogen has been examined, and found to be implausible. Perhaps this conclusion needs to be reexamined.

The question of the mass of white dwarfs currently forming in Messier 4 requires a more subtle examination. At this point, the signal-to-noise ratio in our spectra is insufficient for a robust constraint using the individual spectra, though for our brightest objects, it approaches the requisite quality. While the mass constraint from the composite mass appears to be credible, using the composite spectrum adds another layer of uncertainty. Ideally, we would like to constrain the mass using the individual masses only. It is useful then

to take a step back and reexamine our approach. There are three ways to proceed from here, all with their individual pros and cons.

The first approach is to keep on observing the same field with essentially the same slit mask. Because of the extreme crowding of our field, the final signal-to-noise ratios of our spectra are driven by the signal-to-noise ratio of individual spectra obtained under good observing conditions. If we happen to obtain just several hours of exposure under ideal observing conditions, the signal-to-noise ratio of the brightest spectra would most likely be good enough for our purposes. Considering the instruments presently available, LRIS on Keck seems to have the best combination of blue sensitivity and collecting area. The drawback with Keck is the classical observing mode. Messier 4 is a Southern target, and hence only has a low airmass for several hours a night. It would be far better to get three one hour exposures taken on different nights rather than all in one night, and therefore at smaller air mass.

The second approach would be to design a new slit mask. The location of our first field was driven by the existence of our HST fields. This was done because we had a sample of proper-motion confirmed white dwarfs in this area, and thus did not have to worry about field contamination. While field contamination is a valid concern, preselecting stars from HST photometry is fraught with difficulties. The excellent PSF of HST allows accurate photometry in crowded regions. Because the follow-up spectroscopy will be obtained from the ground, one may not be able to obtain decent spectra from an HST selected source. As has been shown in Anderson et al. [3], good proper motions can be calculated from the ground with only a severalyear temporal baseline. Archival wide field images of Messier 4 exist, and a second epoch would be relatively cheap to obtain. A wide-field ground based image would be a more appropriate source for the selection of spectroscopic targets, and perhaps a field could be located with several bright white dwarfs.

Finally, one could wait for new instrumentation to complete this project. For example, the suite of first-light instruments slated for the TMT include the Wide Field Optical Spectrometer (WFOS). It will provide nearultraviolet and optical (3000–10000 Å) imaging and spectroscopy over a more than 40 square arcminute field-of-view. While WFOS will not be equipped with adaptive optics, it will have an order of magnitude greater collecting area than Keck. Short periods of high-quality natural seeing would enable the collection of sufficient amounts of data. Unfortunately, the TMT is most likely at least a decade away from science operations.

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Appendix A Appendix I



Figure A.1: The astrometry, photometry, and GMOS CCD section of LRIS- $00\,$



Figure A.2: The astrometry, photometry, and GMOS CCD section of LRIS- $01\,$



Figure A.3: The astrometry, photometry, and GMOS CCD section of LRIS- $02\,$



Figure A.4: The astrometry, photometry, and GMOS CCD section of LRIS- 03



Figure A.5: The astrometry, photometry, and GMOS CCD section of LRIS- 04



Figure A.6: The astrometry, photometry, and GMOS CCD section of LRIS- 05



Figure A.7: The astrometry, photometry, and GMOS CCD section of LRIS-06 $\,$


Figure A.8: The astrometry, photometry, and GMOS CCD section of LRIS- $07\,$



Figure A.9: The astrometry, photometry, and GMOS CCD section of LRIS- 08



Figure A.10: The astrometry, photometry, and GMOS CCD section of LRIS-09 $\,$



Figure A.11: The astrometry, photometry, and GMOS CCD section of LRIS-10 $\,$



Figure A.12: The astrometry, photometry, and GMOS CCD section of LRIS-11 $\,$



Figure A.13: The astrometry, photometry, and GMOS CCD section of LRIS-12 $\,$



Figure A.14: The astrometry, photometry, and GMOS CCD section of LRIS-13. This object is close to a chip gap.



Figure A.15: The astrometry, photometry, and GMOS CCD section of LRIS-14 $\,$



Figure A.16: The astrometry, photometry, and GMOS CCD section of LRIS-15 $\,$



Figure A.17: The astrometry, photometry, and GMOS CCD section of LRIS-16 $\,$



Figure A.18: The astrometry, photometry, and GMOS CCD section of LRIS-17 $\,$



Figure A.19: The astrometry, photometry, and GMOS CCD section of LRIS-18 $\,$



Figure A.20: The astrometry, photometry, and GMOS CCD section of LRIS-19 $\,$



Figure A.21: The astrometry, photometry, and GMOS CCD section of LRIS-20 $\,$



Figure A.22: The astrometry, photometry, and GMOS CCD section of LRIS-21 $\,$



Figure A.23: The astrometry, photometry, and GMOS CCD section of LRIS-22 $\,$



Figure A.24: The astrometry, photometry, and GMOS CCD section of LRIS-23 $\,$



Figure A.25: The astrometry, photometry, and GMOS CCD section of LRIS-24 $\,$



Figure A.26: The astrometry, photometry, and GMOS CCD section of LRIS-25 $\,$



Figure A.27: The astrometry, photometry, and GMOS CCD section of LRIS-26 $\,$



Figure A.28: The astrometry, photometry, and GMOS CCD section of LRIS-27 $\,$



Figure A.29: The astrometry, photometry, and GMOS CCD section of LRIS-28 $\,$



Figure A.30: The astrometry, photometry, and GMOS CCD section of LRIS-29 $\,$



Figure A.31: The astrometry, photometry, and GMOS CCD section of LRIS- $30\,$



Figure A.32: The astrometry, photometry, and GMOS CCD section of GMOS-05 $\,$



Figure A.33: The astrometry, photometry, and GMOS CCD section of GMOS-07 $\,$



Figure A.34: The astrometry, photometry, and GMOS CCD section of GMOS-11 $\,$



Figure A.35: The astrometry, photometry, and GMOS CCD section of GMOS-15 $\,$



Figure A.36: The astrometry, photometry, and GMOS CCD section of GMOS-17 $\,$

Appendix B Appendix II



Figure B.1: The spectral line fit of LRIS 00.



Figure B.2: The spectral line fit of LRIS 02.



Figure B.3: The spectral line fit of LRIS 04.



Figure B.4: The spectral line fit of LRIS 05.



Figure B.5: The spectral line fit of LRIS 06.



Figure B.6: The spectral line fit of LRIS 09.


Figure B.7: The spectral line fit of LRIS 15.



Figure B.8: The spectral line fit of LRIS 20.



Figure B.9: The spectral line fit of LRIS 23.



Figure B.10: The spectral line fit of LRIS 24.



Figure B.11: The spectral line fit of LRIS 29.

Appendix C

Notes on Publication

Though this thesis is in the traditional format, Chapters 2 and 3 were based on previously published journal articles.

Chapter 2: For this work, I developed the technique, and wrote the computer code to create models of the binary and single-star sequences. I wrote the code to perform the statistical fits, analyzed the data, and wrote the manuscript. H. Richer provided guidance throughout this project. J. Anderson and J. Brewer provided the photometry and astrometry on the ACS and WFPC2 fields respectively. J. Hurley calculated the N-body models which were extremely helpful in interpreting these results. J. Kalirai, M. Rich, and P. Stetson provided useful comments and suggestions when editing the manuscript.

A version of this chapter was published in The Astronomical Journal, Volume 135, Issue 6, pp. 2155-2162 (2008) as *Deep Advanced Camera for Surveys Imaging in the Globular Cluster NGC 6397: The Binary Fraction* by Davis, D. S.; Richer, H. B.; Anderson, J.; Brewer, J.; Hurley, J.; Kalirai, J. S.; Rich, R. M.; Stetson, P. B. [23]

Chapter 3: For this work, I developed the technique to detect extended radial distributions in close collaboration with I. King and H. Richer. I wrote the computer code to analyze the data, and wrote the manuscript. J. Anderson provided the photometry and astrometry used in this project. J. Coffey, G. Fahlman, J. Hurley, and J. Kalirai provided useful comments and suggestions when editing the manuscript.

A version of this chapter was published in Monthly Notices of the Royal Astronomical Society: Letters, Volume 383, Issue 1, pp. L20-L24, 2008, as On the radial distribution of white dwarfs in the globular cluster NGC 6397 by Davis, D. S.; Richer, H. B.; King, I. R.; Anderson, J.; Coffey, J.; Fahlman, G. G.; Hurley, J.; Kalirai, J. S. [24]

Chapter 4: This project was originally conceived by H. Richer and J. Kalirai, who wrote the original GMOS proposal for this project. I wrote the phase 2 proposal, and analyzed the pre-imaging in order to select candidates. I was principle investigator on a follow-up GMOS proposal to obtain more observing time. The LRIS proposal was written by M. Rich and D. Reitzel. The observations at Keck were collected over several semesters by me (two observing runs), J. Kalirai, M. Rich, and D. Reitzel. I reduced all the spectroscopic data. The spectroscopic fits were calculated by J. Kalirai.