

THE EVOLUTION OF HII REGIONS
IN THE LARGE MAGELLANIC CLOUD

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This thesis is dedicated to my father,
Ian Andrew Wilson and mother, Marion Noreen
Wilson, whose love and sacrifice have made
this thesis possible.

The work described in this thesis is
that of the candidate alone, except where
noted otherwise in the text.

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Ian Wilson

IAN WILSON

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ABSTRACT

A study of the effect of mass-loss from OB stars upon the evolution of HII regions in the Large Magellanic Cloud is presented. This study incorporates spectroscopic observations of 25 OB stars in 10 clusters located within L.M.C. HII regions, together with narrow band hydrogen-alpha surface photometry of an overlapping sample of 15 HII regions.

Empirical calibrations based on prior studies of Galactic OB stars are used to determine the integrated stellar wind mechanical luminosity (L_w) and the integrated Lyman continuum photon luminosity (S_*) of the 10 OB clusters.

The ratio between the momentum carried away from a star by the stellar wind to the momentum carried away by photons is found to be 0.7 ± 0.3 , for OB stars with $\log \dot{M} \geq -6.3$. This fact, together with the scaling law that is found between the stellar wind terminal velocity and the stellar escape velocity, is used to empirically relate the mass-loss rate of an OB star to its luminosity, radius and mass. The dependence of $\dot{M}$ on the stellar parameters found is not predicted by the line-driven wind or Fluctuation theories for mass-loss, indicating that these theories must be modified in order to remain viable.

S_* , L_w , $\dot{M}$ and the stellar wind momentum transfer rates are calculated at intervals of 1.4×10^5 years throughout the

core-hydrogen burning lifetimes for stars with initial masses of 20, 30, 40, 50, 60, 80 and 100 $M_{\odot}$. These calculations show that the magnitudes of L_w and S_* are essentially constant during this phase of evolution, and that they are primarily dependent on the initial mass of the OB star.

These calculations also show that at the beginning of core-hydrogen burning, the ratio L_w/S_* has a constant value equal to $3.2 \pm 0.4 \times 10^{-20}$ (Joules photon $^{-1}$) for 20 to 100 $M_{\odot}$ stars.

A comparison of the integrated values of S_* for the L.M.C. OB clusters with the ionizing photon absorption rate of the surrounding nebulae, via the Zanstra method, shows that the L.M.C. HII regions are in general ionization bounded.

The values of L_w and S_* for the L.M.C. OB clusters, together with the hydrogen-alpha surface photometry, are used to show that the large shell-like HII regions in the L.M.C. are stellar wind bubbles 3 to 5 million years old. In order to reproduce the observed densities in the shells of these HII regions a thick cushion of shocked stellar-wind gas must have been present on the inside of the HII shell for most of the lifetime of the nebula. The stellar wind mechanical luminosity, the age of the central clusters and the radius of the HII shells indicate that the stellar wind bubbles are moving into an I.S.M. with a typical time average density $\geq 2.0 \times 10^7$ particles m^{-3} . Densities this

high are typical of those found in the outer parts of large massive molecular clouds. This raises the possibility that the HII regions are being ram pressure confined by the infalling gas in the molecular cloud.

Average cross-sections of surface photometry maps of HII rings show that the shells are partially filled in, indicating that the hot cushion of gas which has been driving the bubble, has collapsed in the later stages of the nebula evolution.

Acknowledgements	iv
Abstract	vi
List of figures	xiii
List of Tables	xiv
CHAPTER 1 INTRODUCTION	1
CHAPTER 2 - THE EVOLUTION OF MII REGIONS	6
2.1 Evolved MII Regions - The Classical Model	8
2.2 The Blister Model	9
2.3 Stellar Wind Bubbles	20
2.4 The Off-centre Accretion Model	20
CHAPTER 3 - SPECTRAL CLASSIFICATION OF O AND B STARS IN THE L.M.C.	24
3.1 The M.K. Spectral Classification System	24
3.2 The Adopted Method of Spectral Classification for O and B Stars	26
3.2.1 O Stars	27
3.2.2 Luminosity Indicators for O Stars	27
3.2.3 O9 to B5 Stars	29
3.2.4 The Luminosity Class of L.M.C. OB Stars	31
3.2.5 Luminosity Indicators for B0 to B5 Stars	32
3.2.6 The Spectral Classification Scheme	33
3.3 The Observations of L.M.C. OB Stars	35
3.3.1 Constraints on the Observations	35
3.3.2 The Observations	37
3.3.3 The Reduction Procedure	41
3.4 The Spectral Types of L.M.C. OB Stars	43
3.4.1 Spectral Types	43
3.4.2 Uncertainties in the Classification Procedure	44

	PAGE
CHAPTER 4 - THE INTEGRATED LYMAN CONTINUUM PHOTON FLUX OF L.M.C. OB CLUSTERS	46
4.1 Introduction	46
4.2 Theoretical Atmospheric Models for OB Stars	47
4.3 The Lyman Photon Flux (S_L) as a Function of M.K. Spectral Type	50
4.4 Results for L.M.C. Stars	57
CHAPTER 5 - MASS-LOSS IN OB STARS	
5.1 Introduction	64
5.2 Methods for Determining Mass-Loss Rates for OB Stars	66
5.2.1 Mass-Loss Rates Derived from U.V. Resonance Lines	67
5.2.2 Mass-Loss Rates Derived from Radio Observations	69
5.2.3 Mass-Loss Rates Derived from Infra- Red Observations	70
5.2.4 Mass-Loss Rates Derived from Hydrogen Alpha Emission	71
5.3 The Adopted Sample of Mass-Loss Rates	72
5.3.1 Lamers Sample of Mass-Loss Rates	73
5.3.2 Modifications to the Sample of Lamers	76
5.4 The Mass-Loss Rate as a Function of Stellar Parameters	78
5.4.1 $\dot{M}$ as a Function of L_*	78
5.4.2 $E_T \propto L_*$	80
5.4.3 $E_T \propto (t_D/t_K)^{1/2} L_*$	82
5.4.4. $E_T \approx (V_{esc}/c) L_*$	83

CHAPTER 6

6.1	Introduction	92
6.2	The Effective Temperature Scale for OB Stars	94
6.3	Evolutionary Models for OB Stars	100
6.4	Results of the Evolutionary Models	102
6.5	Conclusions	105
6.5.1	The Distinction between Of and non Of Stars in the Adopted Sample	105
6.5.2	The Transition of Of Stars to W-R Stars	107
6.5.3	The Denisty of the Stellar Wind Envelope	109
6.6	The LMC HII Regions	110
6.6.1	Results	110
6.6.2	Discussion and Conclusions	112

CHAPTER 7 - SURFACE PHOTOMETRY AND CONCLUSIONS

7.1	The Instrumentation	115
7.1.1	The S.E.C. Vidicon T.V. 2-D Photometer	115
7.1.2	The Optics	117
7.1.3	The Hydrogen Alpha Filter	118
7.2	The Method of Observation	119
7.3	The Initial Reduction Procedure	122
7.4	The Integrated $H\alpha$ - Flux	124
7.5	The Surface Brightness of the L.M.C. HII Regions	128
7.6	The Conclusions	131

7.6.1	Are the HII Regions Ionization or Density Bounded?	131
7.6.2	The Hot Cushion and Direct Impingement Models	137

REFERENCES	146
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	List of figures	following page
3.1	The apparent visual magnitude and absolute visual magnitude as a function of MK spectral type.	31
4.1	$\log F_L$ as a function of stellar effective temperature.	48
4.2	$\log (N_L / \pi F_V)$ as a function of MK spectral type for dwarf stars.	55
5.1	$\log \dot{M}$ as a function of $\log (L_*/L_\odot)$ for all stars in the adopted sample.	78
5.2	$\log \dot{M}$ as a function of $\log (L_*/L_\odot)$ for supergiant stars.	79
5.3	$\log \dot{M}$ as a function of $\log (L_*/L_\odot)$ for non-supergiant stars.	79
5.4	$\log \dot{M}$ as a function of $\log ((L_*/L_\odot)/V_{esc}^2)$.	81
5.5	$\log \dot{M}$ plotted against the Andriesse formula for mass loss.	83
5.6	$\log \dot{M}$ as a function of $\log ((L_*/L_\odot)/V_\infty)$.	85
5.7	$\log \dot{M}$ as a function of $\log ((L_*/L_\odot)/V_{esc})$.	88
6.1	The effective temperature scale as a function of the MK spectral type for O4 to B3 stars.	97
6.2	The parameters a) $\log \dot{M}$; b) $\log L_{uv}$; c) $\log \dot{M} V_\infty$ and d) $\log L_w$ as functions of stellar age for various initial stellar masses.	102
6.3	HR diagram for OB stars in the adopted sample.	106
6.4	Actual and minimum (necessary to expose enriched core material) percentages of mass loss as a function of stellar initial mass.	107
6.5	$\log \dot{M}$ as a function of $\log (V_\infty R_*^2)$.	109
7.1	The optical train for the TV photometer-Reducer combination.	117
7.2	Transmission profile for the $H\alpha$ filter used for the HII region observations.	118
7.3	Rate at which LMC HII regions absorb ionizing photons (Zanstra method) as a function of the ionizing stellar emission rate.	132
7.4	Dereddened, integrated $H\alpha$ flux as a function of HII region diameter.	136
7.5	Integrated $H\alpha$ flux (derived from stellar Lyman continuum) as a function of HII region diameter.	136
7.6	$\log (t_6 L_w)$ as a function of HII region radius.	138
7.7	L_w for LMC HII regions as a function of their radii.	141

7.8	Lyman continuum photon luminosity (S_{H}) and stellar wind mechanical luminosity (L_{w}) as a function of initial stellar mass.	143
7.9 (a to m)	Absolute $H\alpha$ surface brightness maps for 13 LMC HII regions.	145
7.10	Surface brightness profile for DEM 25.	145

	List of tables	following page
3.1	MK spectral type for O stars based on the $W(4471)/W(4542)$ Ratio W' .	27
3.2	Spectral discriminants for O9 to B5 spectra at MK dispersion.	30
3.3	Standard stars : O9 - B5.	35
3.4	Summary of spectral observations.	39
3.5	LMC HII region bright stars : parameters.	43
3.6	LMC HII region bright stars : equiv. widths.	45
4.1	Adopted relationship between $\text{Log } (N_L/\pi F_V)$ and MK spectral type.	55
4.2	LMC HII regions : physical parameters.	58
5.1	Mass Loss rates : stellar sample.	77
5.2	Table of ϵ , B, V and p for a representative subsample of stars from the adopted data-set.	85
6.1	Adopted temperature scale.	99
6.2	LMC HII region stars : a) astronomical parameters; b) physical parameters.	110
6.3	Calculated errors due to uncertainties in T_{eff} and spectral type.	111
7.1	H α Surface photometry observation summary.	115
7.2	Surface photometry flat-fields.	123
7.3	Surface photometry reduction summary.	124
7.4	Surface photometry: extinction determinations.	125
7.5	Surface photometry: standard observations.	126
7.6	Surface photometry: conversion factors.	128
7.7	LMC HII regions: electron densities.	130
7.8	LMC HII region stellar wind parameters.	137
7.9	LMC HII region densities.	140
7.10	Details of H α surface brightness maps (figures 7.9 a to m).	145

CHAPTER 1

INTRODUCTION

Introduction

The aim of this thesis is to investigate the evolution of HII regions in the large Magellanic Cloud (L.M.C.) through a combined study of the regions and the young OB stars embedded within them.

The L.M.C. was chosen as it is the closest external galaxy to our own, providing us with the opportunity to observe a large number of relatively unobscured HII regions of widely differing morphology and known distance. These properties, proximity and low reddening, are also advantageous for observing the individual OB stars located within the HII regions. It was not until the advent of high sensitivity photon counting detectors though, that a thesis of this nature became viable. Using these modern detectors it is possible to obtain high signal to noise, moderate ($\sim 100\text{\AA}/\text{mm}$) dispersion spectra of bright LMC OB stars in less than 30 minutes.

The HII regions chosen for this study were selected on the basis that they were relatively free of obscuration; that they have a relatively simple geometry with near spherical symmetry; and that they have less than roughly a dozen bright OB stars located within them. The last two criteria were selected against the high surface brightness large HII regions which have tens and sometimes hundreds of OB stars located within them. The choice of simple morphology and spherically symmetric structure was stipulated to minimize complications due to density inhomogeneities in the ISM. The catalogue of L.M.C. HII regions published by Henize (1956) showed that the L.M.C. has many HII regions irregularly distributed over a $6 \times 6^\circ$ region. The HII regions range in size from stellar-like nebulae up to huge objects which are hundreds of parsecs in diameter. The largest nebula in the L.M.C., 30 Doradus, is one of the biggest HII regions known. It dominates a region of

star formation located at the Eastern end of the stellar bar. The other HII regions are either associated with weak spiral arm features, or are found in irregular scattered regions of star formation which cover areas of the order of a kiloparsec. (Meaburn 1979,80)

A more recent catalogue of L.M.C. HII regions is that of Davies, Elliott and Meaburn (1977) (D.E.M.), and is based on deep UK Schmidt Telescope plates taken in conjunction with a mosaic $H\alpha$ filter. This catalogue includes information on the position, size and brightness of over 300 HII regions, and penetrates to a fainter surface brightness than that of the Henize catalogue thus giving a clearer and more comprehensive picture of the distribution of L.M.C. HII regions. The nomenclature used in the catalogue of Davies et. al. has been used in this thesis.

The corresponding Henize number of each HII region can be found in this catalogue.

Both of the above catalogues reveal the presence of large ring-like nebulae which range in size from three to fifteen arc-minutes (roughly 42 to 210 parsecs if a distance modulus of 18.4 is assumed for the L.M.C.). Amongst the L.M.C. HII regions these nebulae have raised most interest, chiefly because some of the brighter rings (eg. D.E.M. 301 (N70), DEM 25 (N185)) are morphologically similar to the optically known supernova remnants in the L.M.C. (Lasker, (1977)). These nebulae have sharp filamentary shells, a high $[SII]$ 6716,32 to $H\alpha$ ratio and supersonic internal motions indicating the

presence of shock phenomena. (Lasker, 1979 , 1981). The shell-like morphology and the fact that the radio fluxes of these nebulae roughly fit on the Σ -D relation for L.M.C. supernovae (Milne et. al. 1980) have led various authors to speculate that they are evolved supernova remnants. Evidence which contradicts this hypothesis includes measured electron temperatures of $\sim 10^4$ K (Lasker 1981, Dopita et. al. 1982), the presence of thermal radio spectra in D.E.M. 301 (N70), (Dopita et. al. 1982) and in D.E.M. 192 (N51D) (McGee and Newton 1972) and D.E.M. 241 (N59A) (Meaburn et. al. 1981), and the presence of OB stars within the HII shells. These have led some people to speculate that these large HII rings are stellar wind bubbles (Meaburn 1978, Meaburn and Blades 1980 and Lasker 1980) formed from the mass-loss from the OB stars. The problem with these objects is that neither the supernova hypothesis nor the classical stellar wind bubble hypothesis can explain all the observational characteristics of those nebulae.

The observations of the dynamics of the ionized gas in these nebula have been so far unable to resolve this dilemma. The velocity fields of nebulae such as D.E.M. 301 (Rosado et. al. 1981, Blades et. al. 1980, Dopita et. al. 1982), D25 (Rosado et. al. 1982), D192 (Lasker 1980, Meaburn and Terrett, 1980, Klaus de Boer and Nash 1981), D241 (Meaburn et. al. 1981) and D229,231 (Meaburn and Blades 1980) are complex and cannot be described in terms of a simple

expanding shell. A number of different hypotheses have been proposed to describe the velocity structure of these nebulae. Lasker (1980) adopted a stellar wind bubble model evolving into an inhomogeneous I.S.M. to fit the velocity structure of D.E.M. 192. Meaburn and Terrett (1980) and Meaburn et.al. (1981) used a model which has an expanding shell, formed by energetic stellar winds or multiple supernova explosions, expanding into an I.S.M. consisting of large sheets of neutral hydrogen. On the other hand Rosado et. al. (1981,1982) preferred a supernova explanation for D301 and D25 to describe the velocity structure of these nebulae.

Dopita (1981) proposed a third model in order to try to explain the confusing observational picture which has developed for the L.M.C. HII rings. Dopita's model has a stellar wind bubble being formed by mass-loss from an OB star (or stars) located off-centre in a large collapsing molecular cloud. The stellar wind bubble quickly establishes a quasi-static configuration where the ram-pressure of the stellar wind just balances the corresponding rampressure of the infalling cloud material. This model was applied to D.E.M. 301 and it was able to produce all of the major observational characteristics of this nebula including the velocity structure. (Dopita et. al. 1982).

A large proportion of the HII regions chosen for study in this thesis are HII ring nebulae. This was deliberately done so as to be able to try and distinguish which of

the models proposed above (if any) apply to the L.M.C. HII regions.

In the following chapter (Chapter 2) the various models which are applicable to evolved HII regions are discussed. This work clearly shows the need to know the ionizing photon luminosity and mass-loss rates ($\dot{M}$) for the OB stars embedded in the L.M.C. HII regions.

In Chapter 3 the M.K. spectral types are determined for the bright OB stars located within the L.M.C. HII regions. Chapter 4 indicates the method which was adopted to determine the integrated Lyman continuum photon flux for the L.M.C. OB clusters, using the M.K. Spectral types derived in Chapter 3. Chapter 5 discusses the various relationships between the mass-loss rates of OB stars and their gross properties, i.e. luminosity (L_*), radius (R_*) and mass (M_*). The relationship between $\dot{M}$ and L_* , R_* and M_* which is adopted in this chapter has been used in Chapter 6 to determine the integrated stellar wind mechanical luminosity ($\frac{1}{2}\dot{M}V_\infty^2$) and momentum transfer rate ($\dot{M}V_\infty$) for the L.M.C. OB clusters (V_∞ is the terminal velocity of the stellar wind).

In the final chapter of this thesis, integrated $H\alpha$ fluxes and surface brightness contour maps of the L.M.C. HII regions are presented. These data enable the integrated stellar parameters of the OB clusters derived in Chapters 3 and 6, to be compared with the properties of the HII regions. This final chapter ends with the discussion and conclusions of the thesis.

CHAPTER 2

THE EVOLUTION OF HII REGIONS

2. THE EVOLUTION OF HII REGIONS

2.1 Evolved HII Regions - The Classical Model

The evolution of a HII region in a uniform low density medium is reasonably well understood (Lasker 1967, Mathews and O'Dell 1969), being characterised by the following features. Soon after the ionizing star (or stars in the case of an OB association) is formed, an R-type ionization front propagates supersonically into the surrounding I.S.M. The R-front continues to expand supersonically until a point is reached where all the available ionizing photons from the star are used up in maintaining the HII region. The radius at which this occurs is designated as the Initial Strömgen Radius R_s , given by the expression

$$R_s = \left(\frac{3 L_{uv}}{4\pi n_{II}^2 \alpha_2} \right)^{1/3} \quad 2-1$$

where L_{uv} is the number of photons/sec emitted by the star, n_{II} is the HII region electron density and α_2 is the recombination coefficient for the excited levels of hydrogen. (Strömgen 1939).

Once the Initial Strömgen Sphere is formed the ionization front (I front) becomes a weak D-type front with an isothermal shock preceeding it. This shock forms a dense neutral hydrogen shell which is in pressure equilibrium with the HII region. The pressure differential between the HII region and the I.S.M., set up by the

establishment of the initial Strömgren Sphere, provides the driving force for this shock and so the weak D-type front will continue to expand until pressure equilibrium is established.

The density of the HII region at the start of the expansion phase is equal to that in the undisturbed I.S.M. The expansion is communicated to the HII region via a rarefaction wave which propagates sonically into the ionized gas. At later phases, when the sound-crossing time scale is short compared to the evolutionary timescale, then the density within the HII region is once again constant, and the constant ionizing flux from the central star is given by

$$\rho_{\text{II}} / \rho_{\text{O}} = (R_{\text{S}}/R)^{3/2} \quad 2-2$$

where ρ_{O} is the density in the I.S.M., ρ_{II} is the density of the HII region and R is the radius.

R is given by the expression

$$R / R_{\text{S}} = \left(1 + \frac{7}{4} \frac{c_{\text{II}}^2 t}{R_{\text{S}}} \right)^{4/7} \quad 2-3$$

Spitzer (1978).

The majority of the L.M.C. HII regions selected for study in this thesis are large nebula with distinct

shell-like structures. It is difficult to imagine how the classical picture discussed above could produce those large shell-like HII regions. At best, the morphology of an evolved HII region would appear as a relatively low density HII region with a very sharp boundary making the ionization front. This model would certainly not produce the large relatively thick shells observed in the L.M.C. HII regions.

2.2 The Blister Model

An alternative theory which appears to fit most of the observations characteristics of Galactic HII regions (Osterbrock 1974, Tenorio-Teagle 1978) is the Champagne (or Blister) Model. This model embraces the fact that regions of active star formation usually take place near large ($\approx 60-80$ pc) molecular clouds of mass 10^4 to $10^5 M_{\odot}$, invalidating the uniform density approximation of the classical theory. In brief outline the evolution described by this model is as follows.

The Champagne model assumes that an ionizing star (or stars in the case of an OB association) switches on near the edge of a large dense molecular cloud and begins to form an Initial Strömgren Sphere. Eventually the I-front breaches the outer boundary of the cloud and rushes into the low density inter-cloud medium (I.C.M.). The large pressure differential between the

ionized cloud material and the ionized I.C.M. results in a strong shock forming at the breach in the molecular cloud boundary. The shock then propagates supersonically into the ionized I.C.M. followed by the ionized cloud gas. The expansion of the HII region out through the breach in the cloud results in the propagation of a rarefaction wave back into the cloud HII region, towards the ionizing stars. This wave decreases the density of the molecular cloud HII region to that of the HII region which is expanding into the I.C.M.

Although this model can reproduce the assymetric blister type HII regions such as the Orion nebula, it does not produce shell structures in the ionized gas. In order to produce these structures we have to consider the effects of stellar winds from the massive OB stars upon the HII region.

2.3 Stellar Wind Bubbles

Models for expanding spherically symmetric stellar wind bubbles have been discussed extensively by Pikel'ner and Shcheglor 1969, Avedisova 1972, Dyson and deVries 1972, Falle 1975 and Weaver et. al. 1977. The following discussion outlines the major characteristics of these bubbles as outlined by Weaver et. al. An evolved stellar wind bubble can be subdivided into four distinct regions. The first and innermost region

is the freely expanding hypersonic wind (zone 1). This stellar wind sweeps up a shell of shocked interstellar gas (or ionized gas if the bubble is expanding into an HII region). The collision of the wind with this shell sets up a reverse "bow" shock in the stellar wind flow which forms the second zone of hot ($\sim 10^6$ K) shocked gas behind the dense shell. The third zone is the shocked interstellar gas (or ionized gas) which forms behind the isothermal shock produced by the piston action of zone 2. Zones 2 and 3 are separated by a contact discontinuity across which a conduction front develops due to the large temperature difference between the two zones. The final zone (zone 4) is the surrounding ambient I.S.M. (or ambient HII region) into which the shell propagates.

The evolutionary history of a stellar wind bubble consist of three stages. In the short formative stage (Falle 1975) the bubble is expanding so fast that the radiative losses in the gas do not have time to affect the gas dynamics of the flow. The second stage begins when the radiative losses in the swept up interstellar gas cause it to collapse into a thin shell. The radiative cooling in the shocked stellar wind is not significant in this stage because of the low density of the material. The major cooling mechanism in this region is by heat conduction across the contact discontinuity between zones 2 and 3. Evaporation of

the cool shell material across this conduction front provides the major source of mass increase in region 2. The final evolutionary stage begins when the radiative cooling becomes important in zone 2, leading to the collapse of this region and allowing the stellar wind to directly impinge on the outer shell. In the following discussion we will only be concerned with the last two stages of the evolution of a stellar wind bubble.

If the radiative losses in zone 2 are ignored Castor et. al. (1975) showed that the radius $R(t)$ and the velocity $V(t)$ of the outer isothermal shock is given by

$$R_s(t) = 427.9 L_{29}^{1/5} n_0^{-1/5} t_6^{3/5} \quad (\text{pc}) \quad 2-4$$

$$\text{and } V_s(t) = 253.6 L_{29}^{1/5} n_0^{-1/5} t_6^{-2/5} \quad (\text{km/sec}) \quad 2-5$$

where $L_{29} (\approx \frac{1}{2} \dot{M} V_\infty^2)$ is the mechanical wind luminosity in units of $1.0 \times 10^{29} \text{ J sec}^{-1}$, n_0 is the ambient I.S.M. density in particles m^{-3} and t_6 is the nebula age in 10^6 years.

This evolution is modified if radiative losses become significant in Zone 2. In the extreme case where the total losses from zone 2 exceed the energy

input from the stellar wind the zone will collapse and the stellar wind will impinge directly onto the outer shell. In this case the time dependence of shell radius (R_{SH}) will follow the formula given by Steigman et. al. (1975).

$$\text{i.e.} \quad R_{SH}(t) = \frac{(3 \dot{M} V_{\infty})^{1/4} t^{1/2}}{4\pi \rho_0} \quad (\text{m}) \quad 2-6$$

with shell having thickness

$$\Delta R/R = \left(\frac{\rho_0 C_{II}^2}{3\dot{M} V_{\infty}} \right) R_{SH}^2 \quad 2-7$$

According to Weaver et. al. (1977) L_b never exceeds L_W because of a regulatory mechanism which governs the losses in zone 2. They claim that if radiative cooling sets in, it will lead to a reduction in the mass flux into zone 2 from the HII shell. The decrease in mass flux across the front causes the density in zone 2 to decrease in such a way as to limit any further increase in the cooling.

This picture is complicated though if we consider situations in which the stellar wind bubble is not spherically symmetric. In this case there will be strong streaming motions in the shock stellar wind material because the wind impinges obliquely on the shell.

Kahn (1980) showed that the interface between the streaming stellar wind bubble and the HII shell, is subject to Kelvin-Helmholtz instability. He also showed that while turbulent mixing is produced by the instabilities, it is limited until the shocked stream has lost most of the thermal energy that it gained in the shock. Kahn finds that the shocked gas would have to cool to roughly 3×10^5 K before the shock layer would be destroyed by turbulent mixing. Thus it is possible the hot cushion of gas in zone 2 could be destroyed at some point during the evolution of the nebula. Without a detailed theoretical model to determine the point at which this occurs, all that we can do is consider the two extreme conditions i.e. the case where zone 2 does not collapse (the hot cushion model) and the case where it does (the direct impingement model). In the following discussion on stellar wind bubbles, both models will be discussed. Due to the fact that ionizing stars within HII regions are also the major source of stellar wind mechanical energy, it is almost certain that the majority of the stellar wind bubbles will begin their evolution within a HII region. In order to gauge the effect this stellar wind bubble would have upon the HII region a number of important time scales need to be considered.

As the ionized shell swept up by the evolving stellar wind bubble increases in mass, it will absorb an ever-increasing number of photons given by

$$S_i' = 4\pi R_{SH}^2 \Delta R \alpha_2 n_{SH}^2 \quad 2-8$$

where ΔR , R_{SH} and n_{SH} are the shell thickness, radius and density respectively.

If the shell were to continue increasing in mass it would eventually absorb all of the photons emitted by the star and the I-front would be trapped. The condition for trapping of the I-front is given by $S_i' = L_{uv}$. The radius at which this event occurs can be determined from the Rankine-Hugoniot conditions, which across the outer isothermal shock give:

$$\frac{n_{SH}}{n_{II}} = \frac{1}{(1 + (V_{SH}/C_{II})^2)} \quad 2-9$$

Combining this with equations 2.1 and 2.8 we have

$$\frac{R_{SH}}{R_S} = 3(\Delta R/R_{SH}) (1 + (V_{SH}/C_{II})^2)^{-1/3} \quad 2-10$$

Now equation 2-10 can be used along with the equations 2-4 and 2.5 describing (V_{SH}/C_{II}) at the point when the I-front is trapped. In the case of the hot-cushion model, (HC model) this is given by the transcendental equation

$$(1 + (V_{SH}/C_{II})^2) = 399.31 (1/(3\Delta R/R_{SH})) S_{48} L_{29}^{-3/2} n_{II}^{-1/2} \times (V_{SH}/C_{II})^{9/2} \quad 2-11$$

where S_{48} is the ionizing photon flux in units of 10^{48} photons sec^{-1} and $C_{\text{II}} = 11.77$ km/sec (the isothermal sound speed at a T_e of 10^4 K). In the case of the direct impingement model (DI model) the corresponding relationship is

$$(1 + (V_{\text{SH}}/C_{\text{II}})^2) = 1.0331 \times 10^5 (1/(3\Delta R/R_{\text{SH}})) S_{48} M_{23}^{-3/2} n_{\text{II}}^{-1/2} \times (V_{\text{SH}}/C_{\text{II}})^3 \quad 2-12$$

where M_{23} is the momentum deposition rate by the stellar wind onto the HII shell, in units of 1.0×10^{23} Newtons.

If we take values of $S_{48} = 40.0$, $L_{29} = 13.0$ and $M_{23} = 5.0$, typical for a small OB association (see Chapter 6) and assume that $\Delta R/R_{\text{SH}} = 0.2$ and $n_{\text{II}} = 10^6$ particles m^{-3} , we find that $(V_{\text{SH}}/C_{\text{II}}) \approx 1.46$ for the HC model and 0.12 for the DI model. These velocities correspond to the I front being trapped at $R_{\text{SH}}/R_S \approx 0.81$ and 1.8 respectively. The stellar wind bubble will reach this radius after 3.0 million years in the case of the HC model and 44.7 million years for the DI model. Clearly, for the assumed outer HII region density, the DI stellar wind shell will not trap the ionization front before it stalls. The HC stellar wind shell on the other hand, will trap the ionization front just prior to stalling.

If we take $V_{SH} > C_{II}$ which is the condition which defines the period in which the bubble has not stalled, then we find from equations 2-11 and 2-12 that the constraint for trapping the I front prior to stalling is

$$n_{II} L_{29}^3 S_{48}^{-2} > \frac{2.26337 \times 10^4}{(3\Delta R/R_{SH})^2} \quad 2-13$$

for the HC model, and

$$n_{II} M_{23}^3 S_{48}^{-2} > \frac{2.66823 \times 10^9}{(3\Delta R/R_{SH})^2} \quad 2-14$$

for the DI model.

The relative thickness of the stellar wind shell is a function of the age of the nebula (assuming a constant density shell) and is given by

$$\frac{\Delta R}{R_{SH}} = \frac{n_{II}}{3n_{SH}} \quad 2-15$$

This is equivalent to

$$\frac{3\Delta R}{R_{SH}} = (1 + (V_{SH}/C_{II})^2)^{-1} \quad 2-16$$

using equation 2-9. The criterion chosen to define times prior to the shell stalling, $V_{SH} \gtrsim C_{II}$ is equivalent to the condition $\frac{3\Delta R}{R_{SH}} \lesssim \frac{1}{2}$. Substituting this into

equations 2-13 and 2-14 we find

$$n_{II} L_{29}^3 S_{48}^{-2} \gtrsim 9.0535 \times 10^4 \quad 2-17$$

for the HC model and

$$n_{II} M_{23}^3 S_{48}^{-2} \gtrsim 1.06729 \times 10^{10} \quad 2-18$$

for the DI model.

Taking the values for L_{29} , M_{23} and S_{48} used previously we see that in order to trap the I-front prior to the bubble stalling $n_{II} \gtrsim 1.24 \times 10^6$ particles m^{-3} for the HC model and $n_{II} \gtrsim 1.37 \times 10^{11}$ particles m^{-3} for the DI model. The values on the right hand side of equations 2-17 and 2-18 are upper bounds, as the shell will not stall until some time after $V_{SH} = C_{II}$

In spite of this we can conclude that if the HC stellar wind model is applicable to evolved HII regions, all except those expanding into very low density media, will eventually have the supply of ionizing photons cut off by the formation of a stellar wind shell.

After the I-front is captured by the expanding shell the nebula would consist of slowly expanding ionization-

limited shell-like HII region surrounded by an outer dense shell of neutral gas. A fossil HII region would exist outside the main shell for a short while until it is swept up by the expanding bubble.

Alternatively, if the DI stellar wind model is correct, then the shell built up by the stellar wind would never cut off the ionizing photons to the surrounding HII region. In this case the stellar wind shell would appear as an intensity enhancement immersed within a larger HII region. This type of morphology is not observed in the HII regions in the L.M.C., suggesting that the DI model is inapplicable.

An additional way in which we can compare the HC and DI models is to compute the density of the ionized gas in the swept up shell. This is because the HII shell is in pressure balance with the shocked stellar wind in the HC model, while in the case of the DI model the shell is in pressure balance with the ram pressure of the stellar wind. The hot cushion of shocked stellar wind material acts to store a considerable fraction of the energy deposited by the wind as thermal energy. The pressure produced by the pad of hot gas is the same as the ram pressure of the material passing through the inner shock, (at radius r_{in}) typically located at 0.3 of the outer radius (at radius r). Hence the ratio of pressures between the H.C. and D.I. models is approximately

$(r/r_{in})^2 \approx 11$. In more detail, the total internal energy in the shocked stellar wind gas of the HC model is given by

$$E_b = \frac{3}{2} PV = \frac{3}{2} \left\{ \frac{4}{3} \pi (R_2^3 - R_1^3) \right\} p \quad 2-19$$

$$\approx 2\pi p R_2^3 \quad \text{since } R_2 \gg R_1$$

Where R_1 and R_2 are the inner and outer boundaries of zone 2 respectively and p is the pressure between these two boundaries. The pressure in this zone can be directly equated to the pressure (both thermal and turbulent) of the HII shell which is separated from zone 2 by a contact discontinuity. The energy balance in zone 2 is given by

$$\frac{dE_b}{dt} = L_w - 4\pi R_2^2 \frac{p dR_2}{dt} - L_b \quad 2-20$$

where the first term on the right hand side is the stellar wind mechanical energy ($= \frac{1}{2} \dot{M} V_\infty^2$), the second term is the work done by zone 2 on the surrounding HII shell, and the last term is the energy lost from this zone (through radiation and conduction).

Differentiating equation 2-19 with respect to time and equating it to equation 2-20 we find that

$$\frac{dp_{II}}{dt} + \frac{5}{R_2} \frac{dR_2}{dt} p_{II} = \frac{\{(1-\gamma) L_w\}}{2\pi R_2^3} \quad 2-21$$

The derivation of equation 2-21 assumed that the cooling in zone 2 could be approximated by $L_b = \gamma L_w$ where $\gamma = \text{const}$, where in fact L_b varies from 0 to roughly L_w during the lifetime of this region. Taking $\gamma = 0.5$ we find from equation 2-21 that

$$\rho_{\text{II}} = \frac{5}{4} \frac{L_w t R_{\text{SH}}^{-3}}{C_{\text{II}}^2} \quad 2-22$$

where t is the age of the nebula. In terms of number density this equation is written as

$$n_{\text{HC}} = 2.80308 \times 10^{10} L_{29} t_6 R_{\text{SH}}(\text{pc})^{-3} \quad 2-23$$

In contrast the number density in the shell of the DI model is given by

$$n_{\text{DI}} = 6.02689 \times 10^7 M_{23} R_{\text{SH}}(\text{pc})^{-3} \quad 2-24$$

2.4 The Off-Centre Accretion Model

Dopita (1981) proposed a third type of model for forming the large HII rings observed in the L.M.C. In brief summary, the evolution described by this model begins with a massive OB Star (or stars) forming in the outer regions of a large molecular cloud ($\sim 10^5 M_{\odot}$) which is undergoing spherical symmetric collapse. The massive star undergoes substantial mass-loss, with the resulting stellar wind interacting with the collapsing cloud. Dopita showed that the collapsing cloud could provide enough ram pressure to confine the mass-loss bubble

formed by the OB star. A quasi-static flow situation is set up where at each point in the bubble, the net ram-pressure of the collapsing cloud (directed towards the mass-loss centre) just balances that due to the stellar wind ram pressure. This situation persists for a considerable period of time due to the fact that the free-fall time for gas located at the outskirts of the collapsing cloud is comparable to hydrogen-core burning lifetime of the star (or stars) producing the mass-loss.

The quasi-static flow which is generated consists of the hypersonic free-flowing stellar wind, together with a thin layer of shocked stellar wind gas mixed to a greater or lesser degree with a surrounding ionization bounded and ram-pressure confined HII region. Outside this HII region is a layer of shocked cloud gas which is accreting into the surface of the bubble. This is separated from the infalling cloud gas by an isothermal accretion shock.

Due to the off-centre location of the bubble there will be considerable streaming of the shocked stellar wind gas away from the accretion centre of the molecular cloud along the inside of the bubble. This is matched by a flow in the opposite direction along the outside of the bubble by accreted cloud gas.

The HII/HI and HII/shocked stellar wind interfaces in this model are both Rayleigh-Taylor and Kelvin-Holmholtz unstable leading to the development of surface irregularities and mixing. The intrusion of cool HII gas into the shocked stellar wind gas will cause it to reach an equilibrium temperature which is only a small fraction of its past shock temperature. It is expected that this would produce a substantial compression and cooling of the shocked stellar wind gas, resulting in the thickness of the hot shocked layer being negligible in this model. This would mean that this model would show the same characteristics we cited previously for the DI model. Thus we should be able to use the discriminants used to differentiate the H.C. model from the D.I. model to see if the off-centre accretion theory is applicable to the L.M.C. HII regions.

An alternative test of this theory would be to compare the predicted velocity structure with that observed for the L.M.C. HII regions. The HII gas on the inner-side of the accreted HI shell is accelerated by the stellar wind impinging on it. This gas would be accelerated up to the same velocity as the streaming shocked stellar wind gas, if it were not for the fact that it is functionally coupled (through boundary instabilities) to the outer HII gas. This has the effect of limiting the HII velocities to a much lower terminal velocities. The main result of this velocity streaming would be to produce

line-splitting of emission lines observed in these HII shells. The splitting would be greatest just inside the brightest HII shell, found on the accretion-centre side of the stellar wind bubble. Moving towards the centre of the bubble the line splitting would disappear only to reappear again just inside the opposite HII shell of the bubble. This velocity structure is quite distinct from that produced by a classical stellar wind bubble which is characterised by uniform expansion about the centre and this can be used to distinguish between these theories (Dopita 1981).

In order to prove or disprove the various stellar wind bubble theories we must have information on the integrated ionizing flux and mass-loss rates from the stars located in the L.M.C. HII regions. These parameters will be determined in the following 4 chapters of this thesis.

CHAPTER 3

SPECTRAL CLASSIFICATION OF O AND B STARS IN THE L.M.C.

3.1 The M.K. Spectral Classification System

The M.K. Classification System (Morgan, Keenan and Kellman, 1943, Johnson and Morgan, 1953) makes use of the fact that the stellar effective temperature (T_{eff}) and surface gravity (g_{eff}) are the two dominating factors which characterise the line spectrum of stellar atmospheres. The system seeks to set up a classification sequence which differentiates stars of different T_{eff} and g_{eff} .

The method of classification is based solely on a comparison between the spectrum of an observed star and a set of standard stellar spectra. Each standard star is differentiated from its adjacent subclass by line strengths and ratios specific to that subclass. A star is spectrally classified by comparing appropriate line strengths and ratios with a grid of standard spectra.

The M.K. Classification system uses widened slit spectra recorded on photographic plates. The spectra are taken with an inverse dispersion of 125 Å/mm (at H γ) and cover the wavelength region 3800–4900 Å. Spectra obtained by the same method can be directly compared with those published in the M.K. catalogue.

The original M.K. spectral system was modified in 1973 by Morgan and Keenan (MK73) and again in 1978 by Morgan, Abt and Tapscott (MK78). Both modifications involved changes in the canonical list of standard stars which define the M.K. system. The MK73 standards were chosen

to increase the internal selfconsistency of the grid of standard stars. The high internal accuracy of this grid enabled precise interpretations of spectral types for 09 stars and later. Only four stars defined spectral types 08 and earlier and there were no luminosity classes in this spectral range.

In the MK78 system each bin of the classification scheme is defined by a single star. This removes the ambiguity which results from two or more stars defining a spectral bin. The grid of standard stars is much more comprehensive and luminosity classes are extended to include the 06 to 08 stars.

If spectra are obtained using a different instrument and/or dispersion, then a grid of standard stars have to be observed using this new configuration, so that a set of standard spectra are available. The stars chosen to make up this grid should preferably be the ones which define the M.K. system. The problem with using these stars is that they are not always visible from the southern hemisphere.

These problems can be circumvented if standard stars can be observed from a spectral catalogue of stars visible in the southern hemisphere. The limitations of classification accuracy are set by the catalogue used. Errors result, from using peculiar or mis-classified stars as spectral standards, although this can be minimised

by observing more than one star for each spectral subtype.

3.2 The Adopted Method of Spectral Classification for O and B Stars

Almost all of the luminous stars associated with the L.M.C. HII regions observed in this thesis, are O or early B stars. The initial method which was adopted to spectrally classify these stars, aimed at creating a grid of standard stars, covering all spectral types and luminosity classes between O4 and B5. Stars in this grid were selected from the Garrison, Hiltner and Schilds' (GHS) Spectral catalogue of 1,113 OB stars, visible from the southern hemisphere (GHS 1977). The stars in this catalogue were classified using the primary spectral standards of the MK73 system. These standards were supplemented with secondary standards in classes for which the only MK standards were too far north for GHS to observe. Some cluster sequences in II Scorpio and N.G.C. 2244/46 were also used.

Unfortunately, not enough standards were observed to form a complete spectral grid. This meant that a modified classification scheme had to be adopted to overcome this difficulty.

The method which was adopted for classifying the L.M.C. OB stars is set out below. Since O stars were classified by a different criterion to B stars, the classification schemes are treated separately.

3.2.1 O Stars

The spectra of O stars are distinguished from B stars by the presence of the HeII 4542 line. The O stars are subdivided into spectral types O3 to O9.5 by using the ratio of the HeI λ 4471 line to the HeII λ 4542 line. HeI λ 4471 progressively decreases in strength in going from O8.5 stars to earlier O stars, disappearing by O3. HeII λ 4542 first appears weakly in O9.5 stars and increases in strength towards O4. Both lines have very similar gravity dependence, as both are broadened by the Stark effect, making the ratios of the two lines a good T_{eff} index.

Conti and Auschuler (1971) showed the quantitative ratio of the equivalent widths of these two lines could be used to determine the spectral type of O stars. Table 3.1 shows the relationship they obtained between $W(4471)/W(4542)$ and MK spectral type. This was adopted as the criterion used to determine the spectral type for O4 to O9 stars.

The weakness of the HeII 4542 line in O9 and O9.5 stars makes it difficult to use the HeI/HeII ratio as a spectral discriminant in low S/N spectra. The method used to classify these stars was the same as used for B stars, and will be discussed in Section 3.2.3.

3.2.2 Luminosity Indicators for O Stars

The difference in M_V between supergiant and main sequence stars is about 3 magnitudes at O9.5, 2 magnitudes

TABLE 3.1.

MK Spectral type for O Stars based on the
W(4471)/W(4542) Ratio W'

<u>SPECIAL TYPE</u>	<u>LOG W' LOWER</u> <u>LIMIT</u>	<u>LOG W' UPPER LIMIT</u>
09.5	+0.45	-
09	+0.30	+0.45
08.5	+0.20	+0.30
08	+0.10	+0.20
07.5	0.00	+0.10
07	-0.10	0.00
06.5	-0.20	-0.10
06	-0.30	-0.20
05.5	-0.45	-0.30
05	-0.60	-0.45
04	-	-0.60
03	Helium I λ 4471 absent	

stars may not be discernable in low S/N or lower dispersion spectra.

Allowing for this problem, Walborn's nomenclature was adopted as a luminosity discriminant for stars earlier than O9. The categories Of, O(f) and O((f)) correspond to the MK luminosity classes I, III and V respectively, for stars in the spectral range O6 to O8.

Conti and Auschuler (1971) suggested that the best luminosity discriminant for late O stars (O8 - O9.5) is the ratio of the SiIV $\lambda 4089$ line to the HeI 4143 line. It was not possible to use these lines, because spectra taken of the L.M.C. OB stars were limited to wave lengths longward of $4100 \text{ \AA}$ (see 3.3.2). Alternative luminosity indicators are, the CIII $\lambda 4650$ /HeII 4686 ratio which can be used for O9 to O9.5 stars, and the HeII 4542/HeI 4387 ratio which is suitable for O8 to O8.5 stars. These ratios make use of the positive luminosity dependence (for a given spectral type) of CIII $\lambda 4650$, the negative luminosity dependence of HeII $\lambda 4686$ and HeI $\lambda 4387$, and the luminosity independence of HeII $\lambda 4542$. These two ratios were tentatively adopted as luminosity indicators for the spectral range O8 to O9.5.

3.2.3 O9 to B5 Stars

For stars later than O9 there is no single line ratio which can be used to unambiguously differentiate spectral classes. In this region the emphasis of classif-

ication must be placed on an intercomparison of line strengths in observed spectra with standard spectra.

As stated previously the grid of standard spectra obtained, was not comprehensive enough to allow accurate interpolation of spectral types. Information from the MK78 catalogue was used to supplement this spectral grid by outlining the general changes in line strengths with spectral type. That is to say, interpolation was carried out with the existing observational grid using the MK78 catalogue as a guide.

The stellar effective temperature in the O9 to B5 spectral region, ranges from $30,000^{\circ}\text{K}$ for an O9 supergiant star, down to $14,000^{\circ}\text{K}$ for a B5 supergiant. Such a large temperature change produces variations in the general excitation and ionization state of atoms in the stellar atmosphere. This is reflected in changes in line intensities as the effective temperature of the star increases.

In the spectra of early B stars there are a number of lines exhibiting the above characteristics, which can be used to limit the spectral type of the stars. Lines which are used in this manner are listed in Table 3.2, together with a description of the spectral range they cover.

TABLE 3.2

Spectral Discriminants for O9 to B5 Spectra at
MK dispersion. (λλ 4100-4900)

SPECIES	WAVELENGTH (Å)	CHARACTERISTIC
1. He II	(4542 (4200	These lines distinguish O from B stars, being found only in O stars
2. He I	4471 4387 4143) 4120)	These lines all peak in intensity at B1 or B2. This line is visible from O4V to B5V This line is visible from O8V to B3V These lines are visible from O9V to B3V The lines have a negative lumi- nosity dependence which is strongest at B1, B2.
3. He II	4686	This line is present in the spectra of O4 to BOV stars. It has a negative luminosity dependence for O9 and BO stars, disappearing in BOIa stars
4. CIII	4650	This line is found in spectral types O7 to B1V, reaching maximum intensity at BOV. Its intensity increases with increasing luminosity
5. CII	4267	This line is found in all luminosity classes at B2. It is also present in all luminosity classes at B3 except V and in the spectra of B5Ia, Ib stars
6. OII	4317-20	This line is present at B1. It decreases in intensity in going from Ia to II-III and is absent at V.

TABLE 3.2 CONT'D.

SPECIES	WAVELENGTH	CHARACTERISTICS
OII	4415-17	This line is present in both B0 and B1 stars being strongest at B1. It decreases in intensity in going from I to V
SiIII	4553	This line first appears weakly at 09.5Ia. It is present at B0 for luminosity classes Ia,Ib. It is also found in B1 and B2 stars for luminosity classes - Ia to III, decreasing in strength with luminosity
SiII	4128-3	This line is found in B3 and B5 stars decreasing in intensity from luminosity class I to V
MgII	4481	This line first appears at B1, increasing in intensity towards B5 stars, where it is comparable in strength to HeI λ 4471

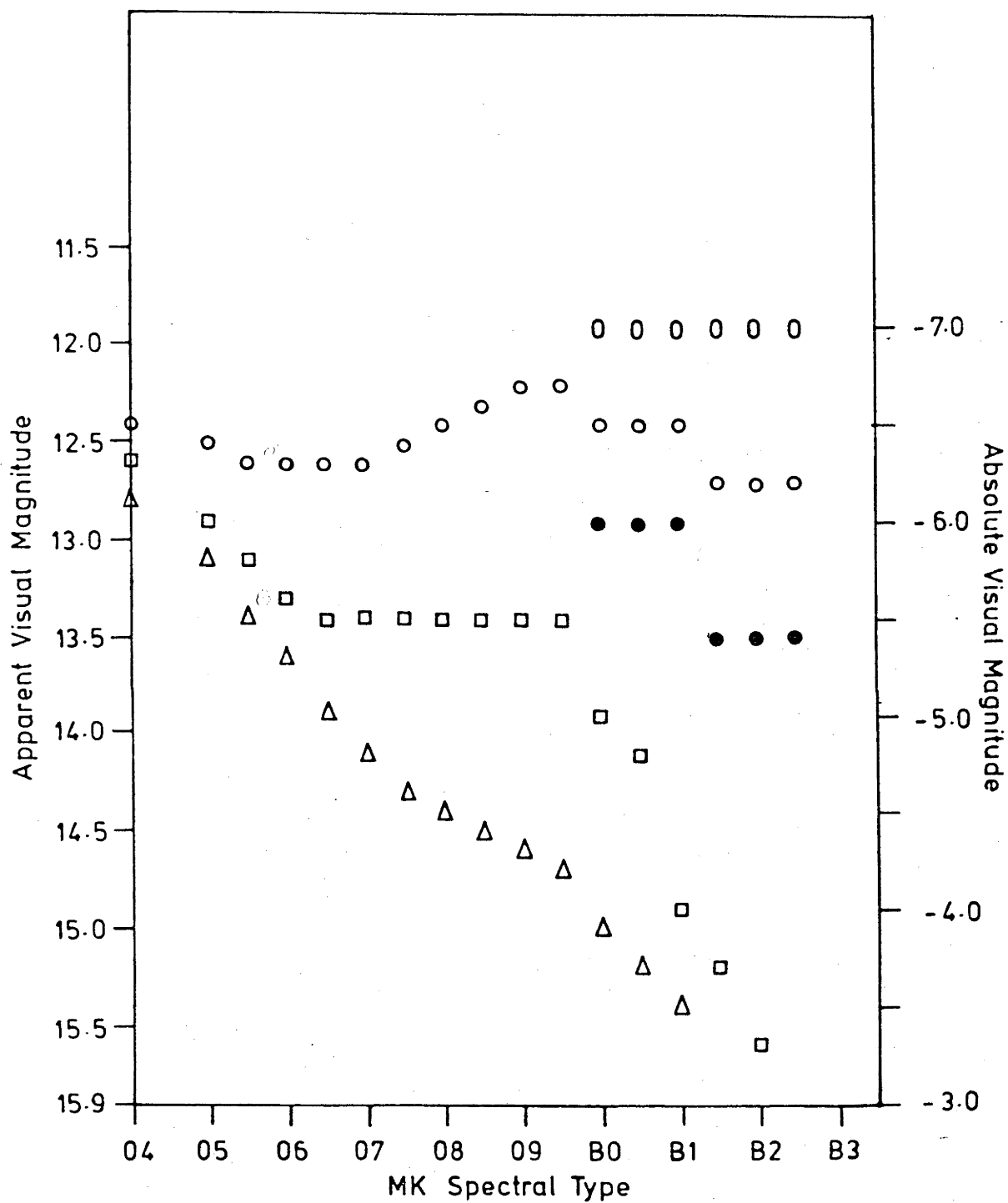
3.2.4 The Luminosity Class of LMC OB Stars

Figure 3.1 shows the relationship between apparent visual magnitude (M_V) and MK spectral type which can be used to determine the luminosity class of LMC OB stars. The graph was obtained by transforming the known relationship between absolute visual magnitude (M_V) and spectral type to the distance of the L.M.C. The transformation was made by taking the true distance modulus (μ_0) to the L.M.C. as 18.4 magnitudes and the mean extinction as 0.5 magnitudes. For stars earlier than B0, the absolute magnitude scale of Panagia (1973) was adopted. Panagia's scale is derived from Conti and Alschuler's relationship between M_V and spectral type (Conti and Alschuler, 1971). This scale is based on absolute magnitudes obtained from OB stars of known spectral type, in clusters of known distance moduli. A comparison of Panagia's scale with that of Walborn (1973), which was obtained using the same technique, shows that the two scales agree very well for luminosity class V and III stars. The scales differ, on average, by 0.1 magnitudes for III stars and 0.3 magnitudes for V stars, with the greatest discrepancy to being 0.6 magnitudes for O4V stars.

Lesh's absolute magnitude scale (Lesh, 1981) was adopted for luminosity classes III and V having spectral types later than O9.5, while Walborn's scale (Walborn 1973) was used for Ia, Iab and Ib supergiants.

FIGURE 3.1

The apparent visual magnitude and absolute visual magnitude as a function of M.K. spectral type.



There is an intrinsic uncertainty of 0.5 magnitudes in these M_V values due to uncertainties in the distance moduli of Galactic OB associations. This uncertainty makes it difficult to distinguish luminosity classes I and V for stars earlier than O6, and classes V and III for stars later than B0.5. For stars outside these spectral regions this method is capable of determining the luminosity of LMC OB stars to within plus or minus a half of a luminosity class.

There is always the possibility that foreground dwarf O9 to B5 stars can masquerade as L.M.C. supergiants (see Section 3.2.6 below). This means that the apparent visual magnitude method described above must be supplemented by luminosity dependent line ratios to confirm luminosity classes.

3.2.5 Luminosity Indicators for B0 to B5 Stars

Listed below are line ratios and strengths which can be used to distinguish luminosity classes from B0 to B5. Considering each spectral type separately, for:

- (1) B5: If $MgII \lambda 4481 \leq HeI \lambda 4471 \Rightarrow$ (luminosity class) I
and If " " $<<$ " " $\Rightarrow$ V
- (2) B3: If $HeI \lambda 4121 \leq HeI \lambda 4144 \Rightarrow$ Ia, Ib
and If " " $<<$ " " $\Rightarrow$ IV, V

(3) B2: i) If HeI λ 4121 < HeI λ 4144 $\Rightarrow$ I
 " " " $\approx$ " " $\Rightarrow$ III
 and " " " $\lesssim$ " " $\Rightarrow$ V

 ii) If HeI λ 4387 > SiIII λ 4553 $\Rightarrow$ I
 " " " >> " " $\Rightarrow$ III
 and " " is the only line present $\Rightarrow$ V

(4) B1: i) If HeI λ 4144 $\lesssim$ OII λ 4415, 17 $\Rightarrow$ I
 " " " > " " $\Rightarrow$ III
 " " " >> " " $\Rightarrow$ V
 and ii) If HeI λ 4378 $\approx$ OII λ 4415, 17 $\Rightarrow$ I
 " " " > " " $\Rightarrow$ III
 " " " >> " " $\Rightarrow$ V

and for

(5) B0: i) If HeII λ 4686 absent, CIII λ 4650 strong $\Rightarrow$ I
 If He 4686 << CIII 4650 $\Rightarrow$ Ib
 If " < " " $\Rightarrow$ V
 and ii) If HeI/SiIV 4116-20 > HeI λ 4144 $\Rightarrow$ I
 " " " " $\approx$ " " $\Rightarrow$ III
 " " " " < " " $\Rightarrow$ V

3.2.6 The Spectral Classification Scheme

The final spectra classification scheme adopted consisted of the following:

- (1) Firstly the HeII λ 4542 line was used to distinguish O and B stars

- (2) If the star was classified as an O type, and HeII $\lambda 4542$ was clearly visible, then the equivalent width of this line and HeI $\lambda 4471$ were measured using a planimeter. This procedure is accurate in spectra obtained with photon-counting instruments, as is the case here. The ratio of the equivalent widths was then used to determine the spectral type of the star from Table 3.1. The luminosity indicators in the spectrum were noted to see if they confirmed the luminosity class given by the apparent visual magnitude of the star.
- (3) If the star was classified as an O star but HeII $\lambda 4542$ was too weak for accurate equivalent widths to be measured, then it was assumed to be O9 or later and classified in the same manner as the B stars.
- (4) If a star was classified as O9 or later then the line strengths in Table 3.2 were used to refine further the classification within this spectral range. The lines in Table 3.2 could usually be used to narrow the classification down to a choice between the two spectral types.

- (5) The final decision on the spectral type was made by taking into account the strength of the lines which were the main spectral indicators, and by comparison of stellar spectra with the appropriate standard spectrum. Care was taken to allow for the effects of the luminosity on line strengths.
- (6) The luminosity class of the star, indicated by the spectrum, had to be consistent with the class indicated by the apparent visual magnitude of the star, for it to be considered a member of the L.M.C. This is necessary as it is possible a foreground galactic B dwarf to have the same magnitude as a super-giant B star in the L.M.C.

Table 3.3 lists the stars which were used to create the grid of standard spectra for O9 to B5 stars. The photoelectric V magnitude is given for each star, together with the M.K. spectral type assigned to it by Garrison, Hiltner and Schild (1977).

3.3 THE OBSERVATIONS OF L.M.C. OB STARS

3.3.1 Constraints on the Observations

A number of constraints were important in determining the form of the observing programme. The primary aim of the observations was to determine the

TABLE 3.3

Standard Stars: 09-B5

HD Number	V Magnitude	MK Spectral Type
1. HD 166832	8.3	B5 II - III
2. HD 164798	8.7	B2 II
3. HD 163899	8.4	B1 II
4. HD 167663	8.4	B1 II
5. HD 156134	8.0	B0Ia
6. HD 73420	9.6	B0III
7. HD 160065	8.4	09.5Iab
8. HD 92850	8.2	09.5Ib
9. HD 165132	8.3	09.5 V
10. HD 149452	9.7	08.5

spectral type of the stars contributing the major share of the photo-ionizing photon flux of the L.M.C. OB clusters. Stars with spectral types later than B0.5 emit a negligible flux in the amounts of Lyman continuum when compared to O stars. This means that we only need to obtain the spectra of stars earlier than B0.5 to determine the integrated ionizing photon flux for the L.M.C. clusters. According to Figure 1, this includes all blue stars which are brighter than 15.2 magnitudes (M_V). The relative contribution of a star brighter than this magnitude limit, to the total ionizing flux of its cluster, depends on the distribution of stars in spectral type and luminosity. For example, if a cluster contains an O6V star, then an O9V star would only contribute an extra 10% to the total U.V. flux. Including O9V stars in the total U.V. flux of the cluster would result in an effective magnitude limit of 14.6. Likewise, if a cluster contains an O7I star, then the faintest star we need to observe would be 14.3 magnitudes, corresponding to an O7.5 star, etc.

In addition to measuring the integrated U.V. flux from the cluster, another aim was to determine the integrated mass-loss for the cluster. The mass-loss rate of a star, to the first approximation, depends on the luminosity of that star. Only OB stars with bolometric magnitudes (M_{bol}) in excess of -6 show significant mass-loss (i.e. $\geq 10^{-7} (M_{\odot})$ Solar masses per year). (Snow and Morton, 1976).

This includes all stars brighter than approximately 15.5 magnitudes in V, corresponding to a B1V star. Again, this magnitude limit will be brighter if the L.M.C. OB cluster is dominated by a more luminous star. For example, if a cluster has an O8If star with a typical mass-loss rate of $1.0 \times 10^{-5} M_{\odot}/\text{year}$, then only the stars brighter than 13.5 in V would have to be observed. This magnitude limit corresponds to an O9.5III star which has a mass-loss rate of approximately $1.0 \times 10^{-6} M_{\odot}/\text{year}$.

3.3.2 The Observations

The objective of the observing programme was to determine the MK spectral types and hence the U.V. Lyman Continuum fluxes and mass-loss rates for all bright blue stars found within the boundaries of the L.M.C. sample HII regions. Typically, this required the observation of between one and a dozen stars in each cluster.

Spectra were obtained using the one-dimensional, red sensitive Photon Counting Array (ID.P.C.A.) (Stapinski, Rodgers and Ellis, 1981). This instrument was used at the cassegrain focus of the Mount Stromlo 1.88 metre telescope together with its Boller and Chivens spectrograph.

During the 1978/80 L.M.C. observing season, i.e. October to February 1979/80, stellar spectra were observed using a 600 line/mm grating blazed to $5,000 \text{ \AA}$

in the second order. This gave spectra with a mean inverse dispersion of $56 \text{ \AA/mm}$ ($0.7 \text{ \AA/channel}$ of the detector), allowing a spectral coverage of approximately $700 \text{ \AA}$. This dispersion was chosen as the highest which could feasibly be used and still meet the constraints of telescope time and S/N requirements. The aim was to be able to obtain a S/N of 15 to 20 for the total counts/channel, in an exposure time of 30 minutes or less. With the instrument configuration indicated above, it was possible to obtain 350 counts/channel (at $4,400 \text{ \AA}$) on a 12.5 (V) magnitude O star, in an exposure time of 30 minutes using a $225 \text{ }\mu$ slit (1.37 arcseconds on the sky).

The effective magnitude limit for these observations was 13.5 magnitudes ($S/N \gtrsim 10$) in the V band. The S/N of all the spectra were increased at the expense of resolution by double binning the data in the reduction procedure. This produced spectra with $S/N \gtrsim 15$ for stars as faint as 13.7 magnitudes in V.

The spectra were centred at $4,550 \text{ \AA}$ and covered the wavelength range $\lambda\lambda 4200\text{--}4900 \text{ \AA}$. This range was chosen to include the lines of HeI 4471, HeII 4542, 4686; CIII 4650 and NIII 4642 $\text{\AA}$, essential for classifying O and early B stars in the MK system. Useful lines for classification which were outside this range included SiIV 4089, 4116, HeII 4200, HeI 4120, 4144 and SiII 4128, 4131. The SiIV and HeI lines can be used to form line ratios which

are useful as luminosity indicators in late O and early B stars (see Section 3.2.5). The SiIII line is useful in distinguishing B3 and B5 stars, while the HeII line can be used as a support for the presence of the HeII 4542 line in late O stars.

During the 1980/81 observing season stellar spectra were obtained using a 600 line/mm grating blazed to 5,000 Å in the first order. This change was made to extend the magnitude limit for spectra to fainter stars and increase the wavelength coverage. The effective spectral range went from 4,100 Å (Hδ) to 5,000 Å, bringing within reach all but the SiIV 4089 of the lines mentioned above.

A summary of the actual instrumental configuration used for the observations is given in Table 3.4 together with the date of observation.

The observations prior to December 15th 1980 were made with the "short" red 1D P.C.A. The detector in this device consisted of two parallel linear reticon arrays physically separated by 2.5mm. Each reticon had 25 x 314 μ pixels which were double binned by the photon event centring logic to give 1024 12.5 μ pixels, in the spectral direction.

After December 15th 1980 the "long" red 1D P.C.A. was used. This had twice as many pixel in the spectral direction as the "short" array.

TABLE 3.4

Summary of Spectral Observations

	Date	Grating	Blaze Wave- Length (1st Order)	Order	(A/mm) Inverse Dispersion	Slit	Arc
1.	12-15/11/79	600	10,000	2	56	225	Helium
2.	14-17/12/79	"	"	"	"	"	"
3.	21-22/01/80	"	"	"	"	"	"
1.	11-12/10/80	600	5,000	1	112	200	Argon ^a
2.	12-17/12/80	"	"	"	"	300	"
3.	31-04/01/81	"	"	"	"	300 31/01 400 02/04	"
4.	02-05/04/81	"	"	"	"	150 - 250	"
5.	02-05/01/81	"	"	"	"	25 - 150	"

The slit aperture of the spectrograph was limited by a dekker producing two windows 0.7mm wide separated by 10.0mm. The dekker allowed the stellar spectrum to be projected onto one of the reticon arrays and the sky spectrum, 144" to the west of the star, to be projected onto the other.

It soon became apparent that (as a result of the electrostatic image tube stack distortion) one of the reticon arrays ran more efficiently than the other. As a result, the stellar spectra were always observed using this reticon array to maximize the S/N. This meant that it was impossible to obtain perfect sky subtraction, but the maximization of S/N was considered more important.

Flat Fields were taken at least once during such observing run to remove the pixel to pixel gain variations, which were of the order of 3% (Rodgers and Stapinski, 1981). The flat fields were obtained by pointing the 1.88 metre telescope towards an evenly illuminated patch of dome and exposing for 3,000 to 5,000 seconds at a count rate of 2-3 photons/pixel/second.

Arcs were also taken at regular intervals during the observation to wavelength calibrate the spectra. The arcs were observed after each second star, if the telescope was set on one cluster, or at each position, if the telescope was moved from cluster to cluster.

3.3.3 The Reduction Procedure

All spectra were reduced on the Vax 11/780 computer at M.S.S.S.O. using the programme SPECT, now incorporated into the PANDORA software suite. Many people have contributed to the development of this package, amongst whom must especially be mentioned: Albert Bosma, Sue Simkin, Andrew Pickles and Bruce Cogan.

Each spectrum was divided by the normalized flat field, taken during its observing run, then wavelength calibrated. The wavelength calibration was done by fitting a polynomial to the wavelength, channel number pairs of unblended arc lines of helium and argon.

To a first approximation, changes in the dispersion produced by the electrostatic pin cushion distortion in the image tube stack of the P.C.A. can be represented by a second order polynomial. This indicates that a third order polynomial should be adequate for fitting wavelength, channel pairs.

The helium arc used in the 1979/80 magellanic season had only 5 to 7 useable lines. This meant that a third order polynomial fit was theoretically possible, but the small number of points made this fit uncertain. This was especially true of wavelength regions far from the arc lines. These regions correspond to wavelengths shortward of 4387 Å and longward of 4921 Å at either end of the spectrum, and between 4471 and 4686 Å where there were

no arc lines. The wavelength uncertainty in the two regions $\lambda 4387$ to $4471 \text{ \AA}$ and $\lambda 4686$ to $4921 \text{ \AA}$ was approximately 0.7 channels (3σ), corresponding to $0.5 \text{ \AA}$. The error outside these regions was larger.

The situation was more satisfactory for the low dispersion data calibrated by the argon arcs, where 9 to 16 evenly distributed arc lines were useable, resulting in a reasonable wavelength fit between 4100 and $4900 \text{ \AA}$. In this wavelength region the uncertainty of the fit was approximately 0.4 channels (3σ) corresponding to $0.6 \text{ \AA}$ (at the dispersion of 1.60 ± 0.01 (1σ) $\text{\AA}/\text{pixel}$ used).

Following the flat fielding and wavelength calibration of the spectra, all spectra observed with a slit width less than 225μ were smoothed with a gaussian filter designed to broaden the instrumental profile to that produced by a 225μ slit. The σ of the instrumental profile (σ_{slit}) was obtained by assuming that it was half the project slit width falling on the reticon array detectors. The σ of the smoothing gaussian (σ_{gauss}) was given by

$$\sigma_{\text{gauss}}^2 = \sigma_{225}^2 - \sigma_{\text{slit}}^2$$

where σ_{225} was the dispersion of the instrumental profile produced by a 225μ slit. Only the standard MK spectral standard stars needed to be treated in this way, so that they could be used as MK spectral comparison stars with the L.M.C. data.

All data taken in the 1979/80 Magellanic observing season was double binned to increase the S/N and make it compatible with that obtained in the 1980/81 Magellanic season.

Finally, all spectra were smoothed by a gaussian with a dispersion of 1.0 detector channel, to reduce the noise in the data.

3.4 THE SPECTRAL TYPES OF L.M.C. OB STARS

3.4.1 Spectral Types

Table 3.5 lists all stars with photoelectric V magnitudes less than 14.5, found within the 12 sample HII regions in the L.M.C. Column 1 gives the DEM number of the HII region (Davies, Elliott and Meaburn, 1976); Column 2 the number of the Lucke OB Association (Lucke 1972); Column 3 the Sanduleak star number (Sanduleak 1969); column 4 the number assigned to the star by Lucke; Column 5 the number of the star used in maps in Appendix A; and Column 6, 7 and 8 the photoelectric V, B-V, and U-B magnitudes and colours (Rousseau et al., 1978).

Column 9 of Table 3.5 lists the spectral type of 25 stars which were classified using the method described in 3.2.6. Eleven of these stars have spectral types determined from the HeI $\lambda 4471$ to HeII $\lambda 4542$ ratio while the remaining 14 have spectral types O9 to B6.

TABLE 3.5

DEM NO.	LUCKE ASSOC	LUCKE NAME	LUCKE NO	STAR NO	V	B-V	U-B	Sp Type	Other	Comment
301	114	SK-67°250	2	1	12.68	-0.17	-1.03	08.5 I(f)	-	
			1	2	13.60	-0.19		05.5V	-	
			7	6	13.66	+0.08		08.5 III+B?	-	
			4	3	13.93	-0.19		09:III-V	-	
			11	10	14.22	-0.25		08:V	-	
			8	12	14.46	-0.25		-	-	
			13	17	12.48	+1.75		-	K5-M5I	?M
			14	18	14.29	+1.56		-	-	F
			17	22	14.27	+0.68		-	-	F
11	3		4	1	13.74	-0.18		08.5III-V	-	
			5	2	13.81	-0.19		07 V	-	
			2	4	14.61	-0.17		09:V	-	
196	55	SK-67°108	4	1	12.56	-0.20	-1.05	06 III	-	
		SK-67°112	-	8	11.90	-0.13	-0.98	09 I	-	
			5	5	14.33	+0.61		-	-	F
			-	2	-	-		09-08:	-	

TABLE 3.5 (cont.)

DEM NO.	LUCKE ASSOC	NAME	LUCKE NO	STAR NO	V	B-V	U-B	Sp Type	Other	Comment
226	-		1ab	12.20	-			08.5I+B?	-	
			4a	13.68	-0.13	-0.94		07V+B?	-	
			2	12.88	+1.25	+1.44		-	G8-K1Ib	F
			3	13.47	+1.03	+0.82		-	-	
			13ab	-	-			-	09.5V	
			4b	-	-			-	BOIII	
25		SK-70 ⁰ 13	1	12.29	-0.12	-0.91		08I	B1	
137	43	SK-65 ⁰ 47	18	12.51	-0.18	-1.06		05.5If	BO	
			15	12.61	+1.67			-	(K4-5I)	M?
		SK-65 ⁰ 50	14	13.19	0.06			BO.5Ib	-	
		SK-65 ⁰ 45	16	14.54	-0.13			-	WN 3	
			1	13.16	+1.09			-	-	F
			13	13.74	+0.07			Early B Star	-	F?
			2	14.19	+0.04			-	-	F?

TABLE 3.5 (cont.)

DEM NO.	LUCKE ASSOC	NAME	LUCKE NO	STAR NO	V	B-V	U-B	Sp Type	Other	Comment
31	6	SK-67°17	21	1	9.69	+0.09	-0.61	-	B9Iae ⁺	-
		SK-67°18	12	2	12.02	-0.20	-0.99	-	06-07I+ WN5-6Var	
			5	11	14.49	-0.33		-	-	
			24	5	14.02	-0.22		-	-	
			27	4ab	14.50	-0.14		-	-	
			4	10	13.82	+0.90		-	-	F
			18	9	13.57	+0.58		-	-	F
106	38	SK-67°67	1	1	11.34	+0.10	-0.45	later than B5Ia	B6Ia	
		SK67°64	3	2	11.83	+0.19	-0.70	-	BI pec	
			11	3a	12.57	-0.15		B2:Iab	B2Iab	
			8	6	13.24	-0.01		B1Ib	B15Ib	
			9	7	13.80	-0.34		-	BQ5Ib	
			-	4	14.6	-		09V	-	
			4	5	12.82	+0.83		-	G2I	F
			7	12	14.35	+0.23		-	-	F

TABLE 3.5 (cont.)

DEM NO.	LUCKE ASSOC	NAME	LUCKE NO	STAR NO	V	B-V	U-B	Sp Type	Other	Comment
231	76	SK-67 ^O 184	1	1	13.13	0.01			WN4+08	
235	-	-	-	1	-	-		08If		
229	76	SK-67 ^O 178a	-	1	9.91	+1.15	+0.60	-	G2I	M
		F 17	67	7	11.53	+0.54		-	(F5-8I)	
		SK-67 ^O 174	-	4	11.67	-0.18	-0.98	09I	-	
		SK-67 ^O 173	51	8	12.04	-0.12	-0.96		(BOIa)	*
		SK-67 ^O 176	-	6	11.82	-0.16	-1.01	08I	-	
		SK-67 ^O 178b	59	5	12.13	-0.16		B1.5Ia	-	
		-	44		13.13	+0.01		-	(B5Ib)*	*
		-	46		13.56	-0.06		-	(B2Ib)*	E _{BV} = 0.10 assumed
			53		13.39	-0.19		-	(09III)*	
			45		14.21	-0.28				
			47		14.21	-0.24				
			49		14.42	-0.34				
			57		14.15	-0.26				
			62		14.16	-0.14				

TABLE 3.5 (cont.)

DEM NO.	LUCKE ASSOC	NAME	LUCKE NO	STAR NO	V	B-V	U-B	Sp Type	Other	Comment
		SK-67 ⁹ 166	1	9	12.27	-0.22	-1.01	-	06If	All these stars are outside main cluster. O-B0
		SK-67 ⁰ 167	21		12.54	-0.19	-1.04	-		
			3		13.78	-0.15				
			5		14.47	-0.08				
			12		14.84	-0.28				
			15		13.17	-0.18				
			16		14.33	+0.02				
			18		13.85	-0.11				
			19		14.10	-0.19				
			20		14.34	+0.10				
			25		14.31	+0.10				
			28		14.09	-0.20				
			35		14.05	-0.16				
			37		14.14	-0.12				

Eighteen additional spectral types are given in Column 9. The source references for these spectral types are denoted by subscripted numbers (in this column) and given below Table 5. An *m* in column 10 indicates that the star is a probable member of the cluster while an "F" indicates that it is a field star. Field stars are distinguished from cluster stars because their ages, as indicated by the position on the H-R diagram, are completely anomalous when compared with the age of the cluster

3.4.2 Uncertainties in the Classification Region

The uncertainty of spectral types of O stars determined from the ratio of the equivalent widths of HeI $\lambda 4471$ to HeII $\lambda 4542$, was set by the errors in the measured equivalent widths of these lines. The major sources of uncertainty in measuring these equivalent widths were the S/N of the spectra and the related problem of determining the true continuum level.

A local continuum level was fitted by eye to the spectra and line profiles were drawn by hand to fit the HeI $\lambda 4471$ and HeII $\lambda 4542$. The equivalent widths of the two lines were then measured by planimeter and the ratio of the line strengths calculated.

The uncertainty of this ratio was estimated by shifting the local continuum up or down by 1σ (in the

noise) for one of the two lines. The noise was assumed to be described by photon statistics. The ratio of HeI to HeII was then redetermined and the change in spectral type was taken as the uncertainty (2σ). The typical uncertainty in spectral type was plus or minus one spectral subclass, with two of the eleven stars having uncertainties of only half a spectral subclass.

Table 3.6 gives the measured equivalent widths of the two helium lines for each star, their ratio and the uncertainty of this ratio.

The uncertainties for the 14 remaining stars, appearing in Column 9 of Table V, were difficult to determine because of the qualitative nature of the method of classification. The average uncertainty should be less than plus or minus 1 spectral class as the luminosity of the stars are well determined due to the known distance modulus to the L.M.C.

The MK spectral types of stars determined in this chapter will be used in Chapter 4 to determine the U.V. Lyman continuum flux and Chapter 5 to determine the mass-loss rate ($\dot{M}$).

TABLE 3.6

Star	W4471 0 (mA)	% W4471	W4542 0 (mA)	% W4542	R	2 σ (R)
SK-67°250	440	-	260	92	1.69	0.81
DEM301/2	300	27	700	-	0.43	0.11
SK-67°174	470	38	180	-	2.61	1.00
SK-67°176	340	-	270	100	1.26	0.63
SK-67°108	330	-	590	32	0.56	0.20
DEM 11/1	570	28	330	-	1.73	0.49
DEM 11/2	510	27	620	-	0.82	0.22
DEM226/4a	340	-	350	71	0.97	0.40
DEM226/1	530	-	250	100	2.12	1.06
SK-70°13	320	50	250	-	1.28	0.61
DEM235/1	310	71	170	-	1.82	1.30

CHAPTER 4

THE INTEGRATED LYMAN CONTINUUM FLUX OF THE L.M.C. OB CLUSTERS

4.1 Introduction

A key parameter in the evolution of HII regions is the total ionizing photon flux (S_L), emitted by the embedded OB stars. However, the evolution theoretically depends on the fraction of this flux which is absorbed in the nebula gas. Fortunately we have methods for both measuring indirectly the Lyman continuum flux from spectral observations of the central stars and measuring directly the flux absorbed in the nebula, from which the above fraction can be derived.

In this chapter the first method for determining S_L is discussed. This method makes use of the Galactic Calibration of S_L with M.K. spectral type. In Chapter 7 the second method will be discussed. This method makes use of observations of the HII region to measure the total Lyman continuum photon flux intercepted by the nebula (N_L) and relies on applying the Zanstra technique to Balmer line surface photometry of L.M.C. HII regions to determine N_L .

If $S_L < N_L$, and all the photo-ionizing stars have been included in the determination of S_L , then some of the Balmer line emission produced by the nebula must come from collisional processes. This would be indicative of the presence of shock phenomena. If $S_L \sim N_L$ then the nebula is ionization bounded, whilst if $S_L > N_L$ the nebula is density bounded. It is also possible for a density bounded nebula to have a collisional contribution to its

Balmer line emission which is not significant enough to make $S_L < N_L$. In this case emission line diagnostics would have to be used to detect the presence of shocks (e.g. Dopita 1977).

Section 4.2 of this chapter discusses whether or not theoretical model atmospheres can be used to determine S_L . This section concludes that the present day input physics to these models is not adequate to predict S_L for OB stars.

Section 4.3 discusses an observational and semi-empirical method for determining S_L as a function of stellar spectral type, which is adopted in Section 4.4 to calculate S_L emitted by the L.M.C. OB clusters.

4.2 THEORETICAL ATMOSPHERIC MODELS FOR OB STARS

The Lyman continuum flux of early type stars cannot be readily observed, due to the dramatic increase in interstellar absorption which occurs at wavelengths shortward of the Lyman jump, resulting from absorption by interstellar hydrogen. This fact dictates that indirect methods be employed to ascertain the magnitude of the photon flux.

Model atmospheres can be used to predict the Lyman continuum flux provided that they take into account all the physical features of the atmospheres of OB stars.

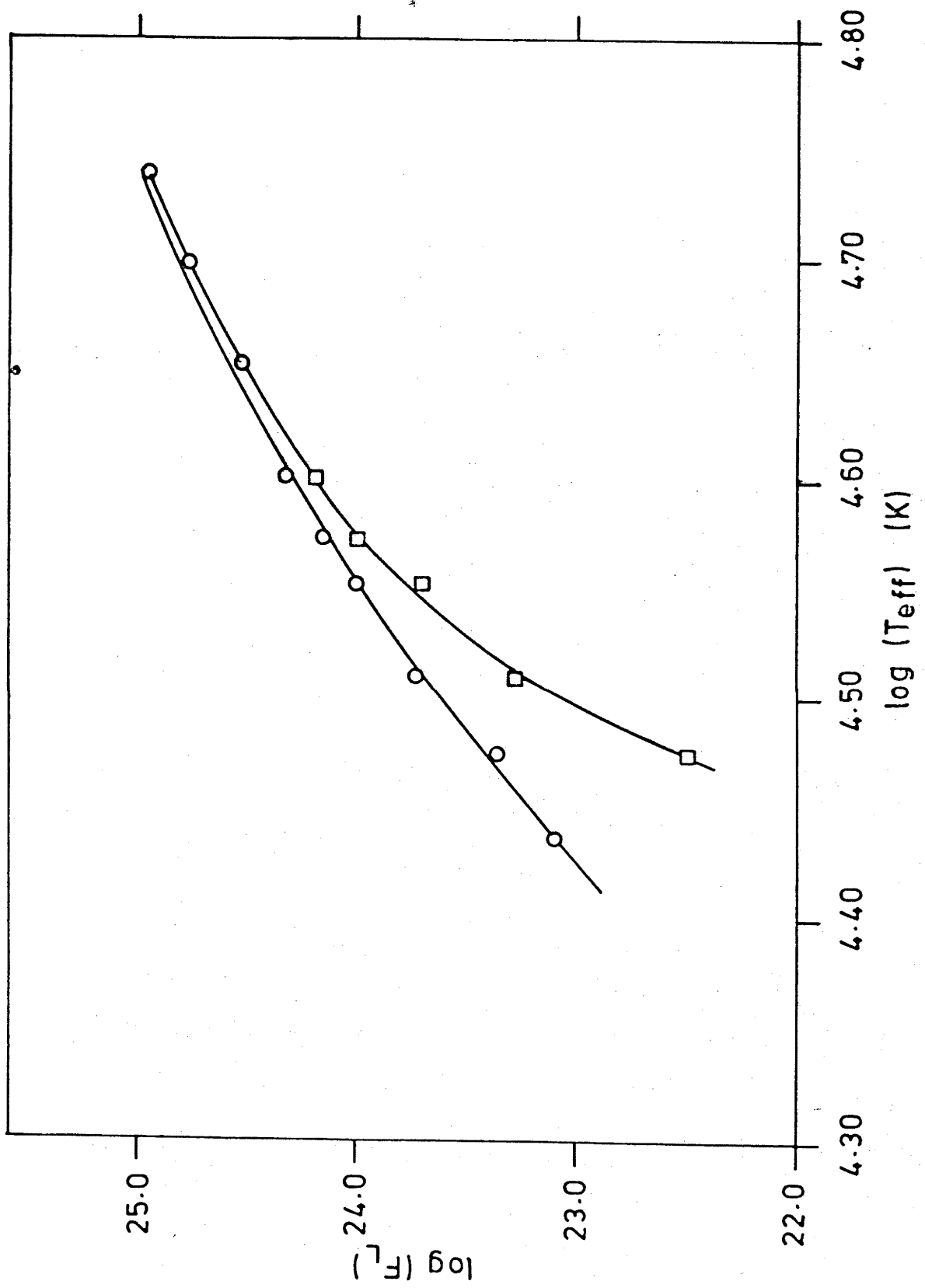
The luminous stars found in the L.M.C. HII regions are mostly O and early B giants and supergiants. These stars are known to possess stellar wind envelopes formed by the ejection of $1.0 \times 10^{-6} - 10^{-5} M_{\odot}/\text{year}$ from their surfaces. The presence of a dense extended envelope around these stars can invalidate the plane-parallel conditions which are assumed to apply to normal stellar atmospheres. It is also possible for NLTE effects to become important in the line and continuum formation processes of these stars.

Thus, to correctly determine S_L , the importance of NLTE conditions and sphericity upon atmospheric models for giant and supergiant OB stars, must be ascertained. In addition, the effects of line blanketing upon S_L would also have to be considered.

Figure 4.1 shows the logarithm of the integrated Lyman continuum photon flux (F_L) as a function of T_{eff} , predicted by the NLTE models of Auer and Mihalas (1972). Panagia (1973) compared the curves shown in Figure 4.1 with the corresponding ones predicted by the blanketed and unblanketed LTE models (Bradley and Morton, 1969, Morton 1969, Mihalas 1969, Van Cutters and Morton 1970 and Hummer and Mihalas 1970). He was able to use this comparison to outline the major effects of NLTE conditions and blanketing upon F_L . The main points which arose from the comparison were:

FIGURE 4.1

The logarithm of the Integrated Lyman continuum photon flux F_L (ergs cm⁻² sec⁻¹) as a function of the logarithm of the stellar effective temperature, predicted by the N.L.T.E. models of Auer and Mihalas (1972). O: Supergiants, □: Dwarfs.



- (1) The value of F_L emitted by a star is a function of g_{eff} as well as T_{eff} , with stars of lower g_{eff} emitting the greater flux for a given T_{eff} (see Figure 4.1).
 - (2) Both NLTE effects and line blanketing increase the value of F_L at a given T_{eff} . The NLTE effects are not as significant as line blanketing in dwarf stars ($\log(g_{\text{eff}}) \sim 4.0$), while the reverse is true for supergiant stars ($\log(g_{\text{eff}})$ 3.3 to 3.5).
- and
- (3) at high T_{eff} , F_L values for all models converge asymptotically on the black body curve (see Figure 4.1).

Cassinelli (1971) showed that it was possible for the near ultraviolet and visible continuum of a star, with an extended atmosphere and a given T_{eff} , to mimic the continuum of a star with a less extended atmosphere and a significantly lower T_{eff} . Thus, the T_{eff} of a star will be underestimated if atmospheric extension is important, leading to an underestimation of F_L (and S_L).

Panagia's and Cassinelli's work indicates that for giant and supergiant stars, NLTE conditions and atmospheric extension can have significant effects on S_L , while line blanketing effects are not so important. The best atmospheric models available for O and B stars are the unblanketed NLTE, plane-parallel atmospheres of Auer and Mihalas (1972) and the blanketed LTE, plane parallel atmospheres of

Kurucz (1979). Neither of these two models addresses the problem of atmospheric extension and only Auer and Mihalas consider NLTE conditions. These latter models could be used to estimate S_L although they would systematically underestimate this parameter.

Another disadvantage in using atmospheric models to determine S_L comes from the fact that the emergent flux is parameterized by the T_{eff} and g_{eff} of the stars. These quantities are not easily determined for O stars and the methods used to obtain them are themselves dependent on the atmospheric models adopted (Rennie and Lamers 1982, Underhill et al. 1979 and Conti and Leep 1974).

The inability of the model atmospheres to take into account the non-planar conditions in the atmospheres and the difficulty in determining T_{eff} and g_{eff} for O stars make it imperative that an observational technique must be found to determine S_L .

4.3 THE LYMAN PHOTON FLUX (S_L) AS A FUNCTION OF MK SPECTRAL TYPE

The Zanstra Principle (Morton 1969) can be applied to HII regions to determine the S_L being emitted by the exciting stars. This method relies on the fact that HII regions are optically thick in the Lyman continuum. In a steady state HII region each photoionization produced by a Lyman photon will be matched by recombination of a

free electron. This electron will recombine to the ground level of hydrogen, either by cascading or by direct recombination. Direct recombination to the ground level will produce a Lyman photon which will only travel a short distance before it is reabsorbed (the on-the-spot approximation). A cascading electron on the other hand, will eventually produce a Balmer line or continuum photon which will escape the HII region. A known fraction of the escaping photons will be hydrogen-alpha emission line photons. Thus, by observing the total dereddened hydrogen-alpha flux from the nebula, it is possible to measure N_L .

It is judicious to express N_L as a ratio $N_L/\pi F_V$ where πF_V is the power emitted in the photometric V-band. This ratio is independent of the distance and absolute magnitude of the stars and only depends on easily observable parameters (see equations 4.1 and 4.2).

Morton (1969) and later Georgelin et al. (1975) (hereafter GLM) used $N_L/\pi F_V$ to determine the total number of Lyman photons (L_{UV}), as a function of the spectral type of the existing star. They chose the relationship between $N_L/\pi F_V$ and spectral type because of the fact that it is almost independent of the luminosity class of the star. Even though $N_L/\pi F_V$ is larger for lower gravity stars at the same T_{eff} , supergiant stars of the same spectral type are roughly 10% cooler than dwarfs (Conti 1973). Thus these two effects nearly cancel making $N_L/\pi F_V$ almost

the same for both supergiant, and dwarf stars of the same spectral type.

Using the formulation given by Morton (1969):

$$\left(\frac{N_L}{\pi F_V}\right)_{H\alpha} = 6.931 \times 10^{26} f_a 10^{0.4(V - A_V + A_a)} \quad (\text{photons } J^{-1}) \quad (4.1)$$

where f_a = the integrated hydrogen-alpha flux at the Earth's surface ($J/m^2/sec$)

V = the photometric V magnitude of the star

A_V = the interstellar extinction in the V band

and A_a = the interstellar extinction at $H\alpha$

(An electron temperature (T_e) of $10,000^\circ K$ was assumed for the HII regions.)

It is also possible to determine $N_L/\pi F_V$ by observing the integrated radio continuum flux from the nebula. Again, following the formulation of Morton (1969):

$$\left(\frac{N_L}{\pi F_V}\right)_{\text{RADIO}} = 3.77 \times 10^{26} \frac{f_v}{g(\nu_1 T_e)} 10^{0.4(V - A_V)} \quad (\text{photons } J^{-1}) \quad (4.2)$$

where f_v = the integrated radio continuum flux
 $J \ m^{-2} \ sec^{-1} \ Hz^{-1}$ and

$$g(\nu, T) = 1 + 0.13 \log(T_e^{3/2}/\nu) \quad (4.3)$$

(The constant of proportionality is again that for a T_e of $10,000^\circ$ for the HII regions).

The HII regions used to determine $N_L/\pi F_V$ must be suited to the application of the Zanstra method.

Conditions which would favour the use of this method are:

- 1) that the nebula contains one exciting star of known spectral type, or at least an exciting star which dominates over others found in the cluster
- 2) that the nebula is not density bounded. This will lead to an underestimation of $N_L / \pi F_V$ as Lyman photons will escape from the nebula without being converted into Balmer photons.
- 3) that all exciting stars have been identified and properly spectrally classified, otherwise an erroneous value of N_L will be attributed to the wrong star, giving rise to error in $N_L / \pi F_V$.
- 4) that the nebula is not irregularly shaped so that the star is set to one side of the nebula. This may make the nebula density bounded on one side while still being ionization bounded on the other
- 5) that the nebula does not have an extended low density halo of material so that the integrated hydrogen-alpha or radio continuum fluxes are under-estimated due to emission being lost in the background
- 6) that the nebula is not confused with nearby sources
- and 7) that dust is not absorbing an appreciable

quantity of the Lyman photon flux.

Morton (1969) used equations 4.1, 4.2 and 4.3, and GLM (1975) equation 4.2 to determine $N_L/\pi F_V$ as a function of spectral type.

Morton did not select HII regions which were necessarily suited to the Zanstra technique. His main criterion was that the nebula have one exciting star or one which dominated over others in the nebula. Morton postulated that $N_L/\pi F_V$ would be overestimated primarily because condition 3 (one of the conditions listed above for applying the Zanstra technique) would not be satisfied, and that $\frac{N_L}{\pi F_V}$ would be under-estimated primarily due to the failure of condition 2). He assumed that for a given spectral type it was equally as likely to overestimate as under-estimate $N_L/\pi F_V$, due to failure of conditions 2) and 3). Thus, Morton adopted the mean of the $N_L/\pi F_V$ values at a given spectral type as the most probable value. This relationship is shown in Figure 4.2 for an assumed T_e of $7,000^\circ\text{K}$ and $10,000^\circ\text{K}$.

G.L.M. confined their sample of HII regions to those which were best suited to the Zanstra Method. Their data gave a reasonably tight relationship between $\log N_L/\pi F_V$ and spectral type, with the average standard deviation being 0.16 in the logarithm. G.L.M. found that by introducing nebulae which did not satisfy their

selection criterion they greatly increased the scatter of the data. This increased scatter was systematic about the mean relationship, suggesting that the assumption adopted by Morton in obtaining his mean relationship was correct.

A comparison of the curves in Figure 4.2 shows that there is excellent agreement between Morton's and G.L.M.'s results in the spectral range 06 to B0. On average Morton's curve is systematically lower than G.L.M.'s by about 0.05 in the logarithm. This is much less than the standard deviation of 0.16 in the logarithm claimed by G.L.M.

G.L.M.'s $N_L/\pi F_V$ relationship was adopted for all except the 05 and B0.5 stars. Morton's value was used in preference to G.L.M.'s for 05 stars, as it was based on more stars. The $N_L/\pi F_V$ value for B0.5 stars given by Morton and G.L.M. were both uncertain and so a logarithmic average was used for these stars. Table 4.1 shows the final relationship between $\log N_L/\pi F_V$ and MK Spectral type which was adopted.

$\frac{N_L}{\pi F_V}$ is independent of the distance and absolute magnitude of the star, and its dependence on spectral type is quite insensitive to g_{eff} and the chosen model atmosphere. These properties make it an ideal parameter for comparing the photon flux emitted by a star with that absorbed by the nebula. The value of $N_L/\pi F_V$ for an HII

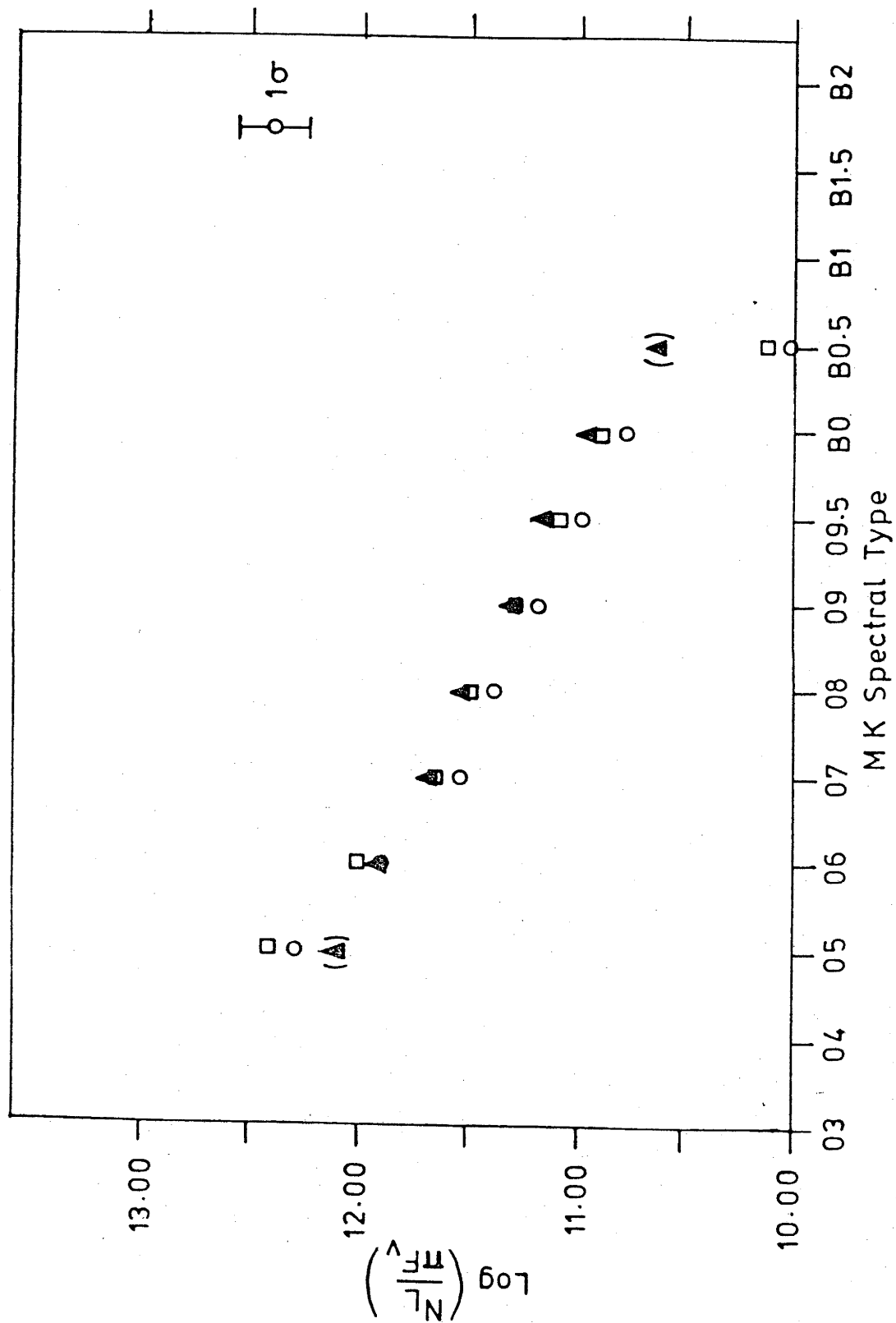
TABLE 4.1

M.K. SPECTRAL TYPE $\log_{10} (N_L / \pi F_V)$ PHOTONS J⁻¹

05	19.41
06	18.92
07	18.72
08	18.54
09	18.32
09.5	18.17
BO	17.95
BO.5	17.39

FIGURE 4.2

Log $(N_L / \pi F_V)$ vs. the M.K. spectral type for dwarf stars. (o) Morton (1969) for $T_e = 10,000\text{K}$; ($\square$) Morton (1969) $T_e = 7,000$; ($\blacktriangle$) Georgelin et. al (1975). $(N_L / \pi F_V)$ is in units of photons erg^{-1} .



region with more than one exciting star is given by:

$$\sum_{\text{All Stars}} \left(\frac{N_L}{\pi F_V} \right) 10^{-0.4(V-A_V)} = 6.931 \times 10^{26} \text{ fa } 10^{0.4A_A} - 4.4$$

(Where the symbols have the same meaning as in equation (4.1)).

The right-hand side of this equation, which is obtained from the integrated hydrogen-alpha flux of the HII region, can be directly compared with the sum on the left-hand side which is derived from the OB stars. The comparison would follow along the lines of the argument which is set out in the introduction.

The ratio $\frac{N_L}{\pi F_V}$ can also be used to determine the total number of Lyman photons emitted per second (L_{uv}) by the OB stars and the excitation parameter (U_L) of the nebula. These quantities are given by:

$$L_{uv} = 1.231 \times 10^{28} 10^{-0.4M_V} \frac{N_L}{\pi F_V} \quad \begin{array}{l} \text{photons/sec} \\ 4-5 \end{array}$$

$$\text{where } L_{uv} = 4 \pi R_*^2 N_L \quad \begin{array}{l} \text{photons sec}^{-1} \\ 4-6 \end{array}$$

$$U_L = 7.282 \times 10^{-2} 10^{-0.4M_V} \frac{N_L}{\pi F_V}^{1/3} \quad \begin{array}{l} \text{pc cm}^{-2} \\ 4-7 \end{array}$$

$$\frac{N_L}{\pi F_V} \text{ is in photons } J^{-1}$$

R_* is the stellar radius and T_e is taken as $10,000^\circ K$.

The total L_{uv} (or U_L) obtained from the OB stars can also be compared with the corresponding L_{uv} (or U_L) absorbed by the HII region, to determine whether or not the HII region is being maintained by photoionization. There is no advantage in using either of these parameters in preference to $N_L/\pi F_V$, because of the additional uncertainties which are introduced by the need to determine absolute magnitudes.

U_L can be used to determine the total mass of gas which could be ionized by the OB stars (in solar units). This is given by:

$$M_{HII} = \frac{M}{M_{\odot}} = 8.612 \times 10^{-8} (U_L)^3 \left(\frac{2.17}{n_e} \right) \quad 4-8$$

where n_e is the electron density in particles m^{-3} and U_L has units of parsecs m^{-2} . The HII region is assumed to have a filling factor of 1.0 and a T_e of $10,000^{\circ}K$. This higher value of T_e is more appropriate for L.M.C. HII regions because of their reduced metal abundance (Stasinska 1982). Hydrogen and helium are taken as being singly-ionized throughout the volume of the nebula and the fraction of Helium in the nebula, by mass and number, are assumed to be those given by Allen (1973).

4.4 RESULTS FOR L.M.C. STARS

Table 4.2 summarises the results for each of the 9 L.M.C. HII regions for which spectral data is available. Only those stars from Table 3.5 which are thought to contribute significantly to the L_{uv} of the cluster are included.

Columns 1 to 5 give respectively the D.E.M. Number (Davies, Elliott and Meaburn 1976) of the HII region, the Lucke association number (Lucke 1972) and the identification number, spectral type and absolute magnitude for each star. The values of $\frac{N_L}{\pi F_V} 10^{-4} (V-A_V)$ (hereafter N_p), L_{uv} and U_L are also listed for each star in columns 6, 7 and 8 respectively. Column 9 gives the percentage contribution of a star to the total L_{uv} of a cluster while column 10 lists the mass of gas which could be ionized by the OB stars. This is calculated using equation 4.8 with n_e equal to 10^6 particles m^{-3} .

The predicted integrated dereddened hydrogen-alpha flux (F_N) for the HII regions are also listed in Column 10 below the nebula Zanstra mass. This predicted flux is calculated using equation 4.4 and will be compared with the observed values of F_N given in Chapter 7.

The dominant source of error for all the results listed in Table 4.2 comes from the uncertainty in $\log N_L/\pi F_V$ and spectral type. The error in spectral type was assumed

TABLE 4.2

DEM.NO.	LUCKE ASS.	STAR	SP.TYP.	-M _V	$N_L 10^{-0.4(V-A_V)}$ $\pi F_V \times 10^{-14}$	Log L _{uv}	U _L	% Total	M _{HII} /M _⊙ x10 ⁻⁵
301	114	1	08.5I(f)	6.2	5.214	49.00	67.9	24.6	
		2	05.5V	5.3	10.696	49.32	86.9	51.5	
		3	09:III-V	5.0	1.280	48.41	43.2	6.3	
		10	08:V	4.7	1.626	48.51	46.6	8.0	
		6	08.5III+B?	5.2	2.114	48.60	50.1	9.8	D
					$2.09^{+0.65} \times 10^{-13}$ $-12.68^{+0.13}$	$49.61^{+0.14}$ $108.5^{+11.5}$			$2.39^{+0.76}$
11	3	1	08.5III-V	5.2	1.964	48.60	50.0	33.1	
		2	07V	5.1	3.591	48.85	60.6	38.9	
		4	09:V	4.0	0.684	48.01	31.8	8.5	
					$6.24^{+2.20} \times 10^{-14}$ $-13.20^{+0.15}$	$49.08^{+0.16}$ $72.4^{+8.8}$			$0.71^{+0.26}$
196	55	1	06III+0B	5.6	9.103	49.25	82.37	52.5	D
		8	09I	7.0	8.302	49.21	79.88	47.9	
		2	09-8:	-	-	-	-	-	
					$1.74^{+0.65} \times 10^{-13}$ $-12.76^{+0.16}$	$49.53^{+0.17}$ $102.2^{+13.0}$			$2.00^{+0.76}$

DEM.NO.	LUCKE ASS.	STAR	SP.TYP.	$-M_V$	$N_L 10^{-0.4(V-A_V)}$ $\pi F_V \times 10^{-14}$	$\log L_{uv}$	U_L	% Total	$M_{HII}/M_\odot$ $\times 10^{-5}$
226	-	1ab	08.5I+B	6.6	8.113	49.16	76.88	63.1	Dom
		4a	07 V+B:	5.2	4.048	48.89	62.49	33.9	Dom
		13ab	09.5V	4.1	-	47.90	29.22	3.5	
		4b	B0.5	-	-				
					$1.22^{+0.48} \times 10^{-13}$	$49.36^{+0.17}$	$89.70^{+11.5}$		$1.35^{+0.52}$
					$-12.92^{+0.17}$				
137	43	11	05.5If	6.4	33.512	49.76	121.82	(94.6)	
		10	WN3	4.4	(1.328)	(48.43)	(43.87)	(4.4)	
		1	B0.5Ib	5.7	0.297	47.76	26.23	(0.4)	
		3b	-	-	-	-	-	-	
		12	-	-	-	-	-	-	
					$2.95^{+1.55} \times 10^{-13}$	$49.76^{+0.23}$	$122.2^{+21.8}$		$3.41^{+1.82}$
					(3.08×10^{-13})	(49.78)	(124.1)		(3.57)
					$-12.53^{+0.23}$				
					(-12.51)				
25	-	1	08I	6.6	9.621	$49.27^{+0.23}$	$83.7^{+5.0}$	-	$1.09^{+0.59}$
					$-13.02^{+0.23}$				

DEM.NO.	LUCKE ASS.	STAR	SP.TYP.	$-M_V$	$N_E 10^{-0.4(V-A_V)}$ $\pi F_V 10^{-14}$	$\log L_{uv}$	U_L	% Total	$M_{HII}/M_{\odot}$ $\times 10^{-5}$
106	38	4	O9V	4.3	0.691	48.13	34.85	(14.6)	
		7	B0.5III	5.0	0.170	47.48	21.17	(3.3)	
		2	(BOIa)	(7.1)	(3.777)	(48.88)	(61.98)	(82.1)	
					$8.61^{+3.77} \times 10^{-15}$	$48.22^{+0.20}$	$37.30^{+7.3}$		$0.096^{+0.057}$
					(4.638×10^{-14})	(48.97)	(66.18)		(0.542)
					$-14.06^{+0.19}$				
					(-13.33)				
31	6	2	(O6.5I (WN6-5	6.6 4.5	23.507 (0.478)	49.55 (47.97)	103.70 (30.83)	(97.7) (2.6)	Dom
					$2.351^{+1.24} \times 10^{-13}$	$49.55^{+0.23}$	$103.70^{+18.7}$		$2.08^{+1.13}$
					(2.355)	(49.56)	(104.6)		(2.14)
					$-12.62^{+0.23}$				
229	76	4	O9I	7.2	10.261	49.29	69.9	32.8	
		6	O8I	7.1	14.832	49.47	97.5	49.7	
		L53	O9III:	5.5	2.105	48.61	50.4	6.9	
		8	BOIa:	6.9	3.113	48.80	58.3	10.6	
					$3.030^{+0.98} \times 10^{-13}$	$49.77^{+0.14}$	$123.1^{+13.5}$		$3.49^{+1.15}$
					$-12.52^{+0.14}$				
235	-	1	O8If	6.5	8.693	$49.23^{+0.23}$	$81.1^{+14.6}$	-	$1.00^{+0.53}$
					$-13.06^{+0.23}$				

to be plus or minus one spectral subclass (2σ), while the error in $\log N_L/\pi F_V$ at a given spectral type was 0.16 (G.L.M. 1975). This resulted in a standard deviation of 0.23 in the final adopted value of $\log N_L/\pi F_V$. This source of error far outweighed the contributions to the error from V, A_V and the L.M.C. distance modulus. A mean A_V of 0.5 ± 0.1 (1σ) magnitudes was assumed for all stars and the true distance modulus to the L.M.C. was taken 18.4 ± 0.1 (1σ) (Wilson 1983).

Individual comments are given below for each of the HII regions, listed in Table 4.2, to indicate whether or not all the exciting stars have been observed.

DEM 301: Except for star N^Q 12, all blue stars brighter than 14.5 in V have been observed in this cluster. N^Q 12 has a V magnitude of 14.46 making it a star of spectral type O8.5 or later if it is a dwarf or B0.5 or later if it is a giant (luminosity Class III). Thus, at most it should only contribute an extra 10% to the total L_{uv} of the cluster.

Star N^Q 6 has a possible B companion which could contribute to an overestimation of N_p for this star.

DEM 11: All blue stars brighter than 14.5 in V have been observed in this cluster.

DEM 196: Except for star N^Q 2, all blue stars brighter than 14.5 in V have been observed in this cluster. Star N^Q 2 is enveloped in its own compact HII region which is

located on the edge of DEM 196. Most of the Lyman flux from this star would be used ionizing the compact HII region and so would not contribute significantly to the total N_p .

Star N^O 1 is double, with the spectral type of the primary being O6III. The luminosity class assigned to the primary is based on the weakness of the HeII 4686 line. The mean absolute magnitude of this spectral type has been used to calculate V , N_p , L_{uv} , U_L and M_{HII} .

DEM 226: All blue stars brighter than 14.5 in V have been observed in this cluster. Stars N^O 1 and N^O 4a have possible companions but their observed absolute magnitudes indicate the primary O star is dominant in each case.

DEM 137: All blue stars brighter than 14.0 in V have been observed in this cluster.

Star N^O 10 (SK-65°45) is a Wolf-Rayet star classified by Breysacher (1981) as WN3.

Barlow, Smith and Willis (1980) have published $\log L_{uv}$ values for Wolf-Rayet stars, which can be used to determine $\log N_L / \pi F_V$. Unfortunately the earliest WN star they consider is WN4. The value they give for a WN4 star is used in Table 4.2 to calculate N_p , U_L and M_{HII} (values which use this estimate of $\log N_L / \pi F_V$ are shown in parenthesis in Table 4.2) even though trends in their data would

indicate that this would lead to a slight overestimate. It is clear from these calculations (see Figure 4.2) that star $N^{\circ} 11$ is the dominant exciting star for this cluster and that the WN3 star only contributes a small percentage to the total L_{uv} .

DEM 25: The brightest star in the photometric band, found within this HII shell, is star $N^{\circ} 1$. This star dominates over all other stars found within this object, making it the most likely source of the majority of Lyman photons.

DEM 106: This cluster is made up of a closely bunched group of 13 OB supergiant stars. $N^{\circ} 4$ and 7 are the only ones with observed spectral types earlier than B1. The photo-ionizing flux from these two stars is insufficient to ionize DEM 106, suggesting that another star is responsible. Star No 2 is the most probable source of these ionizing photons. It is an η Carina-type star with strong FeII, [FeII] and Balmer emission lines (see Figure 4.2). Allen and Glass (1976) found that this star has a dust shell and indicated that its visual spectrum is very similar to the forbidden line B stars of our own galaxy. The spectral features of this star are also characteristic of group 1 Be stars in the classification scheme of Jaschek et al (1980). Jaschek et al found that group 1 Be stars are predominantly of spectral type B0 to B3, but they

could be as late as B6. They also found that the FeII emission is strongest in the stars of earliest spectral type. Star N^Q 2 has particularly strong FeII emission, indicating that it could have a spectral type as early as B0. This classification is adopted in Table 4.2 to estimate the Lyman flux produced by this star (all values which are based on the $\log (N_L/\pi F_V)$ adopted for this star are shown in parentheses in Table 4.2).

DEM 31: This cluster is dominated by two highly luminous OB stars. Star N^Q 1 is a late B supergiant and it does not contribute to the ionization of DEM 31. Star N^Q 2 is classified as O6.5I + WN 5-6 (Var) and is the main source of ionizing photons. The $N_L/\pi F_V$ for the WN 5-6 star is taken from the published data of Barlow, Smith and Willis (1980). (Parameters in Table 4.2 which use the $N_L/\pi F_V$ value for this star are listed in parentheses.) Using this value it can be seen that the Wolf-Rayet star makes a negligible contribution to the Lyman photon flux which is almost exclusively emitted by the O6.5I primary.

DEM 229: The cluster found within this HII region is much richer than others considered in this chapter. There is a relatively compact association of supergiant stars in the brightest part of the nebula which fans out into a large number of lower luminosity stars towards the N.N.W. It was impossible to observe all the stars which could contribute to the photo-ionization of this nebula because of their number. Instead, all blue stars found in the

central cluster, with a V magnitude brighter than 13.4, were observed. The values of N_p , L_{uv} , U_L and M_{HII} in Table 4.2 can only be considered as lower limits because of this.

DEM 235: Star N^o 1 is clearly the dominant exciting star for this nebula.

CHAPTER 5

MASS LOSS IN OB STARS

5.1 INTRODUCTION

The aim of this chapter is, firstly to establish a representative sample of reliable mass-loss rates and stellar parameters (i.e. Luminosity L_* , Mass (M_*), Radius (R_*), T_{eff} , etc.) for OB stars, and secondly to use this sample in order to determine a functional dependency of $\dot{M}$ on these stellar parameters.

Mass-loss has been detected in all O and B stars with Bolometric magnitudes ≥ -6 (Snow and Morton 1976). Typical rates range from 10^{-7} to $10^{-5} M_{\odot} \text{ yr}^{-1}$ with the stellar winds being accelerated up to terminal velocities (V_{∞}) of between 1000 and 4000 km/sec.

Three main models have been proposed to explain the mechanism by which the stellar winds are produced. The first is the line driven wind theory of Castor, Abbott and Klein (1975) (hereafter CAK). In this model the mass-loss is driven by the radiation force, associated with resonance and subordinate line absorption by the ions of the most abundant elements.

This model is by far the most comprehensive of all the models available as it actually makes predictions about $\dot{M}$ and the velocity law for the stellar wind envelopes. However, it has a number of problemsⁱⁿ that it does not predict either the presence of x-ray emission from OB stars or the difference in mass-loss rates between dwarf and supergiant stars of the same luminosity (see Section 5.5).

A further problem is that the flow is unstable to radiative driven instabilities (Nelson and Hearn (1978)) which would produce a break-up of the flow, resulting in either the deposition of energy in a corona or the production of radiatively driven blobs (Lucy and White, 1980).

The second model for describing mass-loss is the Andriesse Fluctuation theory (Andriesse 1980a, 1980b). This theory is based on the concept of thermodynamic fluctuations and partial equilibrium. Andriesse attributes two relaxation timescales for the atmospheres of OB stars. The first is the free fall time scale (t_D) which characterises the dynamical processes at the stellar surface. The second is the Kelvin-Helmholtz (or thermal) time scale (t_K) which characterizes the thermal relaxation time for the star if its nuclear energy source is extinguished.

Andriesse states that the star never entirely reaches thermal equilibrium because of stochastic fluctuations on the time scale t_D . However because $t_D \ll t_K$, it approaches a state of partial equilibrium. These fluctuations are supposed to act as an input of energy to the atmosphere, and the minimum energy condition for the thermodynamic system of the star as a whole is one in which a stochastically steady ejection of matter occurs in the form of a stellar wind.

The fluctuation theory also makes a quantitative prediction about the functional dependence of $\dot{M}$ on stellar

parameters. This theory predicts that

$$\dot{M} \propto L_{\star}^{1.5} (R_{\star}/M_{\star})^{2.25}$$

The third set of models consists of those that predict some form of mechanical input to the stellar wind usually via a corona. These models include the Warm wind coronal model of Lamers and Rogerson (1978), the Cool wind-Thin corona model of Cassinelli et al. (1978) and the Canon and Thomas (1977) Nozzle model.

None of these theories make quantitative predictions about $\dot{M}$ and so cannot be discussed in the context of this chapter. (The presence of x-ray emission from OB stars and NV and OVI in the spectra of these stars argues that some form of mechanical deposition is occurring - possibly in the form of coronal energy deposition.)

5.2 METHODS FOR DETERMINING THE MASS-LOSS RATES FOR OB STARS

The relationship between the mass-loss rate ($\dot{M}$) and the major parameters describing an OB star, Luminosity ($L_{\star}$), mass ($M_{\star}$), radius ($R_{\star}$) and T_{eff} , can only be discussed if reliable measurements of $\dot{M}$ can be obtained, for a representative sample of stars. There are four basic methods for determining $\dot{M}$ for OB stars, each of which is reliant to varying degrees, on various assumptions about the stellar wind envelope.

5.2.1 Mass Loss Rates Derived from U.V. Resonance Lines

The first of these four methods relies on fitting theoretical line profiles to the observed P-Cygni line profiles of U.V. resonance lines. This comparison can be done using the atlas of theoretical profiles published by Castor and Lamers (1979). These profiles were calculated under the assumption of resonance scattering in spherically symmetric expanding envelope, using the Sobolev approximation. The fitting process is dependent on the adopted form of the velocity law ($V(r)$) and the radial optical depth as a function of the flow velocity $\tau_{\text{rad}}(V)$.

The major advantage of the U-V fitting method comes from the fact that it is very sensitive to the presence of mass-loss, with mass-loss rates of $\sim 10^{-8} M_{\odot} \text{yr}^{-1}$ being measurable. This sensitivity has allowed this method to be applied to a large body of OB stars, producing data which is useful in investigating $\dot{M}$ as a function of stellar parameters.

There are two major disadvantages in using this method. Firstly, for supergiant O and B stars it is possible for the absorption component of the P-Cygni profile to be saturated, so that only lower limits to $\dot{M}$ can be obtained. Secondly, the mass-loss rates are dependent on the validity of the adopted form of the $V(r)$ and $\tau_{\text{rad}}(V)$ laws and upon the assumptions used in calculating the ionization fraction of the ion in the wind ($g(n_i)$).

This last quantity is the most uncertain factor in the calculation of $\dot{M}$, due to the present lack of knowledge of the ionization conditions in the wind (Castor 1978).

Olsen (1981) proposed a method which could overcome the first problem, by using the line profiles of excited level transitions such as NIV $\lambda 1718 \text{ \AA}$, OIV $1339,42 \text{ \AA}$ and OV $1371 \text{ \AA}$ to determine $\dot{M}$. O stars which have saturated U.V. resonance lines, show P-Cygni-type profiles for these subordinate lines. The absorption component is unsaturated and thus can be used to calculate $\dot{M}$. Unfortunately the mass-loss rates are dependent on knowing the stellar flux at the exciting wavelengths of the line considered, (i.e. $500\text{--}800 \text{ \AA}$). If the wrong effective temperature or an inappropriate model atmosphere is used to determine the stellar flux, then an erroneous $\dot{M}$ is obtained.

A more satisfactory solution to the problem of determining $g(n_i)$ was given by GHS. They empirically calibrated $\dot{M}$ against a parameter involving $g(n_i)$, using mass-loss rates derived from the more reliable radio and infra-red techniques (see below). This calibration was then used to determine $\dot{M}$ for their program stars. GHS's method assumes that the degree of ionization in the wind is relatively constant with radius and that the absolute values of the ionization fractions of ions in the wind are the same in both evolved and dwarf stars.

Clearly an improvement in the reliability of this

method is dependent on a better understanding of the ionization conditions in the winds of OB stars.

5.2.2 Mass-Loss Rates Derived from Radio Observations

Perhaps the most reliable method for determining $\dot{M}$ is based on radio observations of the free-free bremsstrahlung emitted by the ionized stellar wind envelope. The radio emitting region of the wind is outside the point where the terminal velocity of the flow has been reached. This means that the density structure in the emitting region is inversely proportional to the distance from the star. Thus there is no need to rely on models to obtain the density structure of the envelopes as in the case of the U.V. method.

Wright and Barlow (1975) and Panagia and Fellini (1975) have calculated the relationship between $\dot{M}$ and the observed radio flux for this density structure. They showed that the only parameters required to determine $\dot{M}$ are the distance to the star, the radio flux and the stellar wind terminal velocity. The distance of the star is the major direct source of error in calculating $\dot{M}$ by this method, though there are additional uncertainties due to the poorly known temperature and ionization structure of the wind (H.G.L.M. Lamers, J.P. Cassinelli (1978)).

The radio observations give mass loss rate with formal uncertainties of approximately 50% . These are by far the most reliable because this method requires the least assumptions about the detailed structure of the wind.

The major disadvantage of this method is that it can only be used to measure $\dot{M}$ in stars which have high mass-loss rates. This confines the sample of stars with mass-loss rates determined in this manner, to supergiant O and B stars.

5.2.3 Mass-Loss Rates Derived from Infra-Red Observations

Barlow and Cohen (1977) derived mass-loss rates for 44 luminous OB stars by using the measured free-free infra-red excess of these stars between 2.2 and 10.0 microns. The free-free emission comes from the ionized gas in the accelerating region of the stellar wind. Thus the derived mass-loss rates are dependent on the assumed velocity and temperature structure of the wind.

Barlow and Cohen derived the velocity law appropriate to the star P Cygni by assuming that the temperature of the wind (T_w) was approximately the same as T_{eff} for the star. They then applied this velocity law to the other stars in their sample to obtain mass-loss rates.

Lamers (1981) showed that Barlow and Cohen's mass-loss rates were systematically lower than those derived

by the radio method by a factor of ≈ 2.1 . He attributed this discrepancy to Barlow and Cohen's adoption of P Cygni Velocity law as their standard law. Lamers considered the velocity law of this star as not being representative of OB stars in general, as it was too slow. This would make the measured mass-loss rates lower than the actual values.

The infra-red method suffers from the same disadvantage as the radio, in that it can only be applied to stars with large mass-loss rates. At present, only supergiant O and B stars have had their mass-loss rates measured.

5.2.4 Mass-Loss Rates Derived from Hydrogen Alpha Emission

Mass-loss rates have been determined from the Hydrogen-alpha line which appears in emission in stars with high mass-loss rates. This has been done either through the use of the hydrogen-alpha equivalent widths (Klein and Castor 1978) or by detailed profile fitting in the wings of the hydrogen alpha line (Olsen and Ebbets 1981).

The use of equivalent widths to determine $\dot{M}$ are subject to uncertainties in the adopted velocity law, the predicted contribution of the photospheric absorption component, the effects of rotation on the line profile and the breakdown of the sobolev approximation in the line centre. The large number of sources of uncertainty

make this method the most unreliable of the four methods available for determining $\dot{M}$.

Olsen and Ebbets (1981) attempted to overcome some of the objections to the equivalent width method by comparing the wings of detailed theoretical line profiles with the wings of the observed profiles. They confined their comparison to the wings of the hydrogen-alpha line to avoid the effects of the unknown photospheric profile and the breakdown of the sobolev approximation in the central region of the line. This method is an improvement but it requires very high S/N spectra to obtain accurate line profile information in the wings of the line.

5.3 THE ADOPTED SAMPLE OF MASS-LOSS RATES

Ideally, an homogenous set of mass-loss rates tied to a set of homogenous stellar parameters is required to be able to investigate the dependence of $\dot{M}$ on the gross properties of OB stars, assuming that these parameters are indeed correlated. This data sample must include stars covering the fullest possible range in luminosity and T_{eff} expected for O and B stars, to eliminate biases in the dependency of $\dot{M}$ on stellar parameters.

The best possible way of obtaining an homogenous and reliable set of mass-loss rates would be to observe a representative sample of OB stars using the radio method.

This is not possible because of the limited range in mass-loss rates which can be measured by this technique.

Indeed, none of the four available methods can be used to measure the full range of mass-loss rates which are exhibited by OB stars. Thus the compilation of a representative sample of mass-loss rates for OB stars is only possible if a combination of the various methods available are used.

5.3.1 Lamers Sample of Mass-Loss Rates

Lamers (1981) published mass-loss rates for 53 early-type stars ranging spectral type from O3 to B9, and in luminosity class from V to Ia⁺. He formed this sample from four homogenous sets of mass-loss rates derived by different authors from the radio flux, infra-red excess, U.V. resonance lines and hydrogen-alpha emission. In conjunction with this, Lamers compiled a homogenous set of values for the L_* , M_* , R_* and T_{eff} of each of the stars in his sample. The values that he adopted for these parameters were those that he considered to be the best which were available in the literature at the time. The four sets of mass-loss rates were corrected to be consistent with the new homogenous set of stellar parameters.

Lamers chose the radio mass-loss rates of Abbott et al (1980) as his standard rates. The mass-loss rates which were derived from the other three methods were scaled to the radio results by using stars which were

common to the radio sample. The rationale behind applying this scaling is that it compensates for different models and assumptions adopted by these three methods (Lamers 1981).

The U.V. mass-loss rates adopted by Lamers were those of G.H.S. These required no scaling to the radio results as this had already been done by GHS when they calibrated the mean ionization fraction in the stellar wind using the radio results.

The infra-red mass-loss rates adopted by Lamers were those of Barlow and Cohen (1977). These were scaled up by a factor of 2.1 to correct for the slow velocity law of P Cygni which was adopted by Barlow and Cohen.

The hydrogen-alpha mass-loss rates adopted by Lamers were those of Conti and Frost (1977) and Klein and Castor (1978). These mass-loss rates are derived using the equivalent width method and so are highly uncertain. The scaling factor which brings these mass-loss rates into agreement with the radio, is also uncertain due to the small number of stars which are common to both samples. As a result, all stars with mass-loss rates derived in this manner were tentatively left out of our adopted sample.

There are 36 stars in Lamers sample with spectral types B5 or earlier and with mass-loss rates determined by methods other than the hydrogen-alpha method. 19 of

these stars were derived using the U.V. method and 17 using the radio or scaled infra-red method. The internal uncertainty of these mass-loss rates is approximately 0.20 in the logarithm (Lamers 1981).

The luminosity, radius and mass of a star in Lamers' sample is dependent on the value of T_{eff} that he assigned to the star. Unfortunately, there is still considerable uncertainty as to the true value of T_{eff} for O and B stars (Garmany et al (1982), Rennie and Lamers (1981), Underhill et al (1977) and Conti (1973). It is possible that Lamers' temperatures could be systematically in error by up to 10% for some stars. This is particularly true for O stars earlier than O6, where Lamers has adopted two temperature scales, in recognition of the uncertainty in this value for these stars.

All that can be done, until the uncertainty in the T_{eff} of OB stars is reduced, is to adopt the best available estimates of T_{eff} and then indicate what the effect on the results would be of a change in this quantity.

Lamers' sample is presently the best that is available for investigating the dependency of $\dot{M}$ on stellar parameters and therefore it is adopted as the canonical data base for the purpose of the following discussions.

5.3.2 Modifications to the Sample of Lamers

One shortcoming of this sample is that it is not fully representative of OB stars as a whole. Firstly, there is a distinct bias towards more luminous stars simply because of the limitations of the methods used to determine $\dot{M}$. Secondly there is only a small number of Of stars which have their mass-loss rates measured by techniques other than the hydrogen alpha-method. It is possible to overcome these problems by expanding the sample to include the necessary stars, but care would have to be taken so as not to destroy the homogeneity of Lamer's data.

The most serious flaw in Lamer's sample is the lack of Of stars with reliable mass-loss rates. 3 of the stars with hydrogen-alpha results also have rates determined by Garmany et al. (1981) using the U.V. line fitting technique. The mass-loss rates that Garmany et al. obtained for these three stars agree with the modified hydrogen alpha results of Lamers to within the errors. This would suggest that the rates for these stars are not as uncertain as previously thought. This being the case, the Of stars in Lamers' sample with hydrogen-alpha mass-loss rates were incorporated into the adopted sample, together with 5 additional mass-loss rates from the sample of Garmany et al.

Table 5.1 lists the final sample which was used to investigate the dependency of $\dot{M}$ on stellar parameters.

The following values are listed for each star

- Column 1: The H.D. number
- Column 2: The Bayer number
- Column 3: The M.K. Spectral type of the star
- Column 4: The T_{eff} of the star in $^{\circ}\text{K}$
- Column 5: The logarithm of the luminosity of the star in solar luminosities ($L_{\odot}$)
- Column 6: The radius of the star in Solar radii ($R_{\odot}$)
- Column 7: The mass of the star in solar masses ($M_{\odot}$)
- Column 8: $(1-\Gamma)$ where Γ is the ratio of the luminosity of the star to the Eddington luminosity
- Column 9: The logarithm of the effective gravity (g_{eff}) of the star (c.g.s. units)
- Column 10: the surface escape velocity in km/sec
- Column 11: the terminal velocity of the stellar wind of the star (km/sec)
- and Column 12: the logarithm of the mass-loss rate for stars ($M_{\odot}\text{yr}^{-1}$)

STAR	SP. TYPE	NAME	T _{eff}	log L _* /L _☉	R _* /R _☉	M _* /M _☉	(1-Γ)	log g _{eff}	V _{esc}	V _∞	log M
93129	03f		47,000	6.38	23.4	136	0.53	3.55	1084	-	4.89
269810	03f		48,000	6.07	15.6	90	0.67	3.82	1214	4100:	5.3
15570	04f		42,500	6.18	22.7	93	0.57	3.45	944	3500:	4.93
269698	04f		42,000	5.89	16.6	65	0.70	3.65	1022	2300	5.6
66811	04f	ζP _{up}	42,000	5.93	17.4	70	0.69	3.63	1029	2660	5.45
190429	04f		42,500	5.78	14.3	57	0.72	3.74	1046	-	5.26
164794	04V((f))	9 S _{gr}	44,700	5.98	16.3	75	0.66	3.71	1076	3440	5.52
16691	05I		40,000	5.62	13.5	46	0.76	3.72	994	-	5.39
15558	05(f)		42,000	6.10	21.5	82	0.59	3.46	926	-	5.53
14947	05.5f		36,000	5.78	19.6	51	0.69	3.41	827	3500:	5.28
210839	06ef	λC _{eph}	37,200	5.83	19.7	50	0.66	3.36	799	2500	5.4
152233	06(f)		39,000	5.86	18.7	60	0.68	3.50	912	-	5.73
199579	06V((f))	HR8023	39,000	5.46	11.8	39	0.80	3.79	1004	2980	5.87
163758	06.5Iaf		36,000	5.96	24.4	60	0.62	3.22	762	2600	5.2
47839	07V((f))	15Mon	35,400	5.09	9.3	27	0.88	3.87	987	3060	6.22
24912	07.5III((f))	ξ Per	35,500	5.18	10.3	28	0.86	3.79	944	3000	5.80
151804	08If		33,000	6.09	34.1	70	0.56	2.94	662	2000:	5.1
152408	08If	HR6272	33,000	5.97	29.7	60	0.61	3.04	685	1800:	5.0
188001	08If	Sge	33,000	5.76	23.1	46	0.69	3.20	724	2200:	5.2

TABLE 5.1 (cont..)

STAR	SP. TYPE	NAME	T_{eff}	$\log L_*/L_\odot$	$R_*/R_\odot$	$M_*/M_\odot$	$(1-\Gamma)$	$\log g_{\text{eff}}$	V_{esc}	V_∞	$\log M$
167771	O8I(f)	HR6841	35,000	5.78	21.5	50	0.68	3.30	776	-	5.97
36861	O8III((f))	λ Ori A	35,000	5.13	10.0	28	0.87	3.83	964	2650	6.10
57061	O9II	τ CMa	33,500	5.60	18.8	38	0.72	3.33	745	2260	5.99
37043	O9III	ι Ori	34,000	5.46	15.8	36	0.78	3.49	823	2360	5.94
214680	O9V	10 Lac	35,400	4.99	8.3	29	0.91	4.02	1101	2000:	6.59
30614	O9.5Ia	α Cam	26,300	5.67	32.8	40	0.69	2.85	566	2050	5.29
36486	O9.5II	δ Ori A	27,600	5.17	16.8	25	0.58	3.32	690	2410	6.10
149757	O9.5v	ζ Oph	34,100	5.09	10.0	26	0.87	3.79	929	2000:	6.48
38666	O9.5v	μ Col	32,000	4.58	6.4	19	0.95	4.08	1037	2000:	7.17
14938	O9.7Iab	μ Nor	27,800	5.73	31.5	42	0.66	2.88	579	2190	5.53
37742	O9.7Iab	δ Ori A	25,600	5.23	20.0	25	0.84	3.15	633	2290	5.80
37128	B0Ia	ϵ Ori	26,400	5.65	31.9	38	0.69	2.85	560	2100	5.50
149438	B0V	τ Sco	32,000	4.62	6.7	20	0.94	4.06	1034	2000:	7.13
38771	B0.5Ia	κ Ori	23,100	5.56	37.5	35	0.72	2.69	506	1870	5.53
24760	B0.5III	ϵ Per	27,500	4.86	11.9	20	0.90	3.54	759	1500:	6.90
93030	B0.5Vp	θ Car	29,000	4.30	5.6	15	0.96	4.10	990	2000:	7.25
193237	B1Ia ⁺	P Cyg	21,000	6.02	77.4	62	0.55	2.19	407	600	4.68
169454	B1Ia ⁺		21,000	5.74	56.1	43	0.66	2.39	439	-	5.33
2905	B1Ia	K Cas	21,600	5.39	35.4	29	0.77	2.69	490	1650	5.60
91316	B1Iab	ρ Leo	20,300	5.06	27.4	22	0.86	2.83	513	1580	5.95
24398	B1Ib	ζ Per	19,900	4.98	26.2	21	0.88	2.86	519	1500	6.23

STAR	SP. TYPE	NAME	T _{eff}	log L _* /L _☉	R _* /R _☉	M _* /M _☉	(1-Γ)	log g _{eff}	V _{esc}	V _∞	log M
40111	B1Ib	139Tau	21,400	4.91	20.7	20	0.89	3.06	573	1500:	6.18
157246	B1Ib	γ Ara	20,600	4.95	23.5	21	0.89	2.97	551	1050:	6.14
190603	B1.5Ia ⁺		19,000	5.58	57.0	35	0.71	2.32	408	900	5.26
194279	B1.5Ia		19,000	5.34	43.2	28	0.79	2.51	442	-	5.61
14143	B2Ia		17,500	5.34	50.9	28	0.79	2.37	407	-	5.64
41117	B2Ia	χ ² Ori	17,500	5.47	59.4	32	0.75	2.28	392	-	5.36
14134	B3Ia		14,500	5.14	58.9	24	0.85	2.21	363	-	5.85
53138	E3Ia	0 ² CMa	14,800	5.09	53.2	23	0.86	2.28	377	580:	5.56
198478	B3Ia	55 Cyg	14,300	5.15	61.2	24	0.84	2.17	354	-	5.87
58350	B5Ia	η CMa	13,500	4.91	56.3	20	0.89	2.19	347	590:	6.08
36371	B5Iab	χ Aur	13,500	5.14	67.9	24	0.85	2.08	338	-	5.51
164353	B5Iab	67 Oph	13,000	4.62	37.3	16	0.93	2.47	390	500:	6.58

5.4 $\dot{M}$ AS A FUNCTION OF STELLAR PARAMETERS

5.4.1 $\dot{M}$ as a function of L_*

The first relationship between $\dot{M}$ and stellar parameters was that given by the Radiation-driven wind theory of Castor, Abbott and Klein (1975). Their theory predicted that $\dot{M}$ should vary in direct proportion to L_* . Figure 5.1 shows $\text{Log } \dot{M}$ plotted against $\text{Log } (L_*/L_\odot)$ for all the stars in the adopted sample. There is a considerable degree of scatter in this plot, but there appears to be a general correlation between the two parameters. A linear least squares fit to the data gives

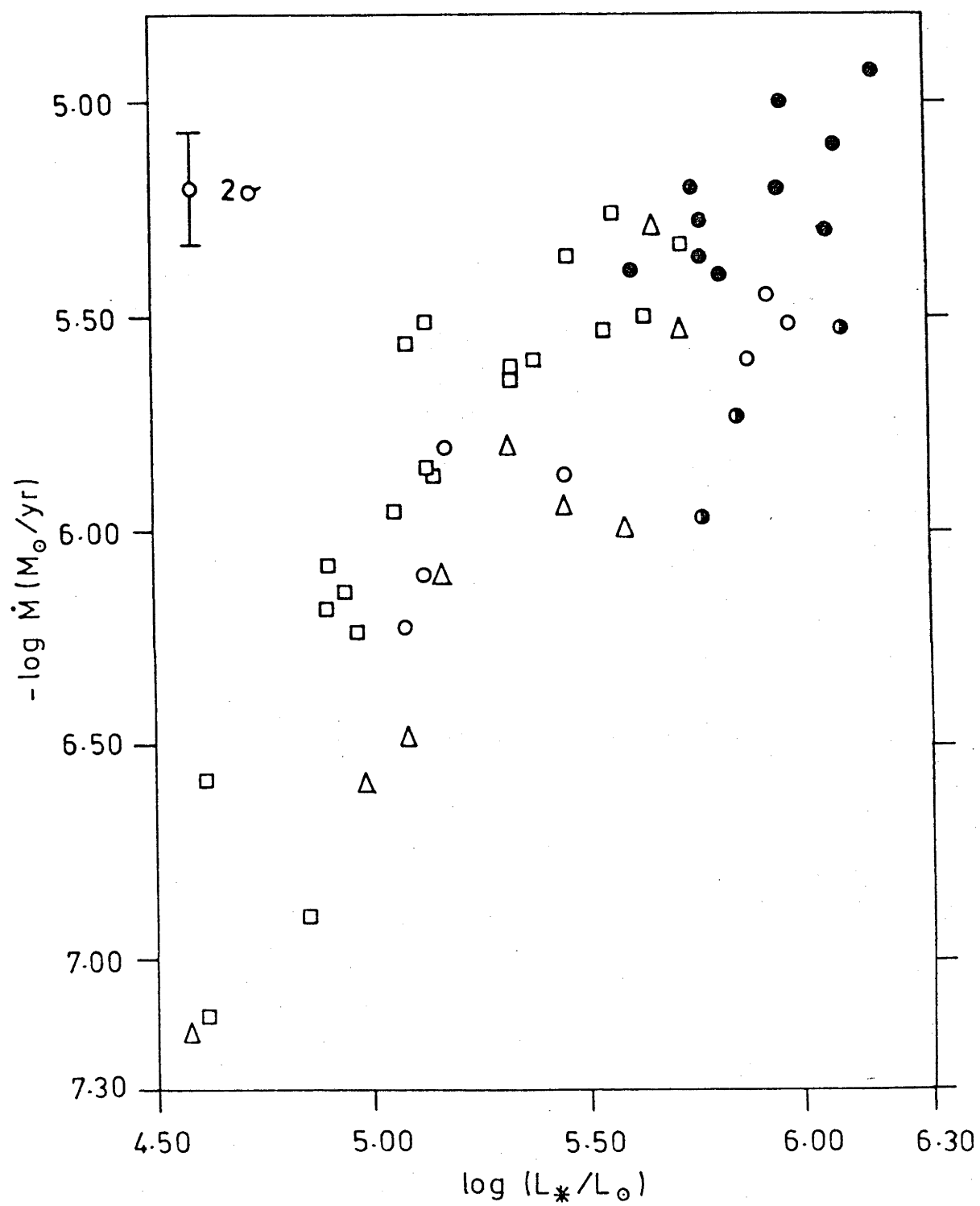
$$\dot{M} = 10^{-11.53} (L_*/L_\odot)^{1.06} (M_\odot \text{ yr}^{-1}) \quad (5-1)$$

with a mean dispersion of 0.30 in $\log \dot{M}$ (about the mean curve). The large amount of scatter in the data results in an uncertainty in the power of $(L_*/L_\odot)$ of approximately ± 0.3 .

Part of the scatter in the data can be attributed to the make-up of the adopted sample. About half of the stars have scaled infra-red mass-loss rates derived from Barlow and Cohen's data, (Barlow and Cohen (1977)), while most of the others are from the U.V. results of GHS. The mass-loss rates derived from the U.V. method are scaled to a sample of rates made up from a combination of the radio results of Abbott (1980) and the infra-red results of Barlow and Cohen (1977). The mass-loss rates should show a dependency, which is the average of the dependency shown by the U.V. and infra-red rates.

Figure 5.1

$\log_{10} \dot{M}$ is plotted against $\log_{10} (L_*/L_\odot)$ for all the stars in the adopted sample. $\bullet$: Of Stars; $\circ$: O(f) stars; $\circ$: O((f)) stars; Δ : O Stars; $\square$: B Stars. $\dot{M}$ is in solar masses per year.



Similarly the mass-loss rates derived from Barlow and Cohen's results should show the same dependency on L_* as the infra-red results. Close inspection of the adopted sample reveals that the stars generally do show the dependency of $\dot{M}$ on L_* , which is characteristic of the method by which they were derived, i.e. the infra-red mass-loss rates scale as $(L_*)^{1.2}$ (for B supergiants) while the radio rates scale as $(L_*)^{1.8}$. If this difference in dependency on L_* were totally an artifact of the two methods used, then a considerable artificial source of error would be present in the adopted sample, particularly at mass-loss rates below $10^{-6} M_\odot \text{ yr}^{-1}$.

Figure 5.2 shows the $\dot{M}$, L_* plot for supergiant stars with mass-loss rates derived from both the infra-red and U.V. methods, and Figure 5.3 shows the same type of plot for other luminosity class stars. The two plots highlight the point that the supergiant stars with mass-loss derived from the U.V. method appear to show the same dependency on L_* as those with rates derived from the infra-red method. This result indicates that at a given luminosity, there is a real difference between the mass-loss rates from supergiants and dwarfs, suggesting that $\dot{M}$ increases as a star evolves away from the main sequence (Conti and Garmany 1980, Chiosi 1980, and Lamers et. al, 1980). This result is not predicted by the line driven wind theory (in its original form) necessitating some form of revision of this theory if it is to remain viable.

Figure 5.2

$\log_{10} \dot{M}$ is plotted against $\log_{10} (L_*/L_\odot)$ for the supergiant stars (Luminosity Class I) in the adopted sample. O: Of Stars; Δ : Non Of Stars; $\square$:B Stars. Shaded symbols represent stars with mass-loss rates obtained from the U.V. method while the unshaded symbols represent stars with mass-loss rates obtained from the Radio Method.

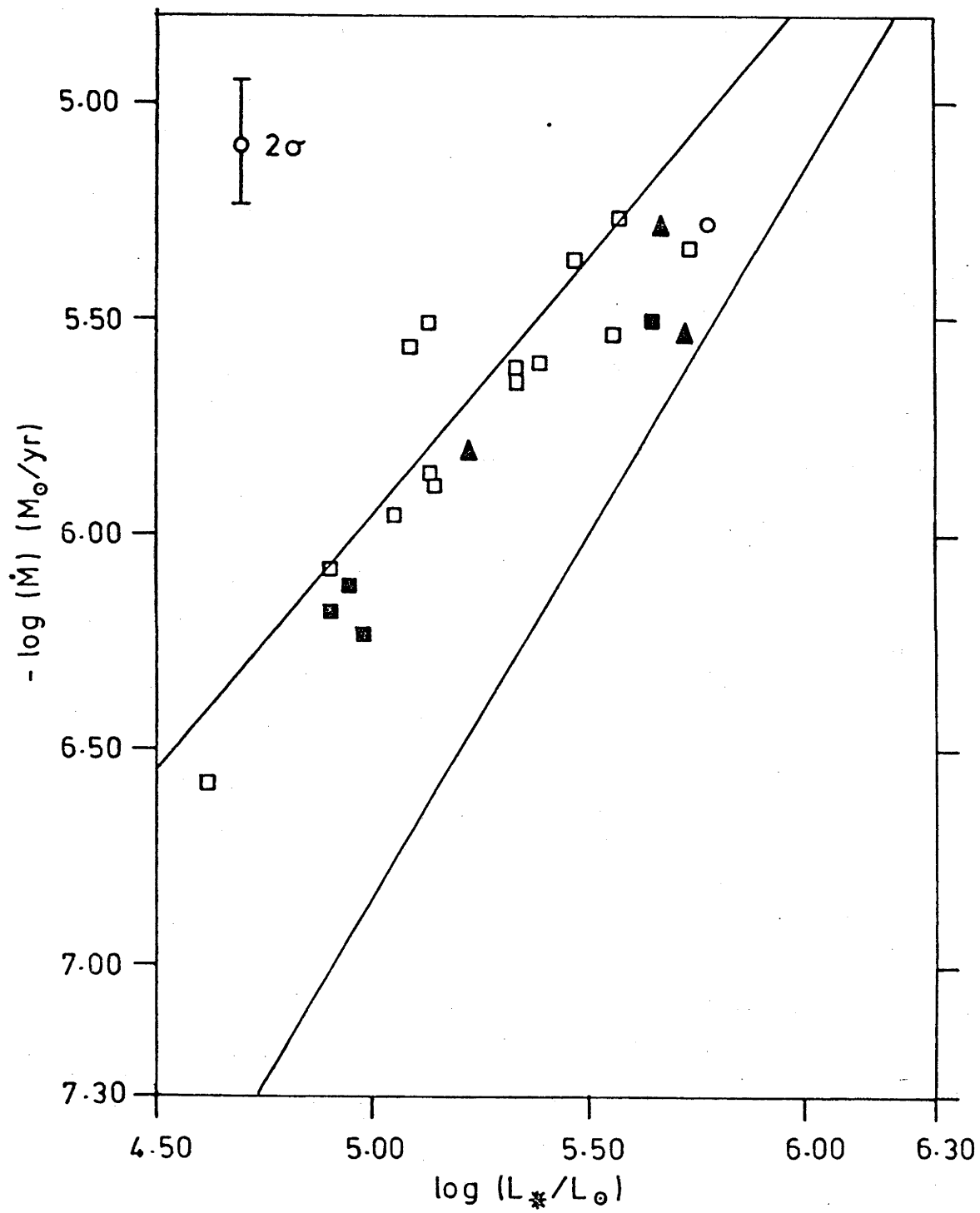
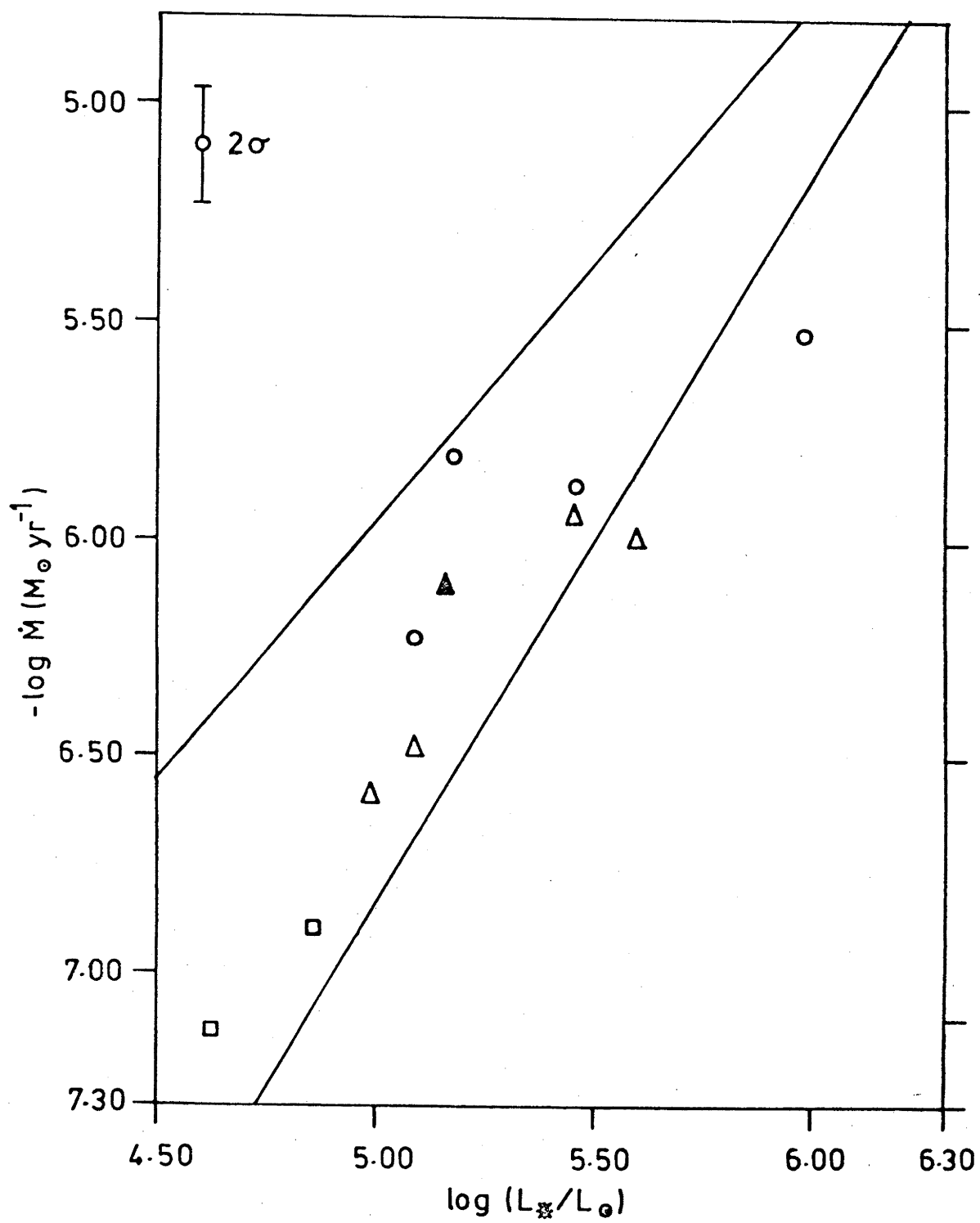


Figure 5.3

$\log_{10} \dot{M}$ is plotted against $\log_{10} (L_*/L_\odot)$ for non-supergiant stars in the adopted sample. The symbols are the same as in figure 5.2.



5.4.2 $E_T \propto L_*$

An alternative method for forming a relationship between stellar parameters and $\dot{M}$ is to consider the formulae for the total power (E_T) required to accelerate the wind up to its terminal velocity. This is given by:

$$E_T = \frac{1}{2} \dot{M} (v_s^2 + v_{esc}^2) \quad (J s^{-1}) \quad 5-2$$

where

$$v_{esc} = 6.1743 \times 10^5 \left(\frac{M_*/M_\odot (1 - \Gamma)}{R_*/R_\odot} \right)^{1/2} \quad (m.s^{-1}) \quad 5-3$$

is the effective escape velocity at the surface of the star and $M_\odot$ and $R_\odot$ are the solar mass and radius respectively.

Equation 5-2 can be used to form an expression relating $\dot{M}$ to stellar parameters provided the dependence of E_T on these parameters is known. Having no a priori knowledge of the dependency of E_T on stellar parameters, it is a reasonable first approximation to assume that $E_T \propto L_*$, which implies that the kinetic energy is derived from the luminous energy by a process with constant efficiency.

In this case $\dot{M}$ would be given by:

$$\dot{M} = \frac{K_1 L_*}{v_s^2 + v_{esc}^2} \quad 5-4$$

Using the result of Abbott (1980), that $v_\infty \propto v_{esc}$, this becomes

$$\dot{M} = \frac{K_2 L_* R_*}{GM_* (1 - \Gamma)} \propto \frac{L_*}{v_{\text{esc}}^2} \quad 5-5$$

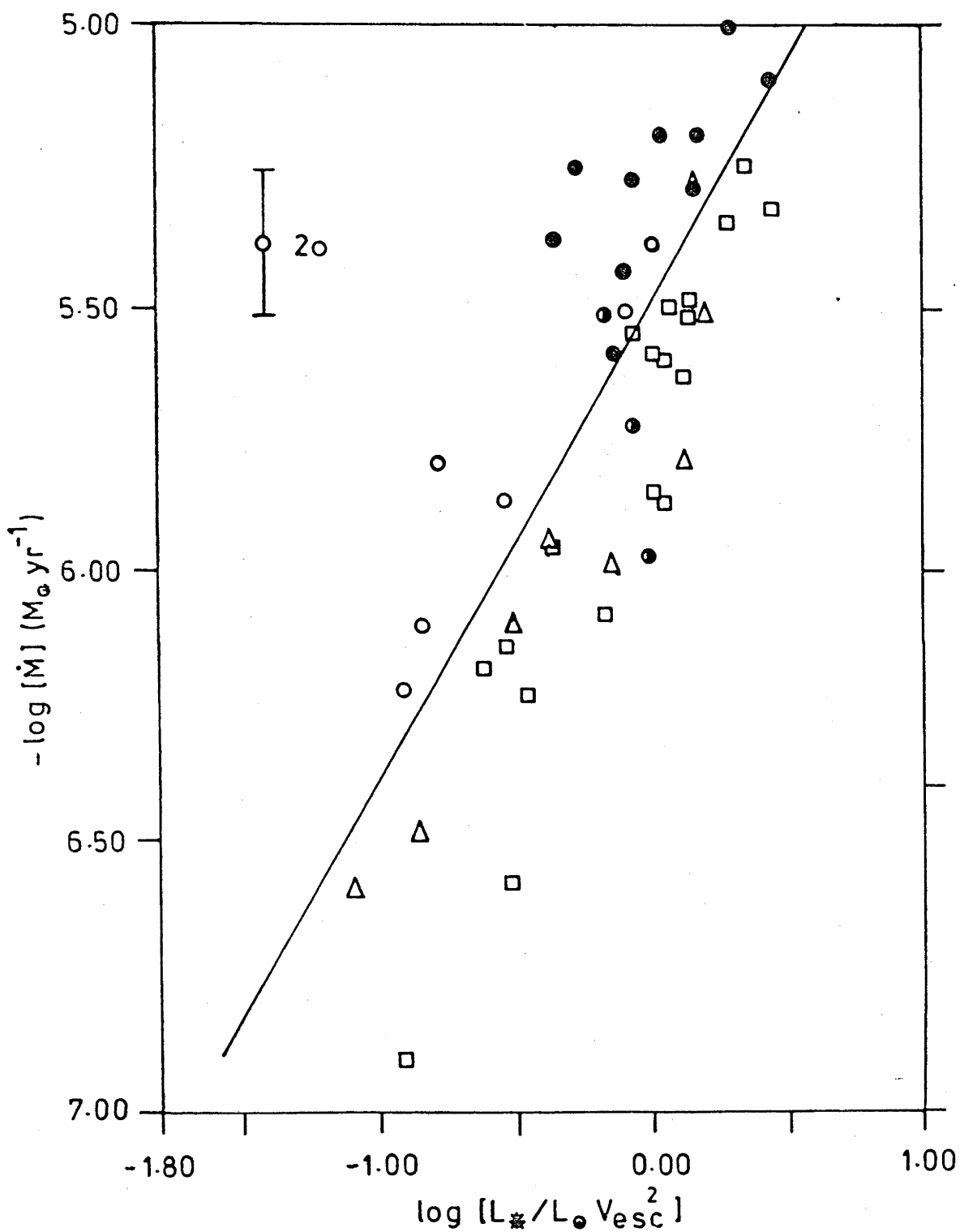
This formula is plotted in Figure 5.4 for all the stars in the adopted sample.

The first thing which becomes immediately obvious is the fact that there is splitting between Of and non Of stars. This is highlighted by a line drawn in this figure, which divides the stars into these two groups. If equation 5-5 is to be used as a mass-loss formula, then Of and non Of stars would have to be distinguished to determine $\dot{M}$. Such a distinction may of course simply imply that the energy partition between wind and luminosity is different in the two classes of objects, that is to say, different efficiency parameters.

It is preferable that one expression be sought which describes the mass-loss rate in all stars, not only because of the convenience of its universability, but because we would expect it to be more fundamentally related to the physics of the mechanism producing mass-loss in OB stars. Care must be taken however, not to overlook the fact that there may be more than one distinct mechanism producing the mass-loss phenomena in OB stars. In this case a universal formula might hide real differences in the mass-loss mechanism. The following discussion is implicitly based on the assumption that there

Figure 5.4

$\log_{10} \dot{M}$ is plotted against $\log_{10} ((L_*/L_\odot)/V_{\text{esc}}^2)$ where V_{esc} is the stellar escape velocity in kilometres per second. The symbols are the same as those used in figure 5.1



is a universal formula describing the mass-loss rates for all OB stars.

The existence of a universal mass-loss mechanism implies that the splitting of the mass-loss rates in Figure 5.4 is the result of the wrong assumption being made about the dependency of E_T on stellar parameters. If the correct (universal) dependence is used then this splitting should disappear.

5.4.3 $E_T \propto (t_D/t_K)^{1/2} L_*$

An alternative dependency on stellar parameters for E_T was proposed empirically by Chiosi (1980) and theoretically by Andriesse (1980a, 1980b). Chiosi suggested a mass-loss formula of the form;

$$\dot{M} = \left(\frac{t_D}{t_K}\right)^{1/2} \frac{L_* R_*}{GM_* (1 - \Gamma)} \quad 5-6$$

where t_D and t_K are the dynamical (free-fall) and Kelvin Helmholtz (thermal) time scales of the star respectively. Except for the factor $(1 - \Gamma)^{-1}$ this formula happens to be the one predicted by the Andriesse Fluctuation Theory. It can also be obtained from equation 5-2 if

$$E_T \propto \left(\frac{t_D}{t_K}\right)^{1/2} L_* \quad 5-7$$

where $t_D \ll t_K$ for OB stars and the ratio t_D/t_K decreases as a star evolves away from the main sequence. Since OB

stars essentially evolve with constant luminosity (Chiosi 1980) this means that E_T decreases during the life time of the star.

Figure 5.5 shows Andriesse's formula plotted for all stars in the adopted sample. The splitting between Of and non-Of stars which is present in Figure 5.4 has become even more distinct, indicating that the functional dependency of E_T is not the universal one we are searching for.

5.4.4 $E_T \propto (V_{esc}/C)L_*$

An alternative expression for $\dot{M}$ in terms of stellar parameters can be obtained by considering the ratio of the momentum carried away by the stellar wind ($\dot{M}V_\infty$), to the total momentum of the photons leaving the star L_*/C , i.e.

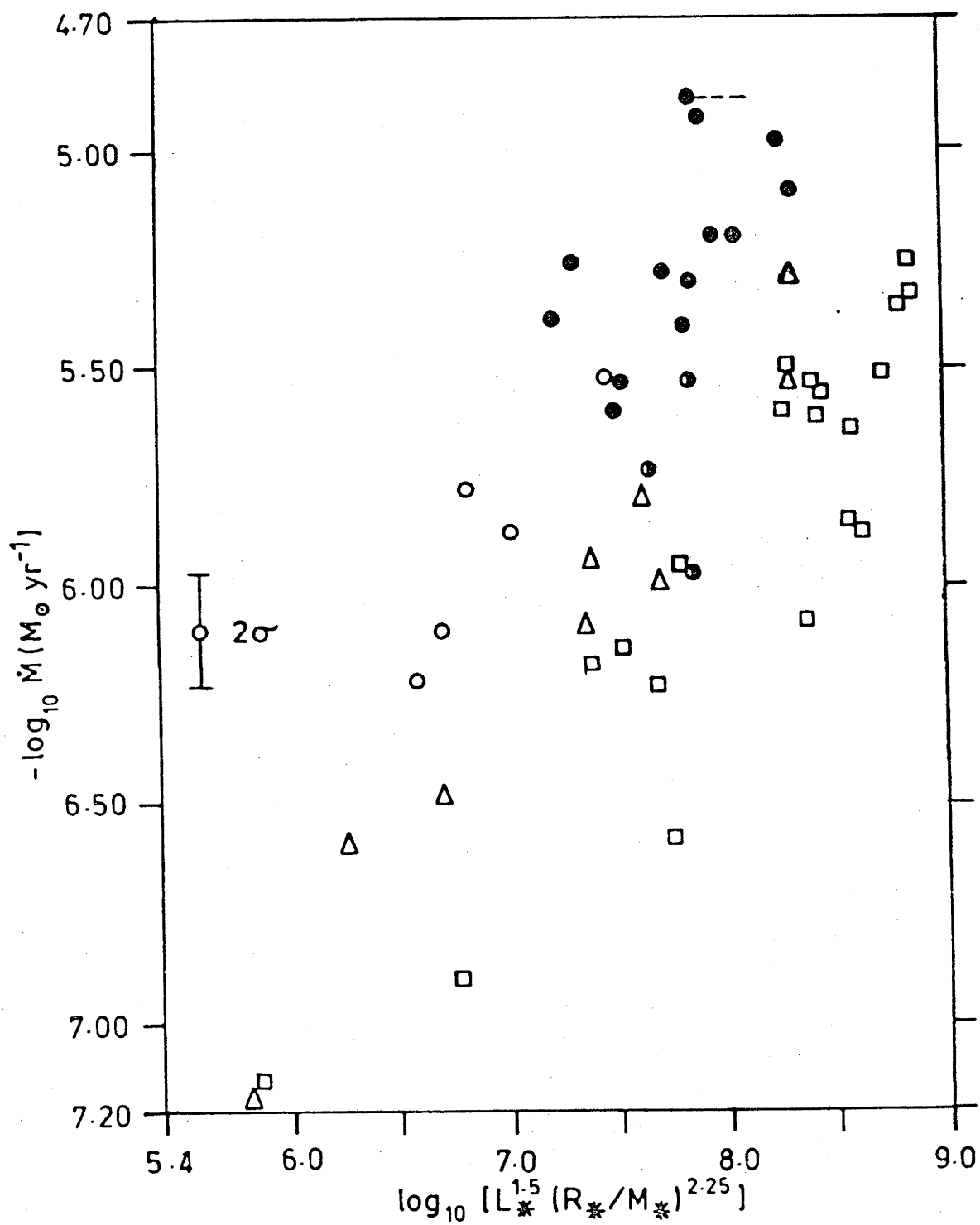
$$\dot{M} = \frac{\epsilon L_*}{V_\infty C} \quad 5-8$$

where C is the velocity of light and ϵ is a constant.

When the major source of energy input for the wind is momentum transfer by line absorption of photons (i.e. the radiation driven wind theory), ϵ can be thought of as a factor expressing the efficiency with which the wind absorbs the momentum from the stellar radiation.

Figure 5.5

$\log_{10} \dot{M}$ is plotted against the Andriesse formula for mass-loss. All parameters are expressed in solar units and the symbols are the same as those used in figure 5-1.



Abbott (1980b) indicated that ϵ must be corrected, for the fraction of the momentum deposition which is used to support the stellar wind envelope against gravity. This can be done by using Abbott's expression for the conservation of momentum integrated over the wind, i.e.

$$\dot{M}V_{\infty} + \frac{L_{*}}{C} \left(\frac{1-\Gamma}{\Gamma} \right) T_H = B L/C \quad 5-9a$$

$$\text{where } T_H = \int_{r_0}^{\infty} \rho \sigma_e dr \quad 5-9b$$

is the electron scattering optical depth of the stellar wind envelope exterior to the sonic point and B represents the fraction of the radiative momentum emitted by the star which is absorbed by the wind (assuming single scattering of photons). T_H can be evaluated if the radial density or velocity structure of the wind is known. Assuming a velocity law of the form

$$V(r) = V_s + V_{\infty} \left(1 - \frac{R_{*}}{r} \right)^{\beta} \quad 5-10$$

where V_s is the velocity of the wind at the sonic point ($V_s \ll V_{\infty}$) and β is a scaling factor describing the steepness of the velocity law, it can be shown that

$$T_H \approx \frac{\dot{M} \sigma_e}{2\pi R_{*} V_{\infty}} \quad \text{for } \beta = \frac{1}{2} \text{ and } \beta = 1 \quad 5-11$$

Detailed studies of the U.V. spectra of ζ P_{up} and τ Sco by Castor and Lamers (1979) showed that $\beta \approx 1$ yields a reasonable approximation to the actual velocity laws for these stars. Accepting this value of β as being representative for all O and B stars, it can be

shown that

$$\dot{M} = \left[\frac{B}{1 + 1/p^2} \right] \frac{L}{V_{\infty} C} \quad 5-12a$$

where
$$p = \frac{V_{\infty}}{V_{\text{esc}}} \quad 5-12b$$

Thus $B \approx \epsilon$ for most stars as $p \approx 3$, though it is possible for $B \approx 2\epsilon$ for stars with $p \approx 1$.

Figure 5-6 shows equation 5-8 plotted for all the stars in the adopted sample. Lines of constant ϵ are drawn on this graph showing that $\epsilon = 0.7 \pm 0.3$ for $\log \dot{M} \geq 6.3$ and that its value quickly decreases towards 0.1 for lower $\dot{M}$. This indicates that for all stars in the sample, except late O and early B stars, the efficiency of momentum transfer is very high, approaching the single scattering momentum limits of $\epsilon = 1.0$.

Table 5.2 highlights this point by listing the values of ϵ, B, V_{∞} and p for a representative subsample of stars from the adopted data set. Also listed in this table are the number of hypothetical lines which would be required to provide the observed momentum in the wind. These hypothetical lines are optically thick, separated by $\Delta v = v \frac{V_{\infty}}{C}$ and found at the peak of the flux distribution of the star.

Figure 5.6

$\log_{10} \dot{M}$ is plotted against $\log_{10} ((L_*/L_\odot)/V_\infty)$ where V_∞ is the terminal velocity of the stellar wind in kilometres per second. Lines of constant ϵ are shown for values of 0.10, 0.40, 0.60 and 1.0. ϵ is the ratio between the momentum carried away by the stellar wind ($\dot{M} V_\infty$) to the momentum carried away by the photons (L_*/c). The symbols are the same as those used in figure 5.1. Points with a colon next to them are uncertain because the terminal velocity for the star is not accurately known.

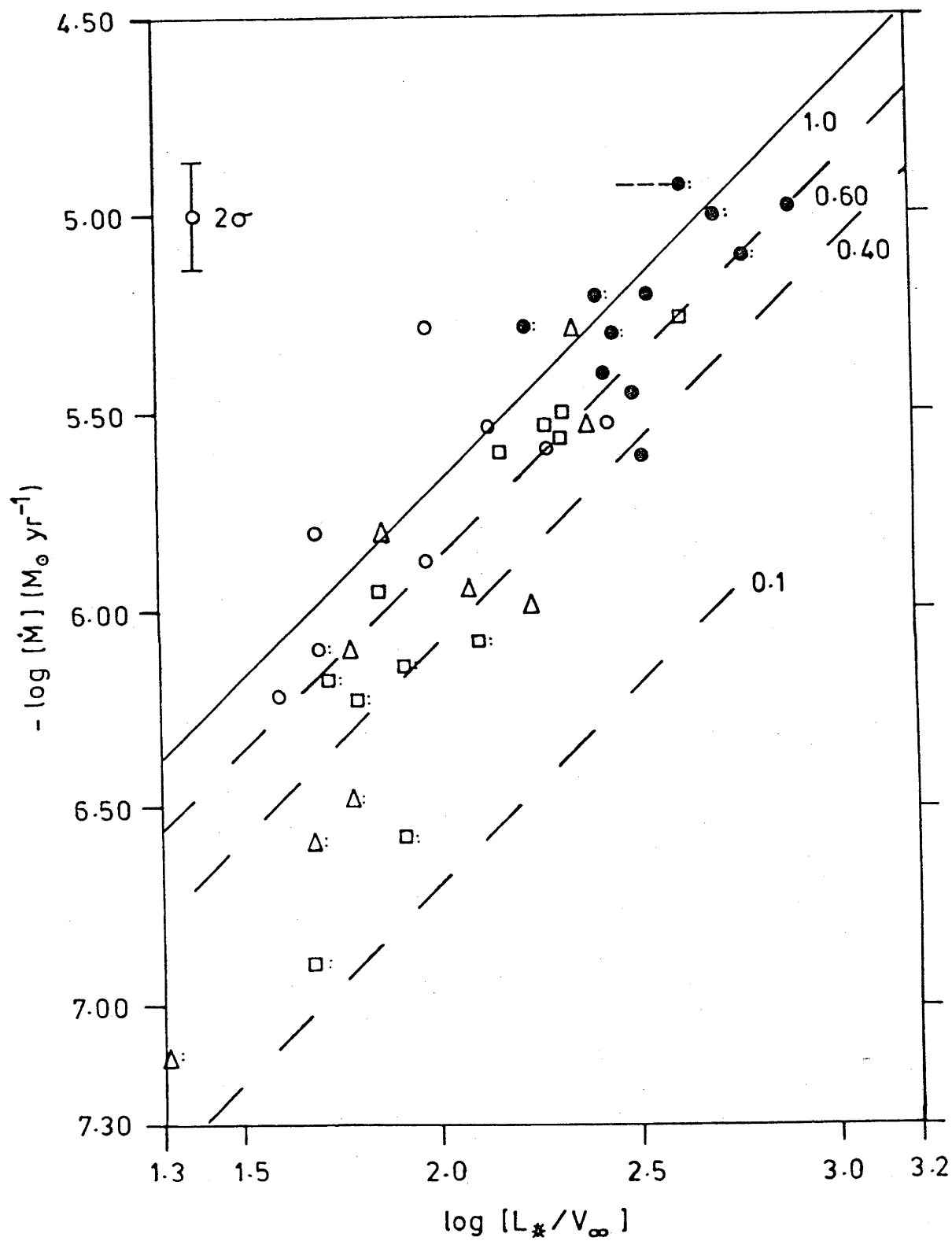


TABLE 5.2

STARS	SP TYPE	ϵ	B	$V_{\infty}/V_{\text{esc}}$	$N=\frac{C}{V}$	V_{∞}	$\frac{C}{V_{\infty}}$
ζ Pup	04f	0.55	0.63	2.585	71.0	2660	112.7
HD152408	08If	0.77	0.88	2.628	146.6	1800	166.5
τ Sco	BoV	0.18	0.26	1.534	38.97	2000:	149.9
ξ Per	BI Ib	0.46	0.51	2.890	101.92	1500	199.8
15 Mon	07V((f))	0.74	0.81	3.100	79.36	3060	97.9
P Cyg	BI Ia ⁺	0.59	0.86	1.474	429.7	600	499.6

Taken at face value Figure 5.6 indicates that the mass-loss formula 5-8 is successful in describing $\dot{M}$ for all O and early B stars simply because there is a small range in the efficiency of momentum transfer from the stellar radiation to the wind. This efficiency also appears to have an upper limit which is given by the single-scattering momentum transfer limit.

This explanation would be sufficient if it were not for a number of problems. Firstly it is reasonable to assume that the major contribution to the momentum in the wind comes from the resonance (and subordinate) lines of the most abundant elements, found near the peak of the stellar flux distribution ($F_{\nu \text{ max}}$). There is a considerable change in the wavelength of $F_{\nu \text{ max}}$ amongst the stars in the adopted sample, varying from $\lambda \sim 1800 \text{ \AA}$ for the early B supergiants ($T_{\text{eff}} \approx 20,000^\circ\text{K}$) to $\lambda \sim 800 \text{ \AA}$ for the hot early O stars ($T_{\text{eff}} \approx 42,000^\circ\text{K}$). If the major line momentum transfer mechanism is produced by single scatterings then Figure 5.6 implies that the net efficiency of this transfer mechanism is the same (to within a factor of 2) for both the cool early B supergiants and the hot early O stars. This fact is hard to reconcile with the claim by Panagia and Macchetto (1980) that the main mode of momentum transfer in stars with $T_{\text{eff}} \geq 30,000^\circ\text{K}$ is multiple scattering, while that in stars with $T_{\text{eff}} \leq 20,000^\circ\text{K}$, is single scattering. They claim that closely spaced optically thick lines

in the far U.V. (λ 200-500 Å) provide the right conditions to produce up to 20 multiple scatterings of photons in the stellar wind of hot stars. This means that the nett efficiency per mean scattering event must be considerably less efficient than the single scattering events in the cool early B supergiants, for ϵ to be roughly the same in the two types of stars. This is difficult to envisage as back scattered photons are the most likely to be involved in multiple scattering processes, raising the mean efficiency of the outward momentum transfer per photon scattering event.

This dilemma could be overcome if multiple scattering was also found in the cool early B supergiants, but this is unlikely because the conditions of having a significant number of closely spaced optically thick lines at the flux peak of the star is not fulfilled (Panagia and Macchetto 1980).

Panagia and Macchetto also claim that the observationally determined values for ϵ (single scattering) do not exceed 0.2 to 0.3. The simple fact that $\epsilon \approx 0.7$ (B, the true line blocking factor is slightly higher) demands that some extra form of momentum transfer must occur in addition to single scattering. This excess momentum could be provided by multiple scattering in hot stars ($T_{\text{eff}} \geq 30,000^\circ\text{K}$) but not the early B supergiants ($T_{\text{eff}} \leq 25,000^\circ\text{K}$). The momentum problem in cool stars

would be overcome if there was some form of mechanical deposition of momentum, in addition to single scattering of photons. The reason why this additional momentum transfer efficiency appears to add to the single scattering transfer efficiency in such a way that the upper bound on ϵ is roughly the single scattering limit is not known. This phenomenon though should provide a clue towards the final understanding of the mode of energy deposition in the wind.

It is possible to obtain equation 5-8 from equation 5-6 by multiplying by $(g_{\text{eff}}/g_{\odot})$, where $g_{\odot}$ is the surface gravity of the sun. It can be shown that $(g_{\text{eff}}/g_{\odot}) * (t_D/t_K)^{1/2}$ is in fact proportional to V_{esc} for the star, reducing equation 5-6 to the simpler form

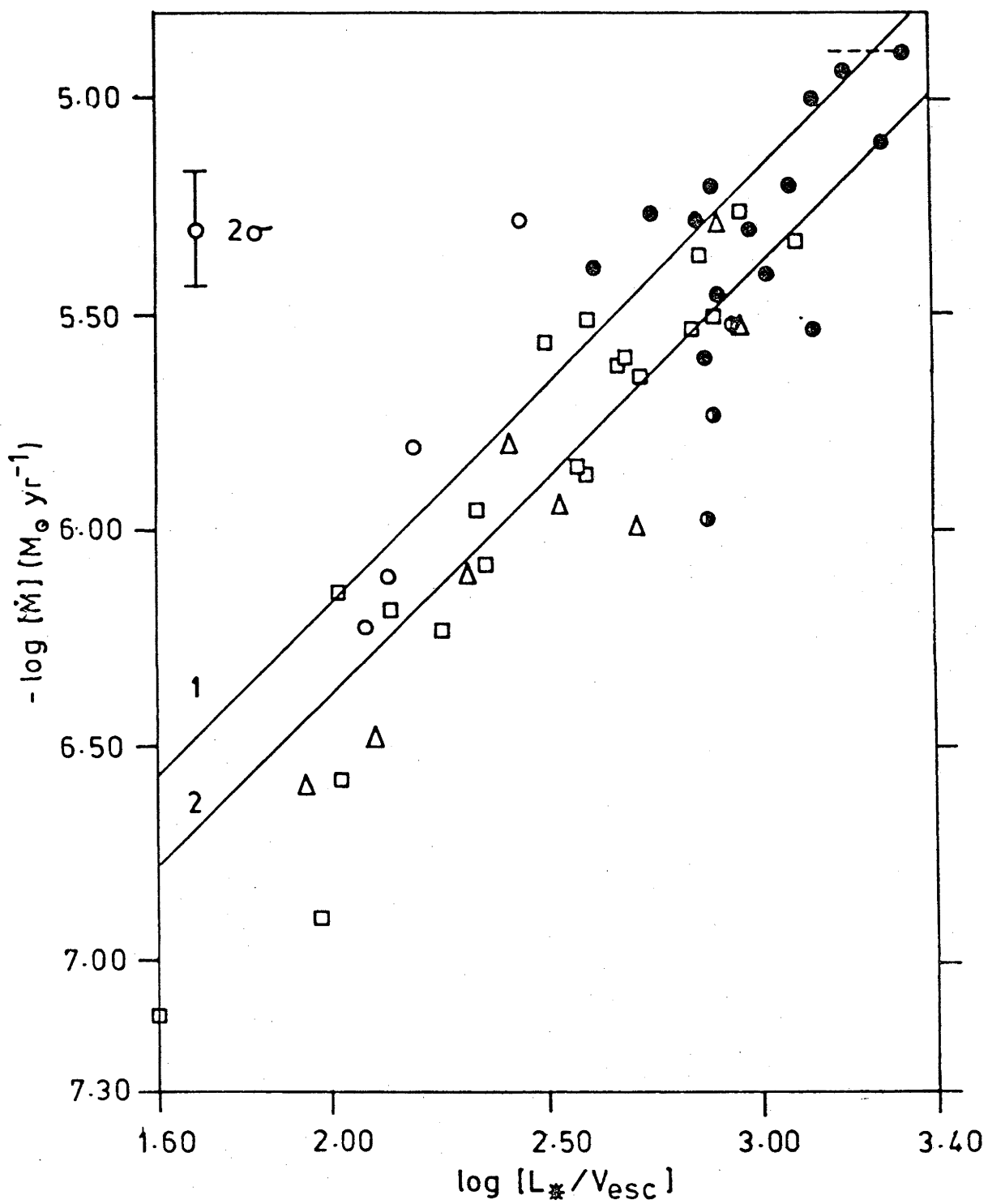
$$\dot{M} = \frac{k L_{*}}{C V_{\text{esc}}} \quad 5-13$$

where the velocity of light C has been incorporated to make both sides of the equation dimensionally equivalent and k is a constant. Equation 5-13 is essentially the same as equation 5-8 since $V_{\infty} \sim 3 V_{\text{esc}}$ for most stars.

Figure 5-7 shows this relationship plotted for all stars in the adopted sample. Again, as in Figure 5-6, there is no splitting between Of and non-Of stars and dwarf and supergiant stars. The scatter in the data of

Figure 5.7

$\log_{10} \dot{M}$ is plotted against $\log_{10} ((L_*/L_\odot)/V_{\text{esc}})$ where V_{esc} is in kilometres per second. The symbols are the same as those used in figure 5.1.



$\sigma_y \approx 0.22$ in the logarithm is comparable to the observational uncertainty of $\dot{M}$ and approximately the same as that in Figure 5.6. Values of k between $1/3$ and $1/6$ encompass the majority of stars, including the late O and early B stars.

Figure 5-6 and 5-7 indicate that it is possible to obtain $\dot{M}$, to within the observational uncertainties, for all O and early B stars, using both equation 5-8 and 5-13, provided $\log \dot{M} \geq -6.3$. Equation 5-13 can be used if the V_∞ for a star is not well determined while equation 5-8 is more appropriate if the mass or radius of the star is uncertain and V_∞ is known.

It is possible to determine $\dot{M}$ for all O and early B stars to within the present uncertainties of measurement using the expressions

$$\dot{M} \approx 0.22 \frac{L_\star}{V_{\text{esc}} C} \quad (\text{kg sec}^{-1}) \quad 5-14a$$

$$\dot{M} \approx 0.7 \frac{L_\star}{V_\infty C} \quad \text{for } \log \dot{M} \geq -6.3 \quad (\text{kg sec}^{-1}) \quad 5-14b$$

This corresponds to assuming that

$$E_T \approx \left(\frac{V_{\text{esc}}}{C} \right) L_\star \quad (\text{J sec}^{-1}) \quad 5-15$$

and indicates that the power driving the wind decreases in proportion to $V_{\text{esc}} \propto (M_*/R_*)^{1/2}$, as massive stars essentially evolve away from the main sequence at constant luminosity. However, the retarding effect of the gravitational potential decreases at a rate $\propto (M_*/R_*)$ and so $\dot{M}$ actually increases at a rate $\propto (M_*/R_*)^{1/2}$.

Neither the line driven wind theory nor the fluctuation theory of Andriesse predicts this form of dependency of $\dot{M}$ on stellar parameters. While this result does not completely rule out these two theories it shows they must be modified to remain viable.

Lamers (1981) showed that it was possible to modify the mass-loss formula predicted by the line driven wind theory to match his data (which is essentially the same as that adopted in this thesis). The line driven wind theory predicts that

$$\dot{M} \propto M_* T_{\text{eff}}^{-1/2} K^{1/\alpha} \Gamma^{1/\alpha} (1 - \Gamma)^{(\alpha-1)/\alpha} \quad 5-16$$

where α and k are adjustable constants and

$$\Gamma = \frac{\sigma_e L_*}{4\pi GCM_*} \quad 5-17$$

(σ_e is the electron scattering opacity/unit mass).

The variables k and α come from the form of the force multiplier used to describe the force produced by line-absorption in terms of the continuum radiation force (CAK). The variable k is a function of density and temperature and α may be a function of radius in the stellar wind envelope. It is still possible considering

the uncertainties which exist in the theory and observation, to obtain values of k and α which make equation 5-16 fit Lamers' data (and thus quantitatively agree with equations 5-14a and 5-14b). Until these parameters are better determined and the errors in measuring $\dot{M}$ are reduced, it is not possible to rule out this theory on the grounds of the functional dependency it predicts for $\dot{M}$ on stellar parameters.

Equation 5-14a will be used in the next chapter to determine $\dot{M}$ and the total kinetic energy in the wind (L_w) for all the stars observed in the LMC HII regions. The integrated value of L_w for these LMC clusters will indicate how important mass-loss is in determining the evolution of these HII regions.

CHAPTER 6

6.1 Introduction

The basic aim of this chapter is to quantitatively determine both the mechanical luminosity of the stellar wind,

$$L_w = \frac{1}{2} \dot{M} V_\infty^2 \quad 6-1$$

and the stellar wind ram pressure

$$P_w(r) = \rho(r) V_\infty^2 = \frac{\dot{M} V_\infty}{4\pi r^2} \quad 6-2$$

for the OB clusters found within the L.M.C. HII regions ($\rho(r)$ is the stellar wind density as a function of radius r). These two parameters were shown in Chapter 2, to be important in determining the role played by mass-loss in the evolution of HII regions.

The three parameters L_* , R_* and M_* are required to calculate $\dot{M}$ and V_∞ . $\dot{M}$ can be obtained from equation 5-14a while V_∞ is calculated from V_{esc} , using the relationship found by Abbott (1980).

The first quantity L_* can be determined for a star, if the Absolute Visual Magnitude (M_V) and the Bolometric correction (B.C.) are known. M_V can be evaluated fairly reliably for stars in the L.M.C. because of the known distance modulus and the low reddening.

Values of M_V used in this chapter are those listed in Table 4.2 and are based on a true distance modulus $\mu_0 = 18.4$ and mean total visual extinction of 0.5 magnitudes (see Chapter 7). The B.C. is dependent on knowing the value

of T_{eff} for the star, which is a relatively uncertain function of its M.K. spectral type (or intrinsic colour).

The second quantity R_* is a function of both L_* and T_{eff} and is given by the relationship (Lamers 1981).

$$\left(\frac{R_*}{R_\odot}\right) = 3.33 \times 10^7 \left(\frac{L}{L_\odot}\right)^{1/2} (T_{\text{eff}})^{-2} \quad 6-3$$

The dependence of L_* on T_{eff} through the B.C. implies that R_* is almost completely dependent on the value of T_{eff} assigned to the star. In view of this dependence, the relationship between M.K. spectral type and T_{eff} for OB stars will be discussed in detail in Section 6.2.

The final parameter M_* , is determined by comparing the position of a star on the H-R diagram with theoretical curves for stars of a given initial mass (mass-loss is assumed). The curves of de Loeres et.al.(1978) for the case $N=100$ (corresponding to moderate rates of mass-loss throughout the core-hydrogen burning lifetime of the star), are used to find the value of M_* .

In section 6.3 current epoch values of T_{eff} and L_* for a star are used, together with the theoretical evolutionary curves of de Loeres et.al. (for $N=100$), to evolve stars backward and forward in time. As the star is evolved, values of L_* , R_* , M_* , M , V_{esc} , V_∞ , L_{uv} (the Lyman photon luminosity of the star), L_w and $MV_\infty (p_w)$ are calculated for 20, 40, 50, 60, 80 and 100 $M_\odot$ stars to indicate how these parameters change throughout the lifetime of a star.

Finally, the values of L_* , R_* , M_* , T_{eff} , M , N_L and V_∞ are tabulated for each star together with the integrated values of N_p , L_w and MV_∞ for the whole cluster.

6.2 The Effective Temperature Scale for OB Stars

Three major methods have been used to construct a T_{eff} scale appropriate to O and early B stars. The first method relies on a comparison of the observed line strengths and line profiles with those predicted by theoretical stellar atmospheric models. Such a comparison was made by Conti (1973, 1975), using the observed HeI $\lambda 4471$ and HeII $\lambda 4542$ line strengths and the predicted line strengths of the unblanketed N.L.T.E. plane-parallel atmospheres of Auer and Mihalas (1972). Conti's temperature scale for dwarfs is shown in Figure 6-1 for the spectral range 04 to 09.5. Conti also found that the effective temperatures of Of and OI stars are about 10% cooler than main sequence stars of the same spectral type. He indicated that this result should be accepted with caution as the effects of atmospheric extension on the line strengths of HeI and HeII are not known.

The second method for determining T_{eff} in OB stars was that used by Underhill et al. (1979) and Rennie and Lamers (1982). Both groups used the method of Blackwell and Shallis (1977) to determine the T_{eff} and angular diameter (θ) of a star. The details of the procedure are summarised in the paper by Rennie and Lamers. In essence,

the method involves the integration, over all wavelengths, of the surface flux emitted by the star F_λ i.e.

$$\frac{\sigma T_{\text{eff}}^4}{\pi} = \int_0^\infty F_\lambda d\lambda \quad 6-4$$

The two groups used published absolute photometric observations of the U.V. and visual continuum of stars, and the method of Blackwell and Shallis, to determine F_λ in the wavelength range $1380 \text{ \AA} \leq \lambda \leq 11084 \text{ \AA}$. Outside this range they used the blanketed plane-parallel L.T.E. models of Kurucz (1979) to determine the unseen flux.

The flux temperatures derived by these two groups are generally 1500K lower than those derived by Conti (1975) for stars in common to both samples. Part of this discrepancy in T_{eff} can be explained by the model atmospheres used by the flux temperature groups. The fact that the Kurucz atmospheric models do not take into account the effects of atmospheric sphericity and N.L.T.E. conditions, both of which lead to increases of the Lyman continuum flux (Panagia 1975, Mihalas and Hummer 1974), means that the flux temperatures will be lower than the true effective temperatures. Atmospheric extension should only be a serious problem for the low gravity hot Of stars (i.e. spectral type O4f to O6f). On the other hand, N.L.T.E. conditions were shown by Panagia (1975) to be an important factor in determining the total Lyman continuum flux for all supergiant OB stars, especially late O and early B supergiants. Since all O and early B stars emit

a considerable fraction of their flux shortward of 1380 Å, the correction factor used by Underhill et al. and Rennie and Lamers to determine T_{eff} will be an underestimate both for the hot O stars and the late O and early B supergiants.

The T_{eff} values given by Underhill et al (1979) for non-supergiant (luminosity classes V, IV, III and II) stars of spectral type O4, O6 and O8 to B3, plus supergiant stars of spectral type B0, B1 and B2, are shown in Figure 6.1. The temperatures for spectral classes O6 and O4 are those for λ Cepheus (O6ef) and ζ Puppis (O4f) respectively. Also shown in Figure 6.1 are the T_{eff} values given by Rennie and Lamers for non-supergiant stars of spectral type B0 to B3 and supergiant stars of spectral type O9.5 to B2.

The final method which has been used to determine T_{eff} in O and early B stars is based on the Zanstra technique described in Section 4.3. The total Lyman continuum photon flux (N_L), obtained from observations of HII regions using the Zanstra method, is compared to that predicted by theoretical stellar atmospheric models at a given T_{eff} .

Georgelin et al. (1975) used this method to determine a T_{eff} scale for dwarf stars. They compared their observed values of $\frac{N_L}{\pi F_V}$ for 45 HII regions with the values predicted by the blanketed plane-parallel L.T.E. models of Morton (1969) for $T_{\text{eff}} \leq 37,500$ K, and the unblanketed, plane-

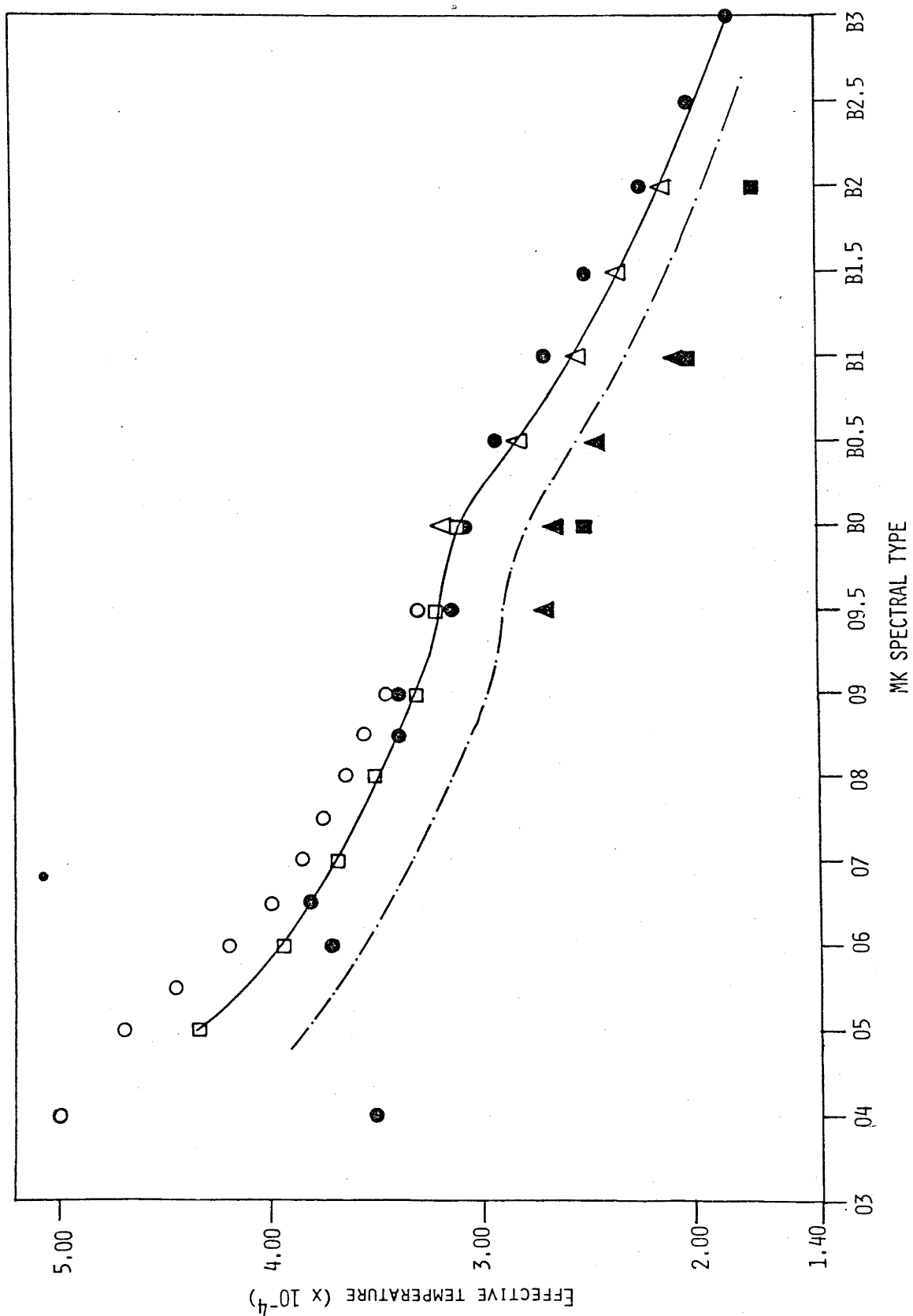
parallel, N.L.T.E. models of Auer and Mihalas for $T_{\text{eff}} \leq 37,500^{\circ}\text{K}$. Panagia (1975) showed that line blanketing has a more important effect upon N_L than N.L.T.E. conditions for dwarf stars. Thus the blanketed L.T.E. models should be adequate for all but the hottest dwarf where N.L.T.E. effects and atmospheric extension may affect N_L . In addition to using the most appropriate atmospheric models, Georgelin et al. also carefully selected their HII regions so that they would be suited to the Zanstra method. This was done to reduce the known sources of error which are associated with this technique.

The T_{eff} values obtained by Georgelin et al. are probably under estimates of the true values of the effective temperature as no allowance is made for absorption of ionizing photons by dust, or for extended low surface brightness components of the HII regions, which were undetected by their radio observations.

The T_{eff} scale given by Georgelin et al. for O5 to B0.5 dwarf stars is also shown in Figure 6.1. It can be seen that the Zanstra temperatures agree remarkably well with the flux temperatures of Underhill et al., except for the stars ζ Pupis and λ Cepheus which have considerably lower values of T_{eff} . The flux temperatures of these two stars are almost certainly depressed by atmospheric extension. The Zanstra scale is also systematically cooler than Conti's line temperature scale, though this difference is only marginally significant because of the

Figure 6.1

The effective temperature scale as a function of M.K. Spectral type for O4 to B3 stars. O: Conti (1975); □: Georgelin et. al. (1977); ●: Underhill et.al. (1979); dwarf stars; Δ, ▲: Rennie and Lamers (1982) dwarf and supergiant stars respectively; ■ Underhill et. al. (1979) supergiant stars. The solid line shows the adopted effective temperature scale for dwarf and the broken line the corresponding scale for supergiants.



large dispersion in the Zanstra scale. The discrepancy may be due to the blanketing by the far U.V. lines in the wind (Rennie and Lamers 1982). The stellar wind envelope becomes opaque at wavelengths below $1200 \text{ \AA}$ because of the crowding of U.V. resonance lines (Lamers and Morton 1976). The opaque scattering stellar wind acts as a blanket, scattering a substantial fraction of the photons with $\lambda \leq 1200 \text{ \AA}$ back into the photosphere. These back scattered photons will raise the boundary temperature of the star where the HeI $\lambda 4471$ and HeII $\lambda 4542$ line forming regions are found. This is probably why the line temperatures are elevated above the Zanstra and flux temperatures for dwarfs.

The preceeding discussions indicate that it is only possible to obtain reliable T_{eff} values for dwarf O and early B stars. Atmospheric extension and N.L.T.E. conditions which are found in the supergiant stars, make any temperature scale based on presently available theoretical models highly uncertain. The good agreement between the Zanstra and flux temperature scales for dwarfs and the probability that the theoretical models adequately describe the atmospheres of these stars, indicates that the temperatures for these stars are reasonably well determined.

The Zanstra temperature scale was adopted for dwarf stars of spectral type O5 to B0, as this scale is more comprehensive than the scale of Underhill et al.

in this spectral range. The temperature scale of Rennie and Lamers was adopted for B0.5 to B2 dwarfs while the scale of Underhill et al was adopted for B2.5 to B6 dwarfs. Rennie and Lamers pointed out that the iterative procedure used to calculate the flux temperatures was not taken far enough by the Underhill group, resulting in their temperatures being too high by about 1500 K for stars with $T_{\text{eff}} \geq 25,000$ K. Recognising this, we have adopted the values of Rennie and Lamers for the temperature range 21,000 K (B2) to 29,000 K (B0.5).

The temperature scale for supergiants was taken as being 10% cooler than that for dwarfs, following the results obtained by Conti (1975), while that for giants (luminosity class III) was taken as being halfway between these two scales. The flux temperature scale for supergiants was not adopted for reasons cited previously.

The final temperature scale for dwarfs is shown in Figure 6-1 by a solid line, and that for supergiants by a dotted line. Table 6.1 lists these temperatures for each spectral type and luminosity class. Also listed in this table are the intrinsic colours associated with each T_{eff} . These are derived from the relationship given by Flower (1977) between T_{eff} and $(B-V)_0$.

The formal uncertainty in the adopted Zanstra temperature scale ranges from $\sigma_T \approx 3500$ K near spectral type 06 to $\sigma_T \approx 800$ near spectral type 09.5 (Georgelin

TABLE 6.1

Spectral Type	V	III	I	V	III	I
	T_{eff}			B-V		
05	43,000	40,600	38,200	-0.36	-0.34	-0.32
05.5	41,200	39,000	36,800	-0.36	-0.33	-0.31
06	39,400	37,400	35,400	-0.35	-0.32	-0.30
06.5	38,100	36,100	34,100	-0.34	-0.31	-0.29
07	36,800	35,100	33,400	-0.33	-0.31	-0.29
07.5	35,800	34,050	32,300	-0.33	-0.30	-0.28
08	34,800	33,050	31,300	-0.32	-0.30	-0.27
08.5	33,900	32,150	30,400	-0.32	-0.29	-0.26
09	33,000	31,250	29,500	-0.31	-0.29	-0.26
09.5	32,100	30,500	29,000	-0.31	-0.28	-0.25
B0	30,900	29,400	27,800	-0.30	-0.27	-0.24
B0.5	28,200	26,800	25,400	-0.28	-0.26	-0.22
B1	25,500	24,200	23,000	-0.26	-0.24	-0.20
B1.5	23,300	22,150	21,000	-0.24	-0.22	-0.18
B2	21,400	20,350	19,250	-0.23	-0.20	-0.16
B2.5	20,100	19,100	18,100	-0.21	-0.19	-0.14
B3	18,100	17,200	16,300	-0.19	-0.17	-0.12
B5	15,000	14,300	13,500	-0.16	-0.14	-0.08
B6	13,500	12,800	12,200	-0.14	-0.11	-0.05

et al. 1975). The uncertainty at 06 is almost certainly an overestimate of the true uncertainty. The difference between the Conti line temperature and the Underhill group's flux temperature at this spectral type should provide a more reasonable estimate of the bounds ($2\sigma_T$) of the true uncertainty, i.e. $\sigma_T = 1500^\circ\text{K}$. Rennie and Lamers claimed an uncertainty σ_T of 1000 to 1500 K for their dwarf temperature scale. We have adopted $\sigma_T = 1500\text{K}$ as the uncertainty, and consider this to be a conservative value for all spectral types.

6.3 Evolutionary Models for OB Stars

The luminosity of a star, derived from its absolute magnitude and bolometric correction, together with its T_{eff} , were used to locate the star in relation to the evolutionary tracks of de Loore, de Greve and Vandeveren (1980). These models consist of evolutionary curves which give T_{eff} , R_* , M_* and L_* as a function of stellar age, for stars with initial (zero-age main sequence) masses of 20, 30, 40, 50, 60, 80 and 100 $M_\odot$. The mass-loss rate during this evolution is of the form $\dot{M} = \frac{NL_*}{c^2}$ where $N=0, 100$ and 300. Once the location of the star is established with respect to the evolutionary curve, linear interpolation in the T_{eff} , L_* plane is then used to determine the current epoch mass and age, as well as the fractional distance of the star between the upper and lower evolutionary curves. The fractional distance between

these two curves is used to locate the initial position of the star on the zero-age main sequence. The star is then evolved from this point in time increments of 0.14×10^6 years and the variables L_* , R_* , M_* and T_{eff} are determined at each time step. These variables are calculated by linearly interpolating in the age, variable plane and they are used to calculate $\dot{M}$, V_{esc} , V_{∞} , $\frac{1}{2} \dot{M} V_{\infty}^2$ and g_{eff} . The evolution is stopped when the age is less than two grid points (the age step used by de Loore et.al.) from the end of the upper evolutionary curve.

The Lyman continuum photon emission rate (L_{uv}) is also calculated at each time step using the relationship

$$L_{\text{uv}} = 1.231 \times 10^{28} 10^{-0.4M_V} N_L / \pi F_V \text{ photons/sec} \quad 6-5$$

M_V values were determined from L_* by using T_{eff} values at each time step to calculate the bolometric correction (B.C.). The relationship between T_{eff} and B.C. given by Lamers (1981) was adopted, i.e.

$$\text{B.C.} = -5.475 \log T_{\text{eff}} + 21,728 \quad \text{for } \log T_{\text{eff}} \leq 4.52 \quad 6-6a$$

and

$$\text{B.C.} = -7.778 \log T_{\text{eff}} + 32.056 \quad \text{for } \log T_{\text{eff}} > 4.52 \quad 6-6b$$

Those expressions give the B.C. to the nearest 1/10 of a magnitude.

A table relating M.K. spectral type with T_{eff} and M_V (the values given in Figure 3.1 were used here) was

used to determine the fractional M.K. spectral type of a star (i.e. 08.7 or 06.3 etc). This was then used to obtain $\log N_L / \pi_{F_V}$ by linear interpolation of the $\log (N_L / \pi_{F_V})$, spectral relationship given in Table 3.1.

If the mass-loss formula used in the evolutionary models of de Loore et al. is compared with the formula given by equation 5-14(b) we find

$$N = 0.7 (\pm 0.3) (C / V_\infty) \quad 6-7$$

Thus the value of N is dependent on the mass of a star and its age. Indeed, as the star evolves N will increase as V_∞ decreases. The values of V_∞ in OB stars range from 1000 to 3500 km/sec with typical values lying around 2,500 km/sec. These values of V_∞ correspond to values of N in the range 220 to 60 with a typical N being $84 (\pm 36)$. Thus the models of de Loore et al. for $N=100$ were chosen to do the evolutionary calculations as they are the closest to reality.

6.4 Results of the Evolutionary Models

Figures 6.2 a, b, c and d show respectively: $\log \dot{M}$ ($M_\odot \text{ year}^{-1}$); $\log L_{UV}$ (photons sec^{-1}); $\log \dot{M} V_\infty$ (Newtons) and $\log L_w = \log (\frac{1}{2} \dot{M} V_\infty^2)$ (Joules sec^{-1}) as a function of stellar age, for initial masses of 20, 30, 40, 50, 60, 80 and 98 $M_\odot$. Also shown in each graph is the curve for a star of initial mass of 98 $M_\odot$, in which the conservative evolutionary models of de Loore are used. The difference

Figure 6.2a

$\log_{10} \dot{M}$ is plotted against stellar age in millions of years for stars with initial masses of 20, 30, 40, 50, 60, 80 and $98 M_{\odot}$. The evolutionary models of de Loore et. al. (1978) for the case $N=100$, are used to calculate the four parameters plotted in figures 6.2 a to d. Solid curves are drawn for each initial mass. The broken curve is that for a $96 M_{\odot}$ star which is evolving with constant mass ($N = 0$). This is shown to exemplify the effect of the choice of evolutionary model.

Figure 6.2b

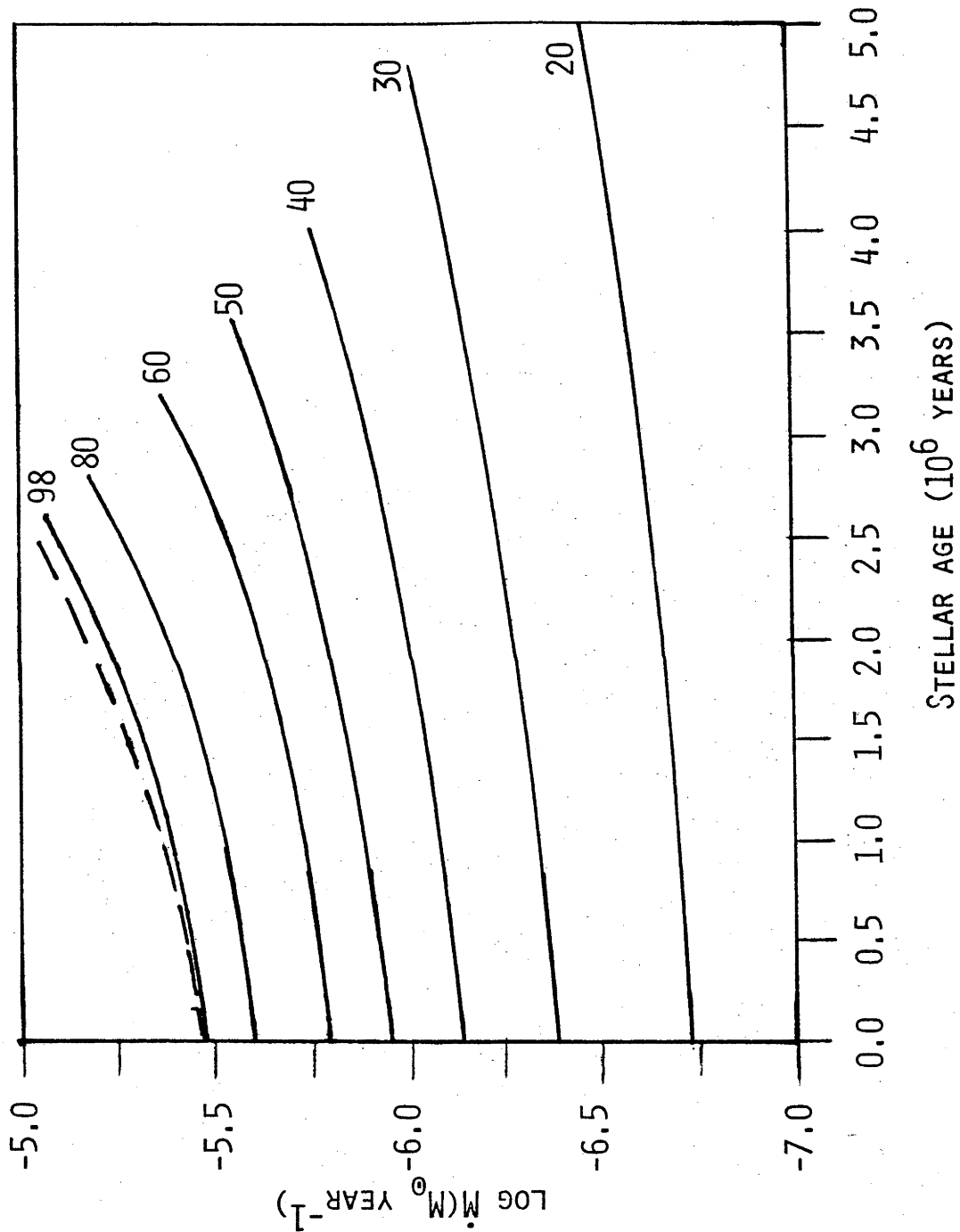
$\log_{10} L_{uv}$ (photons sec^{-1}) is plotted against the stellar age in million of years as in figure 6.2a.

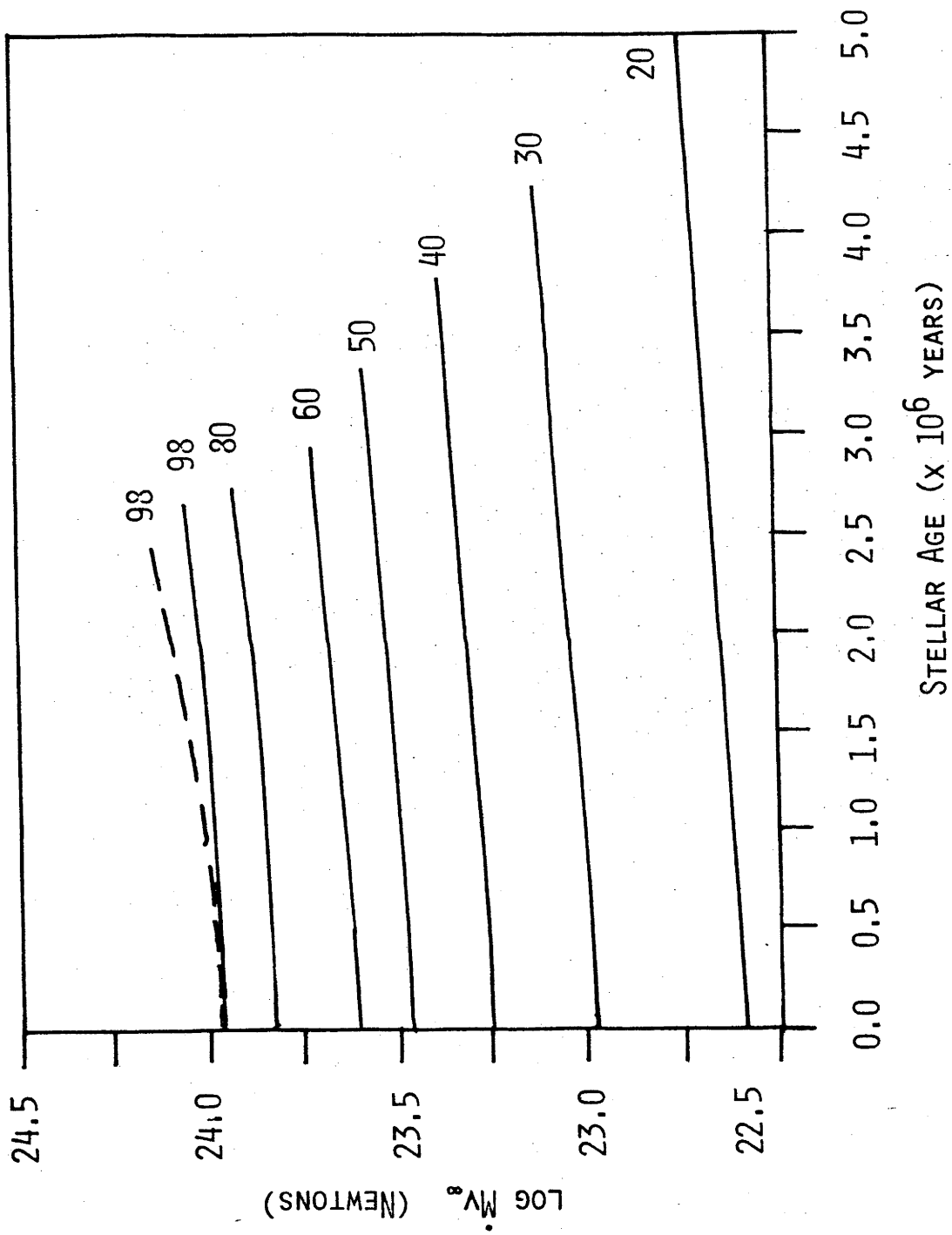
Figure 6.2c

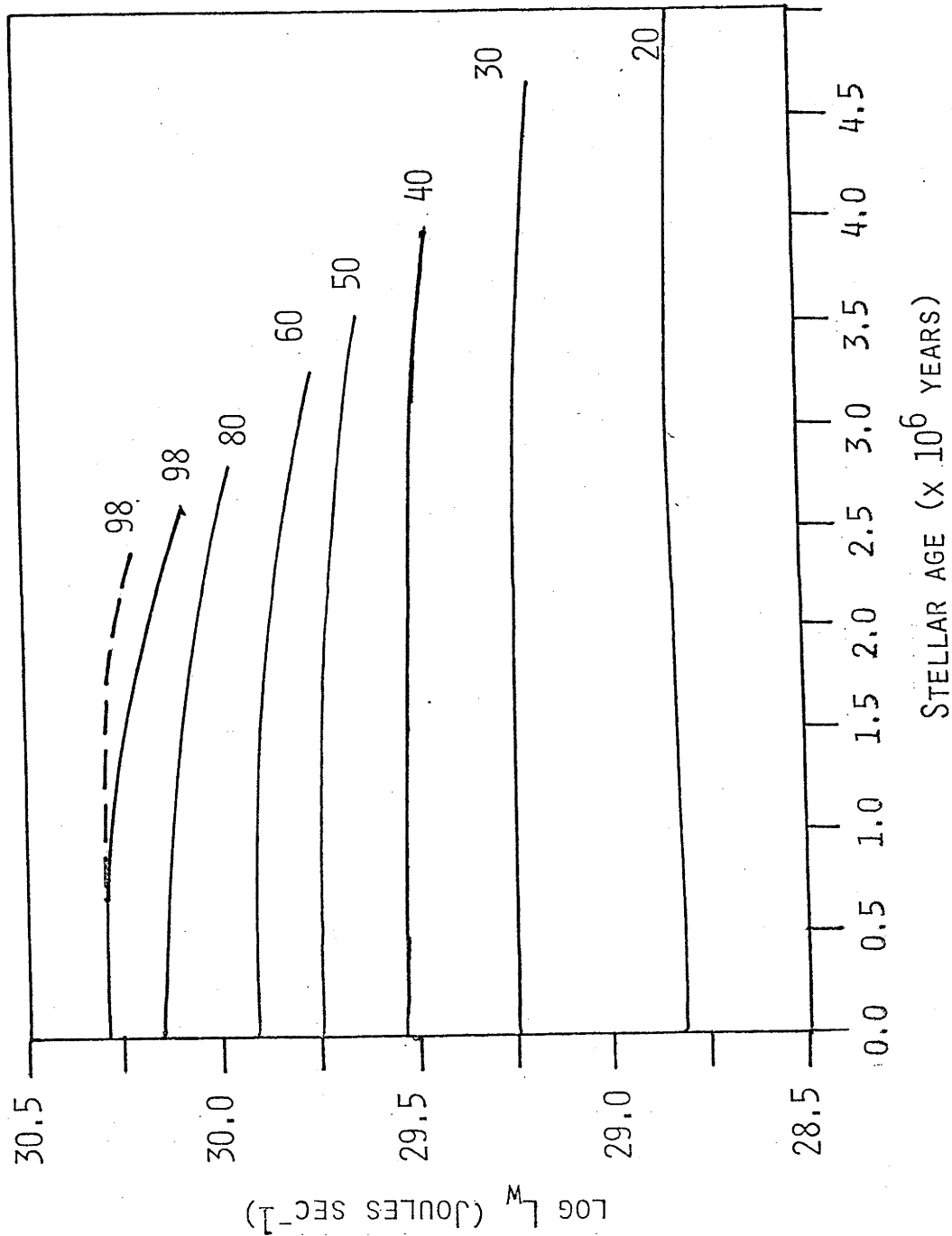
$\log \dot{M} V_{\infty}$ (Newtons) is plotted against the stellar age in millions of years as in figure 6.2a.

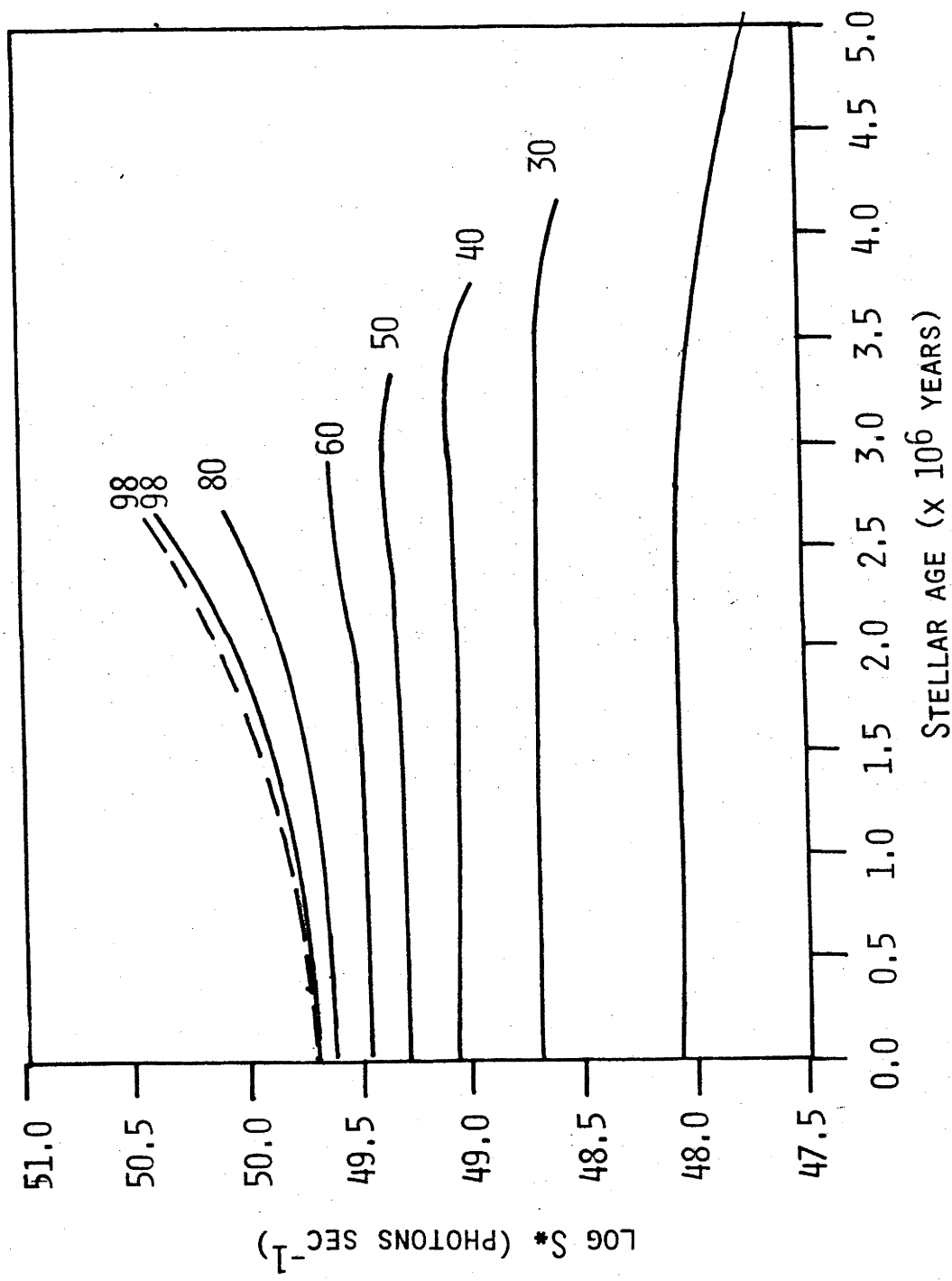
Figure 6.2d

$\log L_w$ (Joules sec^{-1}) is plotted against the stellar age in millions of years as in figure 6.2a.









between the $98 M_{\odot}$ curves for the conservative and non-conservative mass-loss models shows the maximum effect that the choice of evolution model has on the final result.

The main characteristics of mass-loss, during the core-hydrogen burning phase of the star, are brought out in Figures 6.2 a, b and c. Fig. 6.2a shows that $\dot{M}$ increases by a factor of approximately 2.5 during this evolutionary phase. The rate of increase in $\dot{M}$ is thus a function of the hydrogen burning lifetime of a star i.e. its mass. A $100 M_{\odot}$ mass star on the zero-age main sequence is losing mass at a rate which is 18 times faster than a $20 M_{\odot}$ star in the same evolutionary phase. This difference in $\dot{M}$ almost doubles by the end of the $100 M_{\odot}$ stars' hydrogen core-burning lifetime as the change in the $20 M_{\odot}$ star's $\dot{M}$ is small over this time interval.

Figure 6.2b shows that the momentum of the stellar wind ($\dot{M}V_{\infty}$) increases slightly during the core-hydrogen burning lifetime of a star, while Figure 6.2c shows that the mechanical luminosity of the wind decreases as the star evolves. The rate of decrease of L_{uv} is much more pronounced in the star with initial masses $\geq 50 M_{\odot}$.

The trends shown in Figures 6.2a, b and c simply reflect the time dependence of $\dot{M}$ and V_{esc} ($\propto V_{\infty}$) for O and early B stars. The important constraint on these evolutionary trends is the universality of the mass-loss formula 5-14a. This basically says that the momentum carried away by the stellar wind is a constant fraction of the momentum carried

away by the stellar photons.

This fact leads to the following time dependencies for the three parameters shown in Figures 6.2a to c.

$$\dot{M} \propto \frac{L_{\star}}{V_{\infty}} \propto \frac{L_{\star}}{V_{\text{esc}}} \propto L_{\star} (R/M)^{1/2} \quad 6-8a$$

$$\text{while } \dot{M}V_{\infty} \propto \dot{M}V_{\text{esc}} \propto L_{\star} \quad 6-8b$$

$$\text{and } \dot{M}V_{\infty}^2 \propto \dot{M}V_{\text{esc}}^2 \propto L_{\star} (M/R)^{1/2} \quad 6-8c$$

During the immediate post core-hydrogen burning phase a star evolves with roughly constant luminosity (de Loore et al. 1978), and so in this phase $\dot{M}$ will increase as $(R/M)^{1/2}$, $\dot{M}V_{\infty}$ would be constant and $\dot{M}V_{\infty}^2$ would decrease as $(M/R)^{1/2}$. This is not true though for the hydrogen-core burning phase where $L_{\star}$ increases significantly. The increase in $L_{\star}$ is a function of the initial mass of the star, with a $20 M_{\odot}$ star increasing in brightness by 58%, a $40 M_{\odot}$ star by 45% and a $100 M_{\odot}$ star by 26%. These changes in luminosity modify the constant luminosity trends indicated above, producing the results shown in Figures 6.2a to c.

Figure 6.2d shows the Lyman continuum emission rate (L_{uv}) for stars of a given initial mass. L_{uv} is basically constant during the hydrogen core-burning phase of a star with the possibility of a marginal downturn towards the end of this evolutionary stage. The 80 and 98 $M_{\odot}$ curves are exceptions to this rule with L_{uv} increasing substantially

at the end of the main sequence lifetime. This result should be accepted with some caution however, because these stars have effective temperatures which are in the uncertain high temperature region of the adopted T_{eff} scale.

6.5 Conclusions

Two interesting conclusions can be drawn from the evolutionary models described in Section 6.3. The first conclusion deals with the splitting which exists between Of and non-Of stars in Figure 5.4 while the second is concerned about the Conti Scenario (Conti 1976) for the transition of Of stars into Wolf-Rayet stars.

6.5.1 The distinction between Of and non-Of stars in the adopted sample

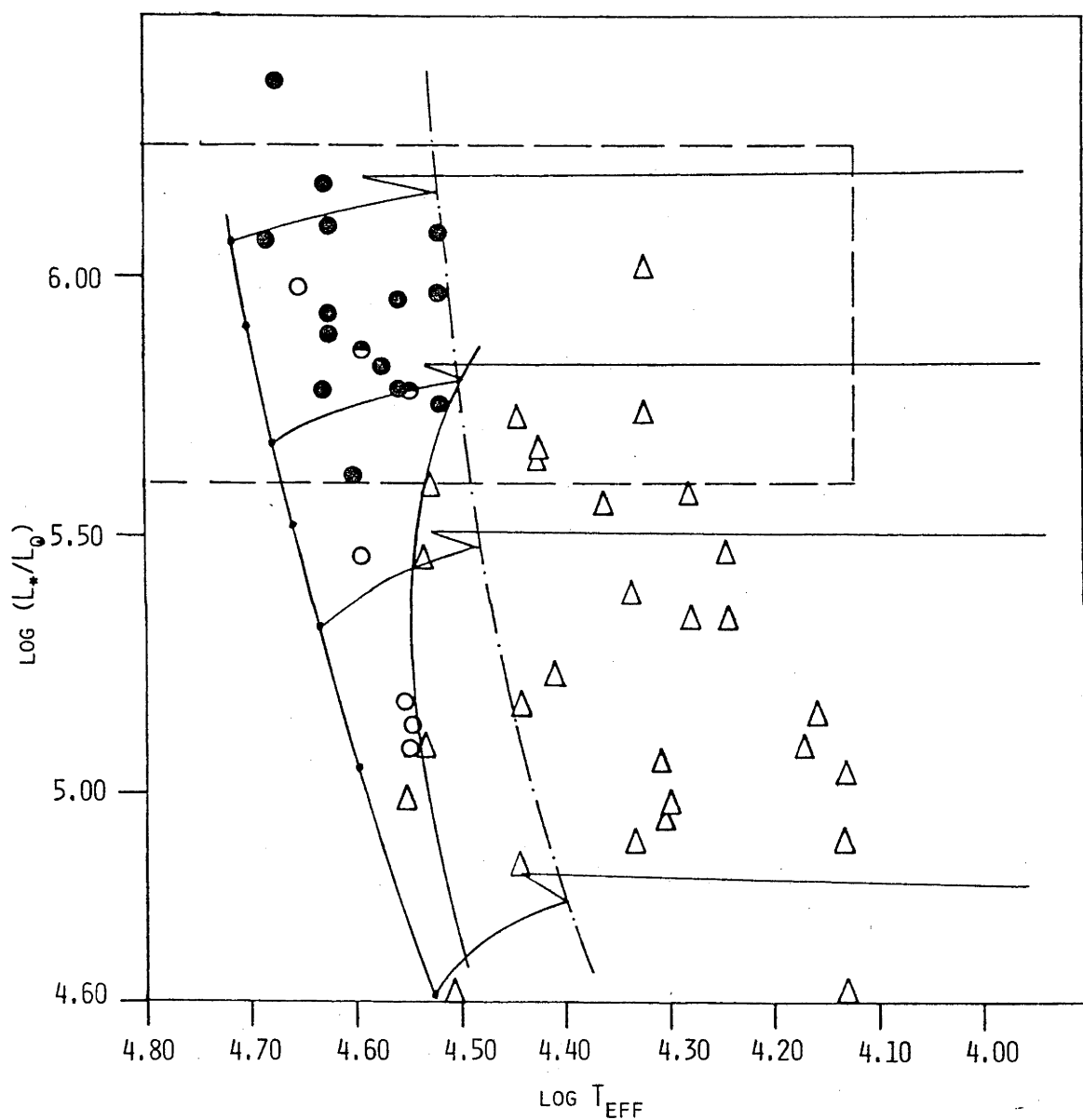
If the mass-loss rate is described by equation 5-14a, the ratio of the mechanical luminosity of the stellar wind to the luminosity of the star (α), decreases as $(M/R)^{1/2}$ irrespective of the stars luminosity (see equation 6-8c). This could be used to explain the splitting which exists between Of and non-Of stars in Figure 5-4 found in the last chapter. This diagram is a plot of $\log \dot{M}$ against $\log (L_*/v_{\text{esc}}^2) \propto \log (\frac{L_*}{v_{\infty}^2})$ and it shows that the ratio is systematically higher for Of stars when compared to non-Of stars. That is to say the driving mechanism producing the winds in Of stars is more efficient than those in non-Of stars at the same $\dot{M}$ ($\propto$ luminosity).

Figure 6.3 is a H-R diagram of the stars found in the adopted mass-loss sample of Chapter 5. The evolutionary curves for stars with initial masses of 20, 40, 60 and 100 $M_{\odot}$, from the models of de Loore et al. (1978) (for $N=100$), are also shown in Figure 6.3. The solid vertical curve at the left of the figure is the zero-age main sequence. Along this line the positions of 20, 30, 40, 50, 60, 80 and 100 $M_{\odot}$ stars are marked. The almost vertical dashed line shows the maximum redward extent of core-hydrogen burning, while the dashed box in the top left-hand corner of the graph shows the expected core-helium burning region for stars with masses $\geq 55M_{\odot}$ (Maeder, 1981, using his case B model with $\langle N \rangle = 65-80$). The solid line to the right of the main sequence curve represents the 3.2 million year isochrone. This isochrone forms a dividing line between Of and non-Of stars for stars with initial masses $\geq 30 M_{\odot}$.

The actual dividing line which determines whether a star is Of or non-Of is a function of the factors contributing to the geometrical extension of the atmosphere, which is thought to produce the Of phenomena. It is only fortuitous that this transition roughly corresponds to the 3.2 million year isochrone in Figure 6.3. What this subdivision does tell us though, is that the splitting in figure 5-4 can be explained by the fact that Of stars are systematically younger than the non-Of stars in the adopted sample. Since the Of

Figure 6.3

This figure shows an H-R diagram for the OB stars in the adopted sample. The evolutionary curves for stars with initial masses of 20, 40, 60 and 100 $M_{\odot}$ are shown for the models of de Loore et. al. (1978) ($N=100$). The solid (near vertical) curve at the left of the figure is the zero age main sequence. Along this line the positions of 20, 30, 40, 50, 60, 80 and 100 $M_{\odot}$ stars are marked. The almost vertical dashed line shows the maximum redward extent of core-hydrogen burning, while the dashed box in the top left hand corner of the graph shows the expected core-helium burning region for the stars with masses $> 55 M_{\odot}$ (Maeder 1981, $\langle N \rangle = 65-80$). The solid line to the right of the main sequence curve represents the 3.2 million year isochrone. $\bullet$: Of stars; $\odot$: O(f); $\circ$: O((f)); Δ : non-Of stars.



stars are younger, the efficiency of the mechanism driving the stellar wind is systematically higher than for non-Of stars at the same $\dot{M}$. This would support the generally accepted view that Of stars are core-hydrogen burning rather than core-helium burning.

6.5.2 The Transition of Of Stars into W-R Stars

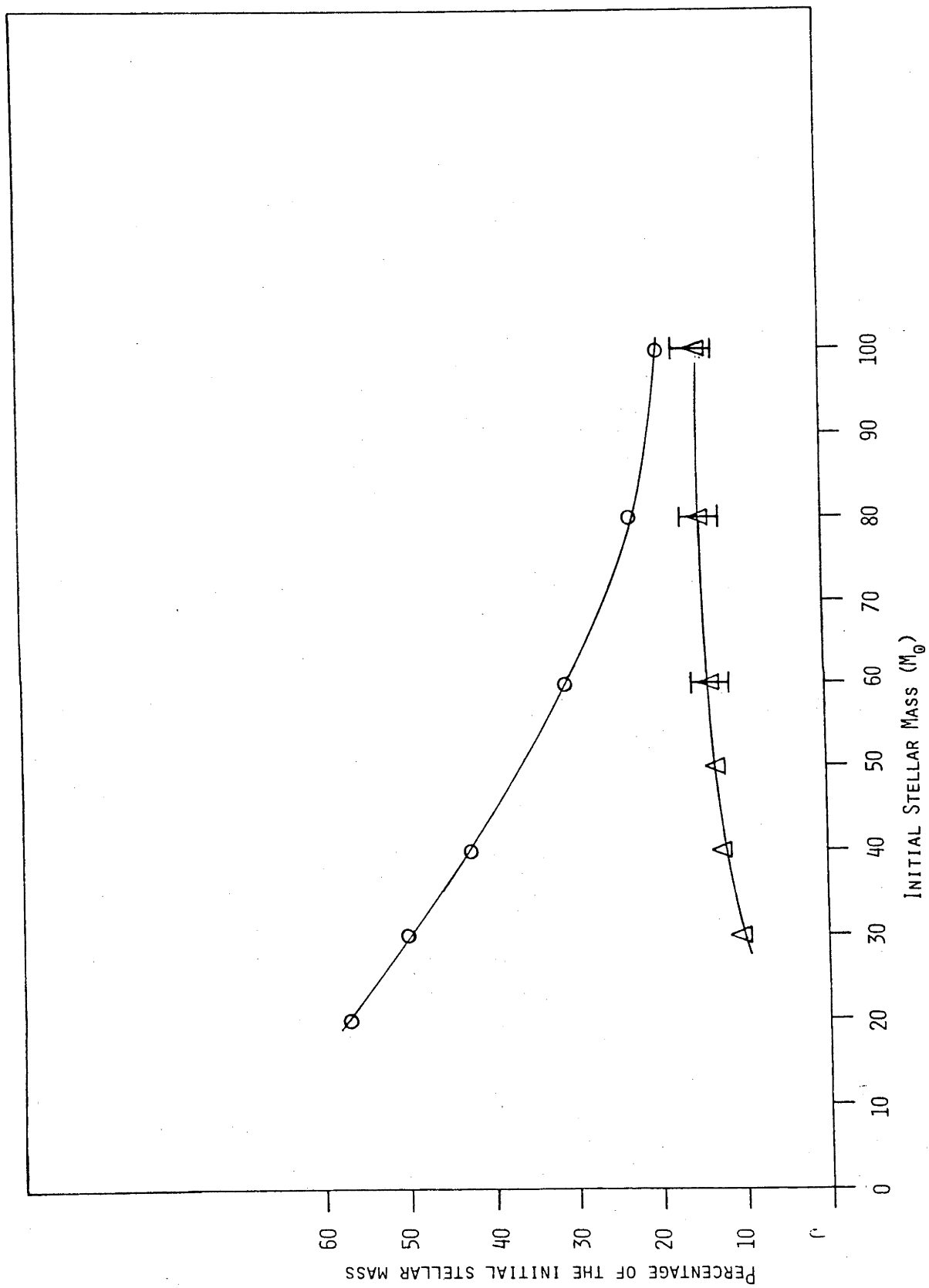
The Conti Scenario for the formation of single W-R stars (Conti 1976) hypothesises that it is possible for single massive Of stars to evolve into late WN stars, if $\dot{M}$ is sufficient to expose the nitrogen enriched material of the core.

Noels et al. (1980) calculated the minimum amount of mass which would have to be removed from a star to expose the core. These masses, expressed as a percentage of the star's initial mass, are shown as a function of the initial mass in Figure 6-4. Also shown in this graph is the mass lost (as a percentage of the initial mass) during the core-hydrogen burning lifetime, predicted by the model given in Section 6.3. The error bars for the 60, 80 and 100 $M_{\odot}$ stars are obtained by assuming that the uncertainty (1 σ) in $\log \dot{M}$ at each time step is ± 0.3 .

It can be seen from Figure 6-4 that to expose its core a star would have to at least have an initial mass in excess of 85 $M_{\odot}$ (allowing for a 3 σ uncertainty in the predicted main sequence mass-loss and for the fact that the given mass-losses required to expose the core are lower bounds).

Figure 6.4

The upper curve in this figure shows the minimum percentage of the initial mass of the star which must be removed according to Noels et.al. 1980, in order to expose the enriched core material. This is plotted against the initial mass of the star. The lower curve shows the actual percentage of the initial mass of the star which is lost during the core-hydrogen burning lifetime, as a function of stellar initial mass. One sigma error bars are shown on the last three points of this curve, showing that only stars with initial masses $\gtrsim 85 M_{\odot}$ can lose enough matter to expose their core material, in their hydrogen-core burning lifetimes.



This result is subject to systematic errors in the mass-loss rates. Since these rates are based on the Radio method for our massive stars, the systematic errors are those of this method. If the mass-loss rates are systematically underestimated by a factor of two then the mass-limit will be reduced to roughly $55 M_{\odot}$, while if they are over estimated by this same factor, then the core would never be exposed in the main sequence lifetime. Since at least 90% of the mass-loss by a massive star ($\geq 40 M_{\odot}$) is lost in the main sequence phase (Brunish and Truran 1981), it would be difficult to imagine that the mass-loss rates are much lower than those given in our models. If the mass-loss rates were a factor of 2 lower for example, then no single W-R stars would form unless some form of catastrophic mass-loss occurred in the post main sequence phase. W-R stars would then be formed by the Maeder process (Maeder 1981). In this process W-R stars are supposed to be formed in the post red supergiant phase, with rapid "mass-ejection" of the star's envelope exposing the core.

Our result indicates that it is possible for a single WN star to form from an Of star while the star is still hydrogen-core burning, provided its initial mass is $\geq 85 M_{\odot}$. Below this threshold mass W-R stars are presumably formed in the post-red supergiant stage of evolution as in the Maeder process.

6.5.3 The Density in the Stellar Wind Envelope of Of Stars

A second distinction which occurs between Of and non-Of stars is highlighted by Figure 6-5. This figure shows $\log \dot{M}$ plotted against $\log (V_{\infty} R_{*}^2)$ for the stars in the adopted mass-loss sample. Both Of and non-Of stars appear to have $\log \dot{M}$ varying in proportion to $\log V_{\infty} R_{*}^2$ but Of stars have a distinctly higher $\dot{M}$ than non-Of stars at the same value of $\log V_{\infty} R_{*}^2$.

Gathier, Lamers and Snow (1981) showed that $\dot{M}$ could be represented by the empirical expression

$$\dot{M} = \text{constant} \times \left(\frac{N(NV)}{R_{*}} \right)^{0.375} \frac{R_{*}^2 V_{\infty}}{A_N} M_{\odot} \text{ year}^{-1}$$

6-9

where $N(NV)$ is the column density of NV between the velocity points $V=0.2V_{\infty}$ and V_{∞} in the wind and A_N is the elemental abundance of nitrogen with respect to hydrogen. This tells us that the splitting in Figure 6-5 is primarily due to the fact that $\left(\frac{N(NV)}{R_{*}} \right)$ is larger in Of stars than in non-Of stars. This ratio is proportional to the stellar wind density and so Of stars appear to have distinctly higher densities in their stellar wind envelopes than non-Of stars. This supports the idea that the Of phenomenon is produced by an extension of the stellar atmosphere.

Figure 6.5

This figure shows a plot of $\log \dot{M} (M_{\odot} \text{ year}^{-1})$ against $\log (V_{\infty} R_{*}^2)$ where R_{*} is the stars' radius in solar radii, and V_{∞} is the terminal velocity of its stellar wind in kilometers per second. The symbols used are the same as those drawn in Figure 5.1. One sigma error bars are drawn in a dummy point at the left of the graph.

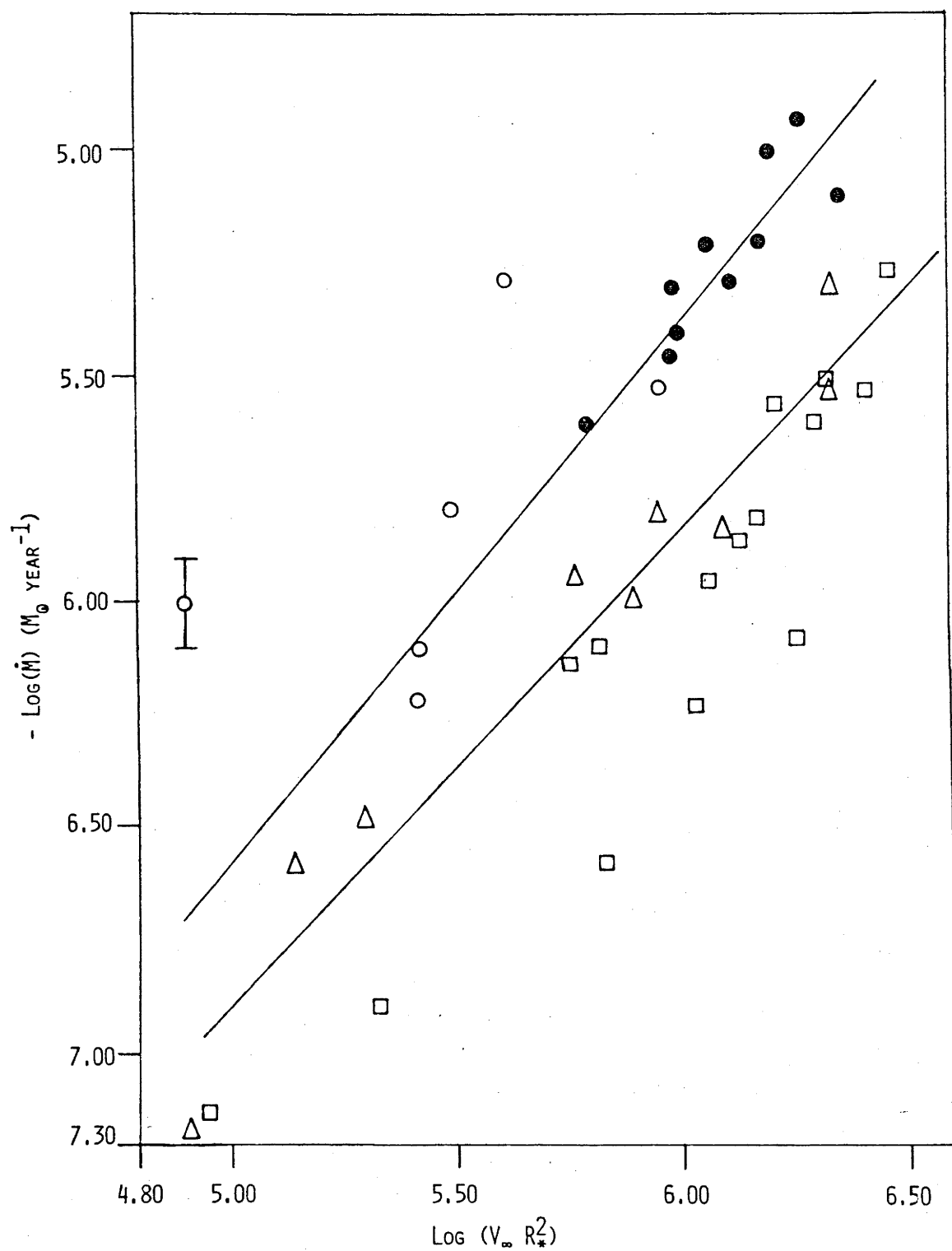


TABLE 6.2a

D.E.M.	STAR	SP TYPE	-M _V	-B.C.	-M _{bol}	T _{eff} (K)	R _* /R _⊙	M _* /M _⊙	log L _* /L _⊙	V _{esc} km/sec	V _∞ km/sec	(1-Γ)
301	1	08.5I(f)	6.2	2.8	9.0	30,400	20.5	35	5.51	701	2100	0.75
	2	05.5V	5.3	3.8	9.1	41,200	11.8	45	5.56	1069	3200	0.79
	3	09:III-V	5.0	2.9	7.9	32,100	11.2	25	5.08	861	2500	0.87
	10	08:V	4.7	3.2	8.0	34,800	9.6	27	5.09	971	2900	0.88
	6	08.5III+B?	5.2	2.9	8.1	32,150	12.2	27	5.16	852	2600	0.86
11	1	08.5III-V	5.2	3.0	8.2	33,000	11.9	27	5.18	858	2600	0.85
	2	07 V	5.1	3.5	8.6	36,800	11.5	34	5.34	967	2900	0.83
	4	09:V	4.0	3.0	7.0	33,000	6.8	21	4.70	1050	3200	0.94
196	1	06III+OB	5.6:	3.5	9.1:	37,400	14.0	41	5.54	930	2800	0.78
	8	09I	7.0	2.7	9.7	29,500	29.7	45	5.78	610	1800	0.64
	3	B1:III-V	3.8	2.5	6.3	24,850	7.7	-	4.41	-	-	-
226	1ab	08.5I+B	6.6:	2.8	9.4:	30,400	24.6	40	5.67:	654	2000	0.69
	4a	07V+B?	5.2	3.5	8.7	36,800	11.8	34	5.36	949	2800	0.82
	13ab	09.5V	4.1:	2.9	7.0:	32,000	7.4	20	4.72:	979	2900	0.93

TABLE 6.2a

[illegible]

TABLE 6.2a (cont.)

D.E.M.	STAR	SP TYPE	-M _V	-B.C.	-M _{bol}	T _{eff} (k)	R _* /R _☉	M _* /M _☉	log L _* /L _☉	V _{esc} km/sec	V _∞ km/sec	(1-Γ)
137	11	05.5If	6.4	3.5	9.9	36,800	20.5	56	5.84	836	2500	0.67
	10	WN3	4.4	-	-	-	-	-	-	-	2000	-
	1	B0.5Ib	5.7	2.5	8.2	25,400	18.5	24	5.17	643	1900	0.84
25	1	08I	6.6	3.0	9.6	33,000	22.7	45	5.74	714	2100	0.68
106	1	B8I:	7.6	0.4	8.0	11,000	97.6	23	5.10	277	800	0.85
	2	B0Ia:	7.1	2.6	9.7	27,800	33.4	45	5.78	575	1700	0.64
	3a	B3:Ib	6.3	1.3	7.6	16,300	37.4	20	4.95	424	1300	0.88
	6	B1Ib	5.7	2.3	8.0	23,000	19.7	23	5.10	617	1900	0.85
	7	B0.5III	5.1	2.6	7.7	26,800	13.4	22	4.98	744	2200	0.88
	4	09V	4.9:	3.0	7.9	33,000	10.4	26	5.06	917	2800	0.88
229	4	09I	7.2	2.7	9.9	29,500	33.3	52	5.88	603	1800	0.61
	8	B0.5Ia:	6.9:	2.6	9.5:	27,800	30.5	41	5.70:	586	1800	0.67
	6	08I	7.1	2.9	10.0	31,300	29.9	54	5.89	648	1900	0.61
	5	B1.5Ia	7.1	2.1	9.2	21,000	40.2	36	5.58	496	1500	0.72
	L53	09III:	5.5:	2.9	8.4	31,250	14.4	28	5.25:	784	2400	0.83
	1	G2Ia	9.0	0.04	9.0	5,500	619.0	-	5.50	-	-	-
	L44	B5Ib:	5.8:	0.9	6.7:	13,500	35.2	-	4.57:	-	-	-
	L46	B2Ib:	5.4	1.8	7.2:	20,200	20.5	-	4.80:	-	-	-

TABLE 6.2a (cont.)

D.E.M.	STAR	SP TYPE	$-M_V$	-B.C.	$-M_{\text{bol}}$	T_{eff} (k)	$R_*/R_{\odot}$	$M_*/M_{\odot}$	$\log L_*/L_{\odot}$	V_{esc} km/sec	V_{∞} km/sec	$(1-\bar{\Gamma})$
31	2	06.5I +	6.3	3.6	9.9	38,100	19.3	60	5.85	902	2700	0.69
	2	WN3-5A	-	-	-	-	-	-	-	-	2000	-
235	1	08.5If	6.6:	2.8	9.4:	30,400	24.6	40	5.67:	654	2000	0.69

TABLE 6.2b

D.E.M.	STAR	$-\log \dot{M}$ $M_{\odot} \text{yr}^{-1}$	$\sigma \log \dot{M}$	$+\log \dot{M}_{\infty}^{+2}$ kgm/sec	$\sigma \log \dot{M}_{\infty}$	$\log \frac{1}{2} \dot{M}_{\infty}^2$ Joules/sec	$\sigma \log \frac{1}{2} \dot{M}_{\infty}^2$	Age 10^6 years
301	1	5.68	0.29	23.17	0.31	29.48	0.35	3.5
	2	5.82	0.31	23.21	0.33	29.71	0.37	2.0
	3	6.20	0.32	22.73	0.34	29.14	0.38	3.8
	10	6.24	0.32	22.77	0.33	29.22	0.38	2.9
	6	6.12	0.33	22.81	0.34	29.21	0.39	3.9
		$\Sigma 5.25$	0.15	$\Sigma 23.67$	0.16	$\Sigma 30.11$	0.19	$\bar{x} = 3.6 \pm 0.4$
11	1	6.10	0.32	22.84	0.33	29.25	0.38	3.7
	2	5.99	0.30	23.01	0.31	29.46	0.36	2.8
	4	6.67	0.32	22.37	0.34	28.86	0.38	2.0
		$\Sigma 5.69$	0.20	$\Sigma 23.29$	0.20	$\Sigma 29.73$	0.24	$\bar{x} = 2.8 \pm 0.9$
196	1	5.78	0.30	23.21	0.32	29.64	0.36	2.7
	8	5.35	0.29	23.45	0.30	29.70	0.35	3.4
		$\Sigma 5.21$	0.23	$\Sigma 23.65$	0.23	$\Sigma 29.97$	0.25	$\bar{x} = 3.2 \pm 0.4$
226	1ab	5.49	0.28	23.33	0.30	29.61	0.35	3.5
	4a	5.96	0.29	23.05	0.31	29.48	0.36	2.8
	13ab	6.62	0.33	22.37	0.34	28.83	0.38	3.0
		$\Sigma 5.34$	0.21	$\Sigma 23.54$	0.21	$\Sigma 29.89$	0.23	$\bar{x} = 3.0 \pm 0.6$

TABLE 6.2b (cont.)

D.E.M.	STAR	$-\log \dot{M}$ $M_{\odot} \text{yr}^{-1}$	$\sigma_{\log \dot{M}}$	$+\log \dot{M}_{\infty}^{+2}$ kgm/sec	$\sigma_{\log \dot{M}_{\infty}}$	$\log \frac{1}{2} \dot{M}_{\infty}^2$ Joules/sec	$\sigma_{\log \frac{1}{2} \dot{M}_{\infty}^2}$	Age 10^6 years
137	11	5.43	0.28	23.53	0.30	29.90	0.35	2.5
	10	(4.49)	-	(24.61)	-	(30.61)	-	
	1	5.98	0.29	22.85	0.31	29.10	0.36	5.4
	-	$\Sigma 4.43$	-	$\Sigma 24.65$	-	$\Sigma 30.70$	-	
				$\Sigma 23.87$		$\Sigma 30.23$		
25	1	$\Sigma 5.46$	0.28	$\Sigma 23.41$	0.30	$\Sigma 29.73$	0.35	3.3
								$\bar{x} = 3.3$
106	1	5.69	0.32	22.77	0.33	28.68	0.38	5.7
	2	(5.33)	0.28	(23.45)	0.30	(29.67)	0.35	3.4
	3a	6.02	0.32	22.61	0.33	28.71	0.38	5.7
	6	6.04	0.29	22.77	0.31	29.00	0.35	5.7
	7	6.24	0.30	22.65	0.32	28.97	0.36	5.1
	4	6.25	0.32	22.73	0.34	29.16	0.38	3.4
		$\Sigma 5.01$	0.16	$\Sigma 23.73$	0.17	$\Sigma 29.96$	0.20	$\bar{x} = 5.6 \pm 0.3$
229	4	5.25	0.29	23.53	0.31	29.78	0.35	3.0
	8	5.41	0.32	23.37	0.34	29.65	0.38	3.4
	6	5.27	0.28	23.57	0.30	29.84	0.35	3.0
	5	5.46	0.27	23.25	0.28	29.38	0.33	3.9
	L53	5.99	0.31	22.93	0.33	29.29	0.37	4.1
		$\Sigma 4.71$	0.13	$\Sigma 24.08$	0.15	$\Sigma 30.34$	0.17	$\bar{x} = 3.4 \pm 0.5$
				$\Sigma 24.20$		$\Sigma 30.46$		

TABLE 6.2b (cont.)

D.E.M.	STAR	$-\log \dot{M}$ $M_{\odot} \text{yr}^{-1}$	$\sigma \log \dot{M}$	$+\log \dot{M V_{\infty}^{+2}}$ kqm/sec	$\sigma \log \dot{M V_{\infty}}$	$\log \frac{1}{2} \dot{M V_{\infty}^2}$ Joules/sec	$\sigma \log \frac{1}{2} \dot{M V_{\infty}^2}$	Age 10^6 years
31	2	5.45	0.29	23.53	0.31	29.94	0.36	2.4
	2	(4.49)	-	(24.61)	-	(30.61)	-	
		$\Sigma 5.75$		$\Sigma 24.64$		$\Sigma 30.69$		$\bar{x} = 2.4$
				$\Sigma 23.83$		$\Sigma 30.24$		
235	1	$\Sigma 5.49$	0.28	$\Sigma 23.33$	0.30	$\Sigma 29.61$	0.34	3.8
								$\bar{x} = 3.8$

TABLE 6.3

M.K. Spectral Type	σ_{sp}	$\sigma_{\Delta sp}$	σ_{tot}
05 - 06	1500	1800	2300
06.5 - 09.5	1500	1100	1900
B0 - B2	1500	2400	2800
B2.5 - B5	1500	1600	2200
B6	1500	750	1700

σ_{sp} = the uncertainty in T_{eff} at a given spectral type

$\sigma_{\Delta sp}$ = the uncertainty introduced by the error in spectral classification

σ_{tot} = the total uncertainty in T_{eff} in the given spectral range

6.6.2 Discussion and Conclusions

The two most important results of this chapter concern the way in which N_L and L_w change during the hydrogen core-burning lifetime of a star. Figure 6 shows that N_L and L_w are both essentially constant throughout this lifetime, for the mass range of the ionizing stars found in the L.M.C. HII regions (i.e. 25 to 70 $M_\odot$ on the zero-age main sequence). Both N_L and L_w vary slightly, toward the end of the hydrogen core-burning phase. N_L shows a tendency to increase if $M_\star$ (initial) $\geq 60 M_\odot$ and decrease if $M_\star$ (initial) $\leq 30 M_\odot$, while L_w decreases slightly if $M_\star$ (initial) $\geq 50 M_\odot$. The largest changes in N_L and L_w only amount to 0.1 in the logarithm, which is considerably less than the formal uncertainties in either of these two parameters (1σ ($\log L_w$) ≈ 0.36 ; 1σ ($\log N_L$) ≈ 0.18). Thus it is reasonable to make the approximation that the present day values of N_L and L_w are applicable throughout the hydrogen core-burning lifetime of a star.

The usefulness of this result is tempered somewhat when we consider a cluster of massive OB stars, because of the rapid evolution of these stars. As a cluster ages, more and more of the massive stars will evolve away from the main sequence and complete their evolution, with almost all of the stars with masses in excess of 30 $M_\odot$ disappearing as the cluster evolves in age from 3.0 to 6.0 million years. The cluster would not have to be much

older than 3.0 million years before the loss of massive stars would make it impossible to extrapolate back to the conditions prior to this date.

Fortunately, in our case, all except one of our clusters are between 3 and 4 million years in age. This fact coupled with the small number of massive stars found in these clusters, means that the probability that more massive stars existed prior to the present epoch, is small.

There are four clusters for which we can only give a lower limit to the values of N_L and L_W . The first of these is D.E.M. 106, which is so old (≈ 5.7 million years) that it has lost almost all of its stars with $M_* \geq 30 M_\odot$. Each of the other three clusters, on the other hand, contains a stellar remanent of a massive progenitor. The clusters D.E.M. 31 and D.E.M. 137 each have a late WN W-R star which is by far the dominant source of present-day mass-loss. If we were to apply the present day values of N_L and L_W to past epochs, we would seriously underestimate the former and overestimate the latter. The stars which formed these W-R stars spent most of their lives as massive hydrogen core-burning stars. In this state the L_W would be considerably lower ($\sim 10X$) and N_L much higher ($\sim 30X$), compared to the W-R phase. Since the W-R phase is significantly shorter than the time spent by a massive star on the main sequence (a few $\times 10^6$ years compared with a few $\times 10^5$ years), we have chosen to

represent N_L and L_W for the two W-R stars by the N_L and L_W of the most massive OB star still present in the cluster. This will mean that our value of N_L will be an upper bound to the true value and our value of L_W a lower bound. As an example, the largest factor by which we could underestimate L_W is roughly two. This can be seen if we compare the typical mass-loss rates of a WN star, $\dot{M} \approx 3 \times 10^{-5} M_\odot \text{ year}^{-1}$, (Barlow, Smith and Willis 1981), with that for $80 M_\odot$ star progenitor, $\dot{M} \approx 3 \times 10^{-6} M_\odot \text{ year}^{-1}$. If the WN star was to contribute the same L_W as the $80 M_\odot$ progenitor then it would have to be a W-R star for 2.7×10^5 years. This age is a lower bound as the typical V_∞ for early WN stars of roughly 2000 km/sec (Barlow, Smith and Willis) is lower than that for an $80 M_\odot$ star, which is closer to 3000 km/sec. Little is known about the relative fraction of time that a massive star spends in the W-R phase, other than that it is thought to be a few by 10^5 years (Maeder 1981). As a result, our approximation is probably as good as any given our present knowledge of W-R stars.

The final HII region D227 contains a G2Ia star which was treated in the same manner as the W-R stars.

CHAPTER 7

SURFACE PHOTOMETRY AND CONCLUSIONS

7.1. The Instrumentation

Hydrogen-alpha surface photometry observations of 17 L.M.C. HII regions were obtained during the 1979/80 and 1980/81 Magellanic Season (October to February). These observations were done with a S.E.C. Vidicon T.V. 2-D photometer/Reducteur Focal combination, placed at the f/18 focus of the Siding Spring Observatories' one metre telescope. A summary of the observations is given in Table 7.1. which lists the date of observation, name, R.A. and Dec and exposure time for each object.

The instrumentation is best described by sub-dividing into three distinct sections, the T.V. photometer, the optics and the filter transmission characteristics.

7.1.1 The S.E.C. Vidicon T.V. 2-D Photometer

The T.V. photometer used for our observations was a 25 mm electrostatically focussed class II Westinghouse S.E.C. Vidicon camera. For a detailed description of the specifications and performance of this instrument the reader is referred to the papers by Rodgers et.al. (1980) and Da Costa et.al. (1977)

The major attributes of this system include; its large field size (a 10' x 10' unvignetted field at f/3.4), low dark current and excellent linearity over the full digitized intensity range of 0 to 256 (each increment corresponds

TABLE 7.1

CURRENT									
DATE	NAME OF OBJECT	R.A.			D.E.C.			EXPOSURE TIME (SEC)	COMMENT
		hr	min	sec	deg	min	sec		
25/10/79	D31	04	55	20.3	-67	13	40	1080	Sum of two exposures
	D199	05	26	35.1	-68	51	21	900	
	D192	05	25	47.9	-67	31	11	600	
	N49	05	26	00	-66	05	59	180	
	N49	05	26	00	-66	05	59	180	
14/01/80	N49	05	25	55.7	-66	05	47	360	Sum of two exposures
	D121	05	16	49.5	-67	21	05	300	
	D138	05	20	29.6	-66	47	41	180	
	IC418	05	26	45.1	-12	42	53	3	
	IC418	05	26	45.1	-12	42	53	3	
	D311	05	45	21.3	-69	46	43	300	
	D281	05	40	44.9	-70	03	04	600	
	IC418	05	26	45.2	-12	42	48	3	
	N49	05	25	55.7	-66	05	47	180	
	N49	05	25	55.7	-66	05	47	180	
	IC418	05	26	45.0	-12	42	53	3	
15/01/80	IC418	05	26	45.0	-12	42	37	3	
	IC418	05	26	45.0	-12	42	37	3.5	
	D229/231	05	33	05.5	-67	43	43	300	
	D145/149	05	21	32.3	-69	41	16	240	
	D301	05	43	14.6	-67	51	00	1500	
	N49	05	25	55.4	-66	05	44	180	
	N49	05	25	55.4	-66	05	44	180	
16/01/80	IC418	05	26	43.9	-12	42	25	150	
	IC418	05	26	43.9	-12	42	25	5	
	N49	05	22	55.1	-66	06	19	180	G.F.
	IC418	05	26	44.0	-12	42	21	2	
	N49	05	25	57.9	-66	05	41	180	G.F.
	N49	05	25	57.6	-66	05	37	180	
	IC418	05	26	45.9	-12	42	31	5	G.F.
	IC418	05	26	45.9	-12	42	31	2	

Cont'd

TABLE 7.1 (cont'd)

<u>CURRENT</u>									
<u>DATE</u>	<u>NAME OF OBJECT</u>	<u>R.A.</u>			<u>D.E.C.</u>			<u>EXPOSURE TIME (SEC)</u>	<u>COMMENT</u>
		hr	min	sec	deg	min	sec		
27/11/80	N1535	04	13	18	-12	48	12	30	
	N1535	04	13	18	-12	48	12	300	G.F.
	N1535	04	13	18	-12	48	12	300	G.F.
	N49	05	25	56	-66	06	64	180	
	D137/South	05	20	36	-65	31	30	1800	
	N49	05	25	56	-66	06	54	160	
	D137/North	05	20	39	-65	28	18	2100	
	N1535	04	13	21	-12	48	00	30	
	N1535	04	13	21	-12	48	00	240	G.F.
	IC418	05	26	32	-12	43	30	2	
	D196	05	25	05	-67	36	42	1200	
	D25	04	53	36.5	-70	02	16	1800	

Notes: G.F. = Galactic Filter

to approximately 11 photoelectron equivalent) (Dopita 1977). These qualities make the T.V. photometer an ideal instrument for doing surface photometry of L.M.C. HII regions.

The target of the T.V. is read over a 15 x 15 mm area, producing a 256 x 256 data array of roughly 60 x 60 pixels. The spatial variations in the gain over this target area consist of three components. The first is a low frequency modulation which results in a 35% reduction in gain in going from the centre to the edge of the target. Superimposed on this is an area of decreased target sensitivity resulting from overexposure of a coude spectrum which cuts diagonally across the centre target (Rodgers (1980)). Thirdly there is a high spatial frequency component produced by pixel to pixel gain variations. In theory, it is possible to remove all of these variations by dividing the raw T.V. data, point by point, by the image read when the camera is exposed to a uniform light source (flat field). In practice, the inefficient magnetic shielding of the tube allows the Earth's magnetic field to both rotate and distort the images, making it impossible to fully remove pixel to pixel gain variations. The majority of the noise left, after a frame has been flat-fielded, is produced by the pre-amplifier in the T.V. (about 30 photoelectron equivalent) except at high signal levels where photon noise becomes important.

Dopita (1977) showed that it is possible to measure

absolute fluxes for planetary nebulae to within an uncertainty of 6% using the T.V. photometer. He also showed that absolute surface brightnesses determined from a properly exposed and calibrated T.V. frame (under good observing conditions) should be reliable at the 10% level.

7.1.2 The Optics

Figure 7.1 shows the whole of the optical train for the T.V. photometer/Reducteur Focal system. The two beams shown in this diagram are those originating at a point on the entrance pupil of the telescope, and those originating at a point at infinity (i.e. a point in the field of view). If we consider the rays originating from the pupil one sees that the role of the wave processor corrector lens is to render the beams telecentric (i.e. all principal axes of ray cones in the field are made parallel). The Aeroekta Imaging lens serves to re-image the pupil. At this pupil is placed the 35mm Leitz Summulux f/1.4 object imager lens, which produces an image of the sky on the 15 x 15 mm T.V. target.

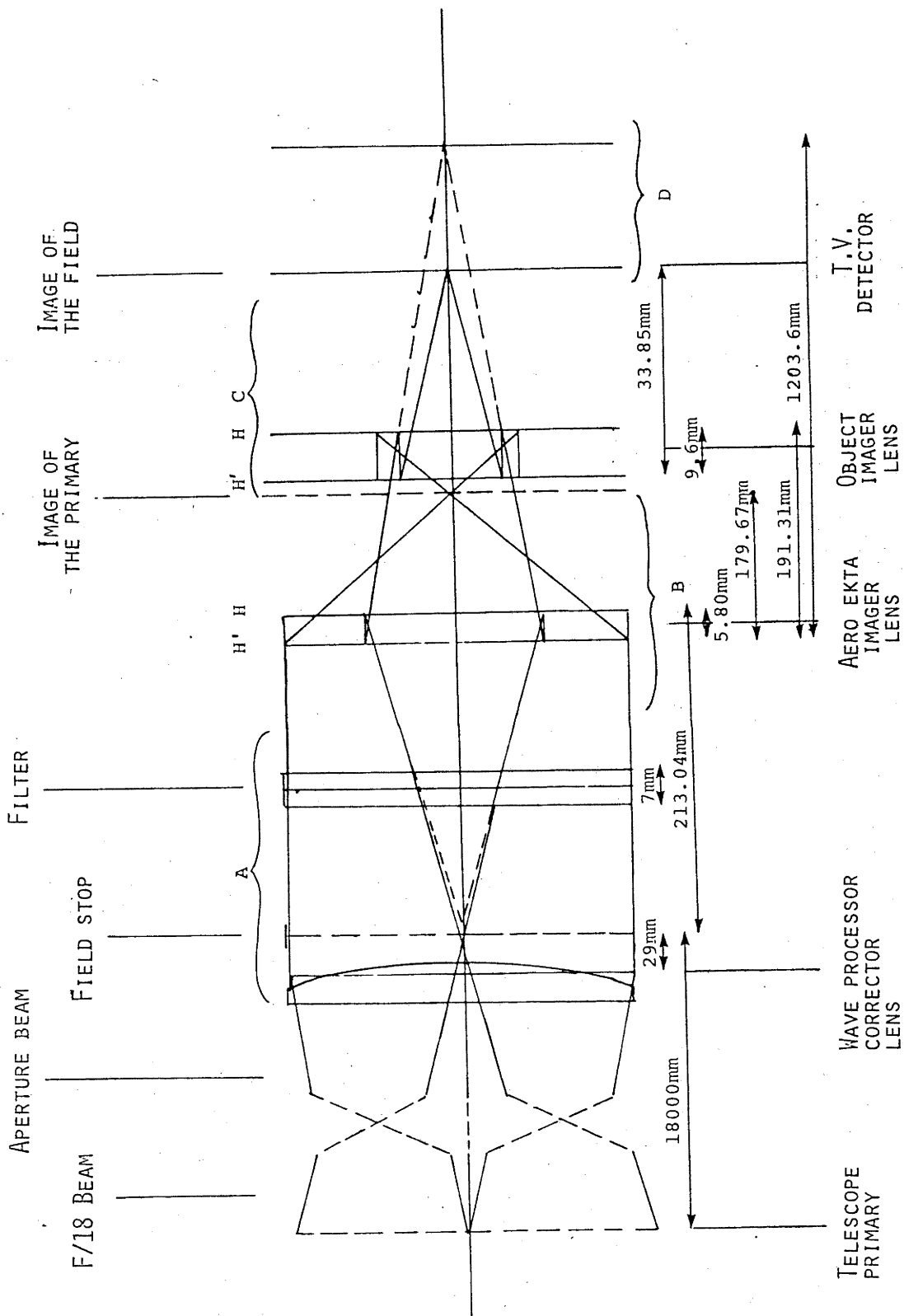
The last two lenses in the optical train, the Aeroekta Imager and the Object Imager lens, combine together to form an inverse teleconverter (or reducteur focale) which changes the incoming f/18 beam to f/3.4, the reduction factor being in the approximate ratio of the focal lengths.

The front corrector lens was removed from this optical system for our observations as strict telecentric operation was unnecessary over the field of view used and

Figure 7.1

This figure shows the optical train of the T.V. Photometer/Focal Reducer combination, used for the hydrogen-alpha surface-photometry observations. Two beams, one corresponding to a point at infinity (the F/18 beam) and the other corresponding to a point on the primary (the Aperture beam), are shown in this diagram. A description of the optical train is given in the text.

THE OPTICAL TRAIN OF THE FOCAL REDUCER/T.V. PHOTOMETER



greater throughput could be achieved without it.

7.1.3 The Hydrogen Alpha Filter

The filter used in almost all of our observations was a 76 mm diameter H-alpha triple cavity Fabry-Perot type interference filter, manufactured by Spectro-Film Inc. The transmission profile for this filter was measured in November 1982 using the Cary 14 scanning spectrophotometer located at the John Curtin School of Research at the Australian National University. Figure 7-2 shows the profile which was obtained from these measurements, using the $\frac{1}{2}$ Å resolution of this instrument. The major characteristics for this filter are;

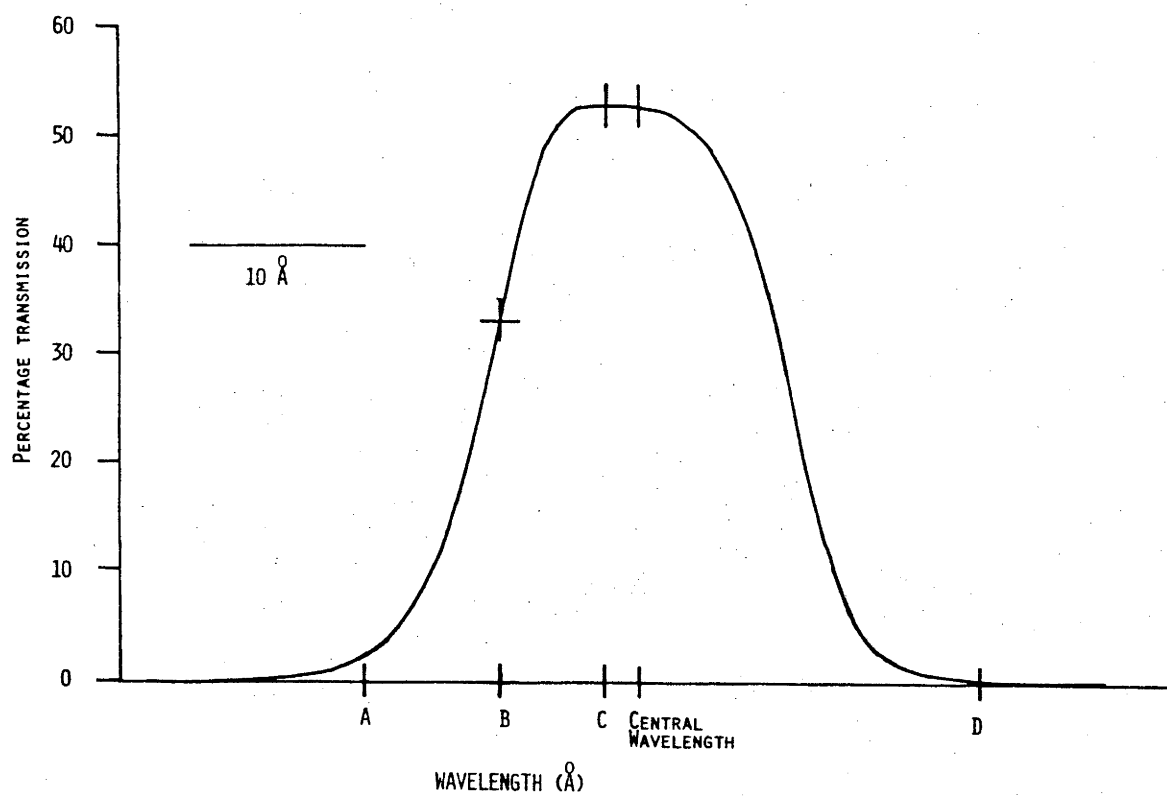
- 1) a Central wavelength $\lambda_c = 6570.6 \pm 0.6$ Å
- 2) a Full width at half maximum $\Delta\lambda = 17.6$ Å
- 3) a Peak transmittance $T = 53.5\%$ (varying by less than 5% across the filter)
- 4) a Temperature drift coefficient of $\Delta\lambda_T = 0.162$ Å/°C

The first three parameters are applicable to the filter when it is used in a collimated beam at a temperature of 20°C.

On the telescope, the filter was used in an f/18 beam at an ambient temperature of roughly 20 ± 5 °C. The effect of the f/18 beam would be to shift the peak transmission wavelength to blue by 1.4 Å and to reduce the peak transmittance slightly. A temperature increase of 5°C on the other hand, would shift the peak of the profile toward the

Figure 7.2

This figure shows the percentage transmission profile for the hydrogen-alpha narrow band filter which was used to observe the L.M.C. HII regions. The horizontal scale is represented by a $10\text{\AA}$ bar near the bottom left hand corner.



red by $0.8 \text{ \AA}$ while a decrease in temperature by 5°C would have the opposite effect. The points marked A,B,C and D on figure 7-2 mark the positions of the $[\text{NII}]$ 6548, H-Alpha (rest velocity), H-Alpha (for a velocity $V=275 \text{ km/sec}$ appropriate to the L.M.C.) and the $[\text{NII}]$ 6584 emission lines respectively (at 20°C), for a collimated beam (changes produced by the f/18 beam and temperature effects can be estimated from the above discussion).

Since, at the velocity of the LMC, $\text{H}\alpha$ emission is located in the flat top of the filter transmission profile, changes in transmittance due to temperature variations are eliminated. It can be also seen that the $[\text{NII}]$ 6548, 6584 emission lines are effectively blocked and so should not contaminate the H - alpha fluxes.

7.2. The Method of Observation

Three main steps are needed to properly calibrate a T.V. frame. The first step is to determine the zero-level of the frame so that it can be removed to avoid zero-level bias. The zero exposure frame of the T.V. photometer shows similar low frequency spatial variations to the flat field and also must be removed from each frame. Throughout our observations the zero level was determined by obtaining the mean of 25 zero exposures at both the beginning and end of each night.

The second step consists of flat-fielding the data to remove any optical variations in gain which are inherent

to the T.V. target. This was achieved by taking multiple frame exposures of either a relatively star-free field of view, or by observing the dome wall with as nearly even illumination as possible. These exposures were made at least once each observing run through a broad-band F filter, in order to maximize the S/N, while still roughly matching the target gain characteristics in a limited wavelength region near hydrogen-alpha. Flat fields taken on the sky could only be used to remove low spatial frequency variations in the T.V. frames because of the low number of total counts. Typically, three five minute frames were averaged to produce these flat fields. The flatfields taken on the dome on the other hand, had sufficient counts to theoretically remove all the inherent spatial frequencies in a T.V. frame. In practice, however, we found that magnetic field shifts meant not only that the majority of the high spatial frequencies could not be removed by flat-fielding but also that amplification of this noise occurred. Thus the dome flat fields were smoothed to remove their high spatial frequencies before they were used for flat-fielding.

Once a frame has had its zero-level subtracted and been flatfielded it must then be calibrated in terms of absolute flux. The best objects for achieving this calibration are the Galactic planetary nebulae, which have well-determined integrated hydrogen-alpha ($H-\alpha$) or hydrogen-beta ($H-\beta$) fluxes. One problem in observing these objects arises from the fact that it is impossible to observe the $H-\alpha$ emission from both the Galactic planetaries and the LMC HII regions with th

same narrow band filter. The emission from Galactic planetaries is located on the blue toe of the transmission profile of the H- α filter used to observe the Magellanic HII regions (the Magellanic filter). The method chosen to overcome this problem was to observe the standard planetary nebulae through a filter with almost identical properties to the Magellanic filter, except that the flat peak in its transmission curve was shifted to encompass the Galactic emission (the Galactic filter). Using the ratio of the percentage transmission at the peaks of these two filters, it is possible to convert the planetary fluxes to the corresponding fluxes which would have been observed in the Magellanic filter. The peak transmission of these two filters were obtained from the original transmission profiles for the two filters (Dopita 1982b). The Galactic filter had a mean transmittance at the peak of $49.5 \pm 1.0\%$ while the peak transmittance of the Magellanic filter was $54.5 \pm 1.0\%$. It was assumed that ageing effects on the peak transmittance in both filters were roughly the same as both were subject to similar environmental conditions since their manufacture.

The standard planetaries chosen for our observations were IC418 and N1535. These were selected by default as they were the only bright and well-observed planetaries which were easily observable during the Magellanic season.

The integrated H β flux of IC418 was obtained by taking the weighted mean of the published fluxes of O'Dell (1962), Capprotti and Daub (1960) and Kohoutek and Martin (1981). This gave a value of $\log F(\text{H}\beta) = -9.57 \pm 0.02$ (ergs cm⁻²/sec).

The values of $\log (F (H\alpha)/F (H\beta))$ published by Osterbrock et. al. 1963, O'Dell 1962 and Kohoutek and Martin (1981) were used to convert the $H\beta$ fluxes to $H\alpha$ fluxes.

Double weight was given to the results of Kohoutek and Martin as they had obtained their results from three or more observations of IC418. The weighted mean of these results gave $\log (F (H\alpha)/F (H\beta)) = 0.57 \pm 0.03$. This implies that the integrated $H\alpha$ flux of IC418 is $\log F (H\alpha) = -9.00 \pm 0.04$ ($\text{erg cm}^{-2}\text{sec}^{-1}$). By comparison, Kohoutek and Martin obtain a value of $\log F (H\alpha) = -9.05 \pm 0.01$ ($\text{ergs cm}^{-2}\text{sec}^{-1}$) by direct observation. Adopting the weighted mean of these two results we obtain $\log F (H\alpha) = -9.03 \pm 0.05$ ($\text{ergs cm}^{-2}\text{sec}^{-1}$).

Similarly, the integrated $H\alpha$ flux for N1535 was obtained from the published $H\beta$ figures of O'Dell (1962) and the $\log (F (H\alpha)/F (H\beta))$ of Osterbrock et. al. (1963). The value that we have adopted for the integrated $H\alpha$ flux for this nebula is $\log F (H\alpha) = -9.86 \pm 0.09$ ($\text{ergs cm}^{-2}\text{sec}^{-1}$).

As a secondary standard, the bright supernova remnant N49 (Henize, 1956) was also observed using the integrated $H\alpha$ flux published by Dopita (1977) to calibrate our T.V. frames. This nebula was observed on at least two occasions each night to monitor any changes in the system gain during the night and from night to night.

7.3. The Initial Reduction Procedure

The first step in the initial reduction procedure

was to remove the zero-level from each frame. If an object covered an area less than 50×50 pixels in size it was found that the zero-level could adequately be represented by a plane. In this case the fitted plane was subtracted from the data rather than the original noisy zero file. Low frequency spatial variations became important in the zero file when areas larger than 50×50 pixels were considered. In this case the original zero-frame was subtracted from the data.

Next, low frequency spatial variations were removed using the flat-field appropriate to the particular observing run. The flatfields used on a given data and the method by which they were obtained is given in table 7.2.

The sky level in each frame was determined by taking the average of the mean value found in anything from one to four box regions located near the object, and this was subtracted from the processed data.

The last step in the initial reduction procedure was to remove the stars from the T.V. frame. It was only possible to do this for the bright stars which were not heavily immersed in nebulosity, as no continuum observations were made of the HII regions. Contamination should not be a problem in all except the smallest nebula, where the stars might contribute a few percent to the total integrated flux.

From this reduced data, it is possible either to

TABLE 7.2

<u>DATE</u>	<u>NUMBER OF EXPOSURES</u>	<u>EXPOSURE TIME/FRAME</u> (SECONDS)	<u>METHOD</u>
25.10.79	3	300	A
14.01.80	5	300	A
15.01.80	5	300	A
16.01.80	3	300	A
27.11.80	25	0.8	B
27.11.80	50	0.8	B

A Flat field taken on the sky

B Flat field taken on the side of the telescope dome

obtain the absolute integrated $H\alpha$ fluxes for the nebulae or to calculate the mean density in the nebulae using absolute surface brightness maps for the HII regions..

7.4. The Integrated $H\alpha$ Flux

The total number of digital units (D.U., approximately 11 photoelectron equivalents) above sky, and the total number of points with intensity greater than 3σ above the sky (where the σ is calculated from the noise in the sky) were determined for each nebula using the output frames from the initial reduction. Each of these values is listed in Table 7-3 (columns 4 and 5 respectively) together with the date of observation (column 1), the D.E.M. number of the HII region (Davies, Elliott and Meaburn 1976) (column 2) and the exposure time in seconds (column 3).

The integrated D.U. for a given HII region must be corrected for atmospheric extinction. This was done using the extinction formula of Hayes and Latham (1975) where the altitude of the Siding Spring Observatory was taken as 1.10 kilometres, the contribution due to ozone was ignored, and the Aerosol constant A_0 (see Hayes and Latham 1975) was 8.7×10^{-2} (Pickles 1982). All observations were corrected to give the total number of detector units which would be observed above the Earth's atmosphere. The atmospheric extinction for each object is listed in column 6 of table 7-3.

Once the atmospheric extinction was removed, the integrated fluxes were corrected for interstellar reddening.

TABLE 7.3

DATE	OBJECT D.E.M. NO.	EXP. TIME (Seconds)	(D.U.) /SEC	TOTAL POINTS	Δm	$F(H\alpha)$ $\times 10^{-14}$ $Jm^{-2} sec^{-1}$	$F_O(H\alpha)$ $\times 10^{-14}$ $Jm^{-2} sec^{-1}$	SIZE 10^{-8} Steradians	SURFACE BRIGHTNESS $10^{-7} Jm^{-2} sec^{-1}$ str ⁻¹
15.01.80	D229	300	842.23	4029	0.130	43.328	54.602	122.05	4.473
	D231	300	140.30	890	0.130	7.218	9.096	26.96	3.374
	D145	240	34.86	208	0.132	1.793	2.490	6.30	3.952
	D149	240	101.34	315	0.132	5.213	7.238	9.54	7.587
	D301	1500	407.26	8054	0.148	20.951	31.816	243.98	1.304
27.11.80	D1375	1800	93.684	1829	0.130	4.885	} 20.247	223.54	0.906
	D137N	2100	190.172	5550	0.126	9.916			
	D196	1200	457.170	8055	0.138	23.837		244.01	1.317
	D25	1800	106.903	1471	0.136	5.574		44.56	1.857
14.1.80	D121	300	62.038	300	0.131	3.284	4.560	9.088	5.018
	D138	180	79.900	217	0.132	4.229	5.872	6.574	8.932
	D311	300	70.568	317	0.134	3.735	5.186	9.603	5.400
	D281	600	76.401	891	0.134	4.044	5.615	26.991	2.080
25.10.79	D31	1080	852.64	11839	0.130	34.033	42.253	358.64	1.178
	D199	900	2095.24	17466	0.132	83.631	116.125	529.10	2.195
	D192	600	2782.15	22698	0.129	111.048	154.195	687.59	2.242

The method chosen to determine the interstellar reddening was based on observed and intrinsic colours of stars found within the L.M.C. HII regions. Only those stars with well determined spectral types and accurate UBV photometry were used to determine the E_{B-V} of the cluster. These stars are listed in Table 7-4. The UBV photometry for the stars with Sanduleak numbers (Sanduleak 1971) was taken from Rousseau et. al. (1978) while the photometry for the other stars in this table comes from the thesis by Lucke (1972).

International Ultra-violet Explorer spectra were taken of the two stars DEM 301/1 and D25/1. These spectra were used to determine the value of E_{B-V} by nulling the $2200\text{\AA}$ features found in the long wavelength spectra. A value of $E_{B-V} = 0.05$ was adopted for the Galactic foreground reddening (Wilson 1983) and the mean Ultraviolet Magellanic reddening curve of Nandy and Morgan (1982) was used to null the $2200\text{\AA}$ feature. Column 6 of table 7-4 lists the spectrographically determined A_v values while column 7 lists the two A_v values determined from the I.U.E. spectra. The final column of table 7-4 lists the adopted A_v for the cluster. The values for D.E.M. 301 and 25 were obtained by double weighting the I.U.E. results. Half-weight was given to the spectrographic extinctions followed by a colon. The halfweighting is due to the uncertainty in the M.K. Spectral type of the star concerned. Where a spectrographic A_v differs significantly from the values obtained from other stars in a cluster, preference was given to the A_v based on the photoelectric photometry of Rousseau et. al. (1978)

TABLE 7.4

NEBULA	SANDULEAK N ^o	STAR N ^o	(B-V)	(B-V) _o	Sp Type	Av _{sp}	Av I.U.E.	Av
301		1	-0.17	-0.26	08.5I	0.29	0.64	
		2	-0.19	-0.36	05.5V	0.54	-	
		3	-0.19	-0.30	09:III-V	0.35:	-	
		6	+0.06	-0.29	08.5III+B?	1.12?	-	0.56
11		1	-0.18	-0.31	08.5III-V	0.42	-	
		2	-0.19	-0.33	07V	0.45	-	0.43
196		1	-0.20	-0.32	06III+OB	0.38	-	
		8	-0.13	-0.26	09I	0.42	-	0.40
226		4a	-0.13	-0.33	07V+OB?	0.64	-	0.64:
25		1	-0.12	-0.27	08I	0.48	0.56	0.53
137		11	-0.18	-0.31	05.5If	0.42	-	
		1	0.06	-0.24	B0Ia	0.96?	-	0.42
31		2	-0.20	-0.29	06.5I ₊	0.29	-	
		1	0.09	0.00	B9Iae	0.29	-	0.29
106		6	-0.01	-0.20	B1Ib	0.61		
		1	+0.10	-0.05	B8:Ia	0.35:		
		3a	-0.15	-0.14	B3:Ib	0.00?		0.52
229		4	-0.18	-0.26	09I	0.26		
		6	-0.16	-0.27	08I	0.35		
		5	-0.16	-0.18	B1.5Ia	0.06?		0.31

A number of HII regions with integrated H α fluxes did not have their central stars spectrally classified. The A_V for these nebulae was taken as the mean of the interstellar extinction of the nebulae listed in table 7-4 i.e.

$$\langle A_V \rangle = 0.44 \pm 0.11.$$

Table 7.5 lists the details of the observations made of the standard planetary nebulae and N49. Column 1 lists the date of observation, column 2 the object; column 3 the integrated number of detector units (D.U.) above sky; column 4 the magnitude correction for atmospheric extinction (already applied); column 5 the average number of detector units for an object on a given night; column 6 the comments and column 7 the exposure time in seconds. The variable R listed in the comments column of table 7-5 is the ratio of the average integrated flux for IC418 to that of N49. The emission from IC418 is in the toe of the transmission profile of the Magellanic filter while the emission from N49 is at the peak. Thus changes in the ratio R should indicate shifts in the transmission profile due to temperature shifts and changes in the system stability. The values of R on the 15th and 16th of January 1980 agree remarkably well indicating that the filter/photometer system forms a very stable unit. The 14th of January was a marginally photometric night and this is reflected in the lower value of R. If the observations from the early part of this night are used, when conditions were photometric, then the value of R approaches the value of the other two nights.

TABLE 7.5

DATE	OBJECT	(D.U.) / SEC	Δ_m	$\langle (D.U.)_O / SEC \rangle$	Comment	EXPOSURE TIME	Seconds
25.10.79	N49	148.995	0.127	} 147.30 $\pm$ 2.40		180	
	N49	145.603	0.128			180	
14.01.80	IC418	574.45	0.114	} 567.30 $\pm$ 16.67		3	
	"	560.93	0.113			"	
	"	547.65	0.120			"	
	"	586.15	0.122			"	
	N49	117.75	0.128	} 111.08 $\pm$ 6.26		180	
	"	105.34	0.166		R=0.196 $\pm$ 0.012	"	
	"	110.15	0.177		(R=0.207)	"	
15.01.80	IC418	531.45	0.112	} 525.19 $\pm$ 8.85		3	
	"	518.93	0.111			3.5	
	N49	113.29	0.154	} 114.28 $\pm$ 1.40		180	
	"	115.27	0.157		R=0.218 $\pm$ 0.04	180	
16.01.80	IC418	506.40	0.109	} 501.75 $\pm$ 6.58		5	
	"	497.09	0.120		Magellanic Filter	5	
	"	1728.05	0.116	} 1728.5	Galactic Filter	2	
	N49	105.24	0.126			180	
	"	112.96	0.133	} 108.36 $\pm$ 4.07		180	
	"	106.89	0.143		R=0.216 $\pm$ 0.09	180	
27.11.80	N1535	23.50	0.120	} 24.14		300	
	"	24.78	0.118			300	
	"	32.66	0.140	} 113.76 $\pm$ 6.20	Magellanic Filter	240	
	N49	109.38	0.135			180	
	"	118.15	0.129			160	
	N1535	221.30	0.129		Galactic Filter	30	

Table 7.5 shows that the planetaries IC418 and N1535 were each observed once through the Galactic filter and then subsequently with the Magellanic filter. Using the adopted fluxes for these two nebula and the ratio between the peak transmittance of the two filters, we find that the observations of IC418 indicate that the flux for N49 is

$$F(H\alpha)_{N49} = 5.314 \times 10^{-14} \text{ J.m}^{-2}.\text{sec}^{-1}$$

while the observations of N1535 give a value of

$$F(H\alpha)_{N49} = 6.444 \times 10^{-14} \text{ J.m}^{-2}.\text{sec}^{-1}$$

The average of these two results was adopted as the final absolute flux for N49

$$F(H\alpha)_{N49} = 5.88 \pm 0.80 \times 10^{-14} \text{ J.m}^{-2}.\text{sec}^{-1}$$

This flux was used to convert the $(D.U.)_0/\text{sec}$ in Table 7.3 into absolute fluxes, by taking the ratio of the number of $(D.U.)_0/\text{sec}$ in the object to that in N49 (the value for N49 used was the average for the night). These fluxes are shown in Column 7 of Table 7.3 together with dereddened fluxes in Column 8.

The absolute (spatial) scale of a T.V. photometer pixel was obtained by astrometry of the data frames. These measurements indicated that the size of one pixel was 3.59 ± 0.08 arc seconds, subtending an area of $3.0293 \pm 0.0015 \times 10^{-10}$ steradians. This scale was used to convert the total number of pixels with intensity

greater than 3σ above sky into a solid angle listed in Column 9 of Table 7.3. The total solid angle of the emission region and the absolute integrated flux were then used to determine the mean surface brightness of the respective HII region (listed in Column 9).

7.5 The Surface Brightness of the L.M.C. HII Regions

The calibration procedure for converting an output frame from the initial reduction procedure into one of absolute surface brightness is based on the determination of the total energy which must be collected by the telescope to register one digital unit (D.U.). This value can be calculated each night using the average integrated D.U. for N49 and the adopted absolute flux for this nebula. Table 7.6 lists the surface brightness equivalent to 1 D.U./pixel/second on each of the observing nights.

TABLE 7.6

Date	Conversion Factor
	$10^{-6} \text{ J.m}^{-2} \text{ .sec}^{-1} \text{ str}^{-1}$
25.10.79	1.318
14.01.80	1.648
15.01.80	1.698
16.01.80	1.791
27.11.80	1.706

The average conversion factor for the four nights in 1980 is $1.711 \pm 0.59 \times 10^{-6} \text{ J.m}^{-2}.\text{sec}^{-1}\text{str}^{-1}$. This differs considerably from the conversion factor for the one night in 1979 i.e. 1.318×10^{-6} . The 50% change in gain between the two years is somewhat of a mystery but must be due to a change in the instrumental configuration or T.V. gain.

Once converted to absolute surface brightness, the frames must be corrected for atmospheric and interstellar extinction. Figures 7.9a to 7.9m show the resulting surface brightness contour plots for the L.M.C. HII regions. The scale and approximate orientation are shown in each figure together with a circle or arc which indicates the adopted dimensions of the HII region. If the HII region had a shell structure then a circle or arc was fitted to the peak of the shell. In all other cases the circle was fitted by eye to the $5.0 \times 10^{-8} \text{ J.m}^{-2}.\text{sec}^{-1}\text{str}^{-1}$ surface brightness contour.

It is possible to determine the average density in L.M.C. HII regions using the surface brightness plots by making the assumption that the nebula temperature and density is constant along the line of sight and by adopting a simple nebula geometry. The nebula surface brightness (B) at any given point is given by

$$B = \frac{2.87}{4\pi} h\nu_{\beta} \alpha_{\text{eff}} \langle n_e^2 \rangle \Delta l$$

where Δl is the path length through the nebula, ν_β is the frequency of the H β emission line and α_{eff} is the effective recombination coefficient for H β ($\alpha_{\text{eff}} = 3.03 \times 10^{-14}$ at $T_e = 10^4 \text{ K}$, Osterbrock(1974)). The factor 2.87 (Osterbrock 1974) allows for the fact that we are dealing with the H- α rather than the H- β line. It follows from equation 7.1 that

$$\frac{\langle n_e \rangle^2}{F} = \langle n_e^2 \rangle = 1.147 \times 10^7 \frac{B}{\Delta l} \quad 7-2$$

where Δl is in parsecs, B in $\text{J.m}^{-2}.\text{sec}^{-1}\text{str}^{-1}$ and F is the filling factor in the HII region.

Column 6 of Table 7.7 lists the parameter $\langle n_e \rangle / F^{1/2}$ for the HII regions for which we have T.V. surface photometry. Where a value of $\Delta R/R$ is shown in column 3 of this table, the nebula geometry was assumed to be that of a constant density thin shell with an outer radius R shown in column 2. In all other cases the geometry was assumed to be that of a constant density sphere of radius R . The HII region radii and shell thicknesses were measured from average surface brightness profiles for the nebula. The shell thickness was taken as being half the width across the base of the surface brightness enhancement corresponding to the shell.

As a double check to densities obtained from the surface brightness plots columns 4 and 5 of Table 7.7 show the densities obtained from the integrated H- α flux of the HII regions. The densities in these two

TABLE 7.7

NEBULA	RADIUS PARSECS	$\Delta R/R$	A. $\frac{\langle n_e \rangle}{F^{1/2}}$ 10 ⁶ particles m ⁻³	B. $\frac{\langle n_e \rangle}{F^{1/2}}$ 10 ⁶ particles m ⁻³	C. $\frac{\langle n_e \rangle}{F^{1/2}}$ 10 ⁶ particles m ⁻³
301	53.0	0.13	3.8	7.7	6.0
311	11.0	-	20.5	15.6	30.6
281	15.5	-	10.7	9.7	24.6
121	9.0	-	21.9	19.8	37.0
138	7.5	-	32.0	29.5	49.7
25	43.5	0.07	5.0	5.5	3.3
137	88.5	0.17	2.4	2.7	2.0
145	9.5	-	18.0	13.5	26.5
149	9.0	-	26.9	25.0	46.3
31	71.5	0.17	3.1	5.3	4.2
231	19.5	-	16.2	8.8	23.5
229	66.1	-	6.2	-	-
196	47.5	0.15	4.0	9.1	7.7
192	73.0	0.19	4.2	6.6	8.5

columns are obtained from the formula

$$\langle n_e^2 \rangle = \frac{n \pi B}{2.87 \alpha_{\text{eff}}} h \nu_{\beta} R \quad 7-3$$

where $n = 3$ for a constant density sphere and $n = 2$ for a shell (provided terms of order $(\Delta R/R)^2$ are ignored).

The densities in column 4 are obtained from surface brightnesses which use the number of points with intensity greater than 3σ above sky as the area of the nebula. The densities in column 5 on the other hand use the dimensions listed in column 1 and 2 to determine the area of the nebula used to calculate the surface brightness.

7.6 The Conclusions

7.6.1 Are the L.M.C. HII Regions Ionization or Density Bounded?

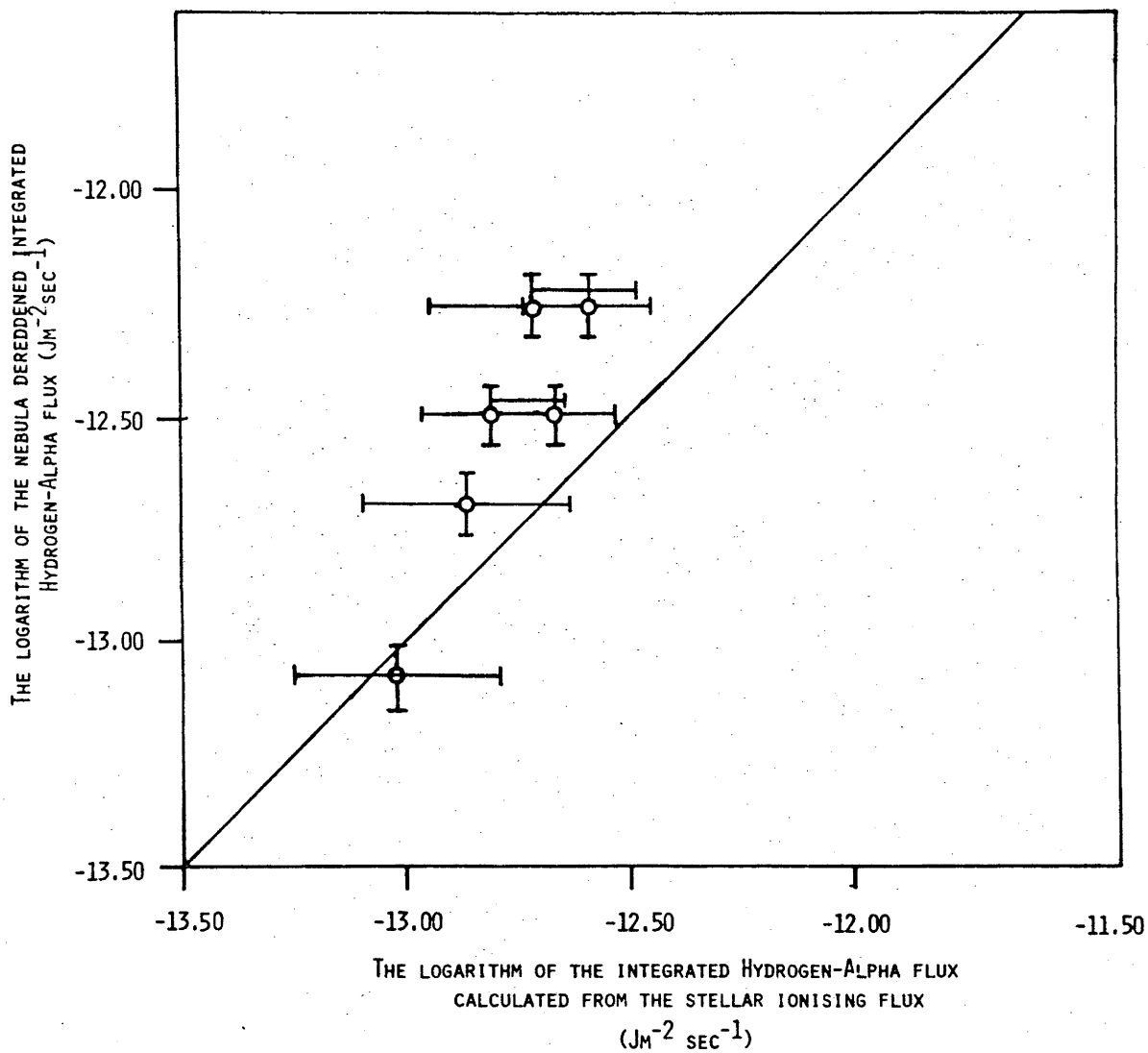
By comparing the number of ionizing photons absorbed by the nebula per unit time (N_L) to that emitted by the stars (S_L), it is possible to determine whether the L.M.C. HII regions are ionization or density bounded. The values of S_L for the OB cluster determined in Chapter 4 can be used to calculate the predicted integrated hydrogen alpha flux for the surrounding nebula. This can then be compared directly to the observed dereddened integrated hydrogen-alpha flux obtained from the T.V. surface photometry.

Figure 7.3 shows these two quantities plotted against each other. This plot indicates that all the HII regions are ionization bounded with all the nebulae (except D.E.M. 25) appearing to absorb more ionizing photons than those provided by the stars. At face value this would imply that these nebula have some form of non-thermal emission but it could also be due to an underestimation of the Lyman continuum flux in the OB stars. The Lyman continuum fluxes for these stars are derived from an empirical calibration between the M.K. spectral type and the ionizing flux, based on Galactic stars (see Chapter 4). Any underestimation of the Lyman flux in the L.M.C. stars must be due to some characteristic difference between these stars and the corresponding stars in the Galaxy. One possible reason for the difference could be due to the fact that the majority of the stars in the Galactic calibration are O dwarf stars while those in the L.M.C. are O giants and supergiants (Georgelin et. al., 1975, Morton, 1969).

This explanation is discounted by the fact that Georgelin et. al. had a small number of supergiants in their sample which did not appear to have significantly different values of N_L / F_V . Another possible reason could be that the helium abundance is lower in the L.M.C. than in the Galaxy. A lower helium abundance would lead to reduced blanketing in the Lyman continuum which

Figure 7.3

This figure shows the rate at which the L.M.C. HII regions absorb ionizing photons (derived from the Zanstra Method) against the rate at which their central OB clusters (or star) emits these photons. A solid line marks the point at which the HII region would be ionization bounded. All points below and to the left of this line correspond to nebulae which are density bounded. Points above and to the right correspond to nebulae which have some form of emission due to collisional excitation.



could explain the under-estimation of the Lyman continuum flux.

In order to confirm the conclusion that the HII regions plotted in Figure 7.3 are ionization bounded, we need to consider each nebula separately in regard to its spatial extent about the ionizing stars. If the HII region does not completely surround the stars then there will be a leakage of photons which will leave the nebula without contributing to the nebula emission. Each HII region is discussed below in order of its position in Figure 7.3, starting from the left most point and moving to the right.

D.E.M. 25: The sharply defined section of the ionized shell in this nebula subtends an angle of roughly 270^0 at star number 1, the dominant ionizing source. The remaining segment of the shell is faint and ill-defined, though leakages through this gap should only be a small fraction of the total ionizing luminosity.

D.E.M. 137: Star number 11, the dominant ionizing source, is found slightly off-centre of this large half shell. As a result only one half of the ionizing flux emitted by this star will be intercepted by the nebula. Accordingly only half the value of S_L was attributed to the stars within this nebula when plotting it

in Figure 7.3. There is a bright compact nebula located on the eastern rim of the HII shell (see Figure 7.9e). The flux contribution from this nebula was removed from the integrated H- α flux since it was reasonable to assume that the small cluster embedded within this HII region is providing the flux to ionize it.

D.E.M. 196: This nebula has two equally dominant ionizing stars which are well separated. The first, star number 1, is located close to the bright north-west rim (see Figure 7.9b), while the second, star number 8, is located close to the south-west corner of a bright knot situated on the eastern edge of the nebula. Star number 1 has almost all of its ionizing flux absorbed by the nebula while star number 8 has less than half absorbed by the nebula. If this were not allowed for it would lead to an over-estimation of the total ionizing flux from the stars in this nebula. Compensating for this effect, the point corresponding to this nebula in Figure 7.3 will move to the left, ensuring that this HII region is definitely ionization bounded.

D.E.M. 301: The major ionizing stars, stars number 1 and 2, are located very near to the centre of an almost complete shell of ionized gas. The shell is only ill-defined over a 90° arc centred about the E.S.E. direction (see Figure 7.9a). Thus only a small fraction of the total ionizing flux emitted by the OB stars should escape.

D.E.M. 31: This nebula consist of a large sharp half shell with a diffuse low surface brightness HII region located on its southern rim. The main ionizing star, star number 2, is located close to the southern side of the HII shell. The extreme off-centre location would make the nebula density bounded on the side nearest the star and ionization bounded on the far side. This leakage of photons is the possible source of ionizing photons for the diffuse southern HII region. A small fraction of the ionizing photons from this star will also leak out through a gap in the shell in the N-E quadrant, though this gap only subtends a small angle at star number 1. These leakages of ionization photons would only shift the point in Figure 7.3 corresponding to this nebula even further to the left of ionization bounded line. Thus it would appear

that this nebula (including the southern diffuse HII region) is ionization bounded.

D.E.M. 229: This nebula is very large and so only the brightest portions of it could be included in the field of view of the T.V. photometer. Extensive low surface brightness ionized gas extends to the north and west of the nebula shown in Figure 7.9g. This nebula also contains a large number of OB stars of which only the most dominant were observed (those nearest the bright shell). If allowance is made for these two effects the point corresponding to this nebula in Figure 7.3 would move upward and to the right and may move into the density bound region. This uncertainty coupled with the complex structure of this nebula makes it difficult to say if this nebula is truly ionization bounded.

Another means by which we can confirm that the L.M.C. HII regions are, in general, ionization bounded is provided by Figure 7.4. This figure shows the observed dereddened integrated H- α flux for the nebulae plotted against their diameters, while Figure 7.5 shows the integrated H- α fluxes, calculated from the stellar ionizing flux, plotted against the same parameter. The

Figure 7.4

This figure shows logarithm of the L.M.C. HII region dereddened, integrated absolute hydrogen - alpha flux ($\text{Jm}^{-2}\text{sec}^{-1}$), plotted against the nebula diameter in arc minutes. The curve, corresponding to a linear least-squares fit to the data points, is drawn in this figure. The two nebulae D.E.M. 25 and 137 were not included in this fit (these two points correspond to the two points > 6 arc minutes in diameter which are significantly below the other points).

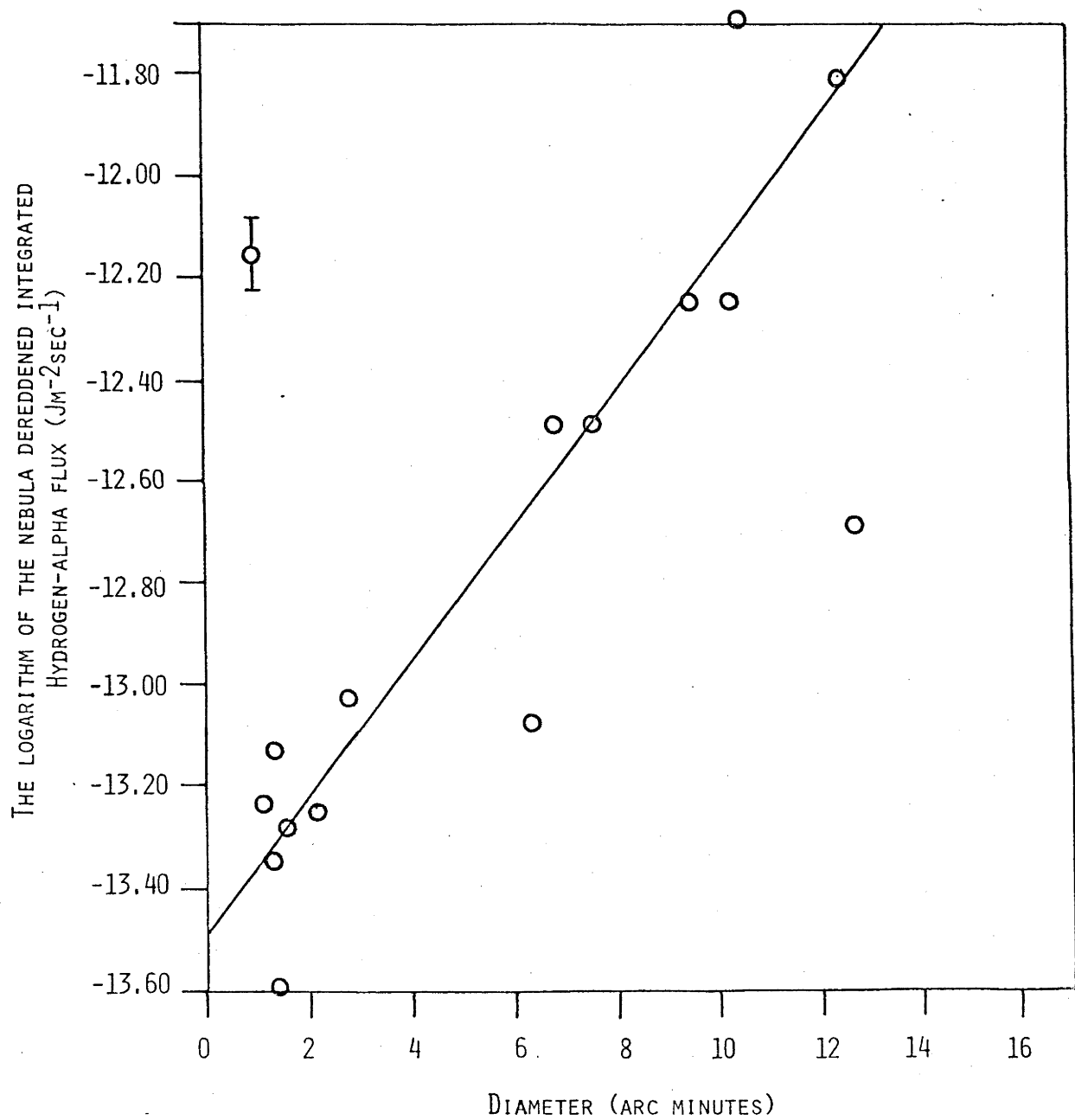
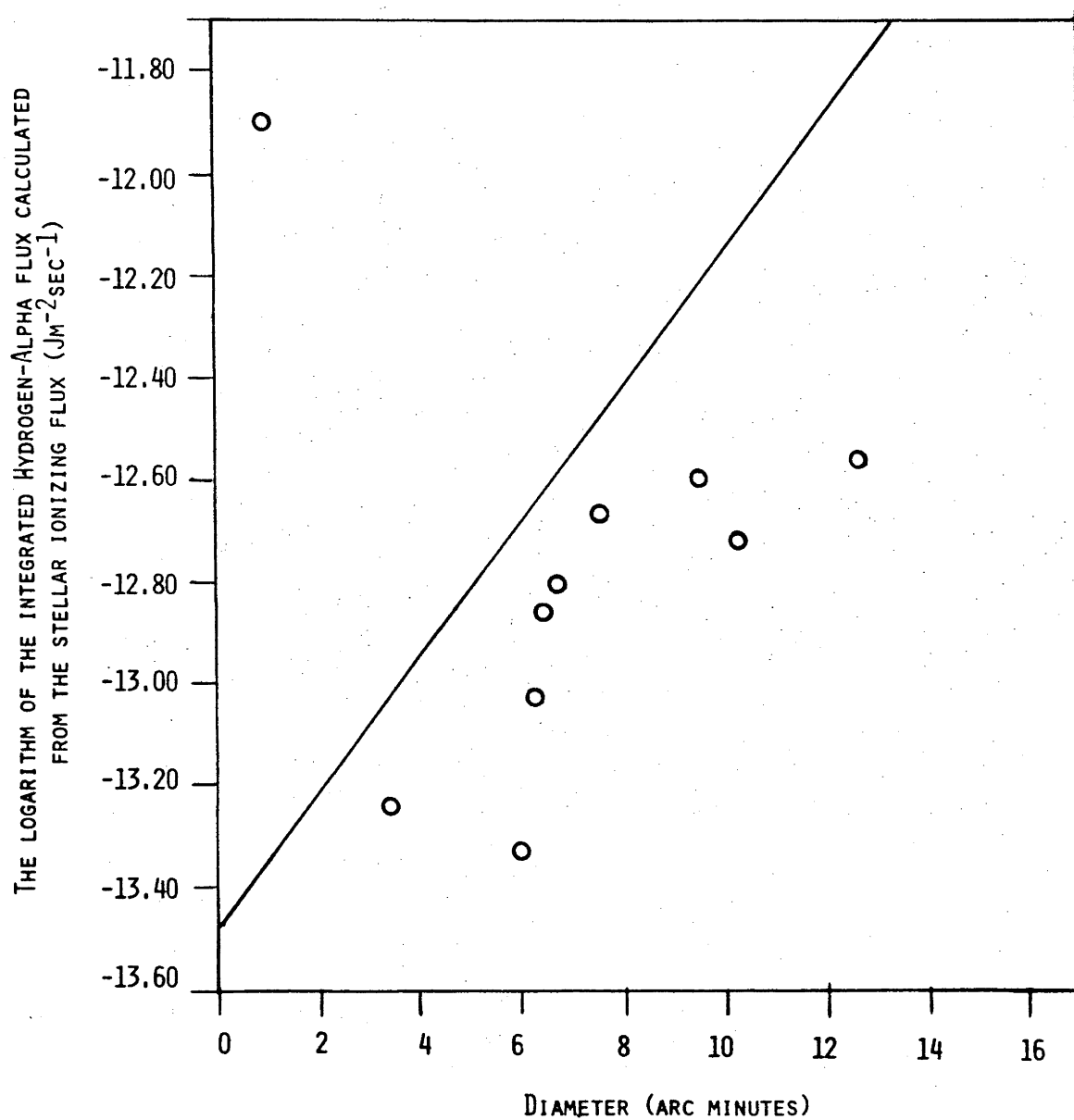


Figure 7.5

This figure shows the logarithm of the integrated hydrogen-alpha flux, calculated from the Lyman continuum photon luminosity of the central OB stars of the L.M.C. HII regions, plotted against the HII region diameter. The curve of best fit from figure 7.4 is also shown in this plot.



first of these two figures indicates that for all except two of the nebulae (D.E.M. 25 and 137) there is a fairly tight linear relationship between the integrated H- α flux and the HII region diameter. This relationship is drawn in Figure 7.5 to compare it with the integrated H- α fluxes predicted by the OB stars. It is obvious from this figure that the H- α fluxes predicted by the ionizing stars are significantly less than the H- α fluxes obtained directly from the T.V. photometry. This result strongly supports the conclusion that the HII regions in the L.M.C. are indeed generally ionization bounded.

7.6.2 The Hot-Cushion and Direct Impingement Stellar Wind Models

As already mentioned in Chapter 2, one way in which we can distinguish between the hot-cushion (H.C.) and the direct impingement stellar wind models (D.I.) is to compare the observed densities within the L.M.C. HII shells with those predicted by these two models.

The 10 HII regions for which mass-loss rates have been determined are shown in Table 7.8. Column 1 lists the D.E.M. number of the HII region (Davies, Elliott and Meaburn 1977); column 2 the nebular radius; column 3 the mechanical luminosity of the wind ($\frac{1}{2} \dot{M} V_{\infty}^2$) in units of $1.0 \times 10^{29} \text{ J.sec}^{-1}$, column 4 the momentum carried away by the stellar wind ($\dot{M} V_{\infty}$).

TABLE 7.8

<u>NEBULA</u>	<u>RDC</u>	<u>L₂₉</u>	<u>M₂₃</u>	<u>ct₆</u>	<u>N_{HC}</u>	<u>n_{DI}</u>	<u>n_{OB}</u>
301	53.0	12.88	4.68	3.6	8.7	0.10	5.8 ± 2.0
25	43.5	5.37	2.57	3.3	6.0	0.08	4.6 ± 1.2
137	88.5	17.15	7.41	-	-	0.06	2.4 ± 0.4
229	66.1	28.84	15.85	3.4	9.5	0.22	6.2
196	47.5	9.33	4.47	3.2	7.8	0.12	6.0 ± 2.6
106	53.0	9.12	5.37	5.6	9.6	0.12	-
11	24.4	5.35	1.95	2.8	28.9	0.20	-
226	31.0	7.76	3.47	3.0	21.9	0.22	-
235	39.0	4.07	2.14	3.8	7.3	0.08	-
31	71.5	26.13	10.17	2.4	4.8	0.12	4.2 ± 1.1

in units of 1.0×10^{23} Newtons; column 5 the age of the stellar association in units of 10^6 years; column 6 the HII region particle density predicted by the H-C model; column 7 the HII region particle density predicted by the D-I model; and column 8 the observed mean particle density ($\langle n_e \rangle / F^{1/2}$) for the HII region.

(F is the filling factor for the ionized gas).

The mean particle density was determined by taking the average of the three different densities listed in Table 7 - 7.

The densities listed in Table 7 - 8 strongly support the H-C model, which predicts values that are only slightly higher than those observed. The densities predicted by the D.I. model fall far short of those observed, indicating that this model cannot produce the L.M.C. HII shells.

Using the classical formula for the expansion of a H-C stellar wind bubble in a constant density medium (Weaver, McCray and Castor 1977)

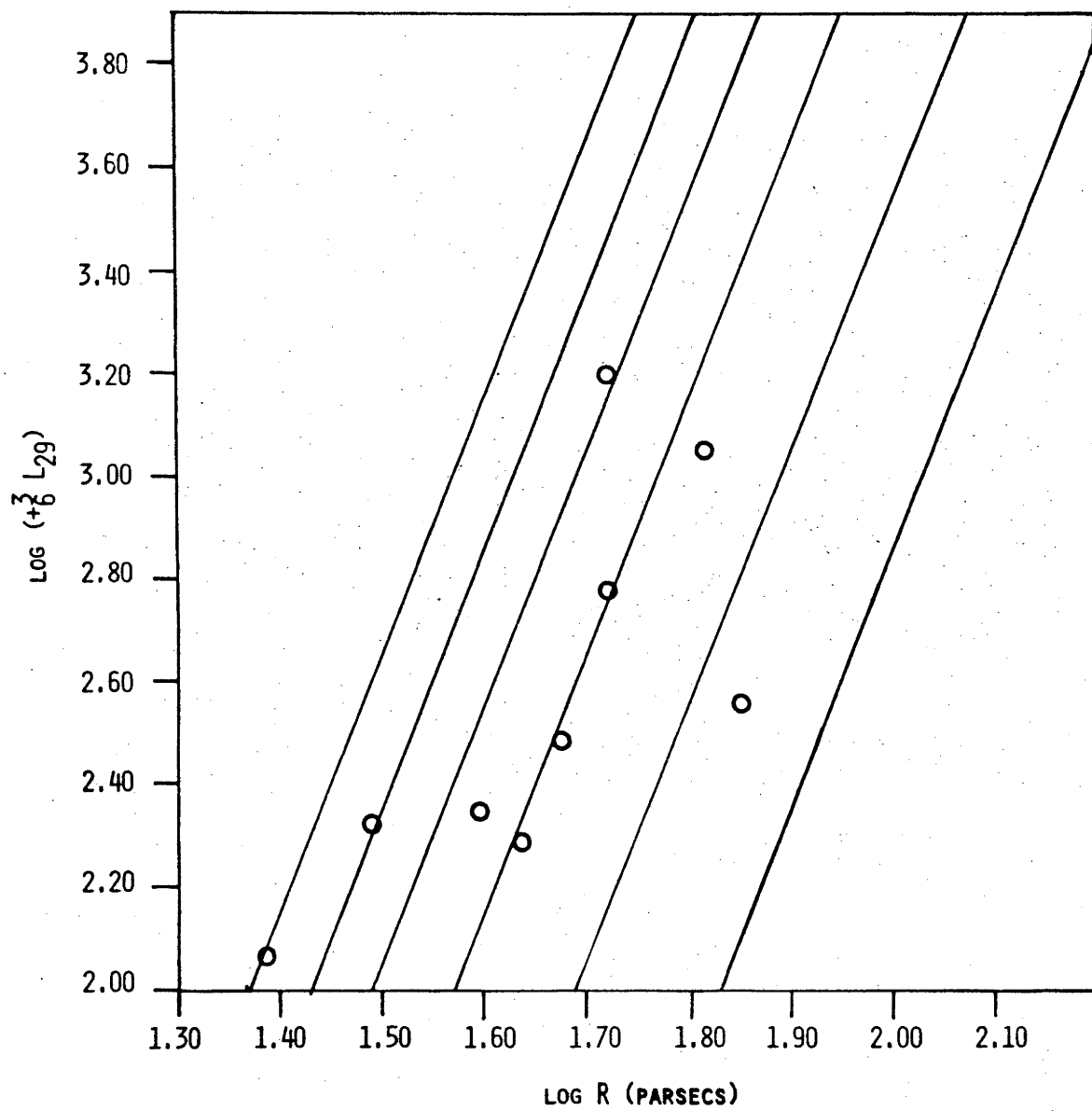
i.e.

$$R = 427.9 L_{29}^{1/5} n_o^{-1/5} t_6^{3/5} \quad (\text{pc}) \quad 7-4$$

we can calculate the density of the I.S.M. into which the HII region is expanding. Figure 7-6 shows a plot of $\log(t_6^3 L_{29})$ versus $\log(R)$ for the 10 HII regions listed in Table 7 - 8. Lines of constant I.S.M. density are drawn on this plot showing that almost all of the HII regions are currently

Figure 7.6

This figure is a plot of $\text{Log}_{10} (t_6 L_w)$ against the logarithm of the radius of the L.M.C. HII region. t_6 is the age of the nebula in units of 10^6 years, L_w is in units of 10^{29} Joules sec^{-1} and the radii are measured in parsecs. Lines of constant density are drawn on this graph corresponding to 1.0×10^6 , 5.0×10^6 , 2.0×10^7 , 5.0×10^7 , 1.0×10^8 and 2.0×10^8 particles m^{-3} . These are labelled a to f respectively in this figure.



expanding in a medium with density $\geq 2.0 \times 10^7$ particles m^{-3} . These densities are much higher than those expected in the general I.S.M. (Long and Helfand 1979) and are more typical of densities found in the peripheries of large massive molecular clouds.

A molecular cloud of mass M_c and radius R_c in the later stages of isothermal collapse will have a radial density distribution of the form

$$\rho(r) = \frac{1}{2} \left(\frac{M_c}{\frac{4}{3}\pi R_c^3} \right) r^{3/2} \quad 7 - 5$$

and a radial velocity distribution of the form

$$V(r) = (2G M_c / r)^{1/2} \quad 7 - 6$$

If the stellar wind bubbles in our sample are located within such large massive collapsing molecular clouds, then it is possible to place a lower limit on the I.S.M. density surrounding the HII region by considering the case in which the HII region radius (r) is comparable to the cloud radius i.e.

$$\rho(R_c) = \frac{1}{2} \left(\frac{M_c}{\frac{4}{3}\pi R_c^3} \right) \quad 7 - 7$$

and

$$V(R_c) = (2GM_c / R_c)^{1/2} \quad 7 - 8$$

In this situation it is possible to determine the total mass of the molecular cloud and thus the I.S.M. density

outside the HII region, by considering the balance between the pressure forces in the outer 3 zones of a stellar wind bubble. The inner ionization bounded HII region is in pressure equilibrium with the pad of accreted HI-gas, which in turn is coupled to the neutral accretion flow through a strong isothermal shock. Ignoring the gas pressure in the neutral accretion flow we obtain the relationship

$$\rho_{\text{II}} c_{\text{II}}^2 = \rho_{\text{S}} c_{\text{I}}^2 = \rho_{\text{O}} v_a^2 \quad 7 - 9$$

where ρ_{II} , ρ_{S} and ρ_{O} are densities in the three zones described above, c_{II} is the isothermal sound speed in the HII shell, c_{I} is the isothermal sound speed in the HI gas and v_a ($\equiv V(R_c)$) is the accretion velocity of the cloud material. Equations 7-7 and 7-8 together with equation 7-9 give the mass of the molecular cloud

$$M_5 = 7.742 \times 10^{-7} n_{\text{II}}^{1/2} R_c^2 \quad 7 - 10$$

and the density of the surrounding I.S.M.

$$n_{\text{O}} = 2.674 \times 10^5 n_{\text{II}}^{1/2} R_c^{-1} \quad 7 - 11$$

where R_c is the radius of the HII region in parsecs, M_5 is the mass of the molecular cloud in units of $10^5 M_{\odot}$ and n_{II} is the density of the HII shell in particles m^{-3} .

Column 4 of Table 7-9 lists the densities plotted in figure 7-6 and compares them with the densities calculated from equation 7-7 (column 3). As indicated above, the latter

TABLE 7.9

NEBULA	n_{II}	n_{O}	$\bar{n}_{\text{O}}$	M_5
301	5.8	13.4	20.6	5.2
25	4.6	13.2	17.8	3.1
137	2.4	4.7	5.7	13.1
229	6.2	10.1	12.9	10.7
196	6.9	14.8	18.1	6.7
106	9.6:	15.6	54.9	8.60
11	28.9:	58.9	194.8	3.5
226	21.9:	40.4	105.0	4.7
235	7.3:	18.5	35.5	5.1
31	4.2	7.7	2.7	11.0

densities are lower bounds since the stellar wind bubble may have not yet reached the outer boundaries of the molecular cloud. Thus it would appear that the HII regions are being rampressure confined by collapsing gas clouds. Thus these observations appear to support a model which combines features of the classical stellar wind model (e.g. Castor McCray and Weaver, 1975) and the ram-pressure confined bubble model of Dopita (1981). This hot cushion, ram-pressure confined model successfully predicts the following observational characteristics;

- 1) the shell structure for the HII regions
- 2) the fact that the HII regions are in general ionization bounded
- 3) the HII region densities
- 4) the ages of the HII regions
- 5) the HII region sizes

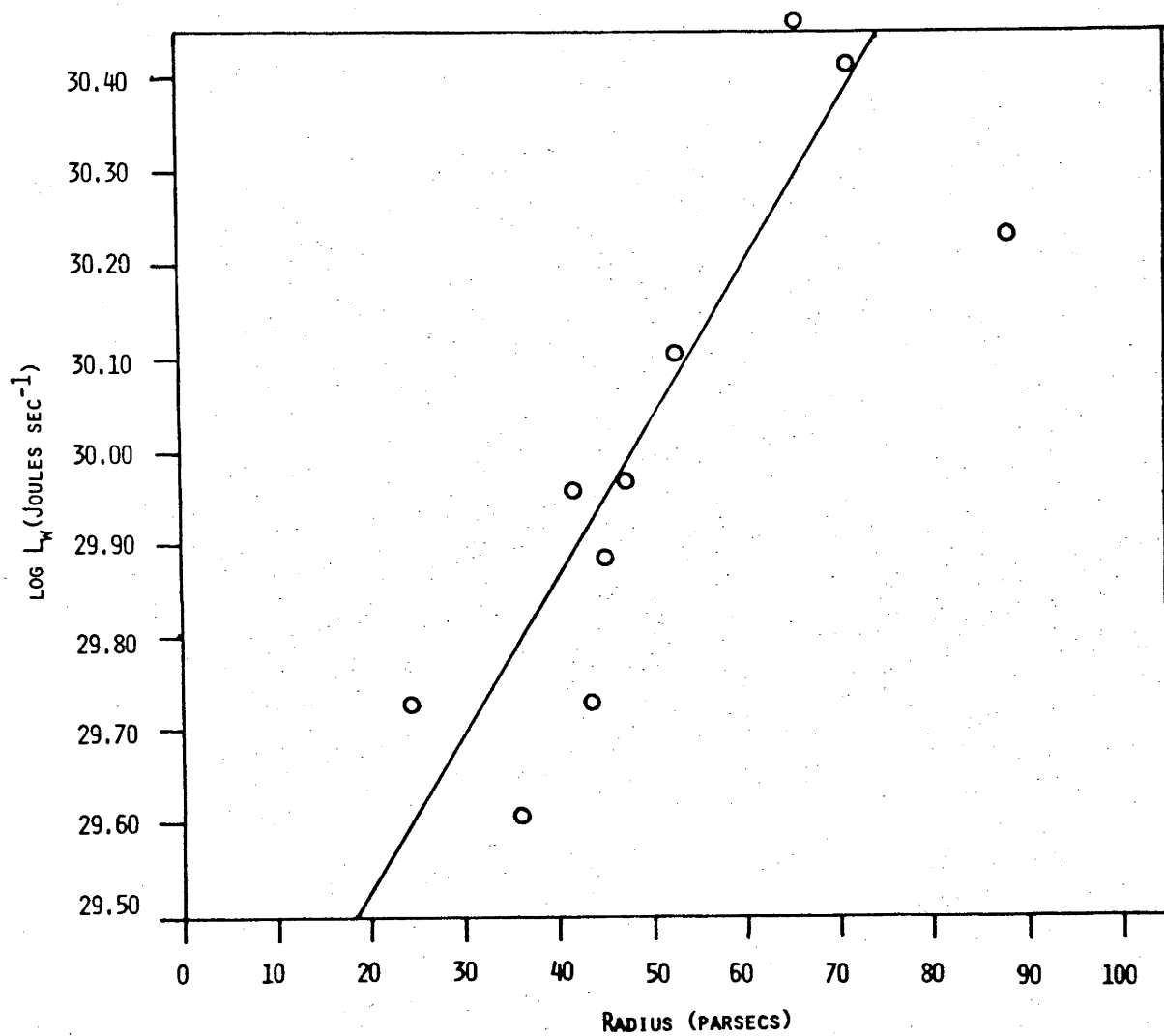
There are two further observational facts which support this thesis. Figure 7-7 shows a plot of $\log L_w$ against the HII region radius in parsecs. This plot shows that there is a reasonably good relationship between these two parameters. The line of least squares drawn in figure 7-7 gives an expression

$$\log L_w = 0.017 R(\text{pc}) + 29.18 \qquad 7 - 12$$

with a dispersion of 0.15 in $\log L_w$ about the mean relation-

Figure 7.7

This figure shows $\text{Log } L_w \text{ (Jsec}^{-1}\text{)}$ for the L.M.C. HII regions plotted against their radii (pc). The curve corresponding to a linear-least-square fit to the data points, is drawn in this figure.



ship. The hot cushion stellar wind model predicts that $R \propto L_w^{1/5}$ which is distinctly different from the observed dependancy. This is taken as further evidence that the radii of the HII regions is determined by the interaction of the stellar wind bubble with a collapsing gas cloud and not a constant density media as assumed in the H-C model.

Larsen (1982) showed that the mass of the most massive star associated with molecular cloud is related to the total mass of the cloud. A more massive cloud will produce a more massive star which, in turn, would provide an increased energy input via stellar winds and ionizing photons. The actual radius of the final stellar wind bubble will thus depend on the interaction between the total cloud mass and the resulting stellar mechanical energy from the stars formed by this cloud. This relationship is constrained by the fact that it must reproduce the observed dependancy of L_w on HII region diameters.

This idea receives support from a comparison of figures 7-4 and 7-7. Figure 7-4 shows that there is a relatively strong correlation between the integrated dereddened H- α flux of a HII region (f_a) and its radius (R) which is given by:

$$\log f_a = 0.019 R \text{ (pc)} - 13.49 \quad 7-13$$

This compares with the corresponding relationship between $\log L_w$ and R given by equation 7-12. The slope of these two relationships are essentially identical within the

errors indicating that

$$L_w \approx 4.7 \times 10^{42} f_a \quad 7 - 14$$

This would imply that

$$L_w \approx 2.4 \times 10^{28} S_{48} \quad 7 - 15$$

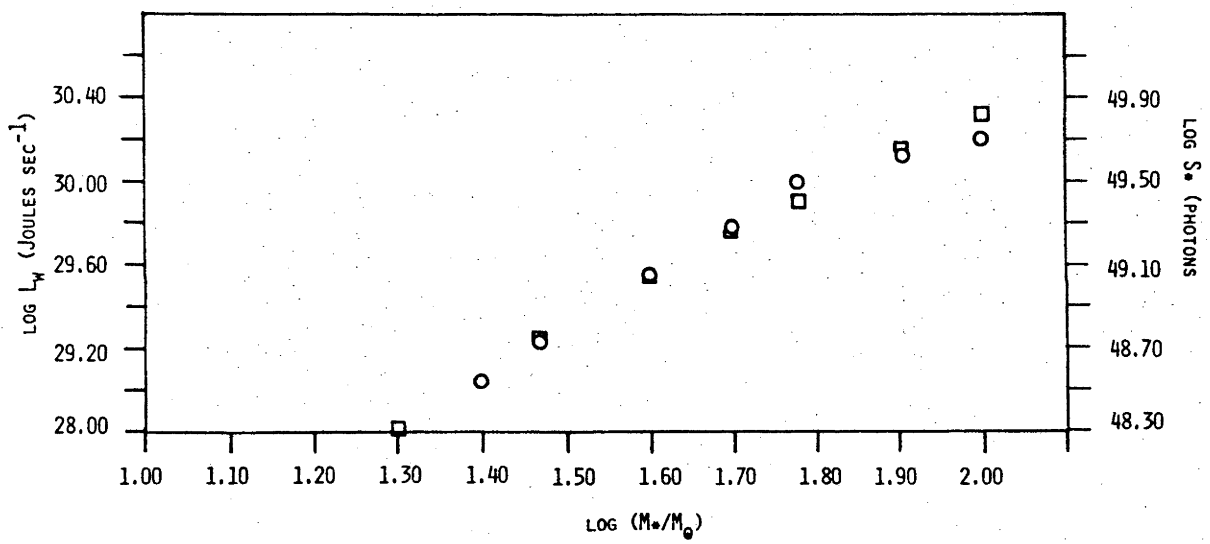
where S_{48} is the Lyman continuum photon emission rate in units of 10^{48} photons sec^{-1} . This is in remarkable agreement with the relationship between L_w and S_{48} which was found in chapter 6 for an OB star of a given mass. Figure 7-8 shows the values of $\log L_w$ (left ordinate) and $\log S_{48}$ (right ordinate) as a function of the stellar mass of the stars on the zero-age main sequence. This shows that

$$L_w = (3.2 \pm 0.4) \times 10^{28} S_{48} \quad 7 - 16$$

Considering the uncertainty of some of the input physics the agreement between these two different approaches is gratifying. It shows that the line drawn in figure 7-4 represents the boundary between ionization and density bounded HII regions, confirming our earlier conclusions on this topic. More importantly though, it shows that the stellar wind mechanical energy is tied to the ionizing photon emission rate. Thus any O and early B star which is a strong ionizing source will also be a source of powerful stellar winds. In this situation it is difficult to envisage a situation in which a HII region will not be affected by stellar winds from OB stars. The only way to avoid this conclusion would require some mechanism which would destroy the hot cushion of gas which would develop as the stellar wind interacts with the surrounding I.S.M.

Figure 7.8

This figure shows the logarithm of both the Lyman continuum photon luminosity (S_* photons sec^{-1}) (left ordinate) and the stellar wind mechanical luminosity (L_w Joules sec^{-1}) (right ordinate), plotted against the logarithm of the stellar initial mass (in solar units). The scales on the two ordinates have been shifted so that the two relationships, relating L_w and S_* to $\text{Log}(M_*/M_\odot)$, correspond. This figure shows that the ratio of L_w to S_* has a value of $3.2 \pm 0.4 \times 10^{-20}$ Joules photon^{-1} which is independent of the initial mass of the star. The relationships shown in this figure correspond to stars on the initial main sequence, but since both L_w and S_* are shown to be roughly constant during the core-hydrogen burning lifetime of the OB stars, this ratio should not change during this evolutionary phase.



A note of caution should be sounded on the ram-pressure confined H-C stellar wind model, however. Figure 7-10 shows a relative intensity plot of an average cross-section of the nebula D.E.M. 25. It clearly shows that ^{the} HII shell is partially filled in, which would not be the case if the H-C model applied to this nebula. This characteristic is also shown by almost all of the other HII shells for which T.V. surface photometry is available. Since the H.C. model is required to produce the observed densities in the HII shells, we are forced to conclude that the hot-cushion of shocked stellar wind gas has recently collapsed in the L.M.C. ring nebulae. The collapse of this pad of gas would reduce the pressure force acting on the inner side of the HII shell resulting in the ionized shell expanding into the low density cavity created by the stellar wind bubble. An alternative hypothesis is that the stellar wind has entrained low density HII material from condensations embedded near the exciting stars.

If the hot cushion has indeed collapsed, this raises a number of questions which must be solved by further study;

- at what stage in the evolution of the stellar wind bubble does the H-C collapse?
- What prevents the break-up of the H-C by Kelvin-Helmholtz and Rayleigh-Taylor instabilities?
- is the collapse of the H-C sudden or gradual?
- if the H-C were to collapse at an earlier point

in the evolution of the stellar wind
bubble would a shell nebula still
be formed?

These intriguing questions represent a good point
to end this thesis, and we hope that future work will be
able to answer them.

Figures 7.9 a to m

Figures 7.9a to m show the absolute hydrogen-alpha surface brightness maps for 13 HII regions in the L.M.C. Table 7.10 lists the order in which each nebula appears (column 1); the scale of the solid bar drawn in each figure (column 2); the contour levels in each figure (column 3); and the correction factor which must be applied to each figure (column 4). This correction factor is necessary, because a mean conversion factor was used to transform the T.V. photometer into absolute units, rather than the conversion factor for each night (see table 7.6). These surface brightness plots must also be corrected for the mean atmospheric extraction of each nebula. This is included in the correction factor in Table 7.10.

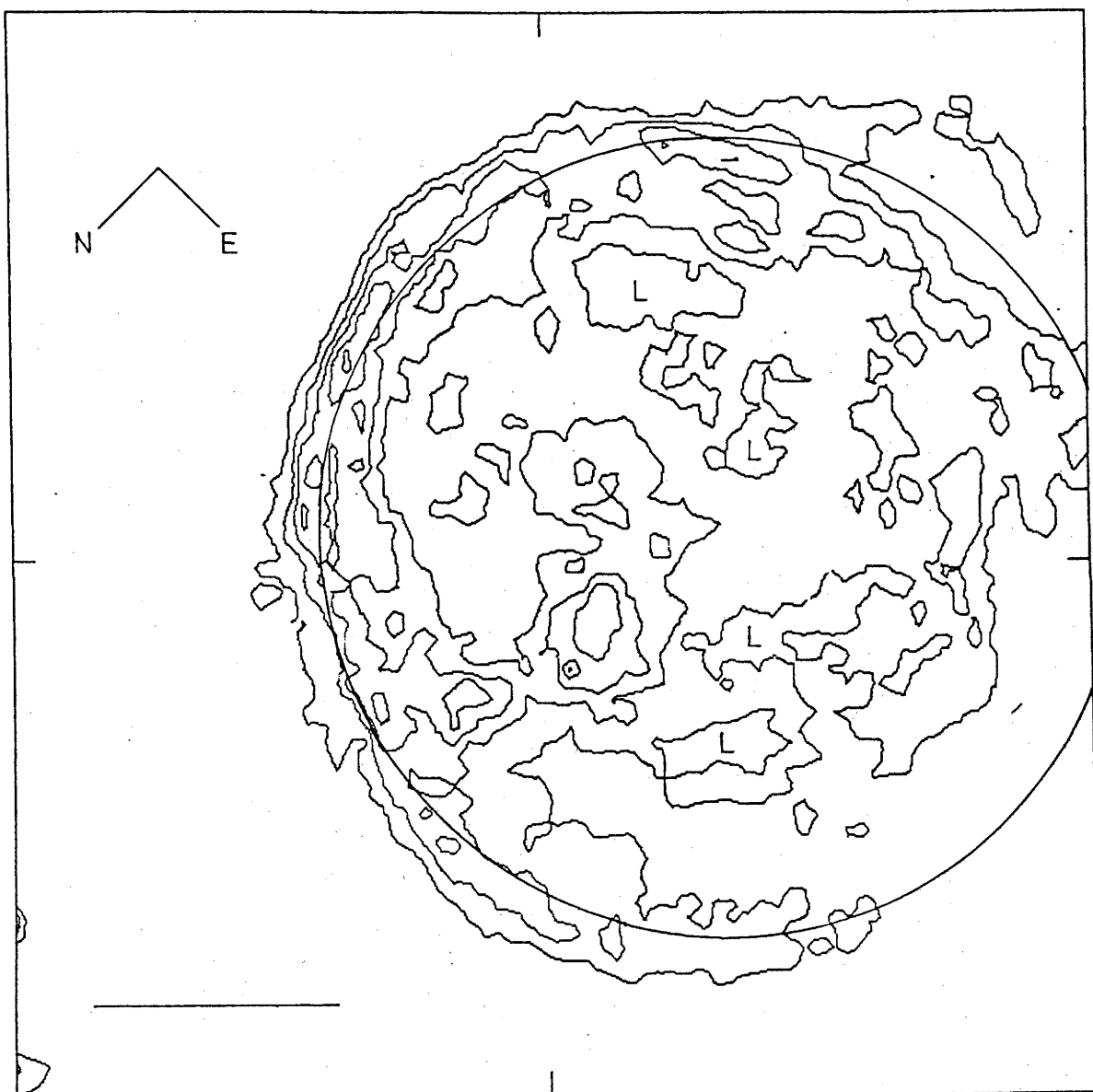
Five additional facts which are necessary for the interpretation of figures 7.9 a to m are:

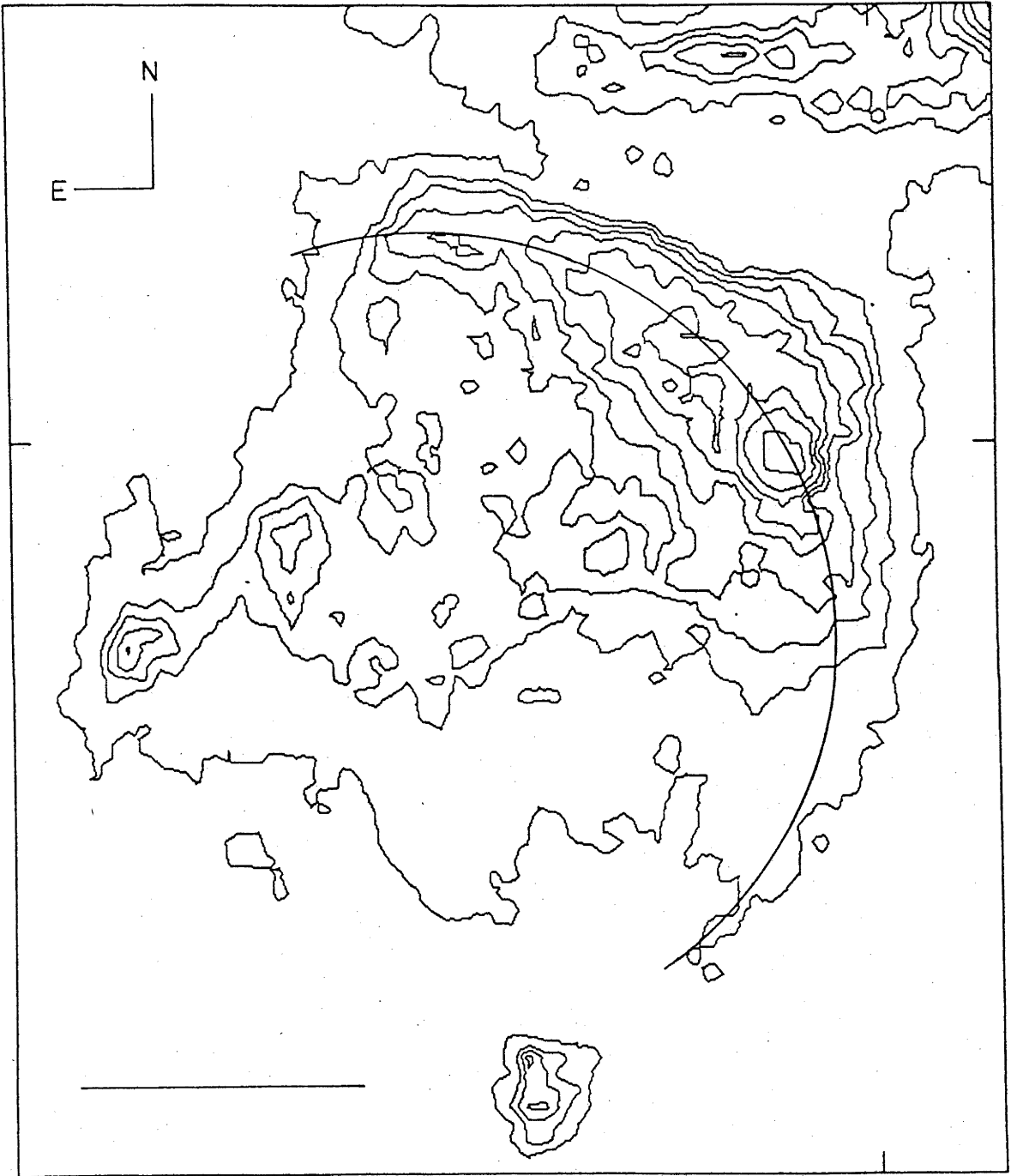
1. The orientation in each plot is shown by arrows which indicate the approximate directions of north and east;
2. An L symbol designates areas which are lower in surface brightness than the surrounding regions;
3. D.E.M. 145 is the nebula in the northwest of figure 7.9.1.,

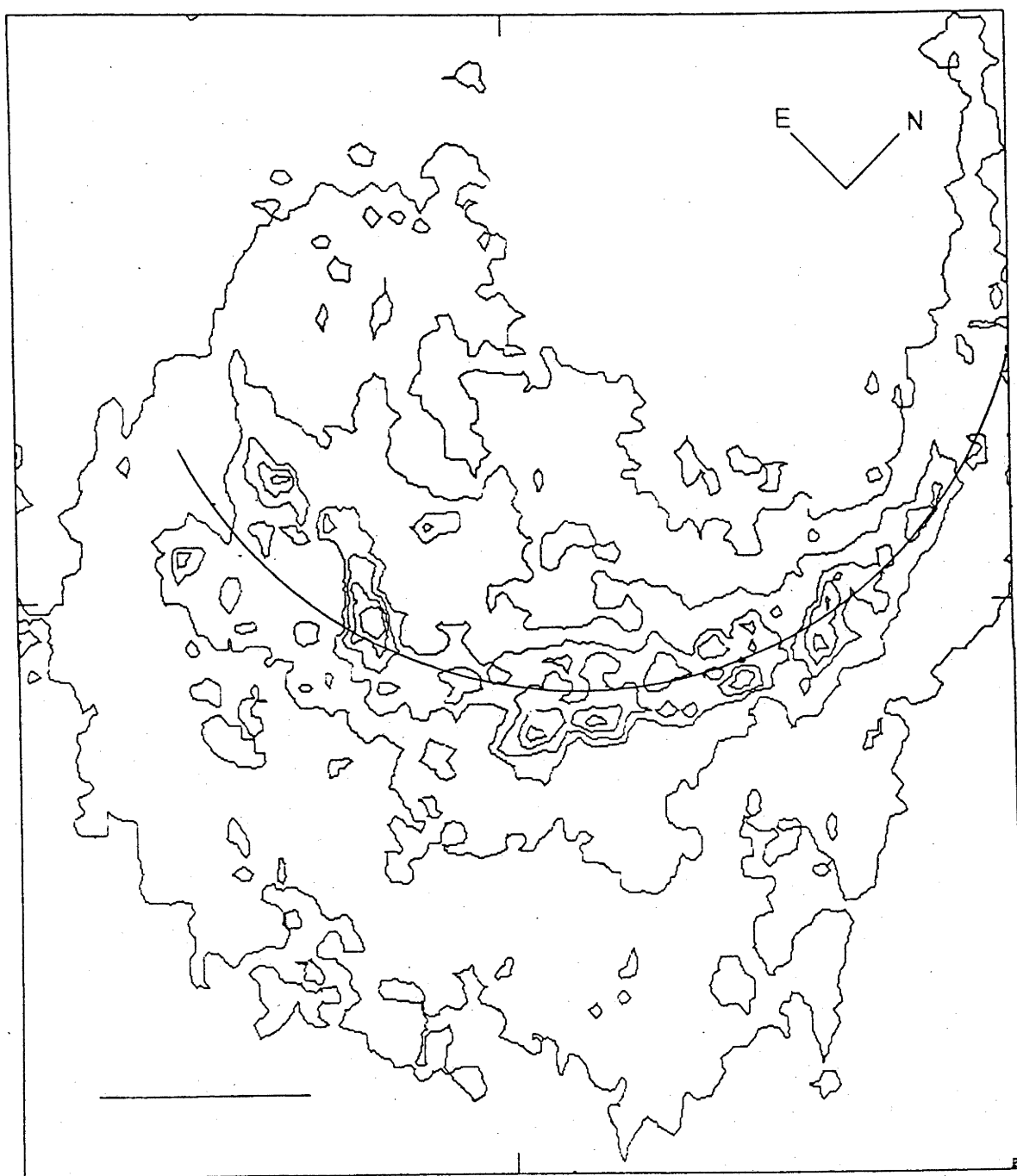
4. Apparent surface brightness enhancements near the edges of the unvignetted field, and
5. The arcs and circles found in each figure are explained in the text.

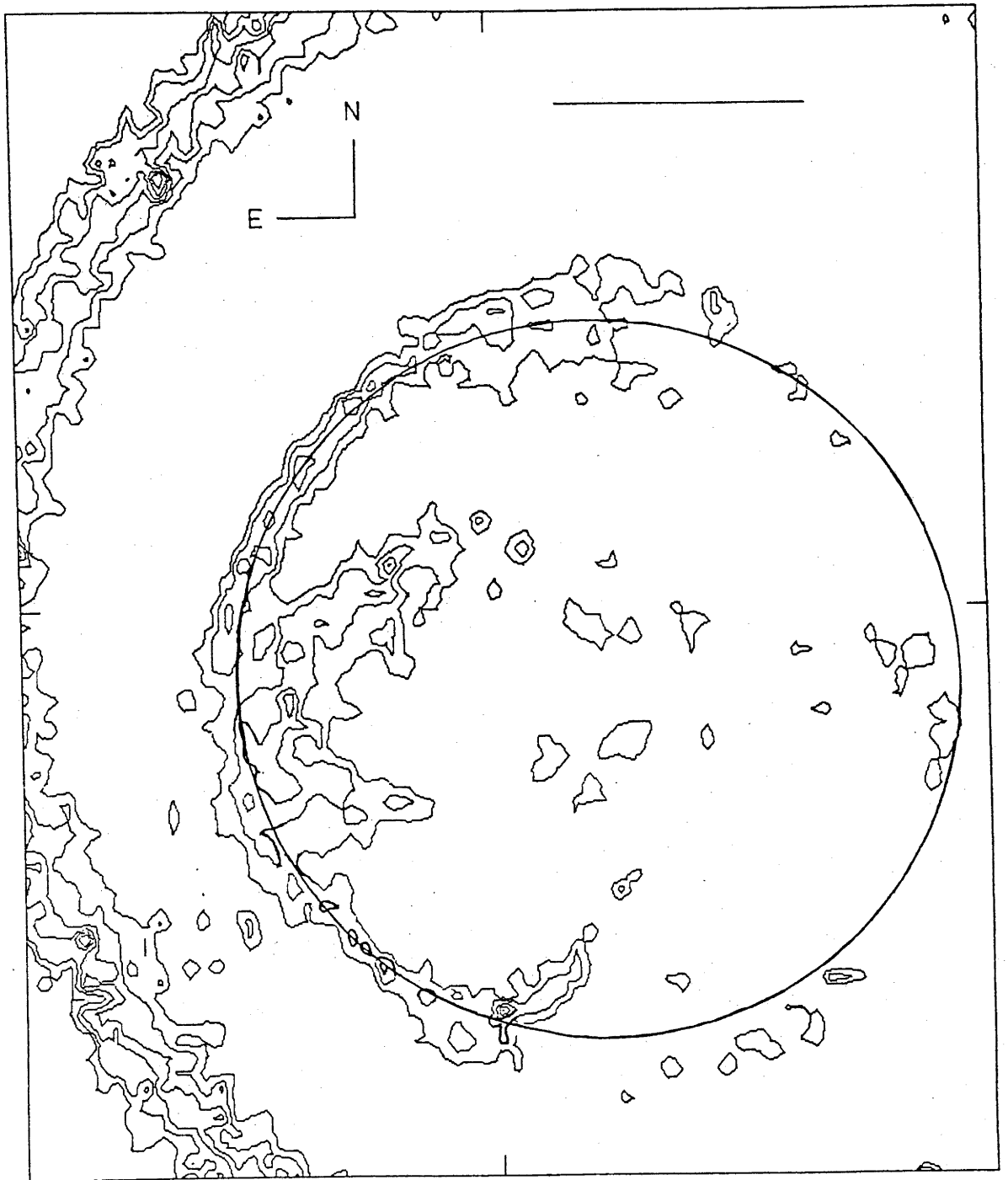
TABLE 7.10

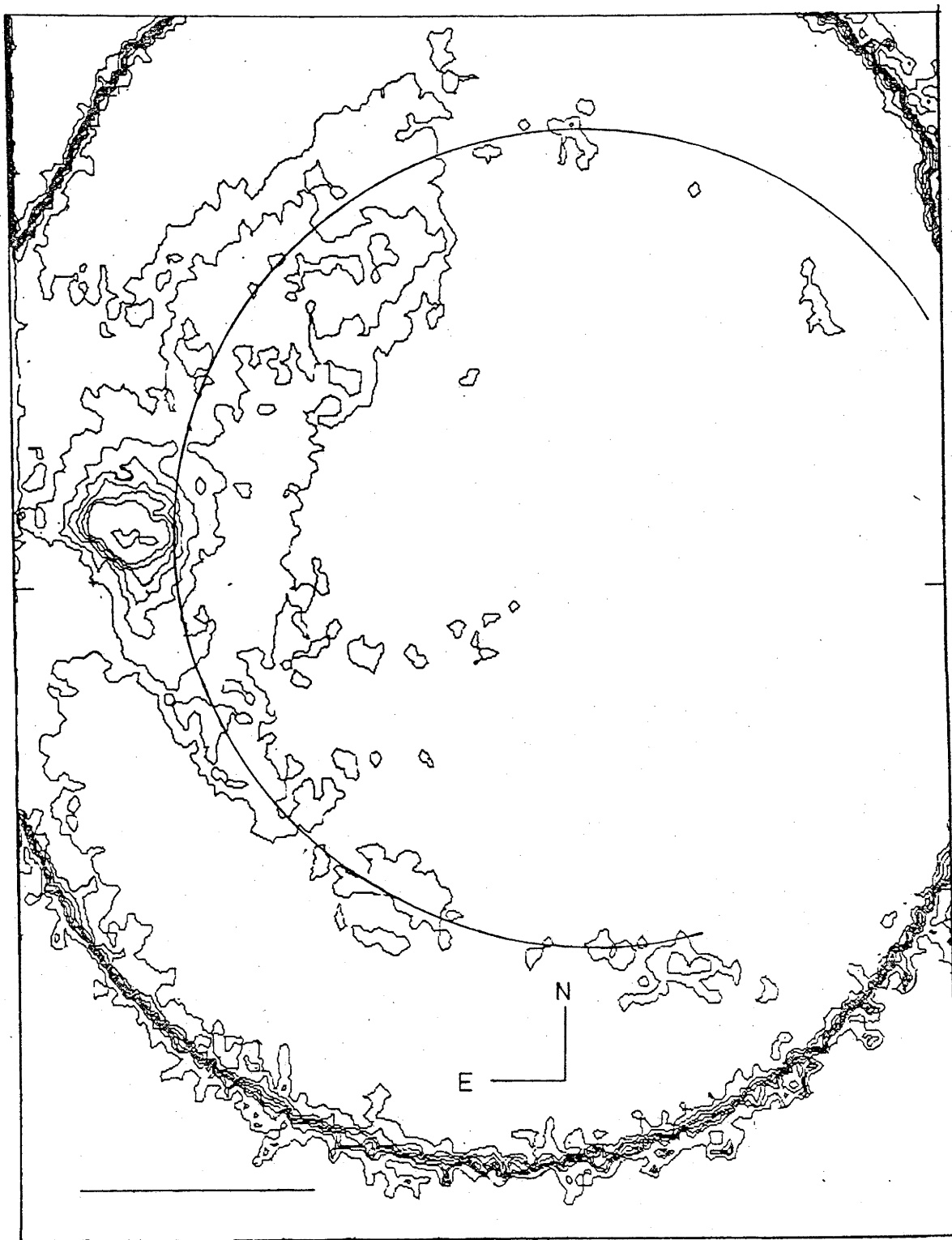
<u>D.E.M.</u> <u>NO.</u>	<u>SCALE</u> <u>ARC MINUTES</u>	<u>CONTOUR INTERVALS</u>	<u>CORRECTION</u> <u>FACTOR</u>
301	2	0.5,1.0,2.0,3.0,4.0	1.115
196	2	0.5,1.0,1.3,1.6,1.9,2.2,2.5,3.0,4.0	1.111
31	2	0.5,1.0,1.5,1.8,2.1,2.4	0.851
25	2	0.5,0.7,0.9,1.1,1.3	1.108
137	3	0.3,0.5,0.7,0.9,1.1,1.3,1.5,2.0	1.100
192	2	0.5,2.0,3.0,4.0,5.0,6.0	0.851
229	2	0.5,2.0,4.0,6.0,8.0,10.0,20.0	1.097
231	1	0.5,1.0,2.0,2.5,3.0,3.5,4.0	1.097
311	1	0.5,1.0,2.0,4.0,8.0	1.069
121	1	0.5,1.0,2.0,4.0,5.0,5.5	1.066
281	1	0.5,1.0,1.5,2.0,2.5,3.0	1.069
145/149	1	0.5,1.0,2.0,3.0,4.0,5.0,10.0	1.099
138	1	0.7,1.0,2.0,4.0,6.0,8.0,10.0,12.0	1.067

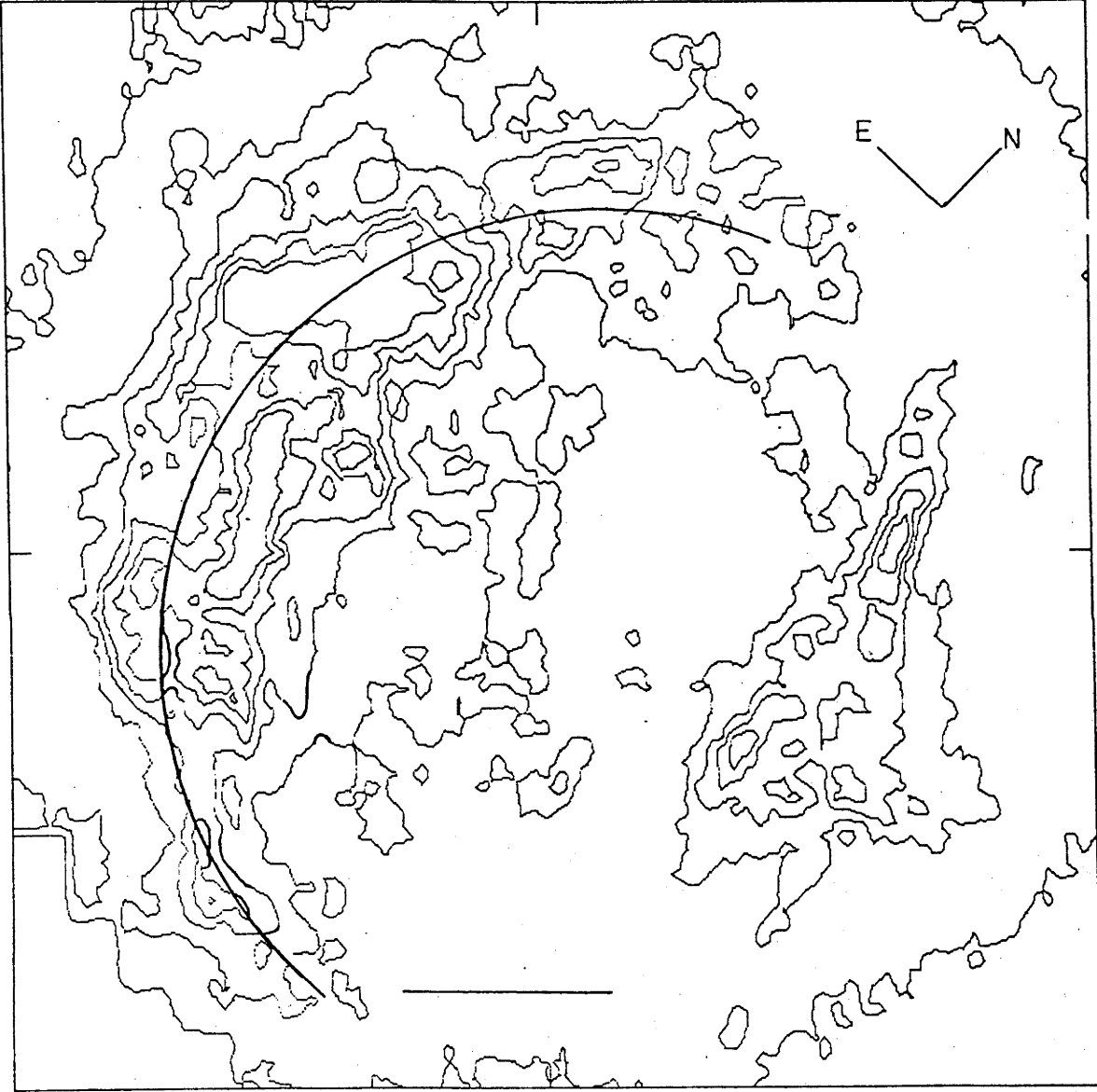


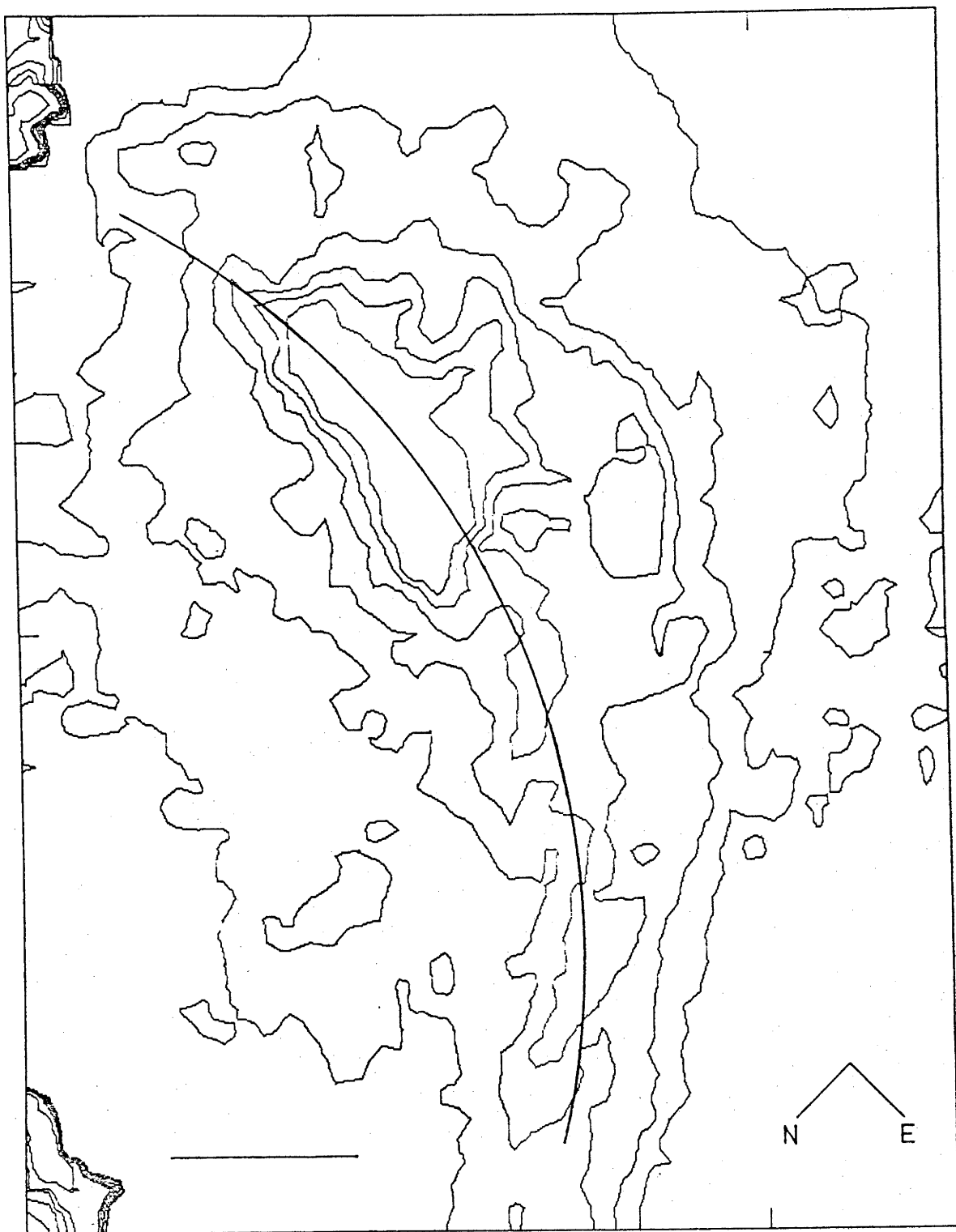


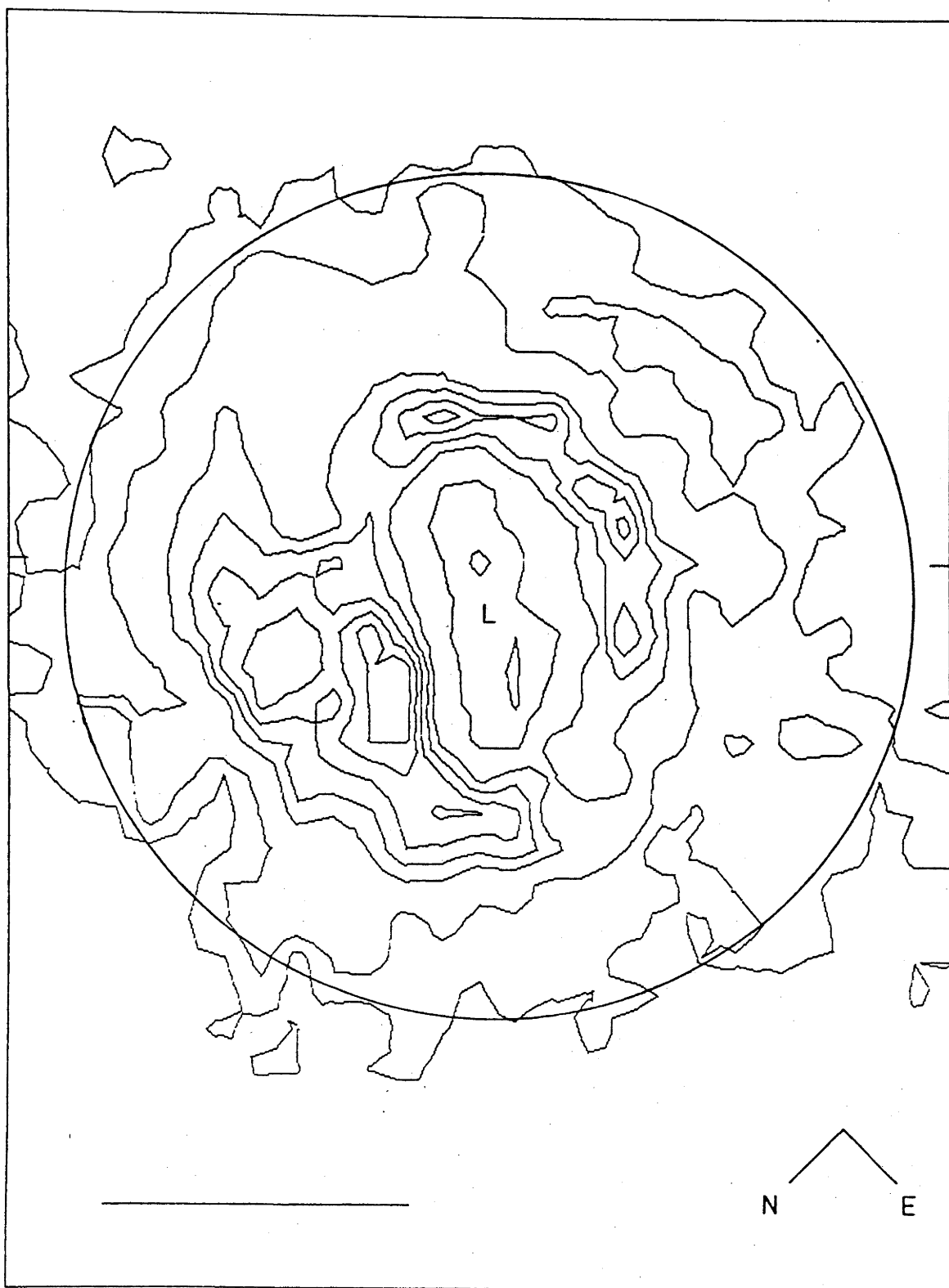


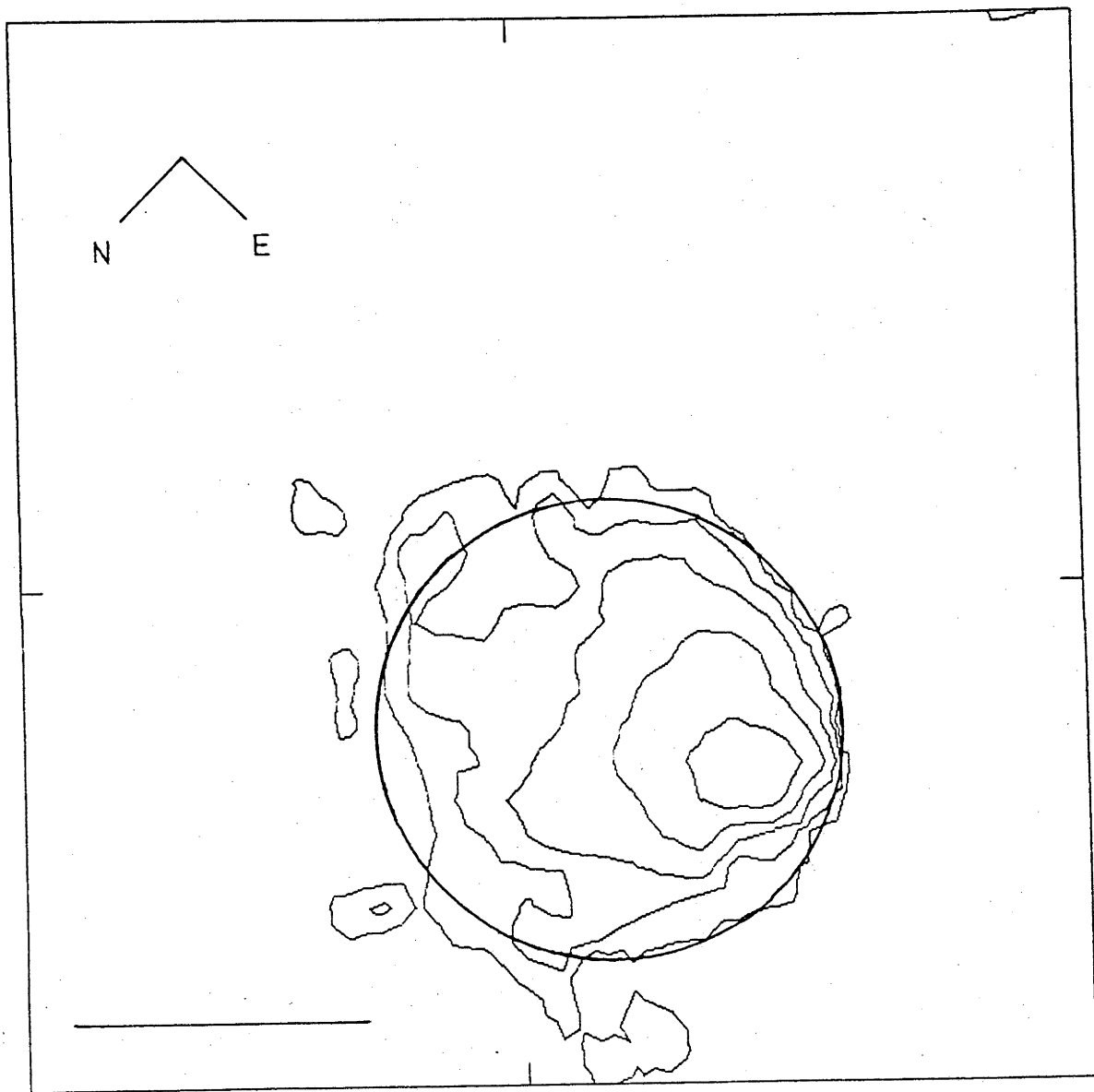


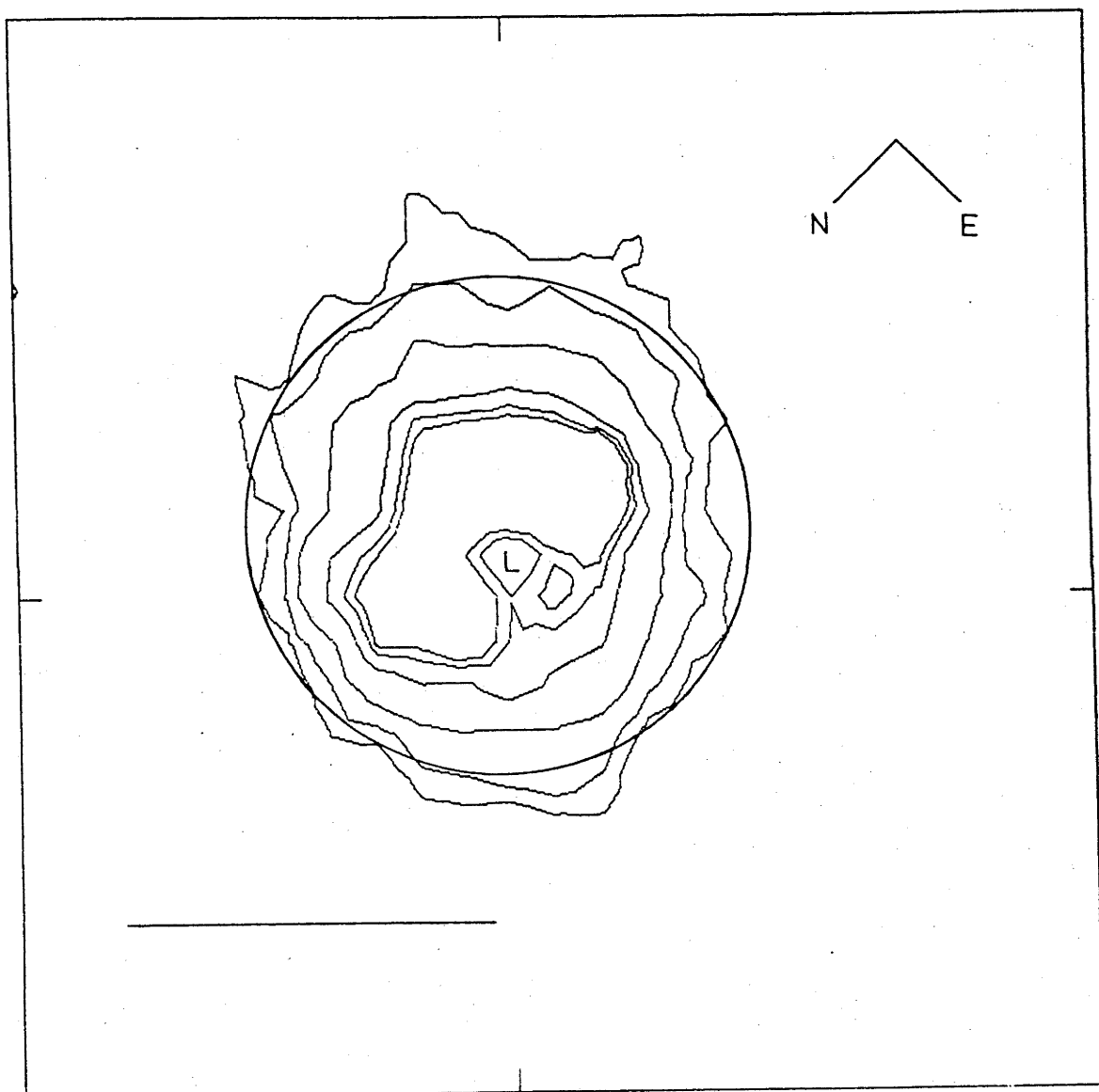


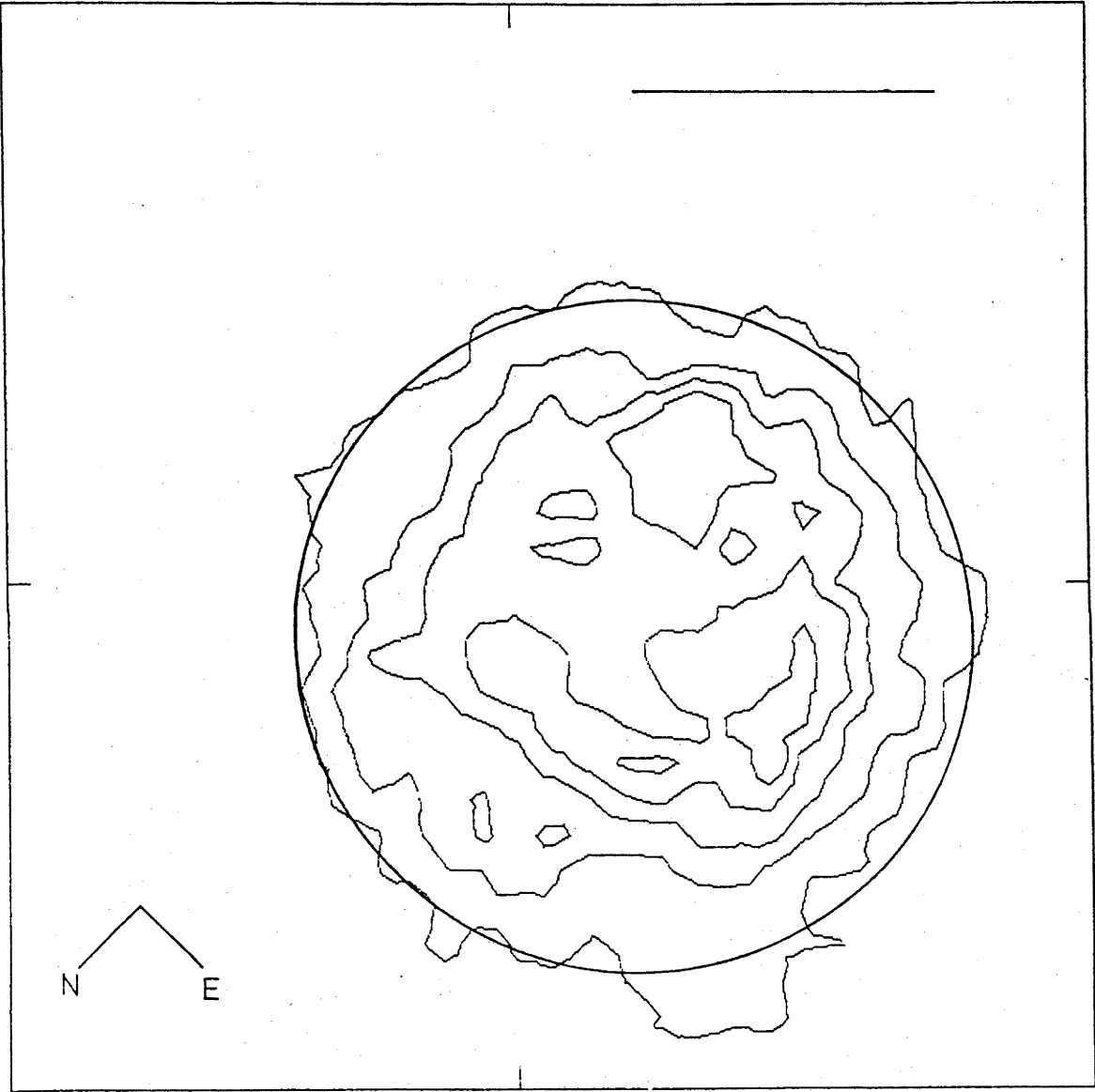


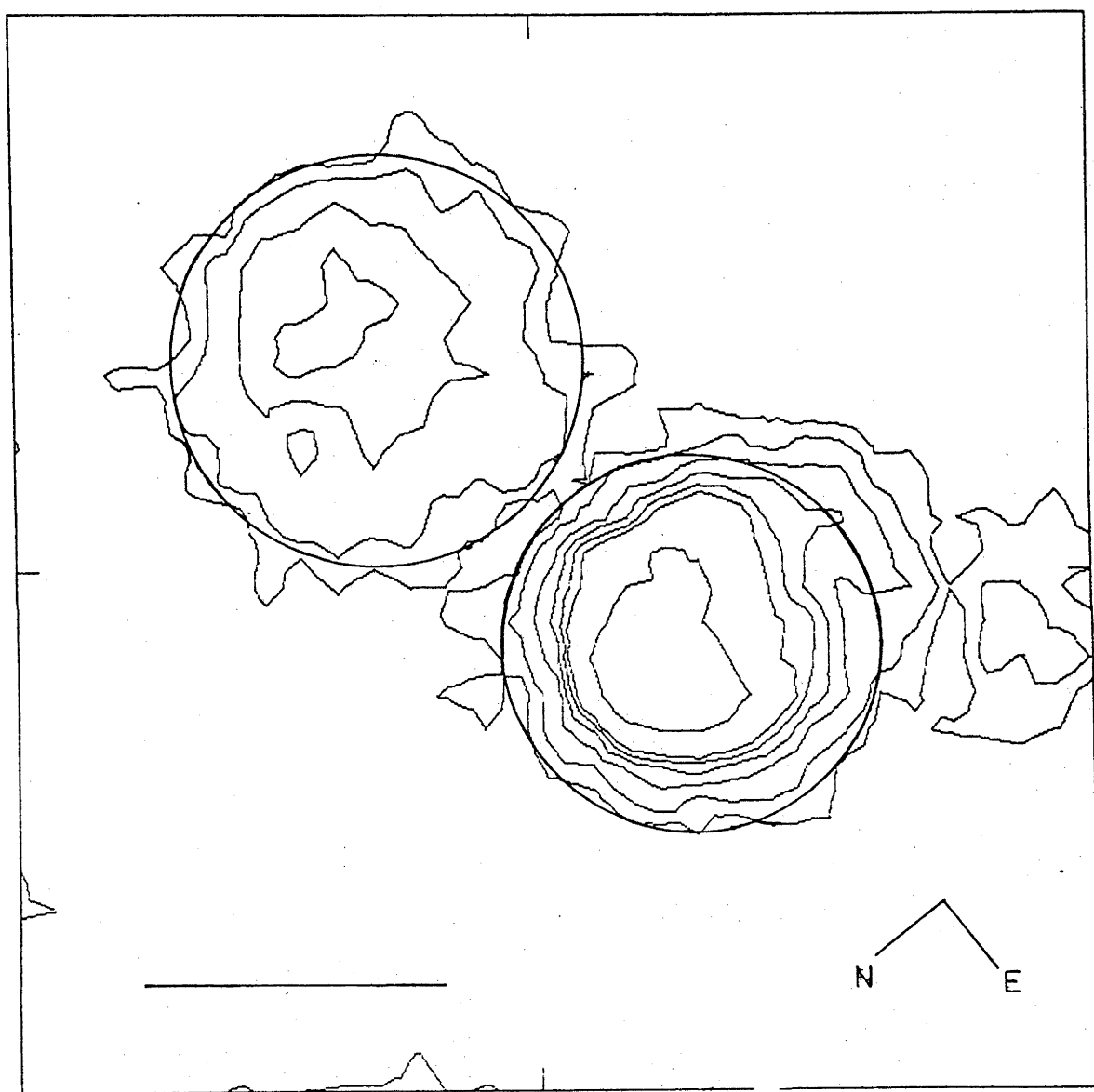












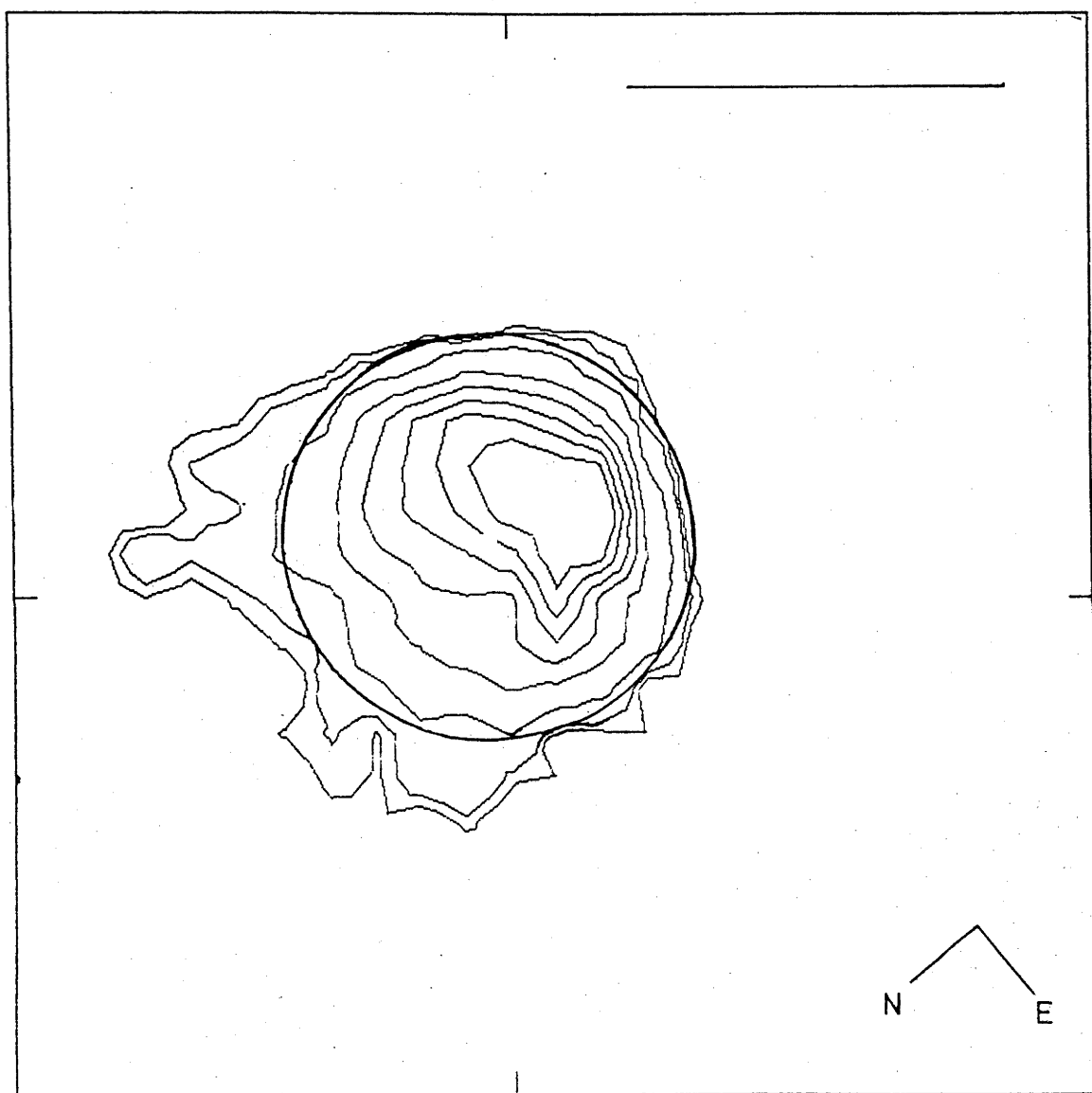
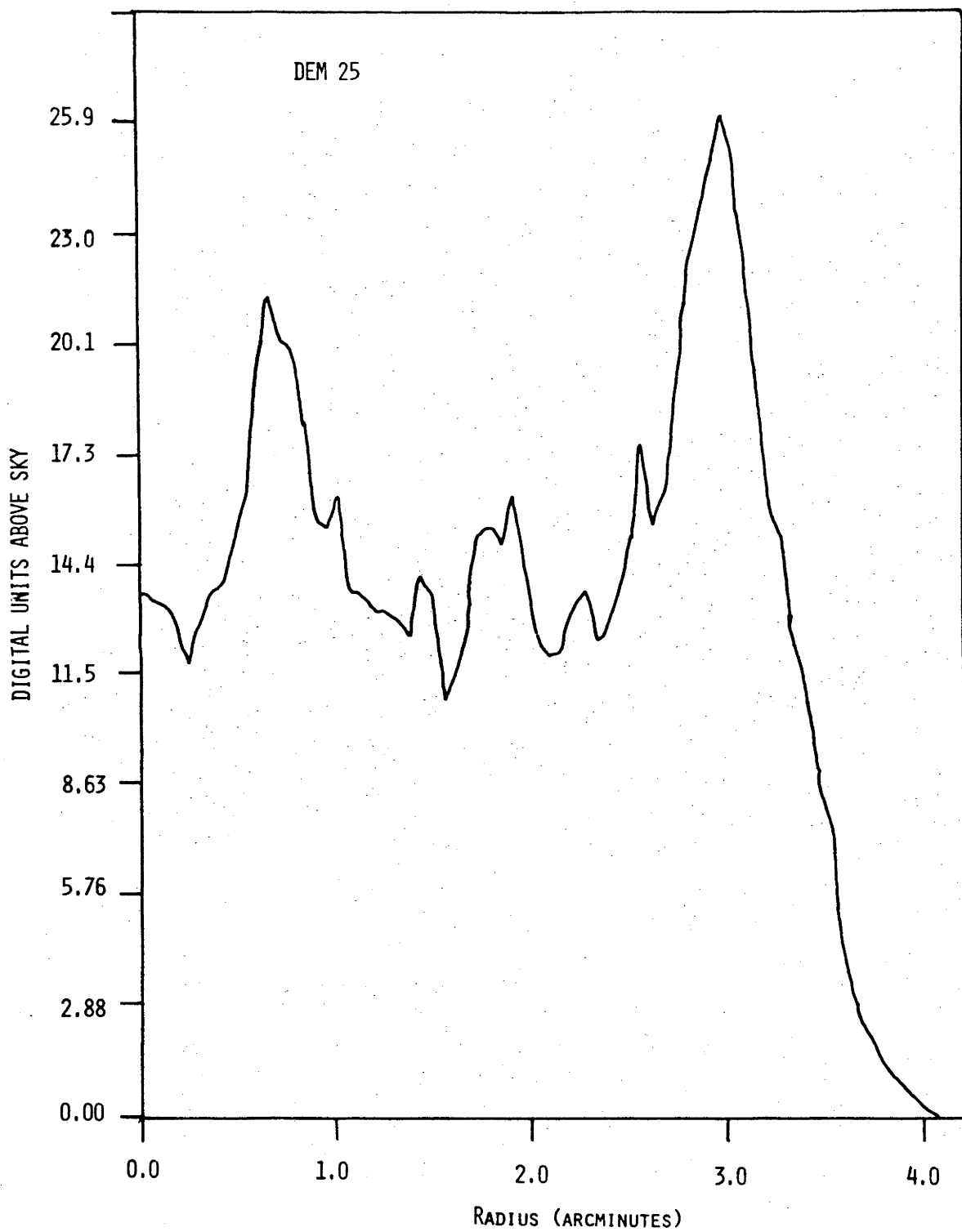


Figure 7.10

This figure shows the average surface brightness profile for the nebula D.E.M. 25. The profile is formed by taking the average of successive cuts through the nebula. The cuts are spaced at 15° intervals between position angles 225° and 45° . The top of figure 7.9d is taken as 0° position angle and all angles are measured from the centre of D.E.M. 25. The position angle and all angles are measured from the centre of D.E.M. 25. The position Angles 75° , 90° and 105° were left out of the final average, as there were local surface brightness enhancements in these directions.

Figure 7.10 clearly shows that the shell structure of D.E.M. 25 has been partially filled in. The numbers on the ordinate are arbitrary while those on the abscissa are in arc minutes.



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