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Exploring the population of compact stellar X-ray sources in our Milky Way

by

Sandra Magdy Kamel Greiss

Thesis

Submitted to the University of Warwick

for the degree of

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Contents

List of	' Table	s vi
List of	Figur	es viii
Ackno	wledgr	nents xv
Declar	ation	xvi
Abstra	act	xvii
Chapt	er 1 I	Motivation 1
Chapt	er 2 S	Stellar evolution, Compact Stars, Accretion discs and
X-r	ay bin	aries 3
2.1	Stellar	$evolution \ldots \ldots \ldots \ldots \ldots 3$
	2.1.1	Star Formation
		2.1.1.1 Protostars
		2.1.1.2 Stellar Mass limits
	2.1.2	Main-sequence stars
	2.1.3	Giant Stars
	2.1.4	Death of stars like the Sun
	2.1.5	Death of massive stars
	2.1.6	The characteristic timescales of stellar evolution $\ldots \ldots \ldots 18$
		2.1.6.1 Dynamic timescale
		2.1.6.2 Thermal timescale \ldots \ldots \ldots \ldots 18
		2.1.6.3 Nuclear timescale
	2.1.7	The evolution of a star on a Hertzsprung-Russell (H-R) di-
		agram
2.2	Compa	act stars
	2.2.1	White dwarfs

	2.2.2	Neutron Stars	23
	2.2.3	Black holes	24
2.3	Binary	stars	26
	2.3.1	Visual, spectroscopic and eclipsing binaries	26
	2.3.2	Roche-lobe overflow (RLO) $\ldots \ldots \ldots \ldots \ldots \ldots \ldots$	27
		2.3.2.1 The orbital angular momentum equation	30
		2.3.2.2 Stable or unstable mass transfer?	31
	2.3.3	Common envelope (CE) evolution	32
	2.3.4	Detached, semi-detached and contact binaries \ldots	32
	2.3.5	Supernova explosions in close binaries	33
2.4	Accreti	ion discs	34
	2.4.1	Disc formation	35
	2.4.2	Thin discs	35
	2.4.3	Alfvén radius	36
	2.4.4	Eddington Limit	38
2.5	X-ray	binaries	39
	2.5.1	Helium stars in binary systems	39
	2.5.2	Cataclysmic variables	40
	2.5.3	Low-mass X-ray binaries (LMXBs)	43
	2.5.4	High-mass X-ray binaries (HMXBs)	46
2.6	Their o	observational properties	48
	2.6.1	Cataclysmic variable	49
	2.6.2	High-mass X-ray binaries	50
	2.6.3	Low-mass X-ray binaries	51
	2.6.4	Intermediate-mass X-ray binaries (IMXBs)	51
	2.6.5	Soft X-ray transients	52
Chapte	er 3 G	Galactic Plane & Bulge surveys	54
3.1	Metho	ds	54
3.2	GBS		59
3.3	UKIDS	SS GPS	61
3.4	2MASS	5	64

3.5	IPHAS	8
3.6	SDSS	
	3.6.1	SEGUE
3.7	VVV	
3.8	UVEX	K
3.9	VPHA	S+
Chapte	er4N	Vear-infrared data of the Chandra GBS sources 80
4.1	GBS v	ersus GPS
	4.1.1	Getting the data from GPS
	4.1.2	2MASS photometry needed
	4.1.3	The contributions to false matches
4.2	Results	s & discussion
	4.2.1	Real matches criteria
	4.2.2	Science results
43	Spectre	oscopic follow-up
т.0	~poot-	
1.0	- T	
Chapte	er 5 I	PHAS calibration 102
4.5 Chapte 5.1	er 5 II IPHAS	PHAS calibration 102 S & SDSS 103
Chapte 5.1	er 5 I IPHAS 5.1.1	PHAS calibration 102 & & SDSS 103 Finding the overlapping regions of both surveys 103
5.1	er 5 I IPHAS 5.1.1 5.1.2	PHAS calibration 102 & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 105
5.1	Er 5 I IPHAS 5.1.1 5.1.2 5.1.3	PHAS calibration 102 S & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 105 Important difference between IPHAS and SDSS magnitude
5.1	er 5 I IPHAS 5.1.1 5.1.2 5.1.3	PHAS calibration 102 S & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 105 Important difference between IPHAS and SDSS magnitude 106
5.2	er 5 I IPHAS 5.1.1 5.1.2 5.1.3 Results	PHAS calibration 102 & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 105 Important difference between IPHAS and SDSS magnitude 106 systems 108
5.2	er 5 I IPHAS 5.1.1 5.1.2 5.1.3 Results 5.2.1	PHAS calibration 102 & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 105 Important difference between IPHAS and SDSS magnitude 106 systems 108 Offsets between IPHAS and SDSS 108
5.2	er 5 I: IPHAS 5.1.1 5.1.2 5.1.3 Result: 5.2.1 5.2.2	PHAS calibration 102 5 & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 103 Important difference between IPHAS and SDSS magnitude 106 systems 108 Offsets between IPHAS and SDSS 108 Is there a gradient in the offsets? 112
5.2	er 5 I IPHAS 5.1.1 5.1.2 5.1.3 Results 5.2.1 5.2.2	PHAS calibration 102 S & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 105 Important difference between IPHAS and SDSS magnitude 106 systems 108 Offsets between IPHAS and SDSS 108 Is there a gradient in the offsets? 112 5.2.2.1 Looking across the entire strip 112
5.1 5.2	er 5 I IPHAS 5.1.1 5.1.2 5.1.3 Results 5.2.1 5.2.2	PHAS calibration 102 & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 105 Important difference between IPHAS and SDSS magnitude 106 systems 108 Offsets between IPHAS and SDSS 108 Is there a gradient in the offsets? 112 5.2.2.1 Looking across the entire strip 112 5.2.2.2 What about across the CCDs? 112
5.1 5.2	er 5 I IPHAS 5.1.1 5.1.2 5.1.3 Resulta 5.2.1 5.2.2 5.2.3	PHAS calibration 102 S & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 103 Important difference between IPHAS and SDSS magnitude 105 systems 106 s 108 Offsets between IPHAS and SDSS 108 Is there a gradient in the offsets? 112 5.2.2.1 Looking across the entire strip 112 S.2.2.2 What about across the CCDs? 112 Why do we find outliers? 114
5.1 5.2	er 5 I IPHAS 5.1.1 5.1.2 5.1.3 Result: 5.2.1 5.2.2 5.2.3	PHAS calibration 102 S & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 103 Important difference between IPHAS and SDSS magnitude 105 systems 106 S 108 Offsets between IPHAS and SDSS 108 Is there a gradient in the offsets? 112 5.2.2.1 Looking across the entire strip 112 S.2.2.2 What about across the CCDs? 112 Why do we find outliers? 114
5.2 5.2	Providence From From <	PHAS calibration 102 & SDSS 103 Finding the overlapping regions of both surveys 103 Getting the data from both surveys 103 Getting the data from both surveys 105 Important difference between IPHAS and SDSS magnitude 106 systems 106 S 108 Offsets between IPHAS and SDSS 108 Is there a gradient in the offsets? 112 5.2.2.1 Looking across the entire strip 112 5.2.2.2 What about across the CCDs? 112 Why do we find outliers? 114 5.2.3.1 Sources with large offsets 114 5.2.3.2 Sources with unusual colours 116

Chapte	er 6 Conclusion and future work	137
6.1	Initial steps	. 137
6.2	Problems	. 139
Refere	nces	143

List of Tables

2-1	Different possible end states of stars according to their initial masses, in	
	the case of a single star and when found in a binary system (Tauris ${\mathfrak E}$	
	van den Heuvel, 2006)	17
2-2	Observational properties of HMXBs and LMXBs (Tauris $\mathfrak E$ van den	
	Heuvel, 2006)	52
3-1	Johnson/Bessel central wavelengths of UBVRI filters	59
3-2	Summary of all five surveys in UKIDSS (Lawrence et al., 2007)	65
3-3	Effective wavelengths of JHK filters used in GPS (Hewett et al., 2006)	65
3-4	Effective wavelengths of JHK_s filters used in 2MASS (Skrutskie et al.,	
	2006)	68
3-5	Effective wavelengths of IPHAS (González-Solares et al., 2008)	71
3-6	The Sloan Digital Sky survey Photometric System (Fukugita et al., 1996)	74
4-1	Results from DR5, DR6 and 2MASS	85
4-2	Numbers and results	87
4-3	Total extinction (A_{λ}) for JHK filters, with $E_{B-V}=1$ (Schlegel et al.,	
	1998)	90
4-4	Numbers and magnitudes of UKIDSS matches < Mean maximum Chan-	
	dra error	95
4-5	Hardness and magnitude of sources	95
5-1	Median of the offsets in the i and r bands in IPHAS fields 3382 and	
	33820 of the December 2006 run, with respect to $SDSS$	109
5-2	List of IPHAS fields, classified as 'outliers' in the (r - H $lpha$) vs (r - i)	
	colour-colour diagram	17

5 - 3	Number of $H\alpha$ emitters found in each $(r - i)$ colour bin. The spectral
	types are those of main-sequence dwarfs with $E_{B-V} = 0$, taken from
	Drew et al. (2005)
5-4	ugriz magnitudes of the H α emitter found. The magnitudes are all given
	in the Vega system. Both reddened and dereddened values are calculated.
	The dereddened value were calculated using the Schlegel dust maps (the
	same method as described in Chapter 4)
5-5	$2MASS JHK_s$ magnitudes of the H α emitter $\ldots \ldots \ldots$

List of Figures

2-1	The mass function of stars in the solar neighbourhood (Rana, 1987).	
	The function is compared to that of Scalo (1986)	7
2-2	Typical H-R diagram (Lecture notes from course PX268, University of	
	Warwick).	20
2-3	$Evolutionary\ tracks\ in\ the\ H-R\ diagram\ of\ six\ stars\ with\ different\ masses,$	
	ranging from 1 M_{\odot} to 50 M_{\odot} (Tauris & van den Heuvel, 2006) \ldots	21
2-4	Diagram of eclipsing binaries and the light curves seen in each case. $\ . \ .$	28
2-5	Equipotential lines in a binary system, also indicating the five Lagrangian	
	points of potential equilibrium (Tauris \mathcal{E} van den Heuvel, 2006)	29
2-6	Different cases of close binary systems (taken from OBSERVATION	
	AND ANALYSIS OF ECLIPSING BINARY STAR, by Derek Wallentin-	
	sen - MSc thesis, B.A., Vassar College, 1985).	33
2-7	$Different\ stages\ of\ the\ formation\ of\ an\ accretion\ disc\ (taken\ from\ Lecture$	
	course PX387, The University of Warwick)	34
2-8	Example of the evolution of a LMXB (Tauris & van den Heuvel, 2006).	45
2-9	Example of the evolution of a HMXB (Tauris & van den Heuvel, 2006).	48
3-1	$Spectrum of the Sun \ (continuum)$ as well as the spectrum of a black body	
	(dotted line) with the same temperature as the Sun (Figure taken from	
	lecture course PX268, The University of Warwick)	56
3-2	UBVRI Johnson filters plotted on top of the spectra of two stars. The	
	I filter is not seen in the wavelength range of this plot but it is one of	
	the Johnson filters. The blue star is the hot one, whereas the red one is	
	the cool star (Figure taken from lecture course PX268, The University	
	of Warwick).	58

- 3-3 Integrated light of what would be detected and measured with the B and
 V Johnson filters in the case of the two stars considered in Figure 3-2
 (Figure taken from lecture course PX268, The University of Warwick).
 59
- 3-4 GBS coverage. Region delimited by black lines is the survey's designated covered area. The magenta dots correspond to the GBS sources detected by the Chandra X-ray satellite. As seen, the survey is not yet complete in the X-ray detections (Figure created using the Chandra X-ray detections and survey coverage area).
- 3-5 Proposed and planned sky coverage of UKIDSS. Cross-hatch : Large Area Survey. Dark Grey : Galactic Plane Survey. Light Grey : Galactic Clusters Survey. Open rectangles : Deep Extragalactic Survey. The dashed line marks the Galactic plane (Lawrence et al., 2007). 64
- 3-6 The top figure is a full-sky distribution of point sources and the bottom figure shows the same distribution of the extended sources. Point sources are presented in Galactic coordinates centered on $b = 0^{\circ}$ and $l = 0^{\circ}$, whereas the extended source map is presented in equatorial coordinates and centered at $\alpha = 180^{\circ}$ and $\delta = 0^{\circ}$. The faint blue band in the bottom figure traces the Galactic Plane, with the intensity proportional to source density. The images are a colour composite of source density in the J (blue), H (green) and K_s (red) bands (Skrutskie et al., 2006). 67
- 3-7 Transmission filters used in IPHAS: Hα, Sloan r and Sloan i (Drew et al., 2005). The dotted lines show the transmission of the narrow-band filter, whereas the continuum lines correspond to the broad-band transmissions.
 71

3-9	The top panel shows the SEGUE Survey footprint in the Equatorial co-	
	ordinates from 360 $^\circ$ to 0 $^\circ$ (left to right) and -26 $^\circ$ to 90 $^\circ$ (bottom to	
	top). The SDSS North Galactic Cap strips are numbered from 9 to 44	
	and the Southern SDSS strips are numbered 76, 82 and 86. SEGUE fills	
	in Southern strips 72 and 79. The lower panel is the SEGUE Survey	
	footprint in Galactic coordinates in Aitoff projection, centered on the	
	Galactic anticenter. The line marking the Southern limit of the tele-	
	scope observing site is indicated in magenta. Red and green filled areas	
	$represent\ South\ and\ North\ SDSS\ and\ SEGUE\ strips\ respectively\ (Yanny$	
	<i>et al., 2009).</i>	75
3-10	VVV coverage. Hard-lined box is the Galactic Bulge area and the	
	dotted-lined box is the adjacent section of the mid-plane (Minniti	
	<i>et al., 2009)</i>	77
4-1	UKIDSS coverage across our sample of sources, in equatorial coordi-	
	nates. The Chandra X-ray sources are plotted in white circles and the	
	red crosses correspond to the ones with UKIDSS GPS counterparts found	
	in the K-band.	82
4-2	Compilation of the source $CX886$ where we can see the images in JHK	
	of the source. The close-up images are within 5 arcseconds of the Chan-	
	dra source. A colour-colour diagram of all the returned matches within 5	
	arcseconds of the CXC source is also plotted. The colours in this plot are	
	dereddened, using the Schlegel dust maps. Note that it is the maximum	
	extinction considered in this case.	84
4-3	Number of matches returned, within 5 arcseconds, for each CXC source.	86
4-4	Histogram of the distance to the closest UKIDSS match to the Chandra	
	source in green, the maximum Chandra radius error (Equation $4-2$) in	
	blue and the Chandra radius error (Equation 4-1) in red	88
4-5	Histogram of the different K magnitudes returned after the searches on	
	UKIDSS of the random positions. The thick line corresponds to the	
	actual Chandra sources we are interested in. We see very similar distri-	
	butions, in line with a significant contribution from false matches. \ldots	89

х

4-6	Colour-colour diagram of the 'Best' UKIDSS counterparts of the Chan-
	dra sources, plotted in white circles. We also added the hardness of the
	sources, an extinction vector for $E_{B-V}=1$ and colours from the stellar
	library of Pickles (in red)
4-7	Colour-magnitude diagram of the best UKIDSS matches, as well as their
	hardness. In red, we see the main-sequence (MS) stars from the Pickles
	library. In general, we notice a similar trend between the MS Pickles
	stars and the UKIDSS distribution in the diagram. However, the Pickles
	sequence can be shifted vertically because we plot the absolute magnitude
	in K and not the apparent one. The extinction vector is plotted for
	$E_{B-V}=1.$
4-8	E_{B-V} of the Chandra sources
4-9	Spectrum of a CV discovered in GBS, taken with Hydra
4-10	Spectrum of a CV discovered in GBS, taken with the NTT. The magenta
	dotted line shows the location of the H α emission line, which can be seen
	<i>in this case.</i>
4-11	Spectrum of a late M-type star, taken with the NTT
4-12	The average normalised optical spectrum of Sco X-1. In the three panels,
	the same normalised flux density is used. In the middle and bottom
	panels, two emission lines, HeII4686 (middle panel) and H eta (bottom
	panel), are off-scale. This was done in order to visualise the weaker
	lines (Steeghs & Casares, 2002)
5-1	Overlapping IPHAS and SDSS fields. The cyan dots are the IPHAS
	fields centers, the ones ones correspond to the SEGUE coverage in the
	Galactic Plane. The magenta strip corresponds to the region which is
	calibrated, using SDSS data. As can be seen, there is a SEGUE strip
	which does not contain IPHAS data (with $l > 210^{\circ}$)
5-2	Offset in the i-band vs offset in the r-band, in all the IPHAS fields.
	The boxes are used to find the fields acting as 'outliers' in this diagram.
	Those same fields were found as 'outliers' in Figure 5-7. Fields with a
	reliable calibration should have $\Delta_i = \Delta_r$

5-3	Offset in the i band of the IPHAS field 3382 from the December 2006
	run, with respect to SDSS
5-4	Offset in the <i>i</i> band of the IPHAS field 33820 from the December 2006
	run, with respect to SDSS
5-5	Gradient of the offset in the i band across the entire strip (per CCD). $.113$
5-6	Gradient of the offset in the i band, in the IPHAS field 3382 of the
	December 2006 run. We notice the L-shape of the CCDs on the WFC.
	Each CCD's contour is colour-coded. CCD1's boundaries are plotted in
	black solid lines, CCD2's are in blue, CCD3's are in red and finally
	CCD4's in green. Moreover, we clearly see the vignetting problem of
	CCD3 mentioned in the text
5-7	Colour-colour diagram of $(r - H\alpha)$ vs $(r - i)$ of all the sources detected
	across the entire strip. The good fields are plotted in blue, whereas the
	bad ones are the green outliers
5-8	Distribution of the outliers across CCD1, in detector coordinates 119
5-9	Distribution of the outliers across CCD2, in detector coordinates 120
5-10	Distribution of the outliers across CCD3, in detector coordinates. As
	mentioned, we notice the vignetting problem in the bottom left corner of
	the detector, as well as the bad columns
5-11	Distribution of the outliers across CCD4, in detector coordinates. We
	can clearly see the bad column in this detector
5-12	Galactic coordinates of all the H α emitters from Witham's catalogue,
	which were detected in the Galactic Plane during his time of search.
	The regions away from the Galactic center had not been observed yet,
	therefore they do not contain $H\alpha$ emitters. The dotted lines show the
	calibrated strip.
5-13	Galactic coordinates of all the sources calibrated, as well as the H α emit-
	ters from Witham's catalogue, which were also found in the strip. The
	large gaps with no sources correspond to the IPHAS fields which were
	observed during non-photometric nights and had to be taken out of the
	calibration. They are scheduled for re-observation

5-14 Colour-colour diagram of $(r - H\alpha)$ vs (r - i) of all the sources in the strip (cyan dots). The magenta stars correspond to all the $H\alpha$ emitters taken from Witham's catalogue, which fall in that same region of the sky. In total, we found 13 emitters from Witham's catalogue. In black is the location of the unreddened main-sequence from O5 to M6, as well as a few spectral types (MS track taken from Drew et al. (2005), in the case 5-15 Distribution of the $(r - H\alpha)$ colour in the (r - i) colour bin of [0.1, 0.3]. In this case, we see a single Gaussian distribution, centered around 0.1 5-16 Distribution of the $(r - H\alpha)$ colour in the (r - i) colour bin of [1.3, 1.5]. In this case, we clearly see two separate Gaussian distributions, each one corresponding to a MS track. \ldots \ldots \ldots \ldots \ldots \ldots \ldots \ldots 1285-17 Colour-colour diagram of $(r - H\alpha)$ vs (r - i). The black dots correspond to all the sources calibrated in the strip. The blue crosses are the $H\alpha$ emitters taken from Witham's catalogue and the magenta stars are the 5-18 Distribution of the $(r - H\alpha)$ colour in the (r - i) colour bin of [0.3, 0.5]. We notice non-perfect Gaussian in this case, but instead a 'bump' in the left tail of the distribution. $\ldots \ldots 130$ 5-19 Density plot of colour-colour diagram of $(r - H\alpha)$ vs (r - i). The colour bar in this case has a linear scale. \ldots \ldots \ldots \ldots \ldots \ldots 1315-20 Density plot of colour-colour diagram of $(r - H\alpha)$ vs (r - i). The colour 5-21 Colour-colour diagram of (u-g) vs (g-r). The black, magenta and cyan dots all correspond to reference stars. The black dots are SDSS main sequence stars, whereas the cyan (DB WDs = helium-rich WDs)and magenta (DA WDs = hydrogen-rich WDs) dots correspond to the location of white dwarfs in this diagram. The yellow star shows the location of the dereddened colours of the H α emitter, whereas the red star corresponds to its reddened colours $\ldots \ldots 134$

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Declaration

This thesis is the sole work of Sandra Magdy Kamel Greiss, all other works and contributions are acknowledged.

This work has not been submitted to any other university or for the purpose of any other degree or qualification.

The research described in Chapter 4 is briefly used in a paper by (Jonker et al., 2011). The paper's authors are: P.G. Jonker, C.G. Bassa, G. Nelemans, D. Steeghs, M.A.P. Torres, T.J. Maccarone, R.I. Hynes, S. Greiss, J. Clem, A. Dieball, R. Wijnands, J.J.M. Int Zand, M. Méndez, N. Rea, E. Kuulkers, J. Grindlay, E.M. Ratti, L. van Haaften, C. Heinke, F. Ozel, F. Verbunt.

Abstract

Two separate projects, with many common goals, are described in this thesis. The main objective of the work accomplished is to develop methods to detect rare and exotic stellar sources. Astronomical surveys are the most important tool used in this project. They provide most of the information required to draw valid scientific conclusions on sources and select the interesting ones for follow-up. The surveys used cover the Galactic Bulge and Plane, a region of star formation. It also contains X-ray binary sources, the main targets of this project. Low-mass X-ray binaries (LMXBs), high-mass X-ray binaries (HMXBs) and cataclysmic variables (CVs), all have interesting properties and are rare. They contain a compact star, accreting from a normal main-sequence star. Such sources are important in order to understand stellar evolution because the compact star is the end point of a star's life. Understanding their behaviour and studying their properties in a statistical way will, in general, help astronomers answer many unresolved questions.

On the one hand, the Chandra Galactic Bulge Survey is cross-matched with the UKIDSS Galactic Plane Survey in order to find the near-infrared counterparts of ~ 1700 Chandra X-ray sources detected in the Bulge. This region on the sky is extremely crowded and suffers from high extinction, which is why multi-wavelength observations of the X-ray sources is required to disentangle the effect of reddening and the selection of real matches can be more accurate. UKIDSS GPS data release 6 was used but not all JHK magnitudes and images were available. The criteria chosen for finding a true counterpart to the Chandra sources was to have a match in UKIDSS within 1.6 arcseconds of the X-ray source and it was found that 50% of the matches were probably unrelated foreground stars. Colour-colour and colour-magnitude diagrams of the near-infrared colours of the ~ 1700 sources were plotted, where the outliers where chosen for follow-up. How-

ever, near-infrared colours are insufficient to find the real exotic sources in the Bulge, which is why optical data of the Chandra Galactic Bulge Survey fields are required to distinguish the X-ray binary systems from the single active stars found in those fields.

On the other hand, another survey of the Galactic Plane, IPHAS, is very useful to find interesting sources such as CVs. Binary stars with accretion discs are $H\alpha$ emitters and IPHAS uses an $H\alpha$ narrow-band filter to detect such sources. However, at the time of this work, the survey did not have a global photometric calibration, therefore it is only possible to work on each field separately, without combining data from different fields in the same diagrams. SDSS, more precisely SEGUE, was used to calibrate IPHAS. There are 15 strips of ~ 20 deg² which overlap between IPHAS and SDSS. Both surveys use the r and i filters, therefore determining their magnitude offsets in IPHAS with respect to SDSS will permit us to calibrate IPHAS. The calibration began in a strip far from the Galactic Center in order to avoid crowding and high extinction, and it was done a CCD by CCD basis.

Moreover a selection method to pick H α emitters was developed. The calibrated data of a 20 deg² strip was used to test the method. Out of ~ 800,000 sources detected in the strip, only 0.35 % of them were found to be H α excess objects. Cross-matching the H α emitters found with Witham's catalogue (Witham et al., 2006) lead to two sources in common. That region of the sky was poorly covered at the time of Witham's search. One of the sources in common is a known and studied H α emitter, whereas the second one has interesting and unusual observational properties. Follow-up on the William Herschel Telescope will confirm its identity.

Ongoing surveys such as VVV, UVEX and VPHAS+ all cover the Galactic Plane region, bringing more recent and additional wavelength coverage of that region. Using colour-colour diagrams, in the near-ultraviolet, optical and near-infrared regions of the electromagnetic spectrum will give us the information required. Binary stars do not have the same colours as single stars, therefore such diagrams will give us a precise list of sources which require spectroscopic follow-up to confirm their identities.

Chapter 1

Motivation

For many years, galactic X-ray point sources have been a mystery to most astronomers due to the extreme difficulty in finding them, the large uncertainty in their positions and the actual origin of the X-rays. Many are found to lie within the Galactic bulge, a region which suffers from very high extinction and crowding ??. An interesting class of X-ray sources are X-ray binaries, consisting of a normal star and a collapsed stellar remnant, which could be a neutron star or a black hole. A binary star system contains two stars which orbit around their common centre of mass. These pairs of stars produce X-rays if the stars are close enough that the material is transferred from the normal star (the donor) to the dense collapsed one, known as the accretor or compact object. This process, called accretion, heats the material to very high temperatures and consequently releases photons of very high energies at very short wavelengths (King, 2003). In this process, potential energy is converted into thermal and kinetic energies; producing very hot plasma near the compact object. The X-rays come from the area around the collapsed star, the so-called accretion disc.

The reason why such interacting binaries are very important to detect and study is because they provide information on the physics of accretion and on the properties of neutron stars and black holes (King, 2003; Carroll & Ostlie, 2006). In fact, it is easier to deduce the mass, radius and velocity of each star when they belong to a binary system rather than as single stars in the sky (Hynes, 2010; Ozel et al., 2010). All of the 23 known stellar mass black holes are in binary systems (Remillard & McClintock, 2006; Ozel et al., 2010), which motivates us to search for more X-ray binaries, in order to broaden our knowledge about black holes. Moreover, such sources allow us to understand accretion discs better, which are important in many other environments such as protostars or AGN (Active Galactic Nuclei). Finally, assessing their properties in a statistical sense should help us analyse their evolution and characteristics.

In Chapter 2 of this thesis, we will explain the physics and properties behind stellar evolution of a single star, in order to understand the evolution of a binary system. We will focus on the characteristics of certain types of X-ray binary systems, which are the main focus of the research methods described in Chapters 4 and 5. Chapter 3 outlines the methods and tools used to find the objects introduced in Chapter 2. The data used by astronomers can be found in surveys. The main regions of the Milky Way, targeted in order to detect the exotic X-ray binaries, are the Galactic Plane and Bulge. Therefore, in Chapter 3, we give a brief description of the different surveys exploited in this project.

It is important to mention that this thesis describes two different projects, which make use of very similar methods and tools. On the one hand, in Chapter 4, we focus on a small area of the Galactic Bulge, and analyse the near-infrared data of X-ray sources in two specific strips of that region. The main focus of this Chapter is the search for X-ray binary candidates. On the other hand, in Chapter 5, we work on calibrating another survey of the Galactic Plane, called IPHAS. The latter does not contain X-ray data, yet it observes a region of the sky which harbours other types of interesting binary systems, such as cataclysmic variables.

The ultimate goal of both projects is to discover exotic binary systems, which will be candidates for spectroscopic follow-up to confirm their identities. Even though the surveys used in both projects are different and do not cover the same regions, we will see in Chapter 6 what the typical steps of an astronomer are and the problems encountered along the way.

Chapter 2

Stellar evolution, Compact Stars, Accretion discs and X-ray binaries

2.1 Stellar Evolution: the birth and death of stars

2.1.1 Star Formation

2.1.1.1 Protostars

Stars are born from huge clouds of interstellar gas and dust, which are found inside galaxies. The gas is mostly composed of hydrogen (71%) and helium (27%), whereas the dust is composed of solid, microscopic particles of silicates, carbon and iron compounds. Interstellar gas is usually very cold, which explains why the pressure and density in such regions are very low. The atoms and molecules in such conditions move too slowly to generate much pressure. If the cloud is disturbed, for instance, by the explosion of a nearby star, or by the collision with a neighbouring cloud, the low pressure will not be able to support the cloud against its own gravity, which leads to its collapse (Arny, 2005). There is a critical mass at which the cloud becomes unstable once it is disturbed and begins to collapse (Jeans, 1902). This mass was discovered by the British physicist, Sir James Jeans in the second half of the nineteenth century. Before mentioning the Jeans mass, it is important to describe the Jeans length which is the limit for stability to gravitational collapse. All scales smaller than the Jeans length will be stable. The

Jeans length is:

$$\lambda_J = \frac{c_s}{\sqrt{G\rho}} \tag{2-1}$$

where c_s is the sound speed of the gas, G is the gravity constant and is equal to $6.67 \times 10^{-11} \text{ m}^3 \text{ kg}^{-1} \text{ s}^{-2}$ and ρ is the gas density.

For a spherical constant density cloud, the Jeans mass can then be written as (Jeans, 1902):

$$M_J = \frac{4\pi}{3}\rho R_J^3 \tag{2-2}$$

where R_J is the radius of the sphere in which the gas is contained.

The studies and observations of interstellar clouds have shown that they are nonuniform regions. In fact, they contain smaller clumps of gas, with higher density than the average. When such a cloud collapses, each clump is also compressed, growing in density and finally collapsing. The rotation of the gas clumps flattens them into discs because the total angular momentum is only conserved in the direction perpendicular to the rotation axis. A few million years later, the discs form hot, small and dense cores at their centres, called protostars. The cores are hot because they accrete from the discs around them and release potential energy (Arny, 2005).

Protostars are hotter than the gas from which they were formed, but they are still much cooler than ordinary stars. Such sources are usually visible in the infrared and radio wavelengths. The visible light emitted by protostars is very weak and immediately absorbed by the dust around them. However, protostars do not remain cool for very long. They continue to accrete from surrounding material, releasing energy and making them hotter. When the temperature in a deuterium reaches a few million Kelvins, nuclear reactions begin. Hydrogen atoms fuse into helium and release energy. This new and powerful energy supply increases the temperature and outward pressure of the star's core, which stops the core from collapsing further.

A protostar passes through these stages and continues to accrete from surrounding material, in a time that is brief by astronomical standards: a few million years for a star like the Sun and even less for a more massive star. However, for all types of protostars, infall creates violent changes in brightness and ultimately creates a strong outflow of gas focused into narrow jets, perpendicular to the disc (Arny, 2005).

The jets are seen though the interaction with the gas that remains around the star. Further out around the jet's directions lie several small bright blobs called Herbig-Haro objects. The jet of gas from the star spreads when it hits surrounding gas, and the impact heats both the jet and the gas around it, making them glow. The jet also pushes gas outward from around the protostar to create bipolar flows, so-called because there are two jets of gas ejected parallel to the star's polar axes. Bipolar flows are important because they clear away gas and dust from around protostars and therefore allow astronomers to see them directly in the visible light. However, even powerful bipolar flows leave some surrounding material behind, thus many young stars are partially immersed in interstellar matter (Arny, 2005). This exact same phenomenon is also found in X-ray binaries systems, where a collapsed stellar remnant accretes material from a normal star, forming an accretion disc around it (Prialnik, 2000; King, 2003). A more detailed description of such systems is found later in this Chapter.

2.1.1.2 Stellar Mass limits

The mass of most stars is found to range between $\sim 0.08 \, M_{\odot}$ and $30 \, M_{\odot} \, (M_{\odot} = \text{solar mass})$. These limits are not absolute, but theory of star formation has proven that very high or very low-mass stars are not very uncommon. Moreover, observations have detected the presence of brown dwarfs and they are believed to exist in abundance, however, they are very difficult to detect.

The reason why we detect a very small number of stars less massive than 0.08 M_{\odot} is because the stage of nuclear burning in the core cannot be reached due to the fact that their mass is too small to be compressed. Such stars, 'brown dwarfs', are

extremely faint and are therefore not easily detected due to the fact that they are very small and dim. This is more of a selection effect because theory predicts their existence and they are very abundant stars, however, even with today's technology and imaging systems, the extremely low-mass sources (M < 0.08 M_{\odot}) are too faint to be detected, unless they are very close by.

Very high-mass stars are, on the other hand, rare for a different reason. In this case, it is a physical explanation rather than a selection effect which limits the number of detected high-mass stars. Having a large amount of material, their gravity compresses them and they rapidly become extremely hot and luminous. Such a high luminosity means that the stars release intense radiation which heats up the gas around them, raising the pressure and preventing additional material from falling onto them. Therefore, it is their high temperature which limits the amount of material they can accumulate.

During galactic evolution, it is believed that massive stars become rarer than low-mass stars because they have shorter life spans and their relative number is correlated with the fractional amount of gas in the considered galaxy. If it is assumed that star formation is independent of the age of the galaxy and its location, one can write the number of stars formed at a given time within a given volume and masses between (M, M + dM) as a function of M (Prialnik, 2000):

$$dN = \Phi(M)dM \tag{2-3}$$

where Φ is the birth function, derived by Salpeter (1955) and can be written as:

$$\Phi \propto M^{-2.35} \tag{2-4}$$

Salpeter studied main-sequence stars in the solar neighbourhood.

The initial mass function (IMF), $\xi(M)$ is the amount of mass locked up in stars with masses in the between (M, M + dM), formed at a given time and within a given volume. In other words, the distribution of stellar masses is referred to as the stellar initial mass function (Rana, 1987). Therefore the following equation



Figure 2-1: The mass function of stars in the solar neighbourhood (Rana, 1987). The function is compared to that of Scalo (1986).

can be deduced:

$$MdN = \xi(M)dM \tag{2-5}$$

Finally, by combining the last two equations, the initial mass function can be written as:

$$\xi(M) \propto \left(\frac{M}{M_{\odot}}\right)^{-1.35} \tag{2-6}$$

By looking at Figure 2-1, the IMF is not consistent with the power-law from Equation 2-6 when considering the lower end of the mass range. In fact, the IMF becomes almost flat on this side of the graph and even decreases with mass. Due to the fact that low-mass stars, more precisely brown dwarfs, are hard to detect because of the fact that they are very faint, the birth function in this mass range is rather uncertain.

2.1.2 Main-sequence stars

A protostar becomes a main-sequence (MS) star when nuclear fuel in its core can supply energy to raise its pressure and stop its collapse. At this stage of its life, it contains a core, where nuclear fusion occurs so hydrogen burns into helium, as well as an envelope of gas around the core to transport energy to the star's surface. The properties of the different layers of the star depend mainly on its mass.

At this stage of a star's life, four physical equations describe stellar structure: hydrostatic equilibrium, mass conservation, energy conservation and energy transport. Hydrostatic equilibrium prevents the star from collapsing (Prialnik, 2000):

$$\frac{dp}{dr} = -\frac{G\rho M}{r^2} \tag{2-7}$$

where p is the pressure, r, M and ρ the radial coordinate, mass and density respectively of the star. G the gravitational constant.

By associating a star to a blackbody, it satisfies the mass, luminosity and temperature relation (Prialnik, 2000):

$$L = 4\pi R^2 \sigma T_{\rm eff}^4 \tag{2-8}$$

where σ is the Stefan-Boltzmann constant and is equal to 5.67 × 10⁻⁸ W m⁻² K⁻⁴, L is the luminosity of the star, R its radius and T_{eff} its effective temperature. Equation 2-8 defines the luminosity of a star, which corresponds to the amount of light radiated by the star per unit of time. As seen, it is determined by the radius and surface temperature of a star.

Mass conservation is given by the following equation (Prialnik, 2000):

$$M_r = \int_0^r 4\pi r^2 \rho dr \qquad \frac{dM_r}{dr} = 4\pi r^2 \rho$$
 (2-9)

where M_r is the mass within the radius r.

Hydrostatic equilibrium is shown in Equation 2-7 and energy conservation can be found by considering the change in the luminosity as a function of radius of the star because of nuclear energy production (Prialnik, 2000):

$$dL = \epsilon_N \rho 4\pi r^2 dr \qquad \frac{dL}{dr} = 4\pi r^2 \epsilon_N \rho \qquad (2-10)$$

where ϵ_N is the nuclear production efficiency which is a function of the density, temperature and chemical composition of the star.

Finally, energy transport by radiation also describes the equilibrium state of the star (Prialnik, 2000):

$$\frac{dT}{dr} = -\frac{3\kappa\rho}{16ac\pi r^2 T^3}L\tag{2-11}$$

where κ is the opacity of the gas and a is the radiation constant and is equal to $4\sigma/c = 7.5657 \times 10^{-15} \text{ erg cm}^{-3} \text{ K}^{-4}$. An erg is a unit of energy and is equivalent to 10^{-7} Joules.

Gravity and pressure must be in balance inside a star, therefore a more massive star will require greater pressure to maintain the balance. According to the ideal gas law, higher pressure can be achieved with higher temperatures. Therefore, a high-mass star (M > 10 M_{\odot}) should have a hotter core than a low-mass star (M < 10 M_{\odot}), which leads to higher luminosity (Prialnik, 2000). The reason why the luminosity has such a high dependance on the mass of the star (see Section 2.1.6 for L \propto M³) is because ϵ_N is a strong function of the temperature.

The nuclear fusion process which goes on in the star's core is different when looking at low and high-mass stars. In fact, in the core of a low-mass star, hydrogen is converted to helium by the proton-proton chain (Prialnik, 2000), which consists of two protons fusing to form heavy hydrogen, to which a third hydrogen is added to make ³He. The ³He then fuses to form ⁴He. All nuclear reactions in stars must follow conservation rules: charge and energy conservation, as well as number of leptons and quarks. For the reaction to be stable, meaning that the formed element does not decay instantly, its mass must be smaller than the total mass of elements producing it.

Proton-proton chain I:

$${}^{1}\mathrm{H} + {}^{1}\mathrm{H} \rightarrow {}^{2}\mathrm{H} + e^{+} + \nu$$
$${}^{2}\mathrm{H} + {}^{1}\mathrm{H} \rightarrow {}^{3}\mathrm{He} + \gamma$$
$${}^{3}\mathrm{He} + {}^{3}\mathrm{He} \rightarrow {}^{4}\mathrm{He} + 2{}^{1}\mathrm{H}$$

Proton-proton chain II:

$${}^{3}\text{He} + {}^{4}\text{He} \rightarrow {}^{7}\text{Be} + \gamma$$

 ${}^{7}\text{Be} + e^{-} \rightarrow {}^{7}\text{Li} + \nu$
 ${}^{7}\text{Li} + {}^{1}\text{H} \rightarrow 2{}^{4}\text{He}$

Proton-proton chain III:

$${}^{7}\text{Be} + {}^{1}\text{H} \rightarrow {}^{8}\text{B} + \gamma$$
$${}^{8}\text{B} \rightarrow {}^{8}\text{Be} + e^{+} + \nu$$
$${}^{8}\text{Be} \rightarrow 2^{4}\text{He}$$

However, in a massive star, the conversion of hydrogen to helium takes place by means of the CNO (Carbon-Nitrogen-Oxygen) cycle (Prialnik, 2000). These 'metals' are already found in the core of all types of stars, however the core temperature of low-mass stars is not high enough to initiate the CNO cycle at a high rate. The CNO cycle is more temperature dependent than the proton-proton chain. Photons are unable to carry this huge amount of energy from small volumes, thus the gas in the core begins to rise in irregular clumps and carries the heat outward by convection. These currents do not extend close enough to the star's surface to allow hydrogen-rich gas from the outer layers to mix into the core and replenish the core's depleted fuel. Only very low-mass stars, with masses less than 0.3 M_{\odot} mix completely.

The CNO cycle:

$${}^{12}C + {}^{1}H \rightarrow {}^{13}N + \gamma$$
$${}^{13}N \rightarrow {}^{13}C + e^{+} + \nu$$
$${}^{13}C + {}^{1}H \rightarrow {}^{14}N + \gamma$$
$${}^{14}N + {}^{1}H \rightarrow {}^{15}O + \gamma$$
$${}^{15}O \rightarrow {}^{14}N + e^{+} + \nu$$
$${}^{15}N + {}^{1}H \rightarrow {}^{12}C + {}^{4}He$$

$$\label{eq:starsest} \begin{split} ^{14}\mathrm{N} + ^{1}\mathrm{H} &\rightarrow ^{15}\mathrm{O} + \gamma \\ ^{15}\mathrm{O} &\rightarrow ^{15}\mathrm{N} + e^{+} + \nu \\ ^{15}\mathrm{N} + ^{1}\mathrm{H} &\rightarrow ^{16}\mathrm{O} + \gamma \\ ^{16}\mathrm{O} + ^{1}\mathrm{H} &\rightarrow ^{17}\mathrm{F} + \gamma \\ ^{16}\mathrm{O} + ^{1}\mathrm{H} &\rightarrow ^{17}\mathrm{F} + \nu \\ ^{17}\mathrm{F} &\rightarrow ^{17}\mathrm{O} + e^{+} + \nu \\ \end{split}$$

A star's MS lifetime depends on its mass and luminosity. Its mass determines how much fuel it has and its luminosity determines how rapidly the fuel is burned. Therefore, in order to calculate a star's MS lifetime, we simply need to divide the amount of fuel by the rate at which it is consumed. In Section 2.1.6, the nuclear timescale of a star, which can also be considered as its MS lifetime, is the nuclear energy of the star divided by its luminosity. The latter is the rate at which the star loses energy. The luminosity L of a star is proportionate to M^3 and its nuclear energy is proportionate to Mc^2 , therefore the MS lifetime of a star is $\tau_{MS} \propto M^{-2}$ (Prialnik, 2000).

2.1.3 Giant Stars

When a MS star has consumed most of the hydrogen in its core, the pressure then begins to drop and gravity compresses the core and heats it. As its temperature rises, hydrogen begins burning outside the core in what is called the shell source (Prialnik, 2000). Core contraction is followed by the expansion of the envelope. Heat from the shell source and the core raises the pressure around the core. That stronger pressure pushes the surrounding gas outward, making the star and its radius expand. The factor by which the star grows depends on its mass. However, this expansion cools the outer layers, thus the star becomes what is called a red giant (Prialnik, 2000). Its name comes from the fact that it cools and therefore radiates more strongly at long (red) wavelengths, and its radius can grow up to several hundred times larger. A MS star becomes a red giant when it uses up the hydrogen in its core, which all fused into helium nuclei.

A giant is very different to a MS star. Even though it has a very large radius, most of its volume is filled by its very low-density, tenuous atmosphere, and most of its mass is found in a tiny, hot and compressed core. The core, or shell source, supplies the star's energy. The huge envelope is relatively cool and the gas in it absorbs the photons flowing from the core towards the star's surface. The absorption of the photons slows them, so the energy is carried by convection rather than radiation. Some of that energy continues to come from burning hydrogen but many stars switch to a new energy supply: helium burning. The core temperature becomes high enough for helium to ignite (Prialnik, 2000).

Helium burning begins very differently in high-mass and low-mass stars. It is important to note that it begins in the core. Helium nuclei can overcome electrical repulsion and fuse only if they collide at enormously high speeds. The required speed is achieved in hot gases because the higher the temperature of the gas, the faster its nuclei move and therefore the more easily they can fuse to form heavier nuclei. Hydrogen can fuse at about 7 million Kelvin; helium, however, must be heated to about 100 million Kelvin.

If a star's core is heated to about 100 million Kelvin, helium nuclei will fuse into carbon. This nuclear reaction combines 3 ⁴He into a ¹²C. Helium nuclei are sometimes called 'alpha particles', which explains why this process is sometimes referred to as 'the triple alpha process', or simply 'helium burning' (Prialnik, 2000):

$${}^{4}\text{He} + {}^{4}\text{He} \rightarrow {}^{8}\text{Be}$$
$${}^{8}\text{Be} + {}^{4}\text{He} \rightarrow {}^{12}\text{C}$$
(2-12)

The reason why this process is called 'the triple alpha process' is because three helium nuclei are needed to produce carbon. In fact, ⁸Be is unstable and decays quickly, unless associated with a ⁴He to produce a ¹²C. ⁸Be decays quickly because the mass of two helium nuclei is smaller than the mass of beryllium, therefore the

reaction does not produce a loss of mass and can be reversed.

Once a star has consumed all its core hydrogen, it must use its helium or contracts. A high-mass star needs to compress its core a little bit before helium begins to fuse because its core is already extremely hot. A low-mass star such as the Sun, however, must compress its core enormously to make it hot enough for helium to begin burning. The compression of such a star packs its gas atoms so closely that they no longer behave like an ordinary gas. Such a gas is called a degenerate gas.

In a degenerate gas, the matter is so densely packed that the electrons act according to the laws of subatomic physics. No more than two electrons of the same energy can be found in the same volume, meaning they can not occupy the same energy level. This law is called the Pauli exclusion law (see Section 2.2.1). In a normal gas, the energy released from nuclear fusion heats the gas and increases its pressure, which leads to its expansion. When the gas expands, it then cools down and reduces the rate of nuclear burning and less energy is therefore released. In normal stars, this mechanism keeps them from collapsing or exploding. In the case of a normal gas, it satisfies the ideal gas law:

$$P = NkT \tag{2-13}$$

where P is the pressure of the gas, N the number of particles in the gas, T its temperature and k the Botlzmann constant equal to 1.38×10^{-23} J.K⁻¹. However, in the case of a degenerate gas, the pressure is written as (Prialnik, 2000):

$$P = \frac{h^2}{20m_e m_p^{5/3}} \left(\frac{3}{\pi}\right)^{2/3} \left(\frac{\rho}{\mu_e}\right)^{5/3}$$
(2-14)

where h is the Planck constant and is equal to 6.626×10^{-34} J.s, m_e is the mass of the electron, m_p is the mass of the proton, ρ is the density and $\mu_e = \frac{N_e}{N_p}$ is the ratio of electron number to proton number. This formula is only valid in nonrelativistic cases at T = 0. When particle energies reach relativistic cases, the degenerate pressure is proportionate to $\left(\frac{\rho}{\mu_e}\right)^{4/3}$.

In this case, the pressure only depends on the density of the gas and not on the

temperature.

When nuclear burning begins in a low-mass star with a degenerate core, the energy released does not raise the pressure because it does not depend on the temperature anymore but only on the density of the star (Equation 2-14). Nuclear burning in degenerate material is thermally unstable (Prialnik, 2000). The gas gets hotter and in only a few minutes it releases several thousand times more energy, which leads to what is termed a helium flash. This happens when the mass of the core reaches $0.5 M_{\odot}$ (Prialnik, 2000). There is no stabilizing expansion and subsequent cooling after the ignition in the case of such stars. This outburst is hidden from our view by the star's outer layers. The energy released heats the core enough for it to become a normal gas once again. With its degeneracy gone, the gas can now expand and adjust the star's outer layers by shrinking, compressing and heating them. The star's reheated surface changes colour from red to yellow and it becomes what is called a yellow giant.

It is important to note a difference in the case of low-mass stars. For stars with M $< 2.3 M_{\odot}$, the helium core becomes degenerate during hydrogen burning and when the mass of helium in the core reaches $0.5 M_{\odot}$, the helium flash occurs (Prialnik, 2000). The fate of stars with $2.3 M_{\odot} < M < 8 M_{\odot}$ is not yet very well understood, but it is well known that low-mass stars (M $< 10 M_{\odot}$) end their lives as white dwarfs.

High-mass stars do not have a helium flash, but instead have a steady helium burning phase (Prialnik, 2000).

2.1.4 Death of stars like the Sun

A star like the Sun will spend about 10 billion years on the MS, before becoming a red giant. However, it will spend less than 1% of that time, about 100 million years, consuming its helium. The evolution will be faster, because helium releases less energy per unit mass than hydrogen when it is burned, making the star brighter. During the triple alpha process, the star's radius will once again shrink, compressing and heating it. The temperature of the core does not reach the limit to burn carbon, but the compression does make it hot enough to increase significantly the rate at which the helium burns. The quicker the fuel is consumed, the more luminous the star gets and its expansion reaches a larger radius than before, making it a supergiant. As the star inflates, the outer layers cool to about 2500 K (Arny, 2005). The temperature becomes cool enough for carbon and silicon atoms to condense and form grains, which do not fall back in the star. The photons released from the star's core push the grains outward, into space, and form a huge shell of dust around the star.

The gas shell then expands and becomes transparent, allowing us to see through to the star's hot core. Because it is so hot, the core's radiation is rich in ultraviolet light, which heats and ionizes the shell around it, making it glow. Such bright objects are called planetary nebulae. They have nothing to do with planets but when they were discovered, due to the size of the telescopes used at that time, they looked like small disks, which were very similar to planets.

A planetary nebula shell contains about one-fourth solar mass of glowing gas but may have as much as several solar masses of cooler, non-luminous gas around it. Shells are typically about one-fourth of a light-year in diameter (1 light-year is the distance travelled by light in one year, which is equivalent to $\sim 10^{16}$ m) and expand at about 20 kilometers per second (Arny, 2005). The shells eventually grow so big that the gas ends up in the interstellar medium. However, the core of the star remains behind as a tiny, glowing star, called a white dwarf.

A very different ending awaits a more massive star.

2.1.5 Death of massive stars

Massive stars do not become planetary nebulae or white dwarfs because their great mass compresses and heats up their cores enough to ignite carbon and allows them to keep burning when their helium is gone. There are many other ways to supply energy to a star by creating heavy elements. Astronomers call the formation of heavy elements by nuclear burning processes nucleosynthesis. The chemical elements in our Universe heavier than hydrogen were made this way.

In massive stars, nucleosynthesis provides enough energy to support the stellar mass from collapsing inwards because of the gravitational force. As each fuel is exhausted, whether hydrogen, helium or carbon, the star's core contracts and heats by compression (Arny, 2005). The higher temperature then allows the star to burn still heavier elements. Oxygen, neon, magnesium, and eventually silicon are formed.

Carbon-Oxygen burning:

$$\label{eq:constraint} \begin{array}{l} ^{12}\mathrm{C} + ^{12}\mathrm{C} \rightarrow ^{24}\mathrm{Mg} + \gamma \\ \rightarrow ^{23}\mathrm{Mg} + n \\ \rightarrow ^{23}\mathrm{Na} + p \\ \rightarrow ^{20}\mathrm{Ne} + \alpha \\ \rightarrow ^{16}\mathrm{O} + 2\alpha \end{array}$$

$$^{16}O + {}^{16}O \rightarrow {}^{32}S + \gamma$$

$$\rightarrow {}^{31}S + n$$

$$\rightarrow {}^{31}P + p$$

$$\rightarrow {}^{28}Si + \alpha$$

$$\rightarrow {}^{24}Mg + 2\alpha$$

The formation of an iron core signals the end of a massive star's life because according to astronomers, the iron nucleus is the most tightly bound of all nuclei, unlike what most physicists think when they claim that it is a nickel isotope which is the most tightly bound element. Once the fuel is entirely consumed, the core of the star starts to shrink and heat up. In the case of massive stars, the core shrinks enough to press the iron nuclei so tightly that protons and electrons merge to form neutrons. Therefore, such stars end up with neutron cores, which leads
Table 2-1: Different possible end states of stars according to their initial masses, in the case of a single star and when found in a binary system (Tauris & van den Heuvel, 2006)

Initial mass	He-core	Single star	Binary star
$< 2.3 \ \mathrm{M}_{\odot}$	$< 0.45 \ \mathrm{M_{\odot}}$	CO white dwarf	He white dwarf
2.3 - $6~{\rm M}_{\odot}$	0.5 - $1.9~{\rm M}_{\odot}$	CO white dwarf	CO white dwarf
6 - $8~{ m M}_{\odot}$	1.9 - $2.1~{\rm M}_{\odot}$	O-Ne white dwarf	O-Ne white dwarf
		or C-deflagration SN?	
8 - $12~M_{\odot}$	2.1 - $2.8~M_{\odot}$	neutron star	O-Ne white dwarf
12 - $25~M_{\odot}$	2.8 - $8~{\rm M}_{\odot}$	neutron star	neutron star
$> 25 \ M_{\odot}$	$> 8 \ {\rm M}_{\odot}$	black hole	black hole

to catastrophic results because most of the pressure which supported the core was supplied by electrons. The electron degenerate pressure is replaced with neutron degenerate pressure. The star's core pressure drops and its interior begins to collapse. The mass of the core of such a star reaches the Chandrasekhar limit (see Section 2.2.1), which is the maximum limit possible for an electron degenerate configuration. That limit is about 1.4 M_{\odot}, which applies to the case of a white dwarf but not a neutron star (Prialnik, 2000). In less than a second, the core is transformed from an iron ball the size of the Earth to a ball of neutrons about 10 kilometers in radius. The outer layers of the star continue to crush the core and heats the infalling gas to billions of degrees. The pressure rises and lifts the outer layers away from the star in a recoil of the infalling material: a supernova.

A supernova explosion marks the death of a massive star. Gas ejected by the supernova explosion travels through the interstellar medium, mixing with other gas it encounters. The cloud of stellar debris, known as a supernova remnant, continues to expand. Eventually, the supernova remnant slows down and cools. However, the remnant's gas is rich in heavy elements, therefore when the remnant mixes with an interstellar cloud, the latter is also enriched in heavy elements (Prialnik, 2000). Such a cloud can collapse and form a new generation of stars, which will contain more heavy elements than the previous generation.

2.1.6 The characteristic timescales of stellar evolution

Three fundamental timescales exist, which help us understand stellar evolution.

2.1.6.1 Dynamic timescale

The dynamical timescale of a star is the rate at which the radius of a star changes, in the absence of pressure. Since gravity is the binding force of a star, the characteristic velocity in a gravitational field, meaning the free fall velocity, gives the typical rate at which the radius R changes. The dynamical timescale of a star is therefore given by (Prialnik, 2000):

$$\tau_{dyn} = \frac{R}{\dot{R}} = \frac{R}{\sqrt{2GM/R}} = \sqrt{\frac{R^3}{2GM}}$$
(2-15)

where \hat{R} is the rate at which the radius changes and is equal to the free fall velocity. The pulsations of a star vary on a dynamical timescale, which corresponds to about 15 minutes in the case of the Sun. As seen, it is extremely short compared to typical stellar ages. A dynamical process is found when the gravitational and pressure forces do not balance one another. If the pressure is too low to counteract gravity, then contraction occurs, whereas if pressure is too high, then the situation develops into expansion. The end result can either be catastrophic with the collapse or explosion of the star, or it can simply lead to the restoration of hydrostatic equilibrium. In either scenario, the end states will occur on a dynamical timescale (Prialnik, 2000).

2.1.6.2 Thermal timescale

When the thermal equilibrium of a star is disturbed, it will restore this equilibrium on a thermal (or Kelvin-Helmholtz) timescale, which is the time it takes to emit all of its thermal energy content at its present luminosity (Prialnik, 2000):

$$\tau_{th} = \frac{E}{\dot{E}} = \frac{E_{th}}{L} = \frac{GM^2}{RL} \tag{2-16}$$

where \dot{E} is the rate at which the star radiates, therefore its luminosity L and E_{th} is the thermal energy of the star, which is equal to half of its potential energy.

The thermal timescale of the Sun is about 10^{15} seconds, which corresponds to about 30 million years. It is much longer than the dynamical timescale, but still small compared to the life span of a star. If a star maintains a constant luminosity, the thermal timescale can be considered as the time it would take to emit its entire reserve of thermal energy while contracting (Prialnik, 2000).

2.1.6.3 Nuclear timescale

The third stellar timescale is the nuclear one, which is the time needed for the star to consume all its nuclear fuel reserve at its present fuel consumption rate:

$$\tau_{nuc} = \frac{\epsilon M c^2}{L} \tag{2-17}$$

where ϵ is the efficiency for which the fuel is exhausted and it equal to 0.007 for hydrogen burning into helium. It is estimated by the typical binding energy of a nucleon divided by the nucleon's rest-mass energy.

In the case of the Sun, the nuclear timescale is much larger than its age therefore stars only consume a small fraction of their available nuclear energy (Prialnik, 2000).

When taking into account all three timescales mentioned above, astronomers were able to deduce a mass-luminosity function where $L \propto M^3$ and a mass-radius function for MS stars where $R \propto M^{0.5}$. These relations are not valid for the lower end of the mass range of stars, meaning for stars with $M < M_{\odot}$ (Prialnik, 2000).

2.1.7 The evolution of a star on a Hertzsprung-Russell (H-R) diagram

The H-R diagram is a plot of the luminosity versus the temperature (or spectral class) of stars. This diagram is useful to understand stellar evolution because as seen in Figure 2-2, stars fall in different areas of the diagram depending on the stage of their lives. Hot luminous stars are usually found at the upper left hand side of the diagram, whereas cool dim stars are in the lower right side. The Sun



Figure 2-2: Typical H-R diagram (Lecture notes from course PX268, University of Warwick).

lies almost in the middle of the big diagonal line in the centre of the diagram, the main-sequence (MS).

An important feature of the diagram, which must be pointed out, is that temperature increases to the left, rather than to the right. When a group of stars is plotted on an H-R diagram, generally about 90% of them will lie along the MS.

Variation of the outer radius

The evolutionary tracks in the H-R diagram of six different stars are shown in Figure 2-3. These stars all have different masses, ranging from 1 M_{\odot} to 50 M_{\odot} (Tauris & van den Heuvel, 2006). The parameters plotted in the diagram are the luminosity L, the effective temperature T_{eff} and the radius R of the stars. These three observable parameters are calculated by using the mass, luminosity



Figure 2-3: Evolutionary tracks in the H-R diagram of six stars with different masses, ranging from 1 M_{\odot} to 50 M_{\odot} (Tauris & van den Heuvel, 2006)

and temperature relation: L = 4 π R² σ T⁴ $_{\rm eff}.$

When looking at the case of a 5 M_{\odot} star in Figure 2-3, we see small numbers indicating different stages of the star's life. Between 1 and 2, the star is in the phase of core hydrogen burning, the MS, which has a nuclear timescale. At point 3, the hydrogen ignites in a shell around the helium core, which itself ignites at point 4. Therefore, between 3 and 4, the star is a red giant, with a dense core and a very large radius. During the helium burning, the star describes a loop in the H-R diagram. Stars with $M \geq 2.3 M_{\odot}$ move from points 2 to 4 on a thermal timescale and after point 4, they go through the helium burning loop on a helium nuclear timescale. Point 4 then corresponds to the expansion of the outer layers during helium burning and then the star becomes a red supergiant on the asymptotic giant branch (AGB). The case is different for stars with $M < 2.3 M_{\odot}$. The helium core becomes degenerate after the hydrogen shell ignition and has a mass of about 0.45 M_{\odot} at helium flash ignition (Tauris & van den Heuvel, 2006).

2.2 Compact stars

There are three kinds of stellar remnants known: white dwarfs, neutron stars and black holes. These three remnants are called compact stars and are the end points of stellar evolution.

2.2.1 White dwarfs

White dwarfs (WDs) have masses comparable to that of the Sun but their diameters are similar to the Earth's, making them very dense and dim because of their small surface area. WDs have a surface temperature which can range between 100,000 K and 4,500 K.

Since WDs are formed from low-mass stars, their cores are mainly composed of oxygen and carbon formed from hydrogen and helium burning. They have a thin layer of hydrogen and helium around their core but they are not hot enough to begin carbon burning. They are in hydrostatic equilibrium, just like normal stars, however it is important to note that their pressure is provided by degeneracy. Their structure, however, differs a lot from that of MS stars. White dwarfs are extremely dense stars, therefore the particles in the core are packed very close to one another (Arny, 2005). The exclusion principle in particle physics, also known as the Pauli principle, shows that for a given volume, only a specific number of electrons can be found. Each electron can be found at a specific energy level from the ground state and only a certain number of electrons can occupy the same level, as long as each electron has a quantum state different to the others. When more than one electron occupies an energy level, its spin will have to be different to the one of the other electron can be squeezed. Bringing the electrons so close together creates a pressure which is not found in normal gas. Such a gas is a degenerate gas, because of the degeneracy pressure. An important property of this pressure is that it only depends on the density of the gas and not on its temperature (see Equation 2-14).

The radius of a WD shrinks if mass is added to it because the added mass increases its gravity and compresses it. The core being degenerate, the pressure is not large enough to stop the increasing gravity to shrink it. Therefore, the more mass added to that star, the stronger the gravity and the less pressure the degenerate gas is able to produce to maintain the star's size. In this situation, the degeneracy pressure becomes relativistic. Therefore, too much mass falling onto the WD will lead to its collapse. As such, exists a limiting mass of WDs, known as the Chandrasekhar mass (it was discovered by the Indian astrophysicist Subrahmanyan Chandrasekhar, who won the Nobel Prize in 1983 for his theoretical calculations on the limiting mass of WDs). For a typical WD, this limiting mass is about 1.4 M_{\odot} (Prialnik, 2000).

2.2.2 Neutron Stars

When the collapsed core of a massive star is of a certain high density, the star's protons and electrons merge together into neutrons. In such conditions, a neutron star is formed. This happens when the pressure of the degenerate electrons is no longer sufficient to sustain the gravitational force, which is one reason why most neutron stars are born with ~ $1.4M_{\odot}$, corresponding to just above the Chandrasekhar mass for WDs. They have radii of about 10 kilometers and masses of 1.4 M_{\odot} , thus they are extremely dense objects. The maximum possible mass of a neutron star has been found to be ~ 3 M_{\odot} , through stellar models and calculations (Prialnik, 2000).

Because neutron stars are formed from the collapse of massive stars, which generally explode as supernovae, the core radius shrinks to a tiny radius. By taking into account the conservation of angular momentum, the neutron star must therefore spin faster when it shrinks. It also creates a very strong magnetic field. Therefore, pulsars have extremely strong magnetic fields. In combination with their rapid rotation, they produce narrow beams of charged particles at their magnetic poles (Carroll & Ostlie, 2006).

In a pulsar, the charges radiate as they accelerate along the magnetic fields. They have a peculiar trajectory around the magnetic field lines, they spiral as they go and radiate electromagnetic energy.

The emission created by the accelerating charges does not depend on the temperature of the heated gas but rather on the magnetic field, therefore it is called nonthermal radiation or synchrotron radiation. For most pulsars, the charges generate low-energy radiation, which is why they were detected at radio wavelengths. This does not exclude the fact that they can also be detected at much higher energy wavelengths such as visible lights and gamma rays.

Stellar models and calculations show that neutron stars have three separate regions: a thin, outer gaseous atmosphere about 1 millimeter thick; a solid crust a few hundred meters thick; and the core of the star composed of neutrons lying below the crust.

Even though neutron stars have been found in binary systems and as radio pulsars, they remain difficult and rare to detect.

2.2.3 Black holes

When a star with an initial mass larger than 30 M_{\odot} collapses, it leaves behind a black hole. In order to understand the properties of black holes, we must remember the concept of escape velocity. This is the speed a mass requires to be free from another object's gravitational attraction. The escape velocity for an object of mass M and radius R is:

$$V = \sqrt{\frac{2GM}{R}} \tag{2-18}$$

where G is the gravitational constant.

This equation shows that for a given mass, the smaller radius it is compressed to,

the larger its escape velocity.

For a body to be a black hole, the escape velocity is equated to the speed of light, because it is known that no light can escape from a black hole (Prialnik, 2000):

$$V = c = \sqrt{\frac{2GM}{R}}$$
 therefore $R = \frac{2GM}{c^2}$ (2-19)

This equation determining the radius of a black hole is called the Schwarzschild radius, named after the German astrophysicist who discovered it.

General relativity shows that gravity is related to the curvature of space and a black hole forms where the curvature is very extreme. According to general relativity, mass creates a curvature of space and as the bodies move along the curvature, gravitational motion occurs. In the case of a black hole, the reason why light can not escape from it is its extreme curvature. Astronomers call its boundary the event horizon. Because we can not see what is inside a black hole, we will never find out its composition. Its mass is one if the few properties of a black hole that can be determined (Carroll & Ostlie, 2006; Arny, 2005).

Since black holes do not emit light or any sort of electromagnetic radiation, it is not easy to observe them. However, they are detected in binary systems where the initial star which evolved had a mass of at least 10 M_{\odot} and exploded as a supernova leaving behind a black hole. As will be seen later in this chapter, the process of accretion draws gas from the companion star to the compact object and forms a ring of gas around the black hole. In this case, the accretion disc can not have an inner radius smaller than a few times the Schwarzschild radius. Since the material in the disc orbits nearly at the speed of light, the dynamics of the gas heats it up to 10 million K, making it emit in X-rays and gamma rays.

Moreover, in the case of binaries, it is possible to calculate the masses of the stars in the system. Therefore, if the mass of the accretor is larger than 3 to 5 M_{\odot} , it is definitely a black hole since it can not be a neutron star (Carroll & Ostlie, 2006).

2.3 Binary Stars and their evolution

It is estimated that about half of all stars that have been detected were found to be in binary systems, systems containing two stars orbiting around their common centre of mass, held together as pairs by their mutual gravitational attraction. Such systems are very interesting because they provide information on stellar mass and evolution. By using Kepler's third law (2-20), we can determine the system's mass (Carroll & Ostlie, 2006).

$$P_{orb}^2 = \frac{4\pi^2 a^3}{G(M_1 + M_2)} \tag{2-20}$$

where M_1 and M_2 are the masses of both stars and a is the binary separation.

2.3.1 Visual, spectroscopic and eclipsing binaries

For some stars, their orbital motion around their common centre of mass can be spatially resolved by comparing images taken several years apart. Such pairs are called visual binaries. However, some binaries are so close that we can not distinguish both stars only by looking at images. In the case of close binary systems, what we detect is the total flux emitted by both stars. A way to find out whether or not the star observed is in a binary system, astronomers study the spectra of the source. If it is indeed a binary system, it is called a spectroscopic binary. From the spectra of the system, astronomers can deduce the radial velocities (RV) of the stars, by calculating the shift in wavelengths of the spectral lines.

As the stars orbit about their common centre of mass, each star alternately moves toward and away from us. Since the observer and emitting object are in movement relative to one another, there is a change in the observed wavelength of radiation. This change is called the Doppler shift (Equation 2-21) in the spectral lines of the system. When the source and the observer move apart, the shift is an increase in the wavelength and it is therefore redshifted, whereas when the source and observer approach, it is blueshifted because the shift decreases (Arny, 2005).

$$\Delta \lambda = \pm \lambda \frac{v}{c} \tag{2-21}$$

where v is the radial velocity of the body moving away from or towards the observer. This equation is only valid in non-relativistic cases, meaning that $v \ll c$.

By observing the system frequently, astronomers can measure the orbital speed of the star, and after an entire cycle they obtain the orbital period. With this information and by using Kepler's third law as seen in Equation 2-20 the size of the orbit is determined, as well as the system's mass. In fact, as Equation 2-20 shows, the total mass $(M_1 + M_2)$ can be calculated. However, it is necessary to know the inclination of the binary system in order to deduce the masses of each star. Therefore, only a limit on the masses of each star can be calculated when using Kepler's third law (Hynes, 2010; Ozel et al., 2010).

In some cases, the orbit of the binary star will be almost edge on as seen from Earth. In such cases, as the stars orbit, one will eclipse the other by transiting between its companion and the Earth. Such systems are called eclipsing binaries (Arny, 2005). When observing such systems, we find that the light detected periodically dims at the time of eclipse because one star covers the other (see Figure 2-4). The system produces a cycle variation in light intensity known as a light curve. An interesting aspect of eclipsing binaries is that astronomers can determine the diameters of the stars from them, as well as the inclination of the system.

2.3.2 Roche-lobe overflow (RLO)

The effective gravitational potential in a binary system, also known as the Roche potential, takes into account the gravitational potential of each star and the centrifugal potential of the system since it is rotating (Tauris & van den Heuvel, 2006):

$$\Phi = -\frac{GM_1}{r_1} - \frac{GM_2}{r_2} - \frac{\Omega^2 r_3^2}{2}$$
(2-22)

where r_1 and r_2 are the distances from each star to the centre of mass, r_3 is the distance to the rotational axis of the binary and Ω is the orbital angular velocity.



Figure 2-4: Diagram of eclipsing binaries and the light curves seen in each case.

We assume that the stars revolve in circular orbits and it is therefore simple to determine fixed equipotential surfaces to the stars. Such surfaces are calculated by taking $\nabla \Phi = 0$, thus Φ a constant. Five solutions to $\nabla \Phi = 0$ are found making equilibrium points and are called the Lagrangian points. In Figure 2-5, the equipotential lines and Lagrangian points are plotted for two stars with masses $M_1 = 15 M_{\odot}$ and $M_2 = 7 M_{\odot}$. L₁, Lagrangian point 1, is in between the two stars and matter can flow freely from one star to another through this point. It also defines the 'pear-shaped' Roche-lobe (RL) of the stars, surfaces which have L₁ as a contact point. L₂ and L₃ are on opposite sides of the secondary and primary stars respectively. Finally, L₄ and L₅ are found in lobes perpendicular to the line joining the binary. It is important to note that L_{1,2,3} are unstable points, which means that a small perturbation will lead to material leaving the L-point, whereas L_{4,5} are stable (Tauris & van den Heuvel, 2006).

In this example, the donor is more massive than the accretor, which is not necessarily the usual case. Since L_1 is unstable, if the more massive star evolves to fill its Roche-lobe mass transfer will occur onto the less massive star. This process is called Roche-lobe overflow (RLO). The size of the RL of the accretor is a function of the orbital separation and the mass ratio of the binary components $q = M_{donor}$ / $M_{accretor}$, given by the following equation (Eggleton, 1983; Tauris & van den



Figure 2-5: Equipotential lines in a binary system, also indicating the five Lagrangian points of potential equilibrium (Tauris & van den Heuvel, 2006).

Heuvel, 2006):

$$\frac{R_{L_{accretor}}}{a} = \frac{0.49q^{2/3}}{0.6q^{2/3} + \ln(1+q^{1/3})}$$
(2-23)

The radius of the donor's RL is (Tauris & van den Heuvel, 2006):

$$R_{L_{donor}} = a \times (0.5 - 0.227 \log q)^4 (1+q)$$
(2-24)

Detached binaries tend to evolve as two single stars, with no effect on one another, but remain interesting in order to measure accurate masses and radii of stars. In the case of close binaries, where the binary separation is comparable to the radii of the stars, a star fills its RL either by expanding during its evolution or because the binary separation shrinks which leads to the RL shrinking as well. The latter scenario happens when orbital angular momentum of the system is lost.

There are three types of RLO, which all depend on the type of binary system

considered. These three types of RLO, cases A, B and C, were defined by Kippenhahn & Weigert in 1967 (Kippenhahn & Weigert, 1967; Tauris & van den Heuvel, 2006). In case A, the system is a contact binary so the donor star begins to fill its RL at the stage of hydrogen burning in the core. The difference between cases B and A of RLO is that in case B, the primary star starts to fill its RL after the end of hydrogen burning but before the helium ignites. Finally, in case C the overflow of the RL begins during or beyond the helium burning stage. The precise orbital period ranges for all three cases depend on the initial mass of the donor star and q. The RLO ends once the donor has lost its hydrogen-rich envelope and can no longer fill its Roche-lobe (Tauris & van den Heuvel, 2006).

2.3.2.1 The orbital angular momentum equation

The orbital angular momentum of a binary system is given by the following equation (Tauris & van den Heuvel, 2006):

$$J_{orb} = \frac{M_1 M_2}{M} \Omega a^2 \sqrt{1 - e^2}$$
 (2-25)

where $M_{1,2}$ are the masses of the donor and accretor, $M = M_1 + M_2$, a is the binary separation, and Ω is the orbital angular velocity and is equal to $\sqrt{GM/a^3}$. As mentioned above, the orbital motion of the close binary star is circular so e = 0.

In order to calculate the rate at which the orbital period changes, the last equation must be differentiated. The following is what results, which constitutes the orbital angular momentum balance equation (Tauris & van den Heuvel, 2006):

$$\frac{\dot{a}}{a} = 2\frac{\dot{J}_{orb}}{J_{orb}} - 2\frac{\dot{M}_1}{M_1} - 2\frac{\dot{M}_2}{M_2} + \frac{\dot{M}_1 + \dot{M}_2}{M}$$
(2-26)

where the total change in orbital angular momentum is given by the following equation:

$$\frac{\dot{J}_{orb}}{J_{orb}} = \frac{\dot{J}_{gwr}}{J_{orb}} + \frac{\dot{J}_{mb}}{J_{orb}} + \frac{\dot{J}_{ls}}{J_{orb}} + \frac{\dot{J}_{ml}}{J_{orb}}$$
(2-27)

The total change in orbital angular momentum (Equation 2-27) has several terms, which will be explained separately. The first one is the change in orbital angular momentum due to gravitational wave radiation (Tauris & van den Heuvel, 2006):

$$\frac{\dot{J}_{gwr}}{J_{orb}} = -\frac{32G^3}{5c^5} \frac{M_1 M_2 M}{a^5}$$
(2-28)

When the orbit of the binary is sufficiently narrow, this first term becomes dominant and will lead to the binary separation to shrink and therefore forcing the stars into contact.

The second term in Equation 2-27 appears because of magnetic braking. It is known that the rotation of low-mass stars decelerates due to magnetic stellar winds. As seen in the section on compact stars, charged particles swirl around the magnetic fields of the stars. Some stars have magnetic field lines which are open, meaning that they do not connect at both their magnetic poles. This structure leads to a deceleration of the spin of the star and in order to conserve angular momentum, the orbital period of the binary system will therefore shrink as well. When the donor star expands or contracts, it can exchange angular momentum with the orbit. This is found in the third term of the equation.

Finally, the last term of the equation describes the change in orbital angular momentum caused by mass loss from the binary system. This term is usually the dominant one in the orbital angular momentum balance equation.

2.3.2.2 Stable or unstable mass transfer?

The stability of the mass transfer depends mainly of the response on the masslosing donor and on the RL. If the mass transfer happens on a short timescale during RLO, such as thermal or dynamical, then the system is most unlikely to be observed. However, if the mass transfer takes longer and occurs during a nuclear timescale then the system can be observed as an X-ray source for a long time because in this case, the accretion rate onto the neutron star or black hole lasts long enough (Tauris & van den Heuvel, 2006).

The donor star gets out of the hydrostatic and thermal equilibrium state when it fills its Roche-lobe. However, it tries to establish the equilibrium once again by either shrinking or growing. The RL changes as well during mass transfer or loss. The mass transfer is stable as long as the RL of the donor star encloses it. If this is not the case, the mass transfer becomes unstable. This usually happens when the donor star expands or shrinks less rapidly than the binary orbit (Tauris & van den Heuvel, 2006).

2.3.3 Common envelope (CE) evolution

During the evolution of a close binary system, a very important stage is the common envelope (CE) phase (Tauris & van den Heuvel, 2006). It usually occurs when the mass transfer is unstable. The donor will fill its Roche-lobe and start mass transfer, which will lead to the orbit shrinking while the star continues to expand. Since the RL volume decreases with the orbital shrinkage and the star is still growing, the mass transfer does not stop but instead accelerates. All this causes both the orbit to shrink and the donor to expand faster, which is a very unstable mass transfer. The name of common envelope (CE) comes from the fact that the donor's envelope expands fast and engulfs the companion star.

The CE phase is usually found in the case where the donor is a giant star, with a large convective envelope, and the accretor is a compact star with a degenerate core. The two stars continue their orbital motion inside the CE. However, this is not the case for long because due to drag forces inside the envelope, the two objects lose energy and are brought closer together. The shrinking of the binary separation inside the CE is known as a spiral-in and it leads to dissipation of orbital angular momentum. The orbital velocities therefore increase but their potential energy decreases more than the kinetic one and this all has a final result of energy loss in the system. The CE evolution is often tidally unstable and all the changes such as the angular momentum transfer, the orbital energy dissipation and the structural changes of the donor happen in very short timescales (Tauris & van den Heuvel, 2006). The CE phase ends when either the envelope is expelled in space, which is often the case, or the two stars inside the envelope merge.

2.3.4 Detached, semi-detached and contact binaries

There are different types of close binary systems: detached, semi-detached and contact binaries. Figure 2-6 shows the obvious difference between each type of binary system. It is clear that in the case of a detached binary system, the stars do not 'touch'. However, with semi-detached ones, the donor star fills its Roche-lobe. Finally, the contact binaries have stars which are so close that they touch each



Figure 2-6: Different cases of close binary systems (taken from OBSERVATION AND ANALYSIS OF ECLIPSING BINARY STAR, by Derek Wallentinsen - MSc thesis, B.A., Vassar College, 1985).

other and share gaseous envelopes (Prialnik, 2000).

2.3.5 Supernova explosions in close binaries

If a star in a close binary system is massive enough, it will collapse and explode in a supernova (SN). This happens after it has consumed all its hydrogen (and possibly helium) during the RLO and/or CE evolution. A neutron star is born after a SN if the minimum mass of the helium star is about 2.8 M_{\odot}. The threshold is lower for a CO star, which is a helium star which has lost its envelope after an additional phase of mass transfer from the helium star to its companion. This critical value of 2.8 M_{\odot} corresponds to an initial mass of ~ 10 M_{\odot} for case C and ~ 12 M_{\odot} for cases A and B of RLO (Tauris & van den Heuvel, 2006). If the mass of the core is below the critical value, the star contracts after a second phase of



Figure 2-7: Different stages of the formation of an accretion disc (taken from Lecture course PX387, The University of Warwick)

RLO, and leaves behind a white dwarf. However, if the mass of the helium core is larger than 8 M_{\odot} , the SN leaves a black hole.

Many different cases can be found and obtained after the evolution of a binary system. The most common X-ray binary systems are the low-mass X-ray binaries (LMXBs) and the high-mass X-ray binaries (HMXBs). More details on those systems can be found in the following section.

2.4 Accretion discs

All of the binary systems mentioned in this thesis accrete via discs, except for HMXBs which accrete via stellar winds. Here we explain the theory behind accretion discs.

2.4.1 Disc formation

The reason why matter accreting onto the secondary star forms a disc and does not fall onto the star directly is because its specific angular momentum J is too large (King, 2003). The circularization radius, defined by the following equation, is where the matter would orbit if it lost energy but no angular momentum (King, 2003):

$$R_{circ} = \frac{J^2}{GM_{accretor}} \tag{2-29}$$

For a disc to form, the circularization radius should be larger than the effective size of the accretor. This size is equal to the radius of a non-magnetic WD or NS if the magnetic field is not dynamically significant. However, if that were the case, then the effective size of the accretor would be of the same order as the magnetospheric radius of the star. In the case of a black hole, the effective size of the accretor is the radius of the last stable circular orbit (King, 2003).

If mass transfer happens through RLO in a compact binary, the formation of an accretion disc is almost always satisfied because the specific angular momentum of the accretor is comparable to that of the binary and R_{circ} is large. The reason why it is said that in the latter case the formation of a disc is *almost* certain is because there are exceptions. For instance, it is the case of some cataclysmic variables where the WD accretor is strongly magnetic (King, 2003).

If we consider the usual case where an accretion disc can be formed, it occurs when energy of the system is lost through dissipation faster than angular momentum is redistributed. For a given angular momentum, the lowest energy for which matter can orbit is circular, therefore, material falling onto the accretor will form a sequence of circular orbits around it. Angular momentum and energy loss are both explained by an additional force, called viscosity.

2.4.2 Thin discs

The initial ring is spread into a disc at R_{circ} from the accretor thanks to the viscosity which helps transport angular momentum. The efficiency with which

the disc can cool should be high enough for it to be considered thin. For a disc to be thin, its scaleheight H should follow the condition (King, 2003):

$$H \simeq \frac{c_s}{v_K} R \ll R \tag{2-30}$$

where R is the radius of the disc, c_s is the local sound speed and v_K is the Kepler velocity defined by (King, 2003):

$$v_K = \left(\frac{GM_{accretor}}{R}\right)^{1/2} \tag{2-31}$$

The conditions of having a thin and efficiently cool disc with a Keplerian velocity are all equivalent and related. The disc can not be thin without having the two other conditions fulfilled as well.

In the case of a thin disc, the vertical and horizontal structures of the disc are decoupled. The vertical structure is almost hydrostatic, whereas the horizontal one can be described by the surface density Σ . If the disc is axisymmetric, Σ obeys a nonlinear diffusion equation because angular momentum and mass are conserved (King, 2003):

$$\frac{\partial \Sigma}{\partial t} = \frac{3}{R} \frac{\partial}{\partial R} \left(R^{1/2} \frac{\partial}{\partial R} [\nu \Sigma R^{1/2}] \right)$$
(2-32)

where ν is the kinematic viscosity and is defined by:

$$\nu = \alpha c_s H \tag{2-33}$$

where α is a dimensionless number. This α ansatz for the viscosity has been a major issue in the understanding of the flow of material in accretion discs. Until today, astronomers do not know how big α is. All we know is that $\alpha < 1$. Shakura & Sunyaev 1973 have the fifth most cited astrophysics paper, directly related to the α Ansatz for the viscosity.

2.4.3 Alfvén radius

The Alfvén radius corresponds to the radius at which the pressure due to the star's magnetic fields is equal to the ram pressure of infalling material. The latter is the

pressure exerted on a body which is moving through a fluid medium. The Alfvén radius is also known as the stopping radius because at this radius, the electromagnetic field of the accreting star will stop the accretion process. The following equations describe the different types of pressure in the case of an accretion disc (Carroll & Ostlie, 2006):

Magnetic pressure:
$$P_m = \frac{B^2}{2\mu_0}$$

Gas pressure: $P_g = NkT$
Ram pressure: $P_{ram} = \rho v^2$
(2-34)

where μ_0 is the permeability of free space and is equal to $4\pi \times 10^{-7}$ NA⁻², k is the Boltzmann constant and is equal to 1.38065×10^{-23} J K⁻¹, T and N are the temperature and the particles density of the gas and v is the velocity of the material.

The Alfvén radius, denoted r_A , corresponds to the radius for which $P_m = P_{ram}$. The magnetic field of the star is (Carroll & Ostlie, 2006):

$$B = \frac{\mu}{r^3} \tag{2-35}$$

where r is the distance to the magnetic field origin and μ is the magnetic moment of the star.

In the case of a spherical accretion, we can write the accretion rate as:

$$\dot{M} = 4\pi r^2 \rho v$$
 therefore $\rho v = \frac{M}{4\pi r^2}$ (2-36)

The velocity v of the material in this case is the escape velocity:

$$v = \sqrt{\frac{2GM}{r}} \tag{2-37}$$

Therefore by considering $P_m = P_{ram}$, we find the Alfvén radius to be equal to:

$$r_A = \left(\frac{2\pi^2 \mu^4}{\mu_0^2 GM\dot{M}}\right)^{1/7} \tag{2-38}$$

2.4.4 Eddington Limit

In the case of typical LMXBs and HMXBs, which have a neutron star accreting at a rate of $10^{-11} - 10^{-8} \text{ M}_{\odot} \text{ yr}^{-1}$, the X-ray luminosities are found in the range of $10^{35} - 10^{38} \text{ erg s}^{-1}$ (Tauris & van den Heuvel, 2006). However, for all accreting binary systems there is a maximum accretion rate at which the gravitational force of the compact star is equal to the radiation pressure force of the accreting matter. This limit is known as the Eddington limit and in the case of a spherical accretion of a hydrogen-rich gas, the mass transfer rate at the limit is $\dot{M}_{Edd} = 1.5 \times 10^{-8}$ $M_{\odot} \text{ yr}^{-1}$. However, it is important to note that the Eddington limit depends on the mass of the accreting compact star.

When material falls onto the compact object, more photons will be produced which will push the infalling material and prevent further accretion. The radiation pressure force can be written as in the following equation (Tauris & van den Heuvel, 2006):

$$F_{rad} = \frac{\kappa F}{c} \tag{2-39}$$

where κ is the absorption coefficient, F is the photon flux and c is the speed of light in vaccum.

The gravitational force is:

$$F_{grav} = \frac{GMm}{R^2} \tag{2-40}$$

where M and R are the mass and radius of the accreting object, m is the mass of the material in the disc and G is the gravity constant.

Considering that plasma is fully ionised, only the Thomson scattering will be taken in effect:

$$F = \frac{L}{4\pi R^2} \tag{2-41}$$

$$F_{rad} = \frac{\sigma_T L}{4\pi R^2 c} \tag{2-42}$$

where σ_T is the Thomson coefficient and is equal to $6.65 \times 10^{-29} \text{ m}^2$, which is the total area a photon can travel before being affected by an electron and L is the luminosity.

In the Eddington limit, $F_{rad} = F_{grav}$, therefore the maximum luminosity that an accreting object can have is the Eddington luminosity (Tauris & van den Heuvel,

2006):

$$L_E = \frac{4\pi G M m c}{\sigma_T} \tag{2-43}$$

Above the Eddington limit, accretion onto the compact object can no longer occur. However, if the accretion rate surpasses the Eddington limit, the excess matter will not form a disc around the compact star but instead will form a cloud which is optically thick to X-rays, and it will therefore be impossible to see the source. Since the Eddington limit seen previously was of $\sim 10^{-8} M_{\odot} \text{ yr}^{-1}$, for LMXBs and HMXBs to be observed, they must have accretion rates of $10^{-11} - 10^{-8} M_{\odot} \text{ yr}^{-1}$ (Tauris & van den Heuvel, 2006). This does not mean that much larger accretion rates can not be found in binary systems. In cases where the mass transfer rate largely exceeds the Eddington limit, the excess material is ejected in a jet.

2.5 X-ray binary systems

Many types of X-ray binary systems have been found. However, the main ones described in this thesis are the low-mass and high-mass X-ray binaries. Binary systems which have a low-mass companion to a neutron star or black hole are low-mass binaries, whereas the systems with high-mass companions to such compact objects are high-mass binary systems (Lewin & van der Klis, 2006).

Before describing the different types of X-ray binaries, it is important to understand the evolution of helium stars in binary systems. In Section 2.1.3, we described the evolution of helium stars in the case of a single star; the outcome is different however, in the case of a binary system.

2.5.1 Helium stars in binary systems

The case of the evolution of helium stars differs from single stars to stars found in binaries only when the helium star's mass is larger than 2.3 M_{\odot} . Naked helium stars are stars which have lost their hydrogen envelope via mass transfer in a close binary system. When helium stars are no longer ongoing RLO in binaries, their cores have tiny hydrogen envelopes with masses less than 0.01 M_{\odot} (Tauris & van den Heuvel, 2006). This is important to realise because it has effective consequences on their radial evolution.

More massive helium stars, also known as Wolf-Rayet stars, should also be considered in the evolution of binaries. It is still not clear how fast these stars transfer mass to their companions by an intensive stellar wind. This uncertainty also affects the determination of the minimal mass required of a core to collapse into a black hole.

Two cases must be considered in order to understand the evolution of helium stars in binaries. On the one hand, if the helium star is naked then it does not lose much mass through stellar wind and can eventually terminate as a black hole after undergoing all burning stages. In the case of single star evolution, this can occur if the helium star has an initial mass > 19 M_{\odot} (Tauris & van den Heuvel, 2006). On the other hand, the helium star must be embedded in its hydrogen envelope for as long as possible to be able to form a black hole in the case of a close binary. This is possible if the binary starts from a wide Case C binary evolution. The CE phase will shrink the binary separation and leave a binary system consisting of an evolved helium core and a low-mass companion. The helium core then collapses and explodes in a SN, leaving behind a low-mass X-ray binary. Once the lowmass companion fills its RL, such systems are called soft X-ray transients (SXTs) (Tauris & van den Heuvel, 2006).

2.5.2 Cataclysmic variables

In binary systems, white dwarfs evolve differently to when they are isolated. A binary system containing a WD accreting from a MS star is known as a cataclysmic variable (CV). When in a semi-detached binary system, the WD could lead the system to evolve into a dwarf nova, a classical nova or a supernova. In the first two end stages of a semi-detached binary system containing a WD, the outburst process recurs, unlike in supernovae where the WD is completely disrupted.

In the case of CVs, the typical mass of the WD is 0.85 M_{\odot} , which is larger than the typical mass of a single WD (0.6 M_{\odot}) (Carroll & Ostlie, 2006). Similarly to all binary systems with accretion discs, the light detected comes from the disc around the compact star. In the case of CVs, some WDs have magnetic fields that are strong enough to prevent the formation of an accretion disc, instead the material follows the magnetic lines, falling onto the magnetic poles of the WD.

In a dwarf nova, the 'outburst' begins in the outer, cooler part of the disc and then spreads to the hotter, inner regions. In fact, outbursts are caused by a sudden increase in the rate at which mass flows down through the accretion disc. According to theoretical models and observations of dwarf novae, the mass transfer rate through the disc, during a long quiescent phase, has been estimated to (Carroll & Ostlie, 2006):

$$\dot{M} \approx 10^{15} - 10^{16} g s^{-1} \approx 10^{-11} - 10^{-10} M_{\odot} y r^{-1}$$
 (2-44)

However, during an outburst, it increases to:

$$\dot{M} \approx 10^{17} - 10^{18} g s^{-1} \approx 10^{-9} - 10^{-8} M_{\odot} y r^{-1}$$
 (2-45)

As seen previously, the luminosity of the disc is proportionate to the mass transfer rate:

$$L_{disc} = G \frac{MM}{2R} \tag{2-46}$$

where R is the radius of the disc.

What astronomers are yet to solve is the mystery of the origin of the increase of the mass transfer rate. There are two theories to why it occurs: either the mass transfer from the secondary to the primary stars becomes unstable, or the accretion disc itself becomes unstable. The latter one is more popular and supported by observations. In this case, the instability is believed to be caused by the hydrogen partial ionisation zone at a temperature of 10,000 K (Carroll & Ostlie, 2006).

The rate at which the material spirals down in the accretion disc depends on its viscosity. If the latter is small then the resistance to the orbital motion of the gases is also low. This leads to a decrease of the inward drift of the material and finally more material in the disc. If the viscosity periodically changes, then it could explain the brightening of the disc in the case of a dwarf nova. The outer part of the disc having a temperature of $\sim 10,000$ K could produce a periodic ionisation

and recombination of hydrogen, which then leads to a periodic switch between low and high viscosity in the disc. When temperatures reach 10,000K or more, the gas is mainly composed of ionised hydrogen which has high opacity, temperature and viscosity, with inefficient cooling and mass released to fall through the disc. When the matter is accumulated in the disc, it slowly heats its outer layer. However, when the matter is released, it leads to a rapid cooling of the disc. This switch and difference in forms of cooling and heating the disc produces its instability. However, this process only occurs for low accretion rates (less than $10^{-11} \text{ M}_{\odot} \text{ yr}^{-1}$) and therefore limit the rate for dwarf novae. This limit is observed, which explains why the theory of the instability of the disc is preferred (Carroll & Ostlie, 2006).

In the case of a WD in a semidetached binary system, the hydrogen-rich gas accumulates on its surface, where it is compressed and heated. At the base of the layer, the gas is mixed with the carbon, nitrogen and oxygen of the WD, where the material is supported by electron degenerate pressure. When the WD has accreted enough hydrogen on its surface for temperatures to reach several million kelvins, a shell of hydrogen-burning ignites, using the CNO cycle (Carroll & Ostlie, 2006). In the case of degenerate matter, the pressure does not depend on the temperature, therefore the reaction rate can not be reduced by expansion and cooling of the shell. This leads to a set of thermonuclear reactions, with temperatures reaching 10^8 K, where electrons lose their degeneracy. When the luminosity surpasses the Eddington limit, the accreted matter is expelled into space by the radiation pressure.

Only about 10% of the hydrogen layer is ejected by the explosion (Carroll & Ostlie, 2006). A hydrostatic hydrogen burning phase is then established and the layer above the shell of CNO burning becomes fully convective and expands by a factor 10 to 100. A months after the hydrogen burning phase begins, the remainder of the accreted surface layer is ejected. The WD begins to cool because it runs out of fuel so the hydrogen burning phase stops. Finally, the binary system return to its quiescent phase and the accretion process begins all over again.

Another type of binary system containing an accreting WD can evolve into a Type Ia supernova. This occurs when a WD in a close binary system accretes enough material from the secondary star to surpass the Chandrasekhar limit. It is important to note that the main difference between Type I and Type II supernovae is found in their spectra near the time of maximum brilliance. When hydrogen spectral lines are found, they are classified as a Type II. Naturally, supernovae which do not have prominent hydrogen lines in their spectra are classified as Type I. The Type I with a strong Si II line at 6150 Å is called a Type Ia. The other Type I supernovae are known as Type Ib or Type Ic, depending on the presence (Ib) or absence (Ic) of strong helium lines (Carroll & Ostlie, 2006).

Type I supernovae have no remnant star because it completely disrupts the white dwarf, unlike a Type II. The latter leave behind a neutron star or black hole and are produced from the collapse of a massive star's iron core.

2.5.3 Low-mass X-ray binaries (LMXBs)

The formation of low-mass X-ray binaries (LMXBs) can also evolve into a millisecond pulsar system. The evolution of both those systems is shown in Figure 2-8 (Tauris & van den Heuvel, 2006). An example of a LMXB system begins with a binary system consisting of a giant star of 15 M_{\odot} and a MS star of about 1.6 M_{\odot} . The orbital period of the initial system is large. The giant star evolves to fill its RL and begins mass transfer. 13.9 Myr later, the system reaches the CE and spiral-in phase described in Section 2.3.3. At the end of the CE phase, the donor star is not massive enough to collapse and explode in a SN. On the contrary, it leaves a helium star orbiting the MS star which is not affected by the previous evolutionary stages of the system. However, as seen previously, the orbital period decreases and binary separation shrinks after the CE phase. About 1Myr later, the helium star evolves, collapses and explodes into a SN leaving behind a neutron star (NS) of about 1.3 M_{\odot} , and the same MS star. The latter being the more massive one and still not having begun helium burning throughout the previous stages, it begins to fills its Roche-lobe and mass-transfer occurs. The NS rotates rapidly while accreting form the evolving MS star. The accretion process spins-up the NS star. The MS star goes through a phase of RLO and finally leaves behind a white dwarf. The final outcome of the evolution of such a system is a millisecond pulsar because the NS acquired enough momentum during accretion to continue to rotate rapidly, orbiting around a white dwarf (Tauris & van den Heuvel, 2006). The X-rays detected in this phase come from the millisecond pulsar, which will slow down due to the absence of accretion which accelerates the rotation of the star.

There are different types of binary millisecond pulsars (BMSPs), all depending on the companion type and orbital period. There are wide-orbit BMSPs with low-mass helium WD companions, which have typical $P_{orb} > 20$ days (Tauris & van den Heuvel, 2006). Moreover, we find close-orbit BMSPs which can have either low-mass helium WD companions or relatively heavy WD companions. The close-orbit BMSPs have typical orbital periods of less than 15 days. The BMSPs with a heavier WD companion did not evolve from a LMXB but rather from an intermediate-mass X-ray binary (IMXB) system, which has a donor star with a mass ranging from 2 to 8 M_☉. Wide-orbit BMSPs are initially formed in LMXBs with $P_{orb} > 2$ days and where the mass transfer happens because the donor star evolves into a giant by the loss of orbital angular momentum. However, in LMXBs with $P_{orb} < 2$ days, the mass transfer is directly conducted by loss of angular momentum due to magnetic braking and gravitational wave radiation (Tauris & van den Heuvel, 2006). In this case, a close-orbit BMSP is formed.

The evolution of close-orbit BMSPs with low-mass helium WD companions can lead to single millisecond pulsars (MSPs). In such cases, it is believed that the companion has been destroyed. Two scenarios could explain such a consequence: either the WD is irradiated by the X-rays during accretion of the NS, or it is destroyed in the form of a pulsar radiation of relativistic particles (Tauris & van den Heuvel, 2006).

There is an orbital period division of LMXBs which can lead to different types of systems. This bifurcation period depends mainly on the strength of the magnetic



Figure 2-8: Example of the evolution of a LMXB (Tauris & van den Heuvel, 2006).

braking torque and is of about 2-3 days. On the one hand, a converging binary system can be formed because the orbital period decreases until the donor becomes degenerate and the system becomes an ultracompact binary. On the other hand, when the orbital period increases, so does the binary separation, which leads to a diverging system. In this case, the orbital period increases until the donor has lost its envelope and the binary separation grows enough to obtain a wide detached binary system (Tauris & van den Heuvel, 2006).

Besides the example of the evolution a LMXB given by Tauris & van den Heuvel (2006), there are cases with LMXBs consisting of a black hole accreting from a non-degenerate secondary star (Remillard & McClintock, 2006). There are 23 confirmed stellar black holes in X-ray binaries (Remillard & McClintock, 2006; Ozel et al., 2010), out of which 8 are transient systems with long orbital periods

(larger than one day), 9 are transient systems with shorter periods and 6 systems have persistent X-ray sources with massive O/B-type secondary stars (Ozel et al., 2010). The latter are classified as HMXBs. Similarly to dwarf novae, transient X-ray sources are sources which appear in the sky for a short time and then disappear. They eventually re-appear after years or decades. However, the source does not really disappear, but only the X-ray emission coming from it. Persistent X-ray sources are objects which have continuous X-ray emission. Besides those 23 known black hole binaries, there are 32 transient black hole candidates (Ozel et al., 2010).

Most black hole X-ray binaries are discovered when they first go into outburst and are then monitored daily by satellites and wide-filed X-ray cameras (Remillard & McClintock, 2006). X-ray outbursts which last between 20 days and many months are usually explained by an instability which occurs in the accretion disc. The same explanation was initially used to explain outbursts in the case of dwarf novae (see Section 2.5.2). The outbursts occur when the accretion rate from the donor is not high enough to maintain a continuous viscous flow to the compact object, which leads to the material accumulating in the outer layer of the disc. The outburst is triggered when a critical surface density is reached. This model shows that outbursts should be recurrent, which is observed in the case of half of the known black hole binaries (Remillard & McClintock, 2006).

2.5.4 High-mass X-ray binaries (HMXBs)

A main difference between LMXBs and HMXBs is not only the fact that the companion star has a high mass in the case of HMXBs, but also the mass-transfer process. In fact, in the case of HMXBs, the compact object often accretes matter from the companion star in the form of a stellar wind (Lewin & van der Klis, 2006). The donor does not need to fill its Roche-lobe in order to transfer mass to the compact star. The companion star in this case has to be massive, with M > 10 M_{\odot}, in order to drive a strong stellar wind (Tauris & van den Heuvel, 2006).

In Figure 2-9, the evolution of a binary system leading to a HMXB is shown. The

formation of a HMXB require two massive stars at their Zero Age Main Sequence (ZAMS), which is their age at birth (once they reach the MS stage). The typical condition of the mass of the stars in a binary leading to a HMXB is to have both masses larger than 12 M_{\odot} . However, if the mass of the secondary ZAMS star is smaller than the threshold mentioned, as seen in Figure 2-9, the system can still evolve into a HMXB. In fact, the secondary ZAMS star needs to be massive enough at birth and also should be able to gain enough material during the mass-transfer process from the primary star to eventually explode in a SN-like the primary star (Tauris & van den Heuvel, 2006).

The first mass transfer stage of the binary evolution of such a system is usually stable. Of course, there are always exceptions: if the mass ratio at the beginning of the RLO phase is too extreme, then the mass-transfer process is unstable. The secondary star accretes material from the primary, which after mass-transfer and evolution becomes a helium star, whereas the secondary star becomes more massive. A first SN explosion occurs after 15Myr, when the helium star evolves. The SN remnant is a neutron star, in a binary system with a massive star. Once the latter has evolved to fill its RL 10 Myr later, the system is a HMXB. What initially was the secondary star becomes the primary and vice-versa. The binary enters CE evolution, followed by the spiral-in phase and the massive star shrinks and loses mass to become a helium star. The latter begins to fills its Roche-lobe and transfers mass to the accreting neutron star. A second SN occurs after the evolution of the helium star. Finally the binary system is left with two pulsars, orbiting one another (Tauris & van den Heuvel, 2006).

It must be pointed out that there are still large uncertainties about the physical conditions which distinguish between the formation of a neutron star or a black hole in such binary systems. The core mass may not be the only effect on the outcome of a SN in these systems, but also the magnetic field and the spin of the collapsing core (Tauris & van den Heuvel, 2006).



Figure 2-9: Example of the evolution of a HMXB (Tauris & van den Heuvel, 2006).

2.6 Observational properties of X-ray binaries

Most of the X-ray sources in our Milky Way can either be a HMXB or a LMXB, without forgetting the nearby single stars and CVs. These groups are different in many ways and their major physical characteristics will be described in this section. Binary pulsars and single MSPs are part of the ultimate fate of an Xray binary system containing an accreting neutron star but are no longer X-ray sources. Table 2-2 shows the general differences between HMXBs and LMXBs. It is important to note that all numbers of systems discovered given in this section are not exact to date. New X-ray binaries are discovered at a quicker rate than their catalogues are being updated. The most recent one is Liu et al. (2007).

2.6.1 Cataclysmic variable

Cataclysmic variables (CVs) are characterised by long quiescent intervals punctuated by outburts in which the brightness of the system increases by a factor between 10 (for dwarf novae) and 10^6 (for classical novae). The mass of the WD is typically 0.85 M_{\odot} and the secondary star, the accretor, is usually a MS star of spectral type G or later and is less massive than the primary star. The period of the binary system ranges from 54 min to 16 hours (Carroll & Ostlie, 2006), even though most of them have periods of less than 8 hours. An interesting observational fact of CVs is that out of the ~ 800 known systems, few of them have periods between 2.25 and 2.83 hours, known as the 'period gap' (Ritter & Kolb, 1992; Carroll & Ostlie, 2006; Gänsicke, 2005).

There are over 300 known dwarf novae. Their outbursts last from about 5 to 20 days, during which their brightness increases by 2 to 6 magnitudes. These small explosions usually reccur every 30 to 300 days.

A nova is characterised by a sudden increase in brightness between 7 and 20 magnitudes. Before reaching its maximum brightness by 2 magnitudes, the rise in luminosity pauses. Once it's reached its peak, a nova could remain extremely bright for ~ 100 days (Carroll & Ostlie, 2006). Depending on the speed class of the nova, meaning whether it is a fast or slow nova, the decline in brightness occurs more slowly and over several months.

The spectra of novae show that outbursts are followed by the ejection of 10^{-5} - 10^{-4} M_{\odot} of hot gases at velocities that can reach several thousand km s⁻¹. The absolute visual magnitude of a nova in its quiescent state is $M_V = 4.5$ (Carroll & Ostlie, 2006).

Classical novae have been observed many times in the X-ray regime. They are weak X-ray emitters, with a strength of 10^{33} - 10^{34} ergs/s, which is a lot weaker than the Eddington luminosity of 10^{38} ergs/s for a 1 M_{\odot} WD. Harder X-ray systems are also detected because they are close enough for sensitive observations. In this

case, the accretion onto the WD is still on-going and explains why the X-rays are present and stronger.

2.6.2 High-mass X-ray binaries

There are over 130 known HMXBs, located in the Galactic Plane, and at least 30% of these systems have well-known orbital periods ranging from 2 to 581 days (Carroll & Ostlie, 2006). In the case of pulsating NS in HMXBs, the entire range of pulse periods starts from 0.069 seconds to 20 minutes. The companion star in HMXBs has special physical features as well. They are found to have optical luminosities larger than $10^5 L_{\odot}$, and their spectral types show that they were born with masses larger than $20 M_{\odot}$. They also have radii ranging from 10 to 30 R_{\odot}, which consequently nearly fill their Roche-lobes (Lewin & van der Klis, 2006). By studying the X-ray spectra of such binary systems, many absorption/emission lines give information on the magnetic fields of the stars. The strength of the fields is B $\simeq 5 \times 10^{12}$ G (Lewin & van der Klis, 2006), even though it can still vary. Two famous HMXBs have an accreting back hole: Cyg X-1 and LMC X-3.

Another class of HMXBs is the Be-star X-ray binaries. They have less eccentric orbits, with orbital periods ranging from 20 to 250 days. In these systems, the companions are rapidly rotating B-emission stars, either on or very close to the MS. Their luminosity class ranges from III to V and the companions have masses ranging from 8 to 20 M_{\odot}. They are also the most numerous class of HMXBs. The X-ray emission from the Be star is not constant; it can go from being completely absent to undergoing huge transient outbursts which could last up to weeks or even months. These outbursts are probably related to the irregular optical outbursts seen in Be-stars. All Be-star binaries are transient sources, therefore they are usually not detected for months or years, whereas the typical HMXB system is a persistent X-ray source (Tauris & van den Heuvel, 2006).

2.6.3 Low-mass X-ray binaries

Over 187 of these systems have been detected in our Galaxy, the Large Magenallic Cloud and the Small Magenallic Cloud (Liu et al., 2007). Their orbital periods are similar to those of cataclysmic variables and range from 11 minutes to 17 days. It is very difficult to obtain the spectrum of the companions in these systems because LMXBs are not wide binaries, however the spectrum usually observed is that of the accretion disc. Their magnetic fields are relatively weak, ranging from 10⁹ to 10¹¹ G, but if their magnetic field strength is larger than 10¹¹ G, X-ray bursts are seen in these systems (Tauris & van den Heuvel, 2006). These bursts happen because of sudden thermonuclear fusion of accreted matter at the surface of the neutron star.

Quasi-periodic oscillations (QPOs) in the X-ray flux of LMXBs have helped astronomers understand these systems. Precise timing signature of the accreting NS or BH has been possible thanks to the discovery of QPOs. Therefore, by observing and studying such systems in detail, it was possible to provide physical evidence of general relativity theory (Tauris & van den Heuvel, 2006).

Most of the LMXBs are found in the Galactic bulge and in globular clusters. About 0.6 kpc away from the Galactic Plane, they have transverse velocities of over 100 km s⁻¹, whereas in globular clusters their velocities are less than 30 km s⁻¹, the escape velocity from the cluster.

2.6.4 Intermediate-mass X-ray binaries (IMXBs)

Intermediate-mass X-ray binaries have companions with masses ranging from 1 to 10 M_{\odot} (Tauris & van den Heuvel, 2006). These systems are very difficult to find for many reasons. Unlike HMXBs, the companion is not massive enough to transfer mass to the neutron star or black hole through stellar wind. Therefore, the system evolves though RLO. Since the companion star is not massive, the mass ratio between the companion and compact object is large enough to imply that the RLO phase will be short and of about a few thousand years. It is either

	HMXBs	LMXBs
X-ray Spectra	$kT \ge 15 \text{ keV} \text{ (hard)}$	$kT \le 10 \text{ keV} \text{ (soft)}$
Type of time variability	regular X-ray pulsations	only a very few pulsars
	no X-ray bursts	often X-ray bursts
Accretion process	stellar wind (or atmospheric RLO)	Roche-love overflow
Timescale of accretion	$10^5 { m yr}$	10^{7} - 10^{9} yr
Accreting compact star	high-magnetic field NS (or BH)	low-magnetic field NS (or BH)
Spatial distribution	Galactic plane	Galactic centre
		and around the plane
Stellar population	young, age $< 10^7$ yr	old, age $> 10^9 \text{ yr}$
Companion stars	luminous, $L_{opt} / L_x > 1$	faint, $L_{opt} / L_x \ll 0.1$
	early-type $O(B)$ stars	blue optical counterparts
	$> 10 M_{\odot}$ (Pop. I)	$\leq 1 \ M_{\odot}$ (Pop. I and II)

Table 2-2: Observational properties of HMXBs and LMXBs (Tauris & van den Heuvel,2006)

on a timescale shorter than the thermal one, or a CE phase occurs.

The mass transfer rate in these systems is much larger than the Eddington limit and could therefore cause the X-rays to be absorbed in the dense cloud around the accreting star. Theory predicts the existence of IMXBs because they are the only systems able to form binary pulsars with heavy CO (carbon, oxygen) or O-Ne-Mg (oxygen, neon, magnesium) WD companions (Tauris & van den Heuvel, 2006). A few IMXBs with NS have been detected, for instance Her X-1 and Cyg X-2. There are also IMXBs with accreting black holes, such as GRO J1655-40, 4U1543-47, LMC X-3 and V4642 Sgr. In these systems, the mass of the companion is smaller than that of the black hole, therefore mass transfer through RLO is stable.

2.6.5 Soft X-ray transients

Soft X-ray transients (SXTs) are binaries with a black hole and a low-mass donor star (usually a K dwarf). They form a class of LMXBs. They are the equivalent of dwarf novae for LMXBs, with recurrent outbursts instead of having a persistent luminosity. They have very strong X-ray luminosities, of about $L_x \sim 10^{38}$ erg s⁻¹, for several weeks (Tauris & van den Heuvel, 2006). During this period, they are also very bright in optical wavelengths. In fact, their optical magnitudes decrease by 6 to 10 magnitudes, making them optical and X-ray novae. Once that phase of
bright X-rays and optical emission is over, it is possible to obtain the spectrum of the donor, usually a K or G star. When the mass of the compact object is larger than 3 M_{\odot} , this indicates the presence of a BH since it is above the limit of the mass of a NS. Such systems have orbital periods ranging between 8 hours to 6.5 days (Tauris & van den Heuvel, 2006).

Chapter 3

Galactic Plane & Bulge surveys

In this chapter, we will introduce the different astronomical tools and databases used to detect the systems described in the previous chapter, where we outlined their physical properties. The beginning of this chapter will explain the imaging methods and how to exploit the data obtained from surveys.

Astronomical surveys use different filters and therefore provide different kind of information on the sources detected. Here, we will describe each astronomical survey of the Galactic Plane and Bulge used for both projects.

3.1 Astronomical imaging and methods

Nowadays, most astronomers use multi-wavelength surveys to understand and classify stellar populations. These surveys provide information on stellar objects, such as their brightness, as well as digital images in different wavebands. Astronomers measure the brightness of a star in magnitudes, derived from the brightness of a star on a logarithmic scale, measured in a specific passband by the use of filters.

It is important to understand the difference between the flux and the luminosity of a star. Stellar flux is the power received per unit area, therefore it is the number of photons per unit area and per second. Therefore it is written in $J m^{-2} s^{-1}$. The flux of a star at its surface depends only on its effective temperature, according to the Stefan-Boltzmann Law:

$$F = \sigma T^4 \tag{3-1}$$

where σ is the Stefan-Boltzmann constant and is equal to 5.67 $\times 10^{-8}$ W m⁻² K⁻⁴.

However, the flux of a star is not the true measure of its energy output. In fact, the total energy output per second of a star is its luminosity. As seen in Equation 2-8, it depends on the flux and surface area of the star.

As the photons leave the star, they travel towards the observer, spreading out and becoming less concentrated as they get farther away from the star. Therefore a star appears to be dimmer the farther away it is. The flux of a star is proportionate to its luminosity divided by the distance squared ($f \propto L / d^2$).

There are several types of magnitudes, such as apparent, absolute and bolometric magnitudes. The apparent magnitude is the brightness of a star with respect to its observed flux:

$$m_1 - m_2 = -2.5 \times \log\left(\frac{f_1}{f_2}\right) \tag{3-2}$$

Since it is a relative definition, meaning a star's magnitude is always calculated with respect to another known star, astronomers had to set a zero-point magnitude, the star Vega, also known as α Lyr. Therefore, $m_{\text{Vega}} = 0$ and the apparent magnitude can therefore be calculated by:

$$m_{\rm star} = -2.5 \times \log\left(\frac{F_{\rm star}}{F_{\rm Vega}}\right) = 2.5 \times (\log F_{\rm Vega} - \log F_{\rm star})$$
 (3-3)

where F_{Vega} and F_{star} are the respective fluxes of Vega and the star in consideration.

The absolute magnitude, M, is directly related to a star's luminosity and is what the apparent magnitude would be if the star were at 10 parsecs (32.6 light-years) from Earth:

$$M = m - 5(\log d - 1) \tag{3-4}$$

where m is the apparent magnitude of the star and d is the distance to the star in parsecs.

The bolometric magnitude takes into account the energy radiated at all wavelengths. There are two important facts about the magnitude system which should be noted. Firstly, the scale is 'backward' in the sense that the brighter the star, the smaller the magnitude. Secondly, the difference between magnitudes of two stars corresponds to the ratio of their fluxes. Hence, the magnitude scales are



Figure 3-1: Spectrum of the Sun (continuum) as well as the spectrum of a black body (dotted line) with the same temperature as the Sun (Figure taken from lecture course PX268, The University of Warwick).

logarithmic units and one magnitude difference is equal to a brightness variation of about 2.512 times (the 5th root of 100).

Astronomers use CCD (charge-coupled device) detectors, attached to big telescopes, in order to image the sky. Before taking the images, a specific filter is chosen. Each filter is designed to transmit only light within a specific wavelength range. A star's spectrum is the plot of its light intensity as a function of wavelength.

Wien's displacement law:

$$\lambda_{\rm max} \times T_{\rm Wien} = 2.897 \times 10^{-3} {\rm m.K} \tag{3-5}$$

where λ_{max} is the peak wavelength of the spectrum of the star. This law considers that a star is a blackbody. In Figure 3-1, we show the spectrum of the Sun with the same spectrum of a blackbody radiating at the same effective temperature as the Sun. It is important to mention that the effective temperature of the Sun is not calculated using Wien's law but rather the definition of the luminosity of a star. It is the temperature that the star would have it it were a blackbody, but no star emits exactly like a blackbody.

There are different types of stellar spectra found, because stars undergo stellar evolution and at each stage of their lives many of their properties change as well (temperature, radius, luminosity, chemical composition, etc). Along the main sequence, the main reason for different spectral appearence is the mass of the star, which translates into its effective temperature. In the spectra of stars, we often find absorption lines which are produced when a continuum travels through a cooler gas before reaching the observer (usually the star's atmosphere). Photons are therefore absorbed by the atoms in the gas and result in a dark feature in the spectrum of the star. Because the temperature in the stellar atmosphere decreases outwards, cooler gas lies in front of the hotter gas, which leads to more absorption lines seen in a typical stellar spectrum. In the absence of a cooler gas in between the hot core and the observer, more emission lines will appear. However, it is interesting to mention that on top of the atmosphere of the star, there is a very low-density inversion layer, the chromosphere, which produces emission lines in the ultraviolet and X-ray regions of the electromagnetic spectrum.

As will be seen in this chapter, astronomical surveys use different filters in order to integrate the total light obtained from stars at different wavelengths. Some stars are hotter than others, therefore their spectra peak at shorter wavelengths.

The colour of a star in astronomy is the difference in magnitude between two filters. A star's colour gives information on its average temperature. All objects release thermal radiation, which is generated when heat from all the movement of electrons and protons is converted to electromagnetic radiation. Thermal radiation from stars is approximated to blackbody radiation, which is a temperature-dependant spectrum of light. As the atoms heat up, they increase the thermal radiation released, thus changing the peak wavelength of that energy changes. When the temperature of the star's surface increases, the star's spectrum peaks at lower wavelengths. All of these properties of stars can be deduced from data obtained



Figure 3-2: UBVRI Johnson filters plotted on top of the spectra of two stars. The I filter is not seen in the wavelength range of this plot but it is one of the Johnson filters. The blue star is the hot one, whereas the red one is the cool star (Figure taken from lecture course PX268, The University of Warwick).

from astronomical survey, thus the importance of their use.

As seen previously, the measure of fluxes is indicated in magnitudes. Therefore, by subtracting the magnitudes of the star obtained in different passbands, astronomers find out their properties. If the star is 'hot', it will emit more light in shorter wavelengths and therefore its magnitudes will be smaller in bluer filters. As can be seen in Figure 3-3, if we observe through the *B* filter, which is bluer than the *V* one, the blue star will be brighter. The opposite result can be found, where the cool star is brighter than the hot one, when observing them with the *V* filter. If we call the hot star Star₁ and the red one Star₂, and B_1 and V_1 the magnitudes of Star₁ through the *B* and *V* filters respectively, then we will find that Star₁ is blue because $B_1 - V_1 < 0$. $B_1 - V_1$ is what astronomers call the colour of the star in this case. Remember that the magnitudes of stars are smaller for brighter sources, therefore $B_1 < B_2$, whereas $V_1 > V_2$.



Figure 3-3: Integrated light of what would be detected and measured with the B and V Johnson filters in the case of the two stars considered in Figure 3-2 (Figure taken from lecture course PX268, The University of Warwick).

Table $3-1$:	Johnson/Bessel	central	wavelengths	of	UBVRI f	filters
---------------	----------------	---------	-------------	----	---------	---------

	U	B	V	R	Ι
	(Ultraviolet)	(Blue)	(Visible)	(Red)	(Near-Infrared)
Central wavelength (nm)	365	440	550	630	900

3.2 The Galactic Bulge Survey (GBS)

X-ray sources have been hard to identify at longer wavelengths due to the lack of optical and near-infrared observations in the Galactic center. This is mainly explained by the fact that the region concerned suffers from high extinction and crowding, which leads to a large uncertainty in the source position.

The Galactic Bulge Survey combines sensitivity for faint X-ray sources, a large area of the sky, the accuracy of the Chandra X-ray Observatory, with a complementary optical survey (Jonker et al., 2011). The three main goals of GBS are to determine the nature of the common envelope in binary evolution by comparing observed number of sources with the predicted ones, to identify rare X-ray binaries and to establish the projected distribution of low-mass X-ray binaries 3.2. GBS

(LMXBs). A large sample of X-ray binaries such as cataclysmic variables (CVs) which are accreting white dwarfs, high-mass X-ray binaries (HMXBs), quiescent eclipsing black hole and neutron star LMXBs and ultra-compact X-ray binaries (UCXBs which have orbital periods of less than 1 hour) form the group of rare X-ray sources searched for. All these sources are important to be detected in order to study neutron star and black hole mass measurements.

The survey team consists of a collaboration between different Astronomical Centers in Europe and The Unites States of America. In Europe, the Netherlands plays a major role with SRON (Stichting RuimteOnderzoek Nederland - Netherlands Institute for Space Research), the University of Amsterdam, Groningen University and the Nijmegen University. Also, in the UK, the University of Warwick and the University of Southampton invest in the work of this survey. Finally, the Harvard-Smithsonian Center of Astrophysics and Louisiana State University are the main collaborators in the States.

The Chandra X-ray Observatory is one of NASA's Great Observatories, launched on July 23, 1999. Chandra is sensitive to X-ray sources 100 times fainter than any previous X-ray telescope, thanks to the high angular resolution of the Chandra mirrors. Because the Earth's atmosphere is opaque to X-rays, Chandra orbits at an altitude of 139,000km above the ground. It is important to point out that Chandra has the best instruments for GBS because of its superb imaging capabilities and its extremely accurate angular resolution. Most sources will have positions accurate to 0.6-1 arcseconds (1 arcsecond is equal to 1/3600 degrees).

The photometric optical follow-up will be done on 4 m-class telescopes down to r, i and K band magnitudes of 24, 24 and 19, respectively, which still allows for Magellan/VLT (Very Large Telescope)/Gemini/SALT (Southern African Large Telescope) spectroscopic observations of discovered counterparts. All the mentioned telescopes have an aperture range from 6.5m to ~ 9.5m. X-Shooter is an efficient optical and near-infrared (u-J band) VLT instrument that will be able

to obtain spectra of objects as faint as 24 magnitudes in 1 hour. Spectroscopic follow-up is important for several science goals and in some cases for individual source classification. At the depth of the optical survey (i=24), it should be possible to detect counterparts to most of the stellar X-ray sources in the GBS. In order to predict the detections, the extinction model of Schlegel et al. (1998) and absolute magnitudes estimates for the different sources were used. The current photometric optical r, i and H α imaging was done on the Blanco Telescope

The area of the sky considered in this survey is $l \times b = 6^{\circ} \times 1^{\circ}$, l and b being the galactic longitude and latitude respectively, centered at 1.5° above and the same region below the Galactic Center. The Galactic coordinate system is centered on the Sun and is aligned with the centre of our galaxy, the Milky Way. The equator is aligned to the Galactic plane. The Galactic longitude l is measured in the plane of the galaxy using an axis pointing from the Sun to the Galactic center. The galactic latitude b is measured from the plane of the galaxy to the object using the Sun as vertex. Due to high extinction and source confusion in the Galactic Bulge, the area with $|b| < 1^{\circ}$ is excluded (Figure 3-4). Although the extinction decreases at higher latitudes, the number of bright X-ray sources drops. To avoid wasting time and money, a compromise between source density and extinction was found where it was decided that it would be unnecessary to observe at higher latitudes.

3.3 UKIDSS Galactic Plane Survey (GPS)

UKIDSS is the UKIRT (United Kingdom Infrared Telescope) Deep Sky Survey, which began in May 2005. It consists of five different surveys, each covering different areas of the sky, with the use of five near-infrared filters (ZYJHK and H_2), and with a total area of 7500 square degrees of the Northern sky (Lucas et al., 2008). Table 3-2 is a summary of all five surveys in UKIDSS. These surveys all use the Wide Field Camera (WFCAM), mounted on the UKIRT. This Cassegrain



Figure 3-4: GBS coverage. Region delimited by black lines is the survey's designated covered area. The magenta dots correspond to the GBS sources detected by the Chandra X-ray satellite. As seen, the survey is not yet complete in the X-ray detections (Figure created using the Chandra X-ray detections and survey coverage area).

telescope, located on Mauna Kea in Hawaii, is a 3.8 metre infrared reflecting telescope and is the largest dedicated infrared telescope in the world.

WFCAM has an unusual design, with an array of infrared detectors inside a long tube mounted above the Cassegrain focus. A field of view as wide as 40 arcminutes (one arcminute is equal to 1/60 degrees) is obtainable thanks to the camera design. The camera has four 2048×2048 spatially separated arrays. The arrays have a projected pixel size of 0.4 arcseconds , which gives an instantaneous exposed field of view of 0.207 deg² per exposure. The arrays are spaced by 0.94 detector widths and a tiling strategy is employed to obtain images of contiguous areas of the sky. The data are reduced and calibrated at the Cambridge Astronomical Survey Unit (CASU) using a dedicated software pipeline. They are then transfered to the WF-

CAM Science Archive in Edinburgh.

The GPS' main goal is to map the Galactic Plane to a latitude of \pm 5°, thus it overlaps with GBS fields. This limit was chosen, mainly, to match other surveys such as the MSX survey (Egan & Price, 1996). In order to make accurate discoveries and draw reliable scientific conclusions, at least three bands, and preferably four, are needed. However, due to high extinction in the Galactic plane the Y band observations are not practical. Therefore, it was decided to observe the northern and equatorial Plane in the J, H and K bands and a ~ 300 deg² area of the Taurus-Auriga-Perseus molecular cloud complex in these three filters and the 2.12 μ m H₂ filter.

Other important goals of the GPS are to measure variability and proper motion. In order to detect variability, it was decided to repeat observations in the K band because the K band can overcome more dust and extinction than the other passbands. The exposure times in JHK are 80 seconds for J and H and 40 seconds for K. The repeats will also have 40 seconds integrations, at intervals of at least 2 years. The central wavelengths of the J, H, K filters are given in Table 3-3.

It is important to mention that the photometric zero-points are bootstrapped from the Two Micron All Sky Survey (2MASS) Point Source Catalogue, which is believed to provide a reliable photometric calibration over the whole sky. 2MASS J, H and K_s filters are different from those used in UKIDSS, thus photometric transformations are used to correct the zero-points. The colours of stars in the Galactic Plane are influenced by interstellar reddening, therefore from Data Release 2 onwards, these transformations include extinction terms for correction.

The area covered by GPS is shown in Figure 3-5, amongst all other UKIDSS surveys. The main area is defined by Galactic latitude range $|b| < 5^{\circ}$, Dec $< 60^{\circ}$ and Dec $> -15^{\circ}$. These limits define two sections of Galactic longitude: $15^{\circ} < l < 107^{\circ}$ and $142^{\circ} < l < 230^{\circ}$. The reason why the region with $107^{\circ} < l < 141^{\circ}$ can not be covered is because of fundamental features of the telescope design. This



Figure 3-5: Proposed and planned sky coverage of UKIDSS. Cross-hatch : Large Area Survey. Dark Grey : Galactic Plane Survey. Light Grey : Galactic Clusters Survey. Open rectangles : Deep Extragalactic Survey. The dashed line marks the Galactic plane (Lawrence et al., 2007).

region lies north of $\text{Dec} = +60^{\circ}$ and the telescope can not observe any area above that declination limit. An additional narrow region, with $|b| < 2^{\circ}$ and $-2^{\circ} < l < 15^{\circ}$, will also be mapped in GPS. The total area covered by GPS is 1800 deg², in JHK to a depth K = 19.0.

UKIDSS GPS data saturates when the magnitudes in JHK reach J < 13.25, H < 12.75 and K < 12, though the typical saturation limits are about half a magnitude brighter than this (Lucas et al., 2008). In this case, it is safer to use 2MASS.

3.4 Two Micron All Sky Survey (2MASS)

The Two Micron All Sky Survey is another near-infrared survey, which began in June 1997 and was completed in February 2001, covering 99.998% of the celestial sphere (Skrutskie et al., 2006). UKIDSS is considered the successor of 2MASS. It

Survey	Area	Filters	Limit	Nights
	(deg^2)			
Large Area Survey (LAS)	4028	Y	20.3	262
	4028	$J \times 2$	19.9	
	4028	H	18.6	
	4028	K	18.2	
Galactic Plane Survey (GPS)	1868	J	19.9	186
	1868	H	19.0	
	1868	$K \times 3$	18.8	
	300	$H_2 \times 3$	-	
Calactic Clusters Survey (CCS)	1067	7	20.4	74
Galactic Clusters Survey (GCS)	1007		20.4	74
	1007 1067	I T	20.3	
	1007	J	19.5 19.6	
	1007	H	18.0	
	1067	$K \times 2$	18.6	
Deep Extragalactic Survey (DXS)	35	J	22.3	118
	5	H	21.8	
	35	K	20.8	
Ultra Deep Survey (UDS)	0.77	J	24.8	296
	0.77	H	23.8	
	0.77	K	22.8	
TOTAL				936

Table 3-2: Summary of all five surveys in UKIDSS (Lawrence et al., 2007)

Area is in square degrees. ' $J \times 2$ ' means that observations in that filter were repeated twice. Limit is the Vega magnitude of a point source predicted to be detected at 5σ with 2 arcseconds aperture. Nights is the estimated number of nights required to complete the observations and calibration of each survey.

Table 3-3:	Effective wavelength	$hs \ of \ JHK$	filters used in	GPS	(Hewett	et al.,	2006)
		Filters	$\overline{\lambda_{\text{eff}}}(\mu m)$				

Filters	$\lambda_{ m eff} \ (\mu m)$	
J	1.2483	
H	1.6313	
K	2.2010	

3.4. 2MASS

produced a Point Source Catalog containing 471 million sources and an Extended Source Catalogue of 1.7 million sources. In order to create such a ground-based near-infrared survey, the choice of wavelength bands was limited due to atmospheric transmission and ambient thermal background. For astronomers to be able to differentiate stellar and extragalactic populations and to distinguish the effects of interstellar extinction, at least three bandpasses are required. The JHKbands were the chosen ones for this survey (see Table 3-4 for effective wavelengths of the filters).

The main goals of 2MASS were to find brown dwarfs and to detect galaxies in the 'zone of avoidance', a strip of the sky which is obscured in visible light by our own Galaxy. Moreover, the discovery of new and rare objects, as well as representing and understanding the Milky Way were other objectives of the survey. Just like most surveys, its final goal was to catalogue all the detected stars and galaxies.

When 2MASS was proposed in the early 1990s, the only near-infrared array format available was 256×256 pixels. For each sky location, 2MASS selected a 2×2 arcseconds² pixel scale, with 7.8 seconds of integration. This total integration time was divided into six 1.3 seconds exposures due to the fact that the background noise would dominate the system read-out noise in exposures longer than 1 second. Dividing the integration time also helps improve the data and find bad pixels or cosmic rays.

In order to map out the entire sky, 2MASS required telescope facilities in both hemispheres. Two identical 1.3 m equatorial telescopes were constructed for the survey's observations. The northern telescope is located at the Whipple Observatory at Mount Hopkins in Arizona (USA) and the southern telescope was constructed at the Cerro Tololo Inter-American Observatory at Cerro Tololo in Chile. An automated software pipeline, the 2MASS Production Pipeline System (2MAPPS), reduced each night's raw data and produced astrometrically and photometrically calibrated images and tables. The entire 2MASS data set was processed twice.

66



Figure 3-6: The top figure is a full-sky distribution of point sources and the bottom figure shows the same distribution of the extended sources. Point sources are presented in Galactic coordinates centered on $b = 0^{\circ}$ and $l = 0^{\circ}$, whereas the extended source map is presented in equatorial coordinates and centered at $\alpha = 180^{\circ}$ and $\delta = 0^{\circ}$. The faint blue band in the bottom figure traces the Galactic Plane, with the intensity proportional to source density. The images are a colour composite of source density in the J (blue), H (green) and K_s (red) bands (Skrutskie et al., 2006).

Filters	$\lambda_{\rm eff} \ (\mu m)$
J	1.25
H	1.65
K_s	2.16

Table 3-4: Effective wavelengths of JHKs filters used in 2MASS (Skrutskie et al., 2006)

This survey is reliable for sources with magnitudes up to 15.8, 15.1 and 14.3 in J, H and K_s respectively (Skrutskie et al., 2006). For this reason, we use 2MASS instead of UKISS in the case of bright sources. The conversions between 2MASS and UKIDSS magnitudes (Lucas et al., 2008) are:

$$J_{\rm WFCAM} = J_{\rm 2MASS} - 0.075 \times (J_{\rm 2MASS} - H_{\rm 2MASS})$$
(3-6)

$$H_{\rm WFCAM} = H_{\rm 2MASS} + 0.04 \times (J_{\rm 2MASS} - H_{\rm 2MASS}) - 0.04 \tag{3-7}$$

$$K_{\rm WFCAM} = K_{\rm 2MASS} - 0.015 \times (J_{\rm 2MASS} - K_{\rm 2MASS})$$
 (3-8)

3.5 The Isaac Newton Telescope (INT) Photometric H α Survey of the Northern Galactic Plane (IPHAS)

The key transition for hydrogen is the H α transition, corresponding to the Balmer series with n=2. The H α spectral line is an obvious feature always found in the spectra of a large group of stars and interacting binaries. In fact, the stars with this dominant spectral line include the evolved massive stars such as supergiants, luminous blue variables, Wolf-Rayet stars and various Be stars. Post-AGB stars, all pre-main-sequence stars and active stars are also included in the group of H α emitters. H α also traces diffused ionized nebulae. Such objects evolve relatively quickly and are therefore hard to find in our Milky Way since it is a spiral galaxy. However, they would be perfect candidates in understanding evolutionary stages of such sources because when 'young' they help explain the maturation of planetary systems, and when 'older' they enable us to determine stellar end states and the recycling of energy and chemically-enriched matter back into the galactic environment.

3.5.

Moreover, H α emission is a very important indicator of interesting sources such as binaries and accretion discs. For instance, the HI recombination emission line is expected in diffuse ionised nebulae. It also appears in the spectra of pre- and post-MS stars, binaries and massive stars. The recombination emission lines come from free electrons which were initially in hydrogen atoms and were hit by photons with much higher energies, but then finally recombined with their atoms. The electrons go down to less excited states, and therefore emit energy after each jump to a lower energy level. This release of energy appears as emission lines in the spectra of stars, where H α emission lines are the strongest in the Balmer series.

Astronomers search for $H\alpha$ emitters in the Galactic Plane and Bulge, where star formation occurs, in order to find more of these sources and expand their knowledge and understanding of these phases of stellar evolution.

Thanks to recent technical developments, astronomers are able to study the entire sky and identify rare objects with great precision by creating large scale astronomical surveys.

In semester B of 2003, the Isaac Newton Telescope (INT) Photometric H α Survey of the Northern Galactic Plane (IPHAS) began taking data with the INT Wide Field Camera, mounted on the 2.5 meter INT in the Roque de los Muchachos Observatory in La Palma. The goal of IPHAS is to survey the entire northern Galactic Plane in the latitude range $-5^{\circ} < b < +5^{\circ}$. A sky area of 1800 deg² is covered by the survey (Drew et al., 2005; González-Solares et al., 2008). The latitude range was limited to \pm 5° and not more because on the one hand, the total telescope time required would increase and on the other hand, the number of sources discovered would not be significantly larger. The total time required to cover the 10 degree-wide strip is 22 weeks of clear time. For point sources, the magnitude range will be 13 < r < 20.

The Wide Field Camera is an imager comprising 4 antireflective-coated, thinned

3.5. IPHAS

4000 × 2000 EEV (English Electric Valve) CCDs (Charge-Coupled Device) arranged in an L - shape. The data obtained by the camera covers an area of the sky of approximately 0.3 of a square degree and has a pixel dimension of 13.5 μ m, which corresponds to an area on the sky of 0.333 × 0.333 arcseconds². The location of the telescope was chosen in order to exploit the sub-arcsecond seeing frequently encountered at the observatory.

A total number of 6000 pointings would be required to cover the entire area of the northern Galactic Plane if we do not take into account the geometric consequences of the L-shaped detector. However, the arrangement chosen for the CCDs has lead to a total number of field centers of 7635. Moreover, each IPHAS field is paired with a second pointing at an offset of 5 arcminutes W and 5 arcminutes S, therefore the total number of pointings at the end of the survey would be 15270 (Drew et al., 2005). Having a paired field for each original one will help cover the gaps from the detectors and enable at least two separate observations for the vast majority of the objects in the IPHAS footprint.

The filters used in IPHAS are two broadband ones, Sloan r and i, and one narrowband filter: H α . Their central wavelengths are given in Table 3-5. During the first season of observations, the exposure times in the three filters were set at 120 seconds for H α and 10 seconds for r and i. However, after working on the initial data obtained from the first season, it was decided to maintain the same exposure time for H α and i, but to increase the r time to 30 seconds starting from the 2004 observing season. It is extremely important to minimise the errors in the r due to the fact that H α excess and broadband colour measurements are linked to r band data. In order to calibrate IPHAS' photometry, standard fields obtained in twilight and every two hours during the night, are included during each observation.

The reason why this survey uses two broadband filters in addition to $H\alpha$ is to be able to compare the $(r - H\alpha)$ excess to another continuum-dominated colour. However, a non-trivial problem occurs with finding genuine $H\alpha$ emitters. They



Figure 3-7: Transmission filters used in IPHAS: $H\alpha$, Sloan r and Sloan i (Drew et al., 2005). The dotted lines show the transmission of the narrow-band filter, whereas the continuum lines correspond to the broad-band transmissions.

Table 3-5: Effective wavelengths of IPHAS (González-Solares et al., 2008)

Filters	λ_{eff} (Å)
r	6254.1
$H\alpha$	7771.9
i	6568.3

can be confused with a molecular band dominated late type stellar spectrum, in which case large $(r - H\alpha)$ colour will correlate with relatively extreme (r - i).

3.6 The Sloan Digital Sky Survey (SDSS)

The Sloan Digital Sky Survey (SDSS), which was formed in 1988 and began observations in 2000, is a multi-colour survey whose main goal was to create a three dimensional map of nearby galaxies, which can then be used to complete models for the formation and evolution of galaxies. The map will consist of 5-colour



Figure 3-8: Status of IPHAS after processing of all 2005 data. It was updated in May 2009. The present quantitative estimate is that 72% of the survey area has been observed/processed to the preferred quality standard (of seeing better than 1.7 arcseconds). 95% of all fields have been observed at least once. (IPHAS Consortium)

imaging, ugriz, to a g-band limiting magnitude of 23 (Loveday, 2002). As seen in Table 3-6, the survey was designed to cover a large area of the electromagnetic spectrum, from the near-UV all the way to the near-IR in five non-overlapping passbands. 10,000 deg² in the Northern sky, as well as three strips in the Southern Sky covering a total area of 740 deg², will be imaged.

The survey is in two parts; one consists of imaging one quarter of the sky and the other part obtains spectra for the sources detected in the first part. From these spectra we obtain the recession velocity v from the Doppler redshift of spectral features. Using the Hubble law: $v = H_o \times d$ ($H_o \simeq 70$ kmsMpc is the Hubble constant), we can determine the distances d to the galaxies and quasars detected. Astronomers often measure extragalactic distances in Mpc where 1 Mpc = 10^6

3.6. SDSS

parsecs, and 1 parsec $\simeq 3.09 \times 10^{16}$ m. The parsec, parallax of one arcsecond, is a unit of length which measures the length of the adjacent side of an imaginary right triangle in space. The two other dimensions of the triangle, which are known, are the parallax angle measured in arcseconds and the opposite side measured in astronomical units (AU). One AU is the distance from the Earth to the Sun and is equal to 149.60 $\times 10^9$ m. Once the two dimensions of the triangle are found, the length of the adjacent side, the parsec, can be calculated.

Until recently, galaxy spectra were measured one-by-one, and it is only in the last few years that optical fibre multiplexing has been used to measure redshifts for many thousands of galaxies.

The main survey telescope, situated at Apache Point Observatory in New Mexico, is a conventional two-mirror Ritchey-Chretien Cassegraine design with a primary aperture of 2.5 meters. The photometric system is defined by the imaging camera at the Monitor telescope and consists of a single ultraviolet-antireflection coated, 54 CCDs and five filters. Furthermore, the survey hardware contains a pair of dual beam spectrographs, each capable of observing 320 fibre-fed spectra. A 10-micron all-sky camera, which provides rapid warning of any cloud cover, is also found on the site.

Out of the 54 CCDs defining the photometric system, thirty of them, containing $2048 \times 2048 \ 24 \ \mu$ pixels, are the main imaging devices. They are arranged in six dewars aligned with the scan direction and holding 5 CCDs each, one CCD for each filter bandpass. The camera operates in scanning mode, therefore the time any part of the sky spends on each detector, known as the effective integration time, is 55 seconds, which results in a limiting magnitude of g $\simeq 23$.

In order to observe the spectra of $\sim 10^6$ galaxies, $\sim 10^5$ quasars and $\sim 10^5$ stars, which is the goal of the spectroscopic survey of SDSS, the overlap of fields will have to be minimised. The solution is to use two identical multi-fibre spectrographs, with two cameras each, one optimised for the red and the other for the

Filters	λ_{eff} (Å)	Full width at half maximum (Å)
u (ultraviolet)	3500	600
$g \; (\text{green})$	4800	1400
$r \pmod{r}$	6250	1400
i (near-infrared)	7700	1500
z (infrared)	9100	1200

 Table 3-6: The Sloan Digital Sky survey Photometric System (Fukugita et al., 1996)

blue. The spectrographs cover the wavelength range of 3900 - 9100 Å at a resolution of 167 km.s⁻¹. To avoid cosmic rays, the exposure time is divided into three times 15 minutes. About 45 minutes exposure times is required to obtain the redshifts for galaxies. Cosmic ray events occur at random locations on the detectors, therefore they are unlikely to appear at the same positions in all three exposures. Spectroscopic observations are carried out whenever observing conditions are not adequate for imaging. Such conditions are found when the seeing exceeds 1.5 arcseconds or when the skies are non-photometric.

3.6.1 SEGUE: A Spectroscopic Survey of 240,000 Stars With g = 14-20

As a result of the productive Galactic science enabled by SDSS, a set of three individual surveys, under the umbrella of SDSS-II, were designed in 2004: Legacy, SN Ia and Sloan Extension for Galactic Understanding and Exploration (SEGUE). SDSS-II began in August 2005 and ended in July 2008, at the same observatory as SDSS, the Apache Point Observatory.

SEGUE is an imaging and spectroscopic survey of the Milky Way and its surrounding halo (Yanny et al., 2009). The data archive from SEGUE was made publicly available in 2008, as part of SDSS-II Data Release 7.

The areas covered by SEGUE include 15 2.5° - wide strips of data along constant Galactic longitude, spaced by approximately 20° around the sky (see Figure 3-9). A wide variety of longitudes is taken in account for SEGUE's imaging, in order to sample the changing relative densities of global Galactic components (thin disk,



-30 -60 -60 86 360 330 300 270 240 210 180 150 120 90 60 30 σ Galactic Langitude Figure 3-9: The top panel shows the SEGUE Survey footprint in the Equatorial coor-

Figure 5-9. The top panel shows the SEGUE Survey footprint in the Equatorial coordinates from 360 ° to 0 ° (left to right) and -26 ° to 90 ° (bottom to top). The SDSS North Galactic Cap strips are numbered from 9 to 44 and the Southern SDSS strips are numbered 76, 82 and 86. SEGUE fills in Southern strips 72 and 79. The lower panel is the SEGUE Survey footprint in Galactic coordinates in Aitoff projection, centered on the Galactic anticenter. The line marking the Southern limit of the telescope observing site is indicated in magenta. Red and green filled areas represent South and North SDSS and SEGUE strips respectively (Yanny et al., 2009).

thick disk and halo). It is important to mention that the precise longitudes of the SEGUE strips are not evenly spaced because of several known open clusters near the galactic plane which could be optically imaged. Two strips of data in the Southern Galactic Cap were added. Since the observatory is at a Northern latitude, essentially no SEGUE data is obtained with the Equatorial coordinates $\delta < -20^{\circ}$. The Galactic anti-centre ($\delta = 29^{\circ}$) is well sampled but it is not the case for the Galactic centre($\delta = -29^{\circ}$). Dust and high stellar densities in the bulge have prevented imaging of that area of the sky.

It is this survey from SDSS which will be used for the work undertaken in this thesis, as it overlaps with IPHAS.

3.7 VISTA Variables in the Via Lactea (VVV)

VISTA (Visible and Infrared Survey Telescope for Astronomy) Variables in the Via Lactea (VVV) survey is a public near-infrared variability ESO (European Southern Observatory) survey. As its name suggests, its main goal is to construct a 3-D map of the surveyed region by using variable stars such as RR Lyrae stars and Cepheids (Minniti et al., 2009). RR Lyrae have comparable mass to that of the Sun, but are about 40 times brighter than the Sun. They are yellow-white giants and have periods of about half a day. They are named RR Lyrae after a star in the Lyra constellation, the harp, which was the first one of this type to be found. Such variables are also perfect indicators for distances in the Milky Way. Astronomers use them to find distances with the standard candles method. Cepheid variables are yellow supergiants, a lot more massive than the Sun. They have a very large luminosity range, starting from several hundred to several tens of thousands that of the Sun. They are named after the star Delta Cephei and not only is their luminosity range wide but their periods are as well. They can have periods that are as short as one day or as long as about 70 days.

The plan is to cover 520 deg^2 of the Galactic bulge and an adjacent section of



Figure 3-10: VVV coverage. Hard-lined box is the Galactic Bulge area and the dotted-lined box is the adjacent section of the mid-plane (Minniti et al., 2009)

the mid-plane (see Figure 3-10 for survey coverage). The Milky Way bulge area which will be covered expands from $l < |10|^{\circ}$ and $-10^{\circ} < b < +5^{\circ}$, thus covering all GBS fields. The entire region contains $\sim 10^{9}$ point sources, out of which \sim 10^{6} are believed to be variables. The way the 3-D map will be done is by using RR Lyrae stars, which are accurate primary distance indicators. They are also very well understood in many ways, such as their chemical abundances and their pulsational and evolutionary properties. The adjacent region of the midplane is located at $-65^{\circ} < l < -10^{\circ}$ and $b < |2|^{\circ}$. The reason why this region was added to the survey was because the star formation rate is high and other optical, mid-infrared and far-infrared surveys have observed it. Therefore, the additional near-infrared information on that region will be practical for additional discoveries.

VISTA is a 4-metre class wide-field telescope, located at the Cerro Paranal Observatory in Chile. Its main purpose is to conduct large-scale surveys of the Southern Sky, in the near-infrared wavelength range. The broad-band filters used are Z Y $J H K_s$, with bandpasses ranging from 0.9 to 2.5 μ m. The camera will also be equipped with an additional narrow-band filter at 1.18 μ m. There are over 10 scientific goals to VVV. Besides creating a high-resolution atlas of our Milky Way bulge and plane region, VVV will aim to find a very large number of RR Lyrae stars in the bulge, identify variables which belong to know stellar clusters and detect rare variable sources. Of interest to us, this survey will also enable us to discover many eclipsing binaries. Furthermore, microlensing events will be searched for, as well as new star clusters of different ages and variable populations. All the work and main goals revolve around variable sources and their properties. However, we saw that X-ray binary systems present variability in their light curves, especially during outbursts and quiescent phases. Therefore, such a survey would provide very interesting data in order to detect binary systems.

As will also be mentioned in Chapter 6, VVV overlaps with GBS and has already covered the Bulge region, therefore it is useful, additional data which will be used in future work.

3.8 The UV-Excess Survey of the Northern Galactic Plane (UVEX)

UVEX is a complementary survey to IPHAS, in shorter bandpasses. The surveyed region is the same one as that of IPHAS but the difference is that for UVEX, the Wide Field Camera on the INT is equipped with three broadband filters U, g and r, as well as an HeI 5875 narrow-band filter (Groot et al., 2009). Together with IPHAS, they form the European Galactic Plane Surveys (EGAPS), which will be the very first map of the Galactic Plane in u, g, r, i, H α and HEI 5875 (Groot, Drew et al, in prep). With the vast range of colours, a different level of the study of stellar evolution models will be achieved.

3.9 The VST Photometric H α and broad-band survey of the Southern Galactic Plane (VPHAS+)

The VLT Survey Telescope (VST) is an 8m telescope at the Cerro Paranal Observatory in Chile. For VPHAS+, the camera mounted on the VST will be equipped with ugri broad-band and H α narrow-band filters. The surveyed area will be the entire Southern Galactic Plane, with a latitude range of $|b| < 5^{\circ}$, to an AB magnitude depth of 21-22. It is therefore a complementary survey to IPHAS in the Southern hemisphere, where the Galactic Bulge is much more visible/reachable. The covered area will also overlap with GBS and VVV fields, adding multi-wavelength photometry to the Chandra X-ray sources. Similarly to IPHAS and UVEX, in order to assure a complete coverage of the area, every field will have an offset. Therefore, each field will be observed twice with a small offset in the second pointing.

Calculations have shown that ~ 50,000 H α emitters should be discovered once VPHAS+ is complete. The VST OmegaCam will acquire much larger spatial resolution and greater depth than any other camera used in Southern Galactic Plane surveys. In this way, a new sample of compact nebulae will be resolved with VPHAS+. Such sources are very interesting because they correspond to stellar evolution phases which are not very well understood and studied.

Chapter 4

Near-infrared data of the Chandra GBS sources

In this Chapter, we will explain the methods used to cross-match the catalogues of two surveys introduced in Chapter 3: the Galactic Bulge Survey (GBS) and UKIDSS Galactic Plane Survey (UKIDSS GPS). We will show how a true match was selected and describe the different problems encountered during the matching process. Our results, consisting of near-infrared data of X-ray sources detected in Bulge, will also be presented at the end of the Chapter. Finally, a few examples of spectroscopic follow-up of GBS sources will be given in order to present the different stages of an astronomer's research methods.

The Chandra X-ray Satellite observed over two-thirds of the GBS fields¹, out of which 1698 X-ray sources were detected. All the X-ray detections were divided in separate tables according to the hardness of the sources. In fact, X-ray sources can be classified as hard or soft X-ray sources, depending on the value of the X-ray energy obtained. In this case, the exact value of the X-ray energy of each source was not given, but instead they were classified in energy ranges. A source was considered soft when it had an X-ray energy between 0.3 keV and 2.5 keV, whereas a hard X-ray source had an energy between 2.5 keV and 8 keV (1 eV = 1.6×10^{-19} J). Usually, the X-ray hardness is defined as a ratio of intensity in a particular range of X-ray energy to the intensity in a softer energy range. It can be compared to the way magnitudes are obtained through filters, which also consist of a number of photons detected per second, which are then converted into a magnitude system.

The final table contained 1698 X-ray sources, all considered 'unique', without

 $^{^{1}}$ www.sron.nl/~peterj/gbs/index.html

distinguishing their X-ray hardness.

Optical and near-infrared observations of the 1698 X-ray sources will help us analyse their properties; hence the necessity to cross-correlate GBS and UKIDSS GPS, the near-infrared survey of the Galactic Plane.

4.1 Cross-matching GBS and GPS

We used the Chandra observations of the X-ray sources, denoted CXC sources, from GBS, and cross-matched them with the UKIDSS GPS. Among the 1698 sources, 171 were found to be detections of the same source within 0.5 arcseconds. Thus we actually have 1527 unique X-ray sources to study. From the GBS, we have the coordinates of each X-ray source, given in right ascension (RA) and declination (Dec), as well as their errors, in degrees. The typical Chandra radius error is 0.47 arcseconds. Unlike the galactic coordinate system mentioned in Section 3.2, the equatorial coordinate system uses the celestial sphere, an imaginary sphere, concentric with the Earth and rotating upon the same axis, to locate positions of stellar objects in the sky. The declination, which measures the angle of an object above or below the celestial equator, corresponds to the latitudinal angle of the equatorial system. The declination of an object can be positive or negative, depending on the celestial hemisphere in which it is observed, and ranging between $+90^{\circ}$ and -90° . The longitudinal angle is called the right ascension (RA) and it is usually measured in hours, minutes and seconds. The Earth takes ~ 24 hours to complete one rotation, therefore there are $\sim 15^{\circ}$ (360°/24 hours) in one hour of RA.

For UKIDSS, the images and magnitudes of the sources were obtained, when available, from the Data Release 6 (DR6) which came out in October 2009.

4.1.1 Getting the data from GPS

The first step was to request 2 arcminute images, centered on the X-ray sources. Images in each filter, JHK, were obtained from the WFCAM website thanks to



Figure 4-1: UKIDSS coverage across our sample of sources, in equatorial coordinates. The Chandra X-ray sources are plotted in white circles and the red crosses correspond to the ones with UKIDSS GPS counterparts found in the K-band.

the 'MultiGetImage' form. The tables uploaded with RA and Dec, in degrees, of the X-ray sources were limited to 500 rows; therefore the search was made 4 times for each filter. Unfortunately, many images were not available: 842 missing in the J filter, 875 in the H filter and 352 in the K filter, as the GPS is not yet complete.

The next step was to acquire the magnitudes of the sources in each filter. The data gathered from the 'CrossID' form of the WFCAM website provides information on the closest UKIDSS match to the X-ray Chandra sources. A table with the RA and Dec, in degrees, of all the X-ray sources was uploaded in the WF-CAM CrossID page and the search is done in the GPS, for the nearest object only within 5 arcseconds. The RA and Dec of the nearest match were returned, as well as the distance to the X-ray sources (in arcseconds), the JHK magnitudes, as well as their errors. Unfortunately, not all the magnitudes for each source were available. In fact, out of the 1698 X-ray sources, 354 of them had no data in any of the filters. 58% of the J magnitudes were missing, 59% of the H ones were not available as well as 26% of the K magnitudes (see Table 4-1). In Figure 4-1 are plotted the positions of the X-ray sources in RA and Dec (in degrees), and we indicate the available UKIDSS data. There are still significant GBS zones that have not been covered in UKIDSS. One may also notice, by looking at Table 4-1, that the coverage in the K-band is reasonable; however, only a limited number of fields have multi-colour information.

Once all the data and images were gathered, all the information for each source was combined in one page (see Figure 4-2): the three images (when available), a close-up of 5 arcseconds centered on the source, and the magnitudes in each filter. Some of the images were too faint and the X-ray sources were not visible unless a certain cut of the brightest and faintest pixels was applied. This process was repeated until most sources were visible.



the Chandra source. A colour-colour diagram of all the returned matches within 5 arcseconds of the CXC source is also plotted. The colours in this

plot are dereddened, using the Schlegel dust maps. Note that it is the maximum extinction considered in this case.

	J	Н	K	J, H & K	No data
DR6 (images)	856	823	1346	818	352
DR5 (photometry)	897	896	1276	805	354
DR6 (photometry)	904	906	1280	817	354
2MASS (photometry)	249	249	249	249	0

Table 4-1: Results from DR5, DR6 and 2MASS

4.1.2 2MASS photometry needed

When investigating the images, one notices many saturated objects. The solution was to use 2MASS data for the brightest sources. The limit which was decided to distinguish these saturated objects was for sources with: J < 13, H < 12.5and K < 12.5. When comparing the values of the UKIDSS saturated magnitudes, the 2MASS values and the WFCAM values calculated from the equations 3-6 in Section 3.4, it is seen that 2MASS and UKIDSS are consistent. Therefore, the 2MASS values were kept for those sources. The number of bright sources searched for in 2MASS is about 250 out of the 1698. UKIDSS images were kept due to the poor spatial quality of 2MASS ones, which enables us to see nearby field source contamination.

The initial search looked at all UKIDSS sources within 5 arcseconds of the Chandra ones. However, it must be pointed out that UKIDSS found more than one match for each source, within that same radius, due to the high density of sources in the Bulge. Figure 4-3 shows a histogram of the number of returned matches for each Chandra position within 5 arcseconds. The matches considered in the end were only the closest ones to our positions. Many sources have more than 5 possible matches, which increases the uncertainty on each match due to the high density of resolved sources.

4.1.3 The contributions to false matches

When all the plots were gathered, inspected and all the information available on each source was combined in one image, the next step was to distinguish the real



Figure 4-3: Number of matches returned, within 5 arcseconds, for each CXC source.

matches from the 'false' ones. As mentioned earlier, the Galactic Bulge is an area which suffers very high crowding. Moreover, using near-infrared photometry permits us to observe through all the interstellar dust existing between us and the Galactic Bulge. The chances of hitting a random source when matching catalogues are high in our case, also because our initial search within 5 arcseconds is much larger than the expected CXC positional errors of about 0.47 arcseconds. In order to overcome this problem, the circular radius error is calculated on the Chandra positions by applying the following equation of spherical trigonometry on the Chandra RA and Dec uncertainties:

$$\gamma \simeq \sqrt{((\alpha_a - \alpha_b)\cos(\delta_a))^2 + (\delta_a - \delta_b)^2}$$
(4-1)

where α corresponds to RA and δ corresponds to Dec, in degrees. γ is then multiplied by 3600 to obtain the radius in arcseconds since RA and Dec are given in degrees. This error is determined by the X-ray source detection software and is a 1σ error, which depends on the brightness of the source, as well as the position within the field of view.

However, there is a second contribution to this error: the boresight correction. The

Number of sources	Percentage
235	13.84%
713	41.99%
396	23.32%
354	20.85%
1009	59.48%
1328	78.21%
-	235 713 396 354 1009 1328

Table 4-2: Numbers and results

This table lists the numbers of the different scenarios and studies considered to find our most reliable matches.

overall 90% uncertainty circle of Chandra X-ray absolute position has a radius of 0.6 arcseconds. The 99% limit on positional accuracy is 0.8 arcseconds and the worst case offset is 1.1 arcseconds (Hong et al., 2005). Since the 99% accuracy is 0.8 arcseconds, which corresponds to a 3σ error, the source position error at 3σ must be added as well. Therefore, the maximum error on each position is calculated by including both the formal centroiding error from the source list, as well as the spacecraft boresight offset. A conservative maximum total error is calculated by combining the 0.8 arcseconds boresight with the 3σ centroiding error in quadrature. For each source, the closest UKIDSS match was considered and compared to these positional errors. To calculate the maximum error on each position, the following equation is used:

$$Maximum_Error = \sqrt{(3 \times Chandra_Error)^2 + (0.8)^2}$$
(4-2)

where the *Chandra_Error*, is the radius error calculated using Equation 4-1. Figure 4-4 plots histograms of the radius error of the Chandra sources as well as the maximum error radius.

4.2 Results and discussion

4.2.1 Real matches criteria

After comparing the Chandra positional errors to the UKIDSS closest match found, a true match was considered when UKIDSS returned distances less than



Figure 4-4: Histogram of the distance to the closest UKIDSS match to the Chandra source in green, the maximum Chandra radius error (Equation 4-2) in blue and the Chandra radius error (Equation 4-1) in red.

the mean value of the maximum Chandra errors. The mean value of the maximum Chandra error radius is 1.66 arcseconds. Out of those ~ 1000 sources, a fair number of them could still be random matches. As the matching radius is increased, the chances of a random match with a field star become significant. In order to investigate this effect, we generate random positions near our X-ray positions and matched them to UKIDSS in the same way. We find similar matching ratios, with ~ 50% of the closest matches that have distances less than 1.66 arcseconds from their sources. This proves that we have a 50% chance of hitting a random match with a field star. Another way of seeing this is by looking at the K-band distributions of our true and random matches (see Figure 4-5). The same conclusion was found when we noticed very similar distributions of the K magnitude values, whether of random sources or of the actual X-ray ones we are interested in. The K magnitudes of the Chandra sources reach a lower minimum value, which corresponds to brighter objects, because of the more accurate and complete search done through 2MASS in that case. However, the trend of the


Figure 4-5: Histogram of the different K magnitudes returned after the searches on UKIDSS of the random positions. The thick line corresponds to the actual Chandra sources we are interested in. We see very similar distributions, in line with a significant contribution from false matches.

histograms is the same in all the cases, with peaks at around 15 mag. The high uncertainty of the accuracy of our matches is an inevitable problem due to the source density, which can not be solved without additional data, such as our own optical imaging.

4.2.2 Science results

Finally, we constructed colour-colour (CCD) and colour-magnitude (CMD) diagrams for UKIDSS matches where we have coverage in more than one filter (Figures 4-6 and 4-7). As a reference, the Pickles stellar library (Pickles, 1998) was used to show the expected location of main-sequence stars. When plotting stellar colours versus brightness, a distinctive and continuous band of stars appears, called the main sequence. The same band is also seen in the H-R diagram, described in Chapter 2. After a star has formed, its main source of energy is created in its core, by nuclear fusion of hydrogen atoms into helium. During this

Table $4-3$:	Total extinction	(A_{λ}) for JE	K filters	, with $E_{B-V}=1$	(Schlegel et	t al., 19	' <u>98</u>).
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Filters	$A_{\lambda} \ (\mathrm{mag})$
J	0.902
H	0.576
K	0.367

stage of the star's lifetime, it is located along the main-sequence. Low-mass, also known as late-type, main-sequence stars (usually K and M stars) have a coronal X-ray activity, which makes them active X-ray stars (Gudel, 1992). Such sources would be nearby stars and they would therefore contribute to false matches. By looking at the near-infrared and optical colours of our X-ray sources, we would be able to see if they fall within the same range as main-sequence stars. We are more interested in the 'exotic' sources, which have colours different to those of 'dwarf' stars. In fact, binary stars should not have the same colour as a single main-sequence star. The Pickles library contains 131 stars, from O to M types. We have their spectral types, a log of the effective temperature, their mass, V - J, V - H, V - K, bolometric magnitudes and bolometric corrections in V (visible) and K.

A reddening vector is also included in the figures, to illustrate the effect of interstellar reddening, an effect related to interstellar extinction. Reddening is caused by the light scattering off dust in the interstellar medium. This particular scattering is known as Rayleigh scattering which consists of the elastic scattering of photons by particles much smaller than the wavelengths of the light, usually by bound electrons. The Rayleigh scattering coefficient σ_R is proportionate to $1 / \lambda^4$, where λ is the wavelength of the photons. This shows that this type of scattering is highly wavelength dependent and according to the Rayleigh scattering coefficient law, we can conclude that short wavelength photons are far more effectively scattered than high wavelength photons. Rayleigh scattering is very common in the interstellar medium, where the dust scatters the light emitted from the stars. Therefore, they appear redder than normal, because blue light is more scattered than red light, leaving an excess of red light. The objects detected seem 'redder'



Figure 4-6: Colour-colour diagram of the 'Best' UKIDSS counterparts of the Chandra sources, plotted in white circles. We also added the hardness of the sources, an extinction vector for $E_{B-V}=1$ and colours from the stellar library of Pickles (in red).



Figure 4-7: Colour-magnitude diagram of the best UKIDSS matches, as well as their hardness. In red, we see the main-sequence (MS) stars from the Pickles library. In general, we notice a similar trend between the MS Pickles stars and the UKIDSS distribution in the diagram. However, the Pickles sequence can be shifted vertically because we plot the absolute magnitude in K and not the apparent one. The extinction vector is plotted for $E_{B-V}=1$.

than their actual intrinsic colours.

In any photometric system, reddening can be described by colour excess. For instance, in the UBV photometric system (UBV stands for ultraviolet, blue and visible magnitudes respectively), the colour excess E_{B-V} is related to the B - V colour (see Section 1):

$$E_{B-V} = (B-V)_{observed} - (B-V)_{intrinsic}$$
(4-3)

where $(B - V)_{intrinsic}$ is the intrinsic (B - V) colour of the star, before being reddened by the dust.

The colour excess is required in order to determine the total extinction of a source. It describes the total absorption and scattering by the dust. It is often measured in magnitudes per kiloparsec of distance. The total extinction depends on the passband used to observe the source and is usually denoted A_{λ} , where λ is the central wavelength of the filter. Interstellar reddening depends on the line of sight of the observer. In our Galaxy, the general extinction curve from ultraviolet to near-infrared wavelengths is well known and characterised by a parameter, R_V :

$$R_V = \frac{A_V}{E_{B-V}} \tag{4-4}$$

In fact, dust maps of our Milky Way have been created in order to take into account the extinction towards observed sources. For our Galaxy, the typical value for R_V is 3.1. The absolute extinction is defined by: A_{λ} / A_V , where A_V is the total extinction in the V band at 555nm. The A_{λ} can be calculated by using the following equation (Cardelli et al., 1989):

$$A_{\lambda} = A_V \left(a(x) + \frac{b(x)}{R_V} \right) \tag{4-5}$$

where $x = \frac{1}{\lambda}$ with λ in μ m⁻¹. This equation is only valid for near-infrared wavelengths, so for 0.7 μ m⁻¹ $\leq x \leq 1.4 \mu$ m⁻¹. Also, a(x) and b(x) are defined by:

$$a(x) = 0.574(x)^{1.61}$$
 $b(x) = -0.527(x)^{1.61}$ (4-6)

In Table 4-3, we give the values of the total extinction in the case of a colour excess of 1, for the UKIDSS GPS JHK (Schlegel et al., 1998).



Figure 4-8: E_{B-V} of the Chandra sources.

For each CXC source, the expected reddening was determined using the Schlegel dust maps, which returned E_{B-V} in each case. By using Equation 4-4, we can calculate A_V for each source and finally obtain A_{λ} thanks to Equation 4-5. Figure 4-8 shows the histogram of the expected E_{B-V} for our sources, indicating a typical E_{B-V} value of 2.9 mag. Once we have E_{B-V} for each source, we can calculate the total extinction in each filter towards each source, using specific programs and equations (Schlegel et al., 1998). The reddening vectors included in the CCDs and CMDs (Figures 4-6 and 4-7) are always calculated for a colour excess of 1. As mentioned previously, the typical value of the colour excess of the CXC sources is ~ 3 mag so the actual length of the vector on the plots has to be multiplied by ~ 3 to have an idea of where the un-reddened star should fall. It is important to note that at optical wavelengths, such a E_{B-V} causes major problems because $A_V \sim 10$ mag, whereas in the K band, A_K is ~ 1 mag.

Some broad differences can be seen between the soft and hard X-ray sources; the soft sources tend to be brighter in the near-infrared and closer to the un-reddened star sequence (see Table 4-5). This can be interpreted in two different ways. On the one hand, this can show that the soft X-ray sources sample, on average, a more

Table 4-4: Numbers and magnitudes of UKIDSS matches < Mean maximum Chandra error

Magnitudes available	Number of sources	Percentage
J, H & K	630	37%
J & K	662	39%
H & K	675	40%

Magnitude	Soft	Hard
Mean(J)	12.49	14.76
$\operatorname{Mean}(H)$	11.78	13.70
Mean (K)	11.48	13.25

Table 4-5: Hardness and magnitude of sources

local population such as active stars since late-type main-sequence stars have coronal X-ray activity. In this case, they are bright and less reddened sources, which are probably closer to us. On the other hand, soft X-ray sources are absorbed by neutral hydrogen in the interstellar medium, therefore such sources are also less reddened, as otherwise we would not detect the soft emission.

Due to missing data from UKIDSS, we were limited to the number of colours we could plot in the diagrams. As seen in Table 4-4, for UKIDSS matches < Mean Maximum Chandra error, there are ~ 610 sources with JHK data at the same time. We plotted a CCD of (J-K) vs (H-K) (see Figure 4-6), with the hardness of the sources as well as the colours of the Pickles stars. Not all sources have a hardness and only 'outliers' were plotted with error bars. Unfortunately, the error bars were too large and it is important to remember that these are 1σ error bars, unlike the maximum Chandra error, used to find our best matches, which is a 3σ error. Outliers were also found in a CMD of (H - K) vs K (see Figure 4-7). They were in common with those from the CCD but the number of outliers found in the CCD is larger than that of the CMD. There were two outliers (CX0886 & CX1495), which have small error bars making them significant.



Figure 4-9: Spectrum of a CV discovered in GBS, taken with Hydra.

4.3 Spectroscopic follow-up of the CXC sources

More than 600 spectra of the CXC sources have been obtained with Hydra at the Blanco Telescope, a 4-m telescope located at the Cerro Telolo Inter-American Observatory in Chile, as well as 56 spectra with the ESO (European Southern Observatory) NTT (New Technology Telescope) in La Silla, Chile. Out of all of these sources, 5 to 10 of them are cataclysmic variables and about 20 of them are active binaries. Because only the optically bright stars have been chosen for spectroscopic follow-up, no neutron star/black hole systems have yet been discovered. The GBS survey is currently still ongoing, therefore more data reduction and spectra are being processed (Andrea Dieball, in prep).

Here, is an example of one of the cataclysmic variables (CVs) discovered in GBS. The spectrum of the CV is plotted in Figure 4-9, where one can clearly see the H α emission line, a particular characteristic found in the spectra of binaries, mainly coming from the accretion discs. The source has a spectrum which peaks closer



Figure 4-10: Spectrum of a CV discovered in GBS, taken with the NTT. The magenta dotted line shows the location of the $H\alpha$ emission line, which can be seen in this case.

to the red end, and has helium and hydrogen emission lines, which confirm its CV identity. In order to confidently identify sources, we mainly need their optical, near-infrared and X-ray data. Thus our next goal is to combine our X-ray and near-infrared matches with optical ones.

Another example of a CV discovered in GBS can be seen in Figure 4-10. This spectrum was taken on the NTT and it is part of the ~ 7 CVs discovered within the 56 spectra taken. Out of those, we also show in Figure 4-11 the spectrum of a late M-type star. As we can see, it is a red star because the spectrum peaks towards the red end of the electromagnetic spectrum, and because of the molecular features visible in the spectrum.

At the moment, all the CXC sources have optical data obtained from the Blanco Telescope, in the H α , r and i bands. We are currently waiting for the optical



Figure 4-11: Spectrum of a late M-type star, taken with the NTT.

photometry of the CXC sources, whilst obtaining spectra of a few hundred of the sources. The chosen sources for spectroscopic follow-up were mainly based on their near-infrared counterparts and provisional optical images obtained by VIMOS. It consists of a visible (360 to 1000 nm) wide field imager and multi-object spectrograph, mounted on one of the VLT in the Atacama desert in Chile. We created combined VIMOS and UKIDSS GPS finder charts that were used to select the brightest sources for spectroscopic follow-up.

Scorpius X-1 After having examined two cases from GBS, it is pertinent to demonstrate a typical LMXB case, in order to see similar sources to what we are looking for in such surveys. Scorpius X-1(Sco X-1) is a low-mass X-ray binary, which consists of a neutron star and a low-mass companion star (Steeghs & Casares, 2002). The compact object is accreting matter via Roche-lobe overflow of the companion. This binary system is one of the brightest persistent X-ray sources

in the sky, therefore it has been studied for many years through all possible wavelength ranges. The system has an orbital period of 18.9 hr and it is at a distance of 2.8 ± 0.3 kpc from us. In order to detect the companion star, Steeghs & Casares used high-resolution spectroscopy, obtained with the 4.2m William Hershel Telescope in La Palma. The instrument used was the ISIS double beam spectrograph.

Figure 4-12 shows the average normalised optical spectrum of Sco X-1. There are strong and broad emission lines from the Balmer series, as well as HeI/HeII. However, at a closer look, a large number of narrow and weak emission lines appear. These lines have double peaks, which are a clear proof of Doppler motions as a function of orbital phase. In fact, the double peaks come from the flow of the accretion disc (Steeghs & Casares, 2002; Hynes, 2010). As was shown in Chapter 2, in binary systems, half of the light detected is blueshifted and the other half is redshifted. Because the half of the material in the disc is moving towards the observer and the other is moving away, the spectral lines of the disc will also be shifted towards shorter wavelengths half of the time. An example of a double-peaked line can be seen in the top panel of Figure 4-12 at the H γ emission line. The single peaks are emission lines from the donor star in the binary system. For instance, the emission lines at around 4650 Å in the middle panel of Figure 4-12 all come from the donor star.

Besides the Balmer lines detected, the spectrum of Sco X-1 shows strong Bowen emission. The strong Bowen lines come from doubly ionised oxygen (OIII). One may notice the broad Bowen emission at 464 nm and several narrow lines from the Bowen transitions. These narrow features come from the irradiated companion star. This binary system was the first one to reveal the direct presence of a mass donor star.

By looking at the spectrum of Scorpius X-1 (Figure 4-12, we can see that it is very different to that of a single star (give example). Therefore, besides obvious colours in colour-colour diagrams, the best way to confirm the identity of a binary system is by obtaining its spectrum.



Figure 4-12: The average normalised optical spectrum of Sco X-1. In the three panels, the same normalised flux density is used. In the middle and bottom panels, two emission lines, HeII4686 (middle panel) and H β (bottom panel), are off-scale. This was done in order to visualise the weaker lines (Steeghs & Casares, 2002).

While spectroscopy clearly identifies X-ray binaries, we would like to select candidates by using astronomical surveys and their photometric properties (Hynes, 2010; Motch et al., 2009). Therefore, a much more powerful analysis can be performed if we match our UKIDSS sources to optical ones. Then effects from reddening can be easily disentangled and a more tailored target follow-up strategy could then be performed. It is clear that a simple spatial matching search is not going to be enough, given the likely presence of many false matches. However, combined UKIDSS and optical colour-colour plots will allow us to assign sensible priorities to specific targets that can then guide spectroscopic follow-up.

Here we described the selection methods of counterparts of X-ray sources detected in GBS by the Chandra X-ray Satellite. Many false matches, due to foreground stars, were found. The next Chapter will explain the second element of the project, which is the calibration of IPHAS, a r, i, H α survey of the Galactic Plane. Finally, in Chapter 6, we will present a more detailed conclusion of this Chapter, as well as future work on the GBS sources, through another survey of the Galactic Plane and Bulge, called VVV (see Chapter 3 for more details on VVV).

Chapter 5

Identifying compact stars by combining IPHAS and SDSS

The purpose of this Chapter is to describe a method developed in order to calibrate the photometric data of IPHAS. In fact, IPHAS data taken from several fields or large areas can not currently be combined and compared in one colour-colour diagram. The survey does not yet have a global calibration therefore the catalogue can not be exploited in a general way. SDSS has a reliable global photometric calibration and many overlapping regions with IPHAS. Both surveys have observed in the r and i bands, therefore cross-matching them and determining the difference in each band between both surveys will give us an estimate of how much the IPHAS data should be shifted in order to match SDSS. The calculated differences between IPHAS and SDSS magnitudes in the r and i filters will be referred to as offsets in the r and i bands throughout this Chapter. They are calculated using the following equations:

$$\Delta_r = r_{\rm SDSS} - r_{\rm IPHAS}$$
$$\Delta_i = i_{\rm SDSS} - i_{\rm IPHAS} \tag{5-1}$$

If Δ_i (or Δ_r) is positive, i_{SDSS} is larger than i_{IPHAS} , thus i_{IPHAS} is brighter than i_{SDSS} . In the same way, if Δ_i (or Δ_r) is negative, i_{IPHAS} is fainter than i_{SDSS} .

Not only will SDSS help us determine the shift required for IPHAS's calibration but it will also provide u and g magnitudes for each source. Therefore, after calibration, each source will have ugriz and H α magnitudes. Interesting science can then be performed with such information on each source, mainly with the use of colour-colour diagrams (see Chapter 4). In our calibration, we only worked with the 'Best' IPHAS fields. Such fields are from good photometric nights, with the main criteria of having an airmass less than 1.3. Light travelling from a celestial source, through the Earth's atmosphere, is attenuated by scattering and absorption. The attenuation increases where the thickness of the atmosphere is greater, which explains why celestial bodies at the horizon appear fainter than when at the zenith. Airmass is the optical path length of this light and it increases as the angle between the source and the zenith increases.

It is important to mention that even within the 'Best' IPHAS fields there could be some with bad photometric data. In one given observation, the weather conditions could have changed within the change of filters and therefore affect the data. Such cases will be shown and studied in this Chapter.

5.1 Cross-matching IPHAS and SDSS

5.1.1 Finding the overlapping regions of both surveys

IPHAS has a total of 7635 field centers, without forgetting that each pointing is paired with a second pointing at an offset of 5 arcmin W and 5 arcmin S, such that the total number of pointings found in the database will be 15270 (Drew et al., 2005), which we downloaded from the CASU website in order to find the overlapping fields with SDSS. As mentioned above, we commence our work with the 'Best' IPHAS fields, found in a separate directory in the survey database.

The SDSS SEGUE regions were requested in Galactic coordinates, whereas the IPHAS fields are all given in equatorial coordinates. In order to find the overlapping regions, we decided to work with Galactic coordinates. Due to the extremely large number of sources detected in each IPHAS field, we calculate the centre of each observed field, as well as its paired field, by taking the mean values of the equatorial coordinates of all the sources detected in each field and we convert these mean values to Galactic coordinates. In Figure 5-1, the centre of each IPHAS field



Figure 5-1: Overlapping IPHAS and SDSS fields. The cyan dots are the IPHAS fields centers, the ones ones correspond to the SEGUE coverage in the Galactic Plane. The magenta strip corresponds to the region which is calibrated, using SDSS data. As can be seen, there is a SEGUE strip which does not contain IPHAS data (with $l > 210^\circ$)

is plotted in Galactic coordinates, as well as the SEGUE sources. In black are the SDSS sources detected in the Northern Milky Way and the cyan dots correspond to the centers of the IPHAS fields overlapping with SDSS. In order to develop our calibration method, we began by working on the strip containing IPHAS fields with Galactic longitude l larger than 200°, which are plotted in magenta in Figure 5-1. In total, we can see 9 strips, of about 20 deg² each, which are detected in both IPHAS and SDSS.

IPHAS fields were considered from the overlapping regions when they were found with a limit of 0.1 arcseconds from SEGUE detections. In total, we found 2248 IPHAS paired fields which overlap with SDSS.

5.1.2 Getting the data from both surveys

The reason why we begin our work with a strip located far away from the Galactic Bulge is because the centre is a region which suffers from high extinction and crowding, which we therefore initially need to avoid. This strip, plotted in magenta in Figure 5-1, is located in the region of Galactic longitude $201^{\circ} < l < 204^{\circ}$. In this overlapping region with SDSS, we find 185 IPHAS fields , out of which 13 are single and 86 are paired $(13 + 86 \times 2 = 185)$. However, out of these 185 fields, 27 are not found to be in the 'Best' observed IPHAS table, therefore we put them aside for the initial calibration and work on 158 fields.

From IPHAS tables, we have the equatorial coordinates of the sources detected in each field, the r, i and $H\alpha$ magnitudes as well as their errors, the position on the detectors in each band, the class of the source, the CCD in which the source was detected, and the difference in coordinates in arcseconds between the matched r, i and r, $H\alpha$ images. The difference in coordinates between matched images will not be used throughout the calibration of IPHAS because one is used to denote the principal coordinate of sources.

In SDSS DR7, we request information from the Table PhotoObjAll, a table containing all photometric information on each source detected in SDSS. The following data is taken: equatorial and Galactic coordinates, u, g, r, i, z magnitudes and their errors and the object ID in the SDSS database (each source detected in SDSS has a unique ID).

In order to calculate the difference between the i and r bands in both surveys, we need to cross-match IPHAS and SDSS. However, it is important to note that a selection criteria was used during the cross-corelation. We only request sources with i magnitudes in SDSS found between 15.5 and 17.5, to stay away from the surveys' magnitude limits and enables us to work on sources with reliable magnitudes. Moreover, we only request SDSS data with reliable photometry by cutting out non-reliable matches using the SDSS 'Clean Photometry' (see Annex). We upload IPHAS tables to the Casjobs interface using Python scripts, SQL queries and a command-line tool. The matching radius chosen for the IPHAS sources is 3 arcseconds. Each match found in SDSS was returned with the magnitudes and errors from PhotoObjAll. Finally, the tables containing all the matches and photometric data on each source were then downloaded from the Casjobs interface.

Due to the matching criteria used while cross-matching IPHAS and SDSS, the final tables downloaded contained many fewer sources than the actual IPHAS lists initially uploaded. On average, ~ 70 % of IPHAS sources were lost during the matching process because they were found to lie outside the magnitude range used for the calibration.

Once all the data are obtained and gathered, we move on to the determination of the offset between the magnitudes in both surveys and eventually work on the calibration of IPHAS.

5.1.3 Important difference between IPHAS and SDSS magnitude systems

Before calculating the required offsets, we need to remember that the magnitude systems in both surveys are different. On one hand, IPHAS uses the Vega system and on the other hand, SDSS uses the AB system.

A photometric system is a set of discrete passbands of filters, with a known sensitivity to incident radiation. A passband is the overall sensitivity of an instrument as a function of wavelength. The sensitivity depends on the optical system, detectors and filters used. For each photometric system, a set of primary standard stars is provided. The first known standardized photometric system is the Johnson-Morgan or UBV photometric system (Morgan et al., 1953). Nowadays, there are more than 200 photometric systems, each based on a particular passband, meaning a particular combination of filter and detector and telescope. One should always remember to specify the system when quoting the magnitude of a star. Photometric standard stars are a series of stars that have had their light output in various passbands of photometric system, measured very carefully. They can be used as reference sources to calibrate surveys.

Most astronomers working in the optical use the UBVRI photometric systems. These are five different passbands which stretch from the blue end of the visible spectrum to beyond the red end.

In the UBVRI systems, the star Vega is defined to have a magnitude of zero. Actually, to be more accurate, the zero point is defined strictly by the mean measurements of a set of bright stars (which may include Vega), rather than by Vega alone. However, since Vega always ends up with a magnitude within a few percent of zero, the simple rule 'Vega's magnitude is zero' suffices for almost all purposes. In the Vega system, the magnitude of an object is defined by the following equation:

$$\operatorname{mag}_{\operatorname{Vega}}(Obj) = -2.5 \times \log(\frac{\int f_{\nu}(Obj) \times R\nu \times d\nu}{\int f_{\nu}(Vega) \times R\nu \times d\nu})$$
(5-2)

where ν is the frequency; $f_{\nu}(\text{Obj})$ is the object flux in ergs s⁻¹ cm⁻² Hz⁻¹; R ν is the instrumental (filter + CCD + telescope) response and $f_{\nu}(\text{Vega})$ is Vega's flux in ergs s⁻¹ cm⁻² Hz⁻¹.

The monochromatic AB magnitudes were defined by Oke & Gunn (1983) as:

$$AB = -2.5 \times \log(f) - 48.60 \tag{5-3}$$

where f is in units of ergs s^{-1} cm⁻² Hz⁻¹ (see also Fukugita et al. 1996). The constant is set so that AB is equal to the V magnitude for a source with a flat spectral energy distribution. We adopt the Vega flux densities recommended by Bohlin & Gilliland (2004). For F expressed in units of Jy, we have:

$$AB = -2.5 \times \log(F) + 8.926 \tag{5-4}$$

A great advantage of AB magnitudes is that the conversion to physical units at all wavelengths can be obtained with a single equation:

$$F = 3720 \times 10^{-0.4AB} \tag{5-5}$$

In order to maintain the H α data that we have from IPHAS, we convert the SDSS magnitudes to the Vega system using the following equations (González-Solares et al, in prep):

$$u_{\rm WFC} = u_{\rm SDSS} - 0.813 - 0.009(u_{\rm SDSS} - g_{\rm SDSS})$$
$$g_{\rm WFC} = g_{\rm SDSS} + 0.106 - 0.136(g_{\rm SDSS} - r_{\rm SDSS})$$
$$r_{\rm WFC} = r_{\rm SDSS} - 0.085 + 0.006(g_{\rm SDSS} - r_{\rm SDSS})$$
$$i_{\rm WFC} = i_{\rm SDSS} - 0.317 - 0.073(r_{\rm SDSS} - i_{\rm SDSS})$$
$$z_{\rm WFC} = z_{\rm SDSS} - 0.450 - 0.013(i_{\rm SDSS} - z_{\rm SDSS})$$

5.2 Results

5.2.1 Offsets between IPHAS and SDSS

We begin by calculating the offsets per IPHAS field, in both r and i bands, for each matched source (see Equation 5-1). In order to analyse the offset, we plot the distribution of $\Delta_{r,i}$ in each field and obtain a normal distribution, not always centered on 0 due to the fact that the IPHAS and SDSS magnitudes are different. Both the offsets and the width of the distributions are interesting to obtain.

Some IPHAS fields have a larger Δ_i than their paired field. Such IPHAS fields with unusually large offsets are mostly explained by 'non-photometric' nights. What we consider a large offset is obtaining $|\Delta_i|$ or $|\Delta_r|$ larger than 0.15 magnitudes. During such nights, clouds could have appeared while changing filters during the observations of a given fields. The change in weather conditions will affect the data in each filter and can then be seen if we plot (Δ_r vs Δ_i) (see Figure 5-2). Large differences between both offsets reflects on a change in weather conditions during the observations. In such cases, the weather conditions changed quicker than the time taken to change filters while observing a given IPHAS field, which lead to problems with calibrating such fields. IPHAS fields observed in such conditions are set for re-observations because no global calibration would fix the data. For instance, we can consider the IPHAS fields dec2006_3382 and dec2006_33820,



Figure 5-2: Offset in the *i*-band vs offset in the *r*-band, in all the IPHAS fields. The boxes are used to find the fields acting as 'outliers' in this diagram. Those same fields were found as 'outliers' in Figure 5-7. Fields with a reliable calibration should have $\Delta_i = \Delta_r$.

Table 5-1: Median of the offsets in the i and r bands in IPHAS fields 3382 and 33820 of the December 2006 run, with respect to SDSS

	3382	33820
r band	0.03 ± 0.05	-0.31 ± 0.05
i band	0.05 ± 0.05	-0.23 ± 0.05

where the offsets in the r and i bands are significantly different. In Table 5-1, we give the offsets in both bands of both fields and Figure 5-3 shows the distribution of the offset in the i band of those two fields. If we compare the magnitudes of the exact same sources detected in both fields, we notice a shift of approximately 0.3 magnitudes between each field.

Moreover, looking at the distributions of Δ_r and Δ_i per CCD in each field, we



Distribution of delta_i (offset between i_sdss and i_iphas) of field: Crossmatching_dec2006_3382 Total number of sources matched: 5724

Figure 5-3: Offset in the i band of the IPHAS field 3382 from the December 2006 run, with respect to SDSS



Distribution of delta_i (offset between i_sdss and i_iphas) of field: Crossmatching_dec2006_33820

Figure 5-4: Offset in the i band of the IPHAS field 33820 from the December 2006 run, with respect to SDSS

realise that there is a difference in offsets even within a given field. In each field, we divide the sources according to the CCD in which it was detected in. We then calculate Δ_i and Δ_r , in the same way as defined in Equations 5-1. We notice that some CCDs have larger offsets than others, and more importantly, the sign of the offsets can be different according to the CCD considered.

For each field, we plot the distribution of Δ_i in each CCD and once again obtain a normal distribution. We fit Gaussians to the distribution in each case, using a least squares method.

As in many scientific cases, we find outliers in the distribution. In general, the Gaussian is centered near zero; however, there are sources in each field which have large offsets between IPHAS and SDSS magnitudes. In fact, these large offsets can be as high as 5 magnitudes. The sources with unusual offsets are put aside while calculating the final offsets required to complete the calibration. Gaussian fits depend on the binning of the data, therefore we work with the median of the offsets in each CCD, instead of the centre of the fitted Gaussians. Thus, to calibrate IPHAS, we apply the median of Δ_i and Δ_r to the sources found in each CCD per field, using the following equations:

$$r_{\rm IPHAS(post-calibration)} = r_{\rm IPHAS(pre-calibration)} + \Delta_r$$

$$i_{\rm IPHAS(post-calibration)} = i_{\rm IPHAS(pre-calibration)} + \Delta_i$$

$$H\alpha_{\rm (post-calibration)} = H\alpha_{\rm (pre-calibration)} + \Delta_r \qquad (5-6)$$

We calibrate H α using the offset in the r band because within IPHAS data, the rand H α zero points are hardwired (Drew et al., 2005). There are no photometric standard stars in H α , therefore the H α magnitudes of all stars are bootstrapped to their r magnitudes. The reason why we calibrate H α using Δ_r and not Δ_i is because the difference between the zero-point magnitudes in r and H α is fixed in IPHAS so by applying the same offset to both magnitudes, we do not affect the colour of the sources. Also, H α lies within the r-band, therefore it seems more natural to apply the same offset to both of them.

In order to test our calibration, we calculate the new offsets between SDSS and

the calibrated IPHAS magnitudes. In each field, we calibrate the *i*-band using Δ_i in each CCD and the *r*-band and H α using Δ_r in each CCD. After calibration, our distributions of Δ_i and Δ_r are centered on 0 and the median of the offsets calculated then are very close, if not equal, to 0.

5.2.2 Is there a gradient in the offsets?

5.2.2.1 Looking across the entire strip

By considering the same example seen in Section 5.2.1, we can conclude that we can not apply a fixed Δ_i and Δ_r to the entire 20 deg² strip but we need to work on a CCD per CCD basis. In Figure 5-5, the distribution of the offsets (in the *i* band) in each IPHAS field is plotted for each CCD. As we can see, there is no obvious trend across the strip.

This result is expected and simply confirms that the IPHAS data need to be globally calibrated.

5.2.2.2 What about across the CCDs?

We can consider looking for a gradient across the CCDs, in each field. In fact, a few problems involving the CCDs of the Wide Field Camera are known. For instance, CCD3 has a vignetting problem which can be seen in Figure 5-10. Moreover, CCD3 and CCD4 both have bad columns (see Figures 5-10 and 5-11). Also, CCD3 is not well aligned with the mirror affects the focusing of the images. If the sources are blurred on the detectors then the calculated Point Spread Function (PSF) for the affected sources is incorrect. This problem is one of the main reasons why IPHAS needs to be calibrated. The PSF describes the impulse response of a focused optical system to a point source or point object. Knowing all these problems could be reflected in a systematic gradient of the offsets across the CCDs.

We plot, for each field, a distribution of the offsets. In Figure 5-6, we show this plot for the field 3382 of the December 2006 run. The colour bar depends on the offset in the i band for each source detected in this field. When comparing this same plot for each field of the strip, we do not find a recurrence of any sort of



Figure 5-5: Gradient of the offset in the *i* band across the entire strip (per CCD).

trend within the CCDs.

We know that the SDSS data becomes less reliable the closer we get to the Galactic Bulge, therefore we would not be able to apply this calibration method in the common regions closer to the Galactic centre. Another reason why the systematic gradient would have been useful is that it would have enabled us to calibrate the IPHAS regions which do not overlap with SDSS. Therefore the calibration method developed can be used when SDSS and IPHAS overlap; however, there is no systematic global correction for the entire Galactic Plane strip.



Figure 5-6: Gradient of the offset in the *i* band, in the IPHAS field 3382 of the December 2006 run. We notice the L-shape of the CCDs on the WFC. Each CCD's contour is colour-coded. CCD1's boundaries are plotted in black solid lines, CCD2's are in blue, CCD3's are in red and finally CCD4's in green. Moreover, we clearly see the vignetting problem of CCD3 mentioned in the text.

5.2.3 Why do we find outliers?

5.2.3.1 Sources with large offsets

As mentioned above, we find outliers in the distributions of Δ_i and Δ_r , where some matches had unusually larger offsets, which could be explained by two main reasons. In this case, we took all the sources with have a Δ_i and Δ_r larger than 0.3 magnitudes. We upload in SDSS a list of coordinates of the outliers and try to understand the significantly large difference between magnitudes in both surveys. On the one hand, the reason comes from the fact that a few fields are crowded and the sources detected have very close neighbours, from which they are not always deblended in SDSS. Sometimes, when two stars are very close to each other on an image, they are hard to distinguish and the PSF measures a combined brightness, which is obviously incorrect. In such cases, the magnitudes in both surveys are not very reliable, therefore we decide not to work with these sources and simply put them in separate files.

However, on the other hand, when looking at the SDSS images of the rest of the outliers which do not fall in the category of sources with deblending issues, we notice single stars, with no obvious reason to explain the large magnitudes difference. In such cases, we realise that SDSS returns different magnitudes for the exact same source. This could have two separate explanations: either the stars considered are variables, or they also have deblending issues in SDSS which are not as obvious in this case. Variable stars are stars which have their apparent brightness fluctuate with time, therefore when they are imaged in epochs, their magnitudes change. These outliers were also put aside in order to investigate their cases further. None of the stars listed as variables in our discoveries were found in Simbad. This could simply imply that either they all have deblending issues or that they are unknown variables.

Moreover, in order to further understand the outliers, we plot them according to their positions on the CCDs, in detector coordinates. As seen in Figures 5-10 and 5-11, there are a few bad columns in CCD3 and CCD4. The edges of all four CCDs are also affected by bad photometry (Figures 5-8, 5-9, 5-10 and 5-11). We also notice that the bad columns and edges in the CCDs are dominated by red dots, which correspond to sources which were always fainter in IPHAS than SDSS. Therefore, all sources which fell on those affected regions of the detectors were not taken in consideration for the calibration. Nonetheless, they give a perfect explanation of the large offset found between both surveys and, in this case, the problem is mainly with the IPHAS data.

5.2.3.2 Sources with unusual colours

Another method of finding outliers is by plotting colour-colour diagrams or colourmagnitude diagrams. In this case, an outlier does not necessarily have a large offset in either filter but it has an unusual colour in colour-colour diagrams. It does not fall in the main-sequence bulk of sources. In order to test our calibrated data, we plot a ccd of $(r - H\alpha)$ vs (r - i) of all the sources detected across the entire strip. From the distribution of the sources on the colour-colour diagram in Figure 5-7, we notice obvious 'blobs' of sources with unusually large shifts in $(r - H\alpha)$, plotted in green and labeled as 'Bad fields', and a large group of sources in blue, labelled as 'Good' fields. By simply plotting this diagram, we were able to pick out IPHAS fields with non reliable data. Table 5-2 is a list of the 'Bad' fields, with their offsets in each band. These fields, even though found in the 'Best' IPHAS table, are taken during 'non-photometric' nights. Such nights are difficult to fix because of unexpectedly low star counts due to poor weather. The transparency changed between different filter observations, which affected the data. Since H α magnitudes are hardwired to the r ones, in such conditions, we get an entire shift of the (r - r) $H\alpha$) colours. The concerned fields are currently scheduled for re-observation. We notice that the 7 fields have large offsets in both bands, which explains why the simple shift of magnitudes after calibration does not 'fix' the data.

A final 'sanity' check of the found 'Bad fields' can be done by plotting Δ_i vs Δ_r of all the IPHAS fields. This plot is shown in Figure 5-2. We clearly see that the fields found in the red box are 'outliers'. The ones lying away from the bulk of fields correspond to the same 'Bad' fields found in the colour-colour diagram. The ones in the green box could also be considered 'outliers' but they do not affect the colour-colour diagram and their offsets are not as large as those calculated in the red-box fields. In order for the calibration to be reliable, we expect to find $\Delta_i = \Delta_r$.

Putting aside all the outliers and bad data from both IPHAS and SDSS, one can still produce enough diagrams with the remaining $\sim 755,000$ sources calibrated, in order to find astrophysically 'interesting' sources. It is important to note that these

Field	Run	Δ_i	Δ_r
3382o	December 2006	-0.232	-0.311
3668	January 2008	-1.295	-0.656
36680	January 2008	-0.311	-0.412
32670	November 2005	-4.723	-3.643
3267	November 2005	-1.160	-1.291
3282	November 2005	-0.292	-0.115
3339	October 2006	0.359	0.134

Table 5-2: List of IPHAS fields, classified as 'outliers' in the $(r - H\alpha)$ vs (r - i) colourcolour diagram

were the objects used for the calibration, meaning that they do not correspond to all the sources detected in IPHAS. Because of the magnitude cut chosen for the calibration process, many IPHAS objects were not cross-matched with SDSS. Moreover, many interesting sources may vary and hence have large offsets with respect to SDSS. Such objects were considered 'outliers' and therefore they were not included in this sample.

5.3 The science which one can do with ugriz and H α magnitudes for a given source

Once all the outliers are put aside and the median of the offsets in each CCD was applied, we plot colour-colour diagrams of the stars detected in each field. Two types of colour-colour diagrams are found to be useful in terms of finding interesting sources: (u - g) vs (g - r) and $(r - H\alpha)$ vs (r - i). The latter helps us find sources with H α excess as seen in Witham et al. (2006), or deficit, which is useful in our case since binary stars with accretion discs are H α emitters. The former allows us to find blue sources, as well as WDs and binaries because they also fall in the top left corner of the diagram. In the case of this colour-colour diagram, the 'y'-axis is usually inverted, in the sense that the (u - g) colour decreases towards the top of the diagram. The 'x'-axis is not inverted, therefore the (g - r) colours increase towards the right-hand side of the diagram. Therefore, the top left-hand corner in the 'bluest' region of the diagram, with negative (u - g) and (g - r)colours. However, in this colour-colour diagram, it is hard to distinguish between



Figure 5-7: Colour-colour diagram of $(r - H\alpha)$ vs (r - i) of all the sources detected across the entire strip. The good fields are plotted in blue, whereas the bad ones are the green outliers.



Figure 5-8: Distribution of the outliers across CCD1, in detector coordinates.



Figure 5-9: Distribution of the outliers across CCD2, in detector coordinates.



Figure 5-10: Distribution of the outliers across CCD3, in detector coordinates. As mentioned, we notice the vignetting problem in the bottom left corner of the detector, as well as the bad columns.



Figure 5-11: Distribution of the outliers across CCD4, in detector coordinates. We can clearly see the bad column in this detector.

quasars and WDs because they fall in the exact same region of the diagram (Groot et al., 2009). Follow-up spectroscopy confirms the identity of such stellar sources once they are discovered.

One of the key aims of the IPHAS survey is to find H α emitters, therefore we will use colour-colour diagrams of $(r - H\alpha)$ vs (r - i). These objects should be found in the top, usually left-hand side, of the colour-colour diagrams used. Similar to the method described in Chapter 4 where the Pickles stellar library was used to determine the colours of MS stars, we use Witham's library of H α emitters in order to find where these sources fall in the diagram (Witham et al., 2006). He used a four-step method in order to find the H α emitters. He begins with an initial selection cut of the sources which consists of three main conditions: the object must be detected in the three photometric bands, without falling on any bad pixels; also their positions in the three bands must match within 1 arcsecond; and they must be classified as 'stellar' in the *i* band and as 'stellar' or 'probably stellar' in the two other bands. Once the objects have passed the initial selection, they are divided into four magnitude bins: r < 16, 16 < r < 17.5, 17.5 < r < 18.5 and 18.5 < r < 19.5. Their colours are then plotted in $(r - H\alpha)$ vs (r - i) colour-colour diagrams, in order to locate the MS locus. An initial straight line least squares fit is applied to the MS. We must point out that in fields with high densities of stars, such as fields in the Galactic Plane, the MS often splits into two mainly because of different reddening values towards the different fields (Drew et al., 2005), but also because of the different stellar populations encountered in the Plane. The upper track of MS stars is usually populated with unreddened stars. Therefore, the straight line least square fit is not enough to set a limit on where to find the $H\alpha$ emitters. The third step is to identify the upper boundary of the MS locus. In order to do so, Witham uses an iterative σ -clipping method, which mainly consists of sources which lie within 3σ above the fit and 'removing' the ones under the fit. The method is repeated four times, until the upper boundary of the MS is found. Finally, the selection of the H α objects takes into account the scatter of the points around the stellar loci and the errors on the colours of each object. He defines the $H\alpha$ excess by:

$$\Delta H\alpha = (r - H\alpha)_{\rm obs} - (r - H\alpha)_{\rm fit}$$
(5-7)

where $(r - H\alpha)_{obs}$ if the observed colour of the source and $(r - H\alpha)_{fit}$ is the value obtained from the fit.

For a source to be an H α emitter, it must follow the selection criterion:

$$\Delta H\alpha > C\sqrt{rms^2 + \sigma_{(r-H\alpha)}^2 + m_{\text{fit}}^2 \sigma_{(r-i)}^2}$$
(5-8)

where rms is the root mean square value of the residuals around the fit, C is a constant and is equal to 4.5 for the initial fits and then 5 for the σ -clipping fits, $m_{\rm fit}$ is the gradient of the fit line and $\sigma_{\rm colour}$ are the errors on the observed colours. His catalogue consists of ~ 5000 sources, detected in a total region of ~ 1500 deg². At the time when the search was done, the region of the sky with $l > 200^{\circ}$ was not complete, as seen in Figure 5-12. Therefore the number of H α emitters which fall in our strip is extremely low (13). Witham worked on a field per field basis because IPHAS was not globally calibrated at the time, but we choose to work on the entire region of 20 deg^2 . In Figure 5-14, we show a colour-colour diagram of $(r - H\alpha)$ vs (r - i) of all the sources in the strip, as well as all the 13 H α emitters from Witham's catalogue, which fall in that region. We notice a scatter of points in the region of the diagram under the main-sequence bulk. We believe that these sources are mainly more 'outliers' with unusual $(r - H\alpha)$ colours. They could also be H α deficit objects, a type of poorly understood characteristic. Such sources are not taken out of the diagram because in the fitting step, all sources found under the MS locus are not included in the next iteration of re-fitting. Therefore, they do not affect the calculations used to find $H\alpha$ emitters.

By making use of the fact that the strip is calibrated, we combine all the data in one colour-colour diagram and use a different method to find the H α emitters. We begin by binning the $(r - H\alpha)$ colours with respect to the (r - i) colour, in equal bins of 0.2 magnitudes. We then plot the distribution of the $(r - H\alpha)$ in each bin. We are simply looking at the $(r - H\alpha)$ colours of the sources in a different dimension and at a smaller scale. As seen in Witham's method description and


Figure 5-12: Galactic coordinates of all the $H\alpha$ emitters from Witham's catalogue, which were detected in the Galactic Plane during his time of search. The regions away from the Galactic center had not been observed yet, therefore they do not contain $H\alpha$ emitters. The dotted lines show the calibrated strip.

in Figure 5-14, at some point the MS locus splits into two. This effect will appear in the distribution of the $(r - H\alpha)$ colour because instead of having a simple Gaussian distribution in some of the bins, we find two peaks, each corresponding to one MS track. In Figure 5-15, we show the case of a (r - i) colour bin where the distribution of $(r - H\alpha)$ is normal, and Figure 5-16 is an example of a bin where the stellar locus is divided. In total, we work on 10 bins, ranging from (r - i) colours of 0.1 to 2.1. Even though we can clearly see that some sources have (r - i) colours outside this chosen range, the total number of sources found in the 0.2 magnitudes bin becomes too small to obtain normal distributions.

Since we are looking for sources with H α excess, their $(r - H\alpha)$ colour must be



Witham's emitters found in the calibrated strip (magenta stars) Total number of emitters found: 13

Figure 5-13: Galactic coordinates of all the sources calibrated, as well as the H α emitters from Witham's catalogue, which were also found in the strip. The large gaps with no sources correspond to the IPHAS fields which were observed during non-photometric nights and had to be taken out of the calibration. They are scheduled for re-observation.

'large' and positive. They must therefore fall on the right-hand side of the distribution of the $(r - H\alpha)$ colours in the different bins. We need to identify a limit defining the selection criterion of H α emitters. In the case of a single normal distribution of $(r - H\alpha)$, we fit a Gaussian to it. In our method, a source is selected if:

$$(r - H\alpha) > 3\sigma_{\rm fit} \tag{5-9}$$

where $\sigma_{\rm fit}$ is the standard deviation of the Gaussian fit.

In Figure 5-17, we show an example of the H α emitters found in the (r - i) bin [1.3,1.5], plotted in the colour-colour diagram, with the H α emitters taken from the Witham catalogue. Once again, not all the H α emitters were plotted here,



Figure 5-14: Colour-colour diagram of $(r - H\alpha)$ vs (r - i) of all the sources in the strip (cyan dots). The magenta stars correspond to all the H α emitters taken from Witham's catalogue, which fall in that same region of the sky. In total, we found 13 emitters from Witham's catalogue. In black is the location of the unreddened main-sequence from O5 to M6, as well as a few spectral types (MS track taken from Drew et al. (2005), in the case of $E_{B-V} = 0$).

but only the ones which were detected in the calibrated strip.

When we consider a 3σ limit, we must remember that the false alarm probability of the source not turning out to have H α excess is 1 in ~ 370. In the case of two peaks in the distributions found in the histograms of $(r - H\alpha)$, we fit them both using Gaussian fits and apply the same 3σ limit to find the H α emitters in the particular bins. In this case of two Gaussian fits, the 3σ limit is taken form the center of the Gaussian on the right of the distributions.Table 5-3 gives the number of H α emitters found, as well as the total number of sources, in each bin. These numbers correspond to the 3σ limit. Out of 751,961 sources calibrated in the strip, we find a total number of 2,656 H α excess objects, which corresponds to 0.35%



Figure 5-15: Distribution of the $(r - H\alpha)$ colour in the (r - i) colour bin of [0.1, 0.3]. In this case, we see a single Gaussian distribution, centered around 0.1 mag.



Figure 5-16: Distribution of the $(r - H\alpha)$ colour in the (r - i) colour bin of [1.3,1.5]. In this case, we clearly see two separate Gaussian distributions, each one corresponding to a MS track.



Figure 5-17: Colour-colour diagram of $(r - H\alpha)$ vs (r - i). The black dots correspond to all the sources calibrated in the strip. The blue crosses are the H α emitters taken from Witham's catalogue and the magenta stars are the sources with H α excess found in that specific (r - i) colour bin.

of all the sources calibrated. Nonetheless, we must remember that not all (r - i) colour bins were considered in the search of H α emitters. The remaining colour ranges did not have sufficient sources to obtain a Gaussian distribution, therefore they were not taken into account when looking for H α emitters. Moreover, we notice that in the (r - i) colour bin of [0.3,0.5], the fit was not perfect because the distribution itself is not very Gaussian (Figure 5-18). This did not affect our results because the centre of the fitted Gaussian does fall very close to that of the obtained distribution.

In Table 5-3, we notice that the number of emitters does not decrease regularly when the sources become redder. This could be worth further investigation in the near future. Furthermore, the (r - i) colour bin of [0.5, 0.7] has the most number



Figure 5-18: Distribution of the $(r - H\alpha)$ colour in the (r - i) colour bin of [0.3, 0.5]. We notice non-perfect Gaussian in this case, but instead a 'bump' in the left tail of the distribution.

(r - i)	Spectral type	Number of $H\alpha$	Total number of	Percentage
colour bin		emitters found	sources in bin	
[0.1, 0.3]	A5 - F5	372	8 813	4.2 %
$\left[0.3, 0.5\right]$	F8 - K2	254	$179\ 157$	0.14~%
[0.5, 0.7]	K2 - K5	$1\ 153$	$388 \ 421$	0.30~%
[0.7, 0.9]	K5 - M0	402	$116 \ 414$	0.35~%
[0.9, 1.1]	M0 - M2	78	30 030	0.25~%
[1.1, 1.3]	M2	113	$13 \ 642$	0.83~%
[1.3, 1.5]	M2 - M3	87	7 489	1.16~%
[1.5, 1.7]	M3 - M4	113	4 611	$2.5 \ \%$
[1.7, 1.9]	M4	47	$2 \ 241$	2.1~%
[1.9, 2.1]	M4	45	843	$5.3 \ \%$
Total		2661	$751 \ 961$	0.35~%

Table 5-3: Number of $H\alpha$ emitters found in each (r - i) colour bin. The spectral types are those of main-sequence dwarfs with $E_{B-V} = 0$, taken from Drew et al. (2005)



Figure 5-19: Density plot of colour-colour diagram of $(r - H\alpha)$ vs (r - i). The colour bar in this case has a linear scale.

of sources and of H α emitters found, but not the highest percentage. By simply looking at the colour-colour diagram in Figure 5-14, we do not notice a larger density of sources in that same colour bin of (r - i) = [0.5, 0.7]. However, by adding a third dimension to the colour-colour diagram, we obtain density plots as seen in Figures 5-19 and 5-20. This is done by dividing the colour-colour diagram in 1000 bins and obtaining the number of sources in each bin. The third dimension is color-coded. One plot has a colour bar with a linear scale, whereas the other colour bar has a logarithmic scale.

Due to the very small number of $H\alpha$ emitters from Witham's catalogue, it was very difficult to find the same sources in both lists of $H\alpha$ emitters. It would be natural to think that since we are both using the same IPHAS data, we would at least find the same $H\alpha$ emitters. However, during Witham's search for such sources, the region of the sky we worked on was poorly covered by IPHAS. The



Figure 5-20: Density plot of colour-colour diagram of $(r - H\alpha)$ vs (r - i). The colour bar in this case has a logarithmic scale.

closest source we find to one of Witham's emitters is ~0.97 arcseconds away from it. We used TOPCAT and our own cross-matching Python scripts to attempt to find the same sources in both catalogues but the same results were obtained in both cases: only one source coexists in the two lists of H α emitters.

Out of the 13 H α emitters from Witham's catalogue, one of them was detected in a paired field not found in our list of IPHAS in the calibrated strip. Two of the emitters were also not found in the IPHAS tables downloaded from the CASU website, even though they were detected in fields which belong to our calibrated strip. Finally, ten of the sources were found in the IPHAS tables of fields before calibrating the data. However, due to our selection criteria, eight sources were either too bright or faint to be selected in the cross-matching criteria with SDSS. Therefore, only two sources were found in our tables, after IPHAS's calibration.

Nonetheless, one of those two sources was an 'outlier', due to a large offset in its magnitudes in the r and i bands with respect to SDSS. Even though it was labeled as an 'outlier' during the calibration process, we decided to further investigate its case. It was detected in IPHAS in November 2005 and in SDSS exactly two years later. The large offset was found in the r-band, where Δ_r was equal to 0.7 magnitudes. It is an H α emitter, which also seems to be variable. Therefore, it shows two of the properties observed in most cataclysmic variables. We plot its (u - g) vs (g - r) colours in the case of a nearby star and in the case of a reddened object (see Figure 5-21). As we can see in the colour-colour diagram of Figure 5-21, the dereddened case falls in the same region of where a CV would be (see Figure 5-22 for an example of where a CV would be found in colour-colour diagrams). In the blue part of the electromagnetic spectrum, this source is quite bright. However, we look it up in 2MASS and find that it is also bright in the red end of the electromagnetic spectrum (see Table 5-5). Given how bright the object is in 2MASS, it does not look like a typical CV, but instead it could be a CV with an early-type donor, or maybe a symbiotic star. Finally, another very interesting aspect of this source is that it was detected in 1901 with 11 magnitudes in the visible range. This final data was found using the Vizier interface, where this source was then detected in the Astrographic Catalogue, +01 to +31 Degrees (Fresneau 1983).

Therefore, this source is very special in many ways: it is an H α emitter, which shows clear signs of variability, and it is bright in both the blue and red ends of the electromagnetic spectrum.

In Figure 5-23, we show a spectrum of this source, taken from IPHAS data. All of Witham's H α emitters were selected for spectroscopic follow-up, therefore we have a spectrum of this object. It has a very strong H α emission line but it does not have a spectrum of a typical CV because it is too flat. Its identity is still unknown and further spectroscopy is required to confirm it. We will make use of time on the WHT (William Hershel Telescope on La Palma-Canary Islands) very shortly to obtain another spectrum of the source.

We are left with one emitter from Witham's catalogue, also found in our list of



Figure 5-21: Colour-colour diagram of (u-g) vs (g-r). The black, magenta and cyan dots all correspond to reference stars. The black dots are SDSS main sequence stars, whereas the cyan (DB WDs = helium-rich WDs) and magenta (DA WDs = hydrogenrich WDs) dots correspond to the location of white dwarfs in this diagram. The yellow star shows the location of the dereddened colours of the H α emitter, whereas the red star corresponds to its reddened colours

Table 5-4: ugriz magnitudes of the $H\alpha$ emitter found. The magnitudes are all given in the Vega system. Both reddened and dereddened values are calculated. The dereddened value were calculated using the Schlegel dust maps (the same method as described in Chapter 4).

Case	u	g	r	i	\overline{z}
Reddened	17.44	18.20	17.36	16.45	15.63
Dereddened	15.46	16.46	16.30	15.67	15.04

Table 5-5: 2MASS JHK_s magnitudes of the $H\alpha$ emitter

Filter	Magnitude
J	14.31
H	13.71
K_s	13.32



Figure 5-22: Spectrum of a CV found in SDSS, with its colours in different colourcolour diagrams. This confirmed cataclysmic variable gives us information on where to find such systems in the diagrams.

emitters. This one H α source found in both lists is, in fact, a known H α emitter since 1984 (Ogura, 1984; Walsh et al., 1992). Through Simbad, we found this object to be known as KHA (Kiso observations detected the H α object). The source lies in the vicinoty of NGC 2264 (New General Catalogue, an astronomical catalogue containing many deep-sky objects), and was found to have H α emission on the objective-prism plates used at the time of discovery (Ogura, 1984).

It is interesting to notice that we only find 0.35% of all our calibrated sources to be H α excess objects. We were expecting to find a much larger number of emitters because we know that low-mass stars, M and K dwarfs in particular, are active stars and usually indicate the presence of H α in their spectra. Perhaps such sources have very weak H α emission to be selected in the 3 σ limit chosen in our



Figure 5-23: FAST spectrum of the H α emitter, taken from the IPHAS spectral database. As we can see, the H α emission line is very strong.

method.

It is clear that a much more powerful conclusion and catalogue can be produced with further investigations with the data. We are currently waiting for the next IPHAS data release, which will contain all the calibrated photometry of the Galactic Plane sources. These data will enable us to verify our calibration method and also add many more sources in our region of the sky, particularly all the ones which were flagged as 'outliers'. Furthermore, with more reliable IPHAS data, we will be able to produce a more updated and complete catalogue of H α emitters.

Chapter 6

Conclusion and future work on more Galactic Plane and Bulge Surveys

6.1 A typical astronomer's initial steps...

In Chapter 3.2, we went through the common steps an astronomer takes in order to find rare sources among many non-directly 'interesting' ones. We begin by choosing a strategic region of the sky, which will contain a large number and diverse sample of stellar objects. After observing that region in a specific wavelength range, usually chosen according to the type of sources searched for, we cross-match our catalogue with other known catalogues in different bandpasses. The ultimate tool an astronomer uses to find peculiar sources is the colour-colour diagram (or colour-magnitude diagram). The odd objects usually stand out in such plots and the selection for spectroscopic follow-up is then more natural.

We cross-matched a list of ~ 1700 X-ray sources from GBS, detected by the Chandra X-ray satellite, with the near-infrared catalogue of UKIDSS' GPS. The Galactic Bugle and Plane are highly populated due to the fact that the bulk of stars, dust and gas in the Milky Way are confined to those regions. Therefore, they suffer from high extinction and crowding. We find that ~ 50% of the sources could have false matches, where in that case, the Chandra source could be matched to a non-related foreground star in the field. We set a limit to the cross-matching radius for which the given sources would be reliable matches. This radius was chosen at ~ 1.6 arcseconds. We also encountered another unavoidable problem

where a lot of UKIDSS GPS data was missing. We found that from DR6, 354 sources were not found in the UKIDSS GPS database, thus they must be in areas which have not yet been observed by GPS. Moreover, 58% of the J magnitudes were missing, 59% of the H ones were not available as well as 26% of the K magnitudes.

We plot a colour-colour diagram of (H - K) vs (J - K) as well as a colourmagnitude diagram of K vs (J - K) of all the sources which were considered reliable matches. A few 'outliers' were found in the diagrams and were chosen for spectroscopic follow-up, amongst many other bright optical sources in the survey region.

Finally, r, i and H α magnitudes of sources detected in the GBS fields were obtained with the Blanco Telescope. Such data cross-matched with the Chandra X-ray sources and UKIDSS near-infrared counterparts would provide enough information to distinguish the true matches from the false ones, as well as separate single low-mass active stars from compact binaries.

Besides UKIDSS GPS, there is more recent near-infrared Galactic Plane and Bulge survey: VISTA Variables in the Via Lactea (VVV) (see Chapter 3 for a description of the survey). VVV can be considered the successor of UKIDSS GPS, but at a much larger scale. The survey has just begun and has not had any data releases so far but we may have access to their data in the very near future. In effect, we plan on cross-matching the Chandra GBS sources (CXC) with VVV as soon as this will be possible. This data would be more recent, more reliable with better image quality and more complete than that of UKIDSS GPS. We believe that the GBS coverage areas have already been observed by VVV and are at the data reduction stage. We will soon be able to compare our current results with the VVV data and hopefully obtain data on the CXC sources which were neither found in UKIDSS GPS, nor had all the photometry required in order to reach final conclusions on all the sources.

6.2 When things go wrong

IPHAS is a scientifically interesting survey to make many discoveries, and expand our understanding of the Galactic population. However, its current photometric calibration is not, to absolute standards, very reliable. Our goal was to use another survey, SDSS, to compare their data and potentially calibrate IPHAS's. By cross-matching both surveys and calculating the differences in magnitudes, in the r and i bands, between the same sources detected in both SDSS and IPHAS, we find an initial calibration of the IPHAS data.

Further research and understanding of the data enabled us to realise that calibrating the data on a CCD per CCD basis was more accurate and efficient than on a IPHAS field per field basis. We do not live in a perfect world, therefore we encountered problems related to the technical issues of the camera such as bad columns on the detectors (mainly CCD3 and CCD4), or simply data taken during cloudy nights.

Nonetheless, even though we had to discard many sources while calibrating, due to the fact that they were flagged as 'outliers' and affected the calibration, we finally worked on ~ 755,000 sources, located close to the Galactic anti-centre. We chose that region of the sky because the SDSS data is reliable away from the Galactic Bulge and it suffers from less extinction and crowding than the Galactic center. After the calibration of the large sample of sources detected in the strip, we tested a new method to select H α emitters.

A next step, which would be of practical use to the astronomical community, could be to produce an up-to-date H α emitters catalogue of sources in the Galactic Plane and Bulge. The current and nearly complete IPHAS data is being calibrated in Hertfordshire by a group lead by J. Drew. Their global calibration data is planned to be released before the end of 2010. Comparing our calibration with theirs would be an interesting piece of work, as making use of the data on a much larger scale would enable us to select a much larger sample of H α emitters. By mid-2011, we should be able to access data from both UVEX (see Chapter 3 for more details on UVEX) and IPHAS. An interesting mini-project would be to search for the H α emitters found with our method in UVEX and look for their positions on a colour-colour diagram of (U - g) vs (g - r). It would be practical to get their UV/'bluer' colours and select the hot ones for spectroscopic follow-up. Mainly, the understanding of binary evolution and phases will be possible by combining the different colour-colour diagrams which could be created in this case. As seen in Chapter 5, blue colours of H α emitters are very useful for selecting the exotic and interesting sources. Therefore, cross-matching IPHAS and UVEX would enable us to find more CVs and rare objects.

Finally, we can expand our search to the Southern Galactic Plane with yet another survey, VPHAS+ (Drew et al, in prep) (see Chapter 3 for more details).

The combination of all the colours obtained by VPHAS+ will enable us to gather and classify intrinsically faint, blue objects. Such stars could be single or binary WDs, subdwarf B stars and close binaries. Many questions will be answered once a large sample of such stars is found. Astronomers will be able to understand how close binaries evolve, more specifically the common envelope phase. Moreover, obtaining u and H α data at the same time would play an important role in the study of accretion-powered binaries.

Conclusion

With data from all the mentioned surveys, we can continue our work in the search of H α emitters and produce larger and more reliable catalogues of these sources. We also work hard on finding close binaries in order to understand their evolution and properties. The reason why so many surveys are ongoing is because the larger the sample of sources, the more accurate the results and models will be. A large sample of binary systems is essential to develop our knowledge and understanding of stellar evolution.

Two projects were described in this thesis, where in one case, the counterparts of Chandra X-ray sources in a crowded region of the sky were obtained from UKIDSS GPS, and in the other case a survey, IPHAS, was calibrated to an absolute scale using data from SDSS.

Both methods used are very useful in the near-future as several directly related surveys will begin. Many interesting discoveries can be made by combining all the tools and strategies used in both projects.

Annex : SDSS clean photometry SQL query

We give the SDSS SQL query used to pick the reliable SDSS matches for the IPHAS sources:

SELECT TOP 10 u,g,r,i,z,ra,dec, flags_r

FROM Star

WHERE

ra BETWEEN 180 and 181 AND dec BETWEEN -0.5 and 0.5

AND ((flags_r & 0x10000000) ! = 0)

– detected in BINNED1

AND ((flags_r & 0x8100000c00a4) = 0)

- not EDGE, NOPROFILE, PEAKCENTER, NOTCHECKED, PSF_FLUX_INTERP,

- SATURATED, or BAD_COUNTS_ERROR

AND (((flags_r & 0x4000000000) = 0) or (psfmagerr_r <= 0.2))

– not DEBLEND_NOPEAK or small PSF error

- (substitute psfmagerr in other band as appropriate)

AND (((flags_r & 0x1000000000) = 0) or (flags_r & 0x1000) = 0)

- not INTERP_CENTER or not COSMIC_RAY

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