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MASSIVE STARS IN THE LOCAL GROUP: Implications for Stellar Evolution and Star Formation

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■ **Abstract** The galaxies of the Local Group serve as important laboratories for understanding the physics of massive stars. Here I discuss what is involved in identifying various kinds of massive stars in nearby galaxies: the hydrogen-burning O-type stars and their evolved He-burning evolutionary descendants, the luminous blue variables, red supergiants, and Wolf-Rayet stars. Primarily I review what our knowledge of the massive star population in nearby galaxies has taught us about stellar evolution and star formation. I show that the current generation of stellar evolutionary models do well at matching some of the observed features and provide a look at the sort of new observational data that will provide a benchmark against which new models can be evaluated.

1. INTRODUCTION

Massive stars are extremely rare: For every 20 $\mathcal{M}_\odot$ star in the the Milky Way there are roughly a hundred thousand solar-type stars; for every 100 $\mathcal{M}_\odot$ star there should be over a million solar-type stars. However, their importance is considerably out of proportion to their scant numbers. Through the actions of their strong stellar winds and eventual disruption as supernovae, they provide most of the mechanical energy input into the interstellar medium (Abbott 1982b). They also generate most of the ultraviolet (UV) ionizing radiation in galaxies, and power the far-infrared luminosities through the heating of dust. And, massive stars serve as the primary source of carbon, nitrogen, and oxygen (CNO) enrichment of the interstellar medium (ISM) (Maeder 1981). Massive stars are believed by many to be the source of the most energetic phenomenon yet found, emitting gamma-ray bursts as they collapse into black holes (Woosley 1993, Bloom et al. 2002, Price et al. 2002).

The evolution of massive stars is difficult to model, owing primarily to the complications of mass loss. A very massive star might lose half of its mass during its core H-burning main-sequence lifetime. Such mass loss has a profound effect on the evolution of massive stars, as shown by de Loore et al. (1977, 1978, 1979), Chiosi et al. (1978, 1979), and Brunish & Truran (1982) and the studies that followed. The parameterization of mass loss on physical parameters (such as luminosity, effective temperature, and metallicity) is still somewhat uncertain even during the main-sequence phase (Chiosi & Maeder 1986, de Jager et al. 1988, Lamers & Leitherer 1993), although there has been significant improvement in our understanding in recent years (Kudritzki & Puls 2000, Kudritzki 2002, Puls et al. 2000, Vink et al. 2001). Beyond the main sequence, the characterization of mass loss is even more problematic. For instance, the luminous blue variables (LBVs) are thought to be a short-lived phase that the most massive stars enter after core H-burning. These stars suffer huge but episodic mass-loss events, the origins of which are poorly understood (Lamers 1987, Humphreys & Davidson 1994). Similarly the mass-loss rates for red supergiants (RSGs) may be much higher than previously recognized (Salasnich et al. 1999), with large effects on the subsequent stellar evolutionary tracks. Wolf-Rayet stars (WRs) are the final evolutionary stage of high-mass stars, and they have the highest mass-loss rates of all, but their deduced physical properties are highly sensitive to the details of atmosphere models such as whether the stellar winds are assumed to be "clumped" (Crowther et al. 2002b). Because mass loss is driven by radiation pressure acting through highly ionized metal lines, the mass-loss rates on the main sequence (before the surface layers are enriched) will depend on an unknown power of the initial metallicity Z. It is uncertain to what extent the mass-loss rates of evolved stars will depend on Z, as the surface abundances of WRs are primarily the results of their own nuclear burning, although native elements such as iron may still be important in accelerating the wind (Crowther et al. 2002b).

In addition to the problems of modeling mass loss, other physics is also uncertain, and so observations are an integral part of testing and tweaking the models. Among these are the treatment of mixing and convection (Maeder & Conti 1994). In this regard, the inclusion of rotation into the models may prove to have as profound an effect as the inclusion of mass loss. It is well known that the spectra of unevolved massive stars (i.e., O-type) show highly Doppler-broadened lines (Conti & Ebbets 1977, Penny 1996), with typical rotational velocities $v \sin i \sim 100$–300 km s^{-1}. Recent calculations have shown that this rotation is an important transport mechanism of evolved material from the stellar core (Maeder & Meynet 2000a,b; Meynet & Maeder 2000).

Observations of massive stars in the Local Group provide a key diagnostic tool for furthering our understanding of massive star evolution. Given the theoretical uncertainties, progress is largely dependent upon good feedback between the observers and theorists, with critical observations serving to reveal the successes and failures of the models. (See figure 5 of Conti 1982 for one observer's amusing take on this process.) The star-forming galaxies of the Local Group provide an ideal laboratory for such studies, as they span a range of $\sim$15 in Z (WLM to M31). And this is a wonderful time for such studies, with improved stellar models becoming available, and the advent of large field-of-view imagers on 4-m telescopes along with multi-object spectroscopic capabilities on 8-m telescopes.

Such studies are also of value for what we can learn about the physical parameters of massive stars as a function of metallicity. Does the conversion between spectral type and effective temperature depend upon metallicity? Does the observed dependence of mass-loss rates on metallicity match what is expected from radiatively driven wind theory? Similarly, these studies can answer fundamental questions about star-formation processes: What can we infer about the initial mass functions (IMFs) and upper mass cutoffs as a function of metallicity?

In this review I summarize what is known about massive stars in the star-forming galaxies of the Local Group, describing what we have learned about stellar evolution and star formation as a result. Complementary reviews on massive stars are given by Conti (1988), Maeder & Conti (1994), Maeder & Meynet (2000b), and Kudritzki & Puls (2000). Section 2 covers the main-sequence (hydrogen-burning) massive stars, emphasizing that the most massive stars are not the visually brightest in a galaxy (Section 2.1) and describing searches for these stars in the Milky Way, Magellanic Clouds, and beyond (Section 2.2). Section 2.3 describes what the resulting Hertzsprung-Russell (H-R) diagrams have taught us about star formation, the initial mass function, and stellar evolution. Section 2.4 concludes with a brief discussion of mass loss. In Section 3 I describe the evolved (He-burning) massive stars, beginning with the evolutionary link between them (Section 3.1) and then describing what is known about luminous blue variables (Section 3.2), red supergiants (Section 3.3), and Wolf-Rayet stars (Section 3.4). That section concludes by discussing the implications for stellar evolution and making some critical comparisons to the predictions of evolutionary models (Section 3.5). The review ends with a short summary and list of remaining questions (Section 4). First, though, I introduce the galaxies of the Local Group.

1.1. Local Group Galaxies Known to Contain Massive Stars

In Table 1 I list the oxygen abundances, distances, and reddenings of Local Group galaxies thought to contain massive stars. Galaxies are included if they are known to have H II regions and/or luminous blue stars. Additional information about these systems can be found in recent reviews by van den Bergh (2000) and Mateo (1998). The "metallicity" Z is actually the mass fraction of all elements heavier than H and He, but the values for Z in this table are scaled linearly from the oxygen abundances rather than, say, iron (compare Mateo 1998), as oxygen, along with carbon and nitrogen, is one of the primary accelerators of the stellar wind at the high effective temperatures characteristic of main-sequence massive stars (Abbott 1982a). Similarly, in selecting what values of the color excess $E(B-V)$ to add to the list, the selection is biased in favor of the young massive star population and is based on comparing charge-coupled device (CCD) photometry with the intrinsic colors expected on the basis of spectral types, when this information is available.

2. MAIN-SEQUENCE MASSIVE STARS

Massive stars begin their main-sequence lives as O-type stars, with effective temperatures of 30,000–50,000°K. During the main-sequence phase, they convert hydrogen to helium via the CNO cycle, during which the effective temperatures

TABLE 1 Star-forming galaxies of the Local Group

	Metallicity			Distance		Reddening	
Galaxy	**log(O/H) + 12 (dex)**	**References[a]**	**Z[b]**	**$(m - M)_o$**	**References[a]**	**$E(B - V)$**	**References[a]**
MW (d < 3 kpc)	8.70	1	0.018	—	—	—	—
SMC	8.13	2	0.005	18.85	3	0.09	4
LMC	8.37	2	0.008	18.50	3	0.13	4
M31	9.00	5	0.036	24.4	3	0.16	6
M33 ($\rho^c = 0.0$)	8.85	7	0.013	24.5	3	0.13	6
($\rho^c = 1.0$)	8.12	7	0.005				
NGC 6822	8.25	8	0.006	23.48	3	0.38	6
IC 10	8.25	9	0.006	24.10	3	0.80	10
IC 1613	7.85	11	0.003	24.3	3	0.09	3
WLM	7.77	12	0.002	24.83	3	0.02	3
Pegasus	7.93	13	0.003	24.4	3	0.02	3
Phoenix	—	—	—	22.99	3	0.02	3

Notes: [a]References: 1. Esteban & Peimbert (1995); 2. Russell & Dopita (1990); 3. van den Bergh (2000); 4. Massey et al. (1995c); 5. Zaritsky et al. (1994); 6. Massey et al. (1995a); 7. Garnett et al. (1977); 8. Pagel et al. (1980); 9. Garnett (1990); 10. Massey & Armandroff (1995); 11. Talent (1980); 12. Hodge & Miller (1995); 13. Skillman et al. (1997).

[b]Z is scaled from the oxygen abundances (with $Z = 0.018$ for the solar neighborhood) and, hence, is characteristic of the young stellar population.

[c]The quantity ρ is the distance from the nucleus within the plane of M33, normalized to a Holmberg radius of 25 arcminutes (5.8 kpc), following Kwitter & Aller (1981).

decrease slightly, while the bolometric luminosity stays nearly the same. By the end of core H-burning these stars are typically B-type supergiants. Table 2 shows the expected change in spectral type during the main-sequence phase, based upon the $Z = 0.02$ (solar neighborhood metallicity) evolutionary tracks of Schaller et al. (1992) and the effective temperature scale of Vacca et al. (1996).

It is misleading to think in terms of a spectral–subtype-to-mass relation for massive stars. For instance, in Table 2 we find that a star of spectral type "O4 V" may be a zero-age $60\mathcal{M}_\odot$ star, or a slightly older (0.5 Myr) $85\mathcal{M}_\odot$ star. Instead, the spectral types are stages through which stars of different masses pass.

2.1. The Most Luminous Stars are not the Brightest

One of the other things we can infer from Table 2 is that although massive stars evolve at nearly constant bolometric luminosity (at least while on the main sequence), as they grow cooler with time, they become considerably brighter visually, as indicated by the absolute visual magnitude M_V. For instance, a $120\mathcal{M}_\odot$ star remains at nearly constant bolometric magnitude -11, but during its short 2.6 Myr main-sequence life, its absolute visual magnitude changes by 2.4 mag, nearly a factor of 10 in brightness. The reason for this is that much of the bolometric luminosity occurs outside of the visible region for stars with these effective temperatures; even a slight cooling shifts a significantly greater fraction of light into the visible.

When we look at stars in a nearby galaxy, we are looking at a mixed-age population: the stars all have the same distance, and possibly similar reddenings, but we usually have a heterogeneous mix of ages. In such a population, the most luminous stars are not the brightest stars. We illustrate this in Figure 1, where we show lines of constant visual magnitude superposed upon the $Z = 0.004$ (SMC-like) evolutionary tracks from the Geneva group (Charbonnel et al. 1993).

The visually brightest stars in Figure 1 are not the most bolometrically luminous, and hence are not the most massive. Instead, at $M_V < -8$ ($V < 11$ in the SMC) we find A-type ($\log T_{\rm eff} \sim 4.0$) supergiants with masses of 25 $\mathcal{M}_\odot$. The zero-age main-sequence (ZAMS) stars of 60 $\mathcal{M}_\odot$ are 3 mag fainter than this visually; i.e., a factor of 16 less in brightness, despite being a magnitude or so more luminous bolometrically. This faintness is simply caused by the changing bolometric correction with effective temperature.

One of the implications of this effect is that magnitude-limited catalogs may be incomplete for the youngest massive stars. A catalog that goes deep enough to include A-type 7 $\mathcal{M}_\odot$ stars will nevertheless be incomplete for ZAMS stars below $\sim 30\mathcal{M}_\odot$, as can be seen by following the $V = 15$ curve in Figure 1(lower). A 25 $\mathcal{M}_\odot$ star on the ZAMS will have $V \sim 15.5$, whereas a 15 $\mathcal{M}_\odot$ star on the ZAMS will have $V \sim 16$, although both are more bolometrically luminous than the evolved A-type 7 $\mathcal{M}_\odot$ star.

2.2. Searches for O Stars

A well-known spectroscopist is said to have told her students, "You can't do astrophysics by just taking images through little colored pieces of glass!" This adage applies particularly to massive main-sequence stars, whose high effective temperatures result in spectral energy distributions far on the tail of the Rayleigh-Jeans law at optical wavelengths. There is little temperature sensitivity in the colors of these stars, and stars of similar intrinsic colors may have very different bolometric corrections. The presence of reddening is an additional complication, requiring the construction of reddening free indices. Massey (1998a) uses model atmospheres to compare the sensitivity of the colometric correction (BC) to different reddening-free indices, including the extension into the UV wavelengths for which photometry is becoming increasingly commonplace, thanks to *Hubble Space Telescope* (*HST*) and the *Ultraviolet Imaging Telescope* (*UIT*) (e.g., Massey et al. 1996, Parker et al. 1998).

Spectroscopy resolves this issue neatly: a single classification spectrum provides an accurate assessment of the effective temperature and hence the BC. However, one issue that remains uncertain is to what extent the relationship of effective temperature to spectral classification depends upon metallicity. The classification of OB stars depends primarily on the relative strengths of different ionization states of the same element (He I to He II for O-type stars, Si III to Si IV for early B-type stars), and hence should be relatively metallicity independent. Indeed, fully blanketed, non-LTE models computed for solar metallicity result in effective temperatures that are less than 10% lower than those found from pure H and He models (Martins et al. 2002). With increased telescope size it is now possible to perform

TABLE 2A Evolution of massive stars at galactic metallicity (60–120 $\mathcal{M}_\odot$)

	Age (Myr)						
	0.0	**0.5**	**1.0**	**1.5**	**2.0**	**2.5**	**3.0**
120 $\mathcal{M}_\odot$ (τ_{ms} = 2.56 Myr)							
Log T_{eff}	4.727	4.706	4.683	4.648	4.635	4.420	—
M_{bol}	−10.9	−10.9	−11.0	−11.0	−11.1	−11.1	—
log g [cgs]	4.1	4.0	3.9	3.7	3.6	2.6	—
M_V	−6.2	−6.4	−6.6	−6.9	−7.0	−8.6	—
Sp Type	O3 V	O3 V	O4 III	O5.5 III	O5 If	WNL	—
85 $\mathcal{M}_\odot$ (τ_{ms} = 2.82 Myr)							
Log T_{eff}	4.705	4.691	4.673	4.645	4.577	4.449	—
M_{bol}	−10.3	−10.3	−10.4	−10.5	−10.6	−10.6	—
log g [cgs]	4.1	4.0	3.9	3.8	3.5	2.9	—
M_V	−5.7	−5.9	−6.1	−6.4	−6.9	−7.9	—
Sp Type	O3 V	O4 V	O4 III	O5.5 III	O7 If	B0 I	—
60 $\mathcal{M}_\odot$ (τ_{ms} = 3.45 Myr)							
Log T_{eff}	4.683	4.671	4.660	4.645	4.620	4.576	4.464
M_{bol}	−9.6	−9.6	−9.7	−9.8	−9.9	−10.0	−10.1
log g [cgs]	4.2	4.1	4.0	3.9	3.8	3.6	3.1
M_V	−5.2	−5.4	−5.5	−5.7	−5.9	−6.3	−7.2
Sp Type	O4 V	O5 V	O5 V	O5.5 III	O6.5 III	O7.5 If	B0 I

TABLE 2B Evolution of massive stars at galactic metallicity (20–40 $\mathcal{M}_\odot$)

	Age (Myr)							
	0.0	**1.0**	**2.0**	**3.0**	**4.0**	**5.0**	**6.0**	**8.0**
40 $\mathcal{M}_\odot$ (τ_{ms} = 4.30 Myr)								
Log T_{eff}	4.640	4.626	4.609	4.570	4.443	—	—	—
M_{bol}	−8.7	−8.8	−9.0	−9.1	−9.3	—	—	—
log g [cgs]	4.2	4.1	3.9	3.7	3.1	—	—	—
M_V	−4.6	−4.8	−5.1	−5.5	−6.6	—	—	—
Sp Type	O6 V	O6.5 V	O7 III	O8 III	B0.5 I	—	—	—
25 $\mathcal{M}_\odot$ (τ_{ms} = 6.40 Myr)								
Log T_{eff}	4.579	4.570	4.563	4.552	4.535	4.504	4.438	—
M_{bol}	−7.5	−7.6	−7.7	−7.8	−7.9	−8.1	−8.3	—
log g [cgs]	4.2	4.1	4.1	4.0	3.8	3.7	3.3	—
M_V	−3.8	−4.0	−4.1	−4.3	−4.6	−4.9	−5.6	—
Sp Type	O8 V	O8 V	O9 V	O9 V	O9.5 III	O9.5 III	B0.5 I	—
20 $\mathcal{M}_\odot$ (τ_{ms} = 8.14 Myr)								
Log T_{eff}	4.640	—	4.626	—	4.609	—	4.570	4.443
M_{bol}	−6.9	—	−7.0	—	−7.2	—	−7.4	−7.8
log g [cgs]	4.2	—	4.1	—	4.0	—	3.8	3.3
M_V	−3.5	—	−3.7	—	−4.0	—	−4.4	−5.3
Sp Type	O9.5 V	—	O9.5 V	—	B0 V	—	B0 III	B1 I

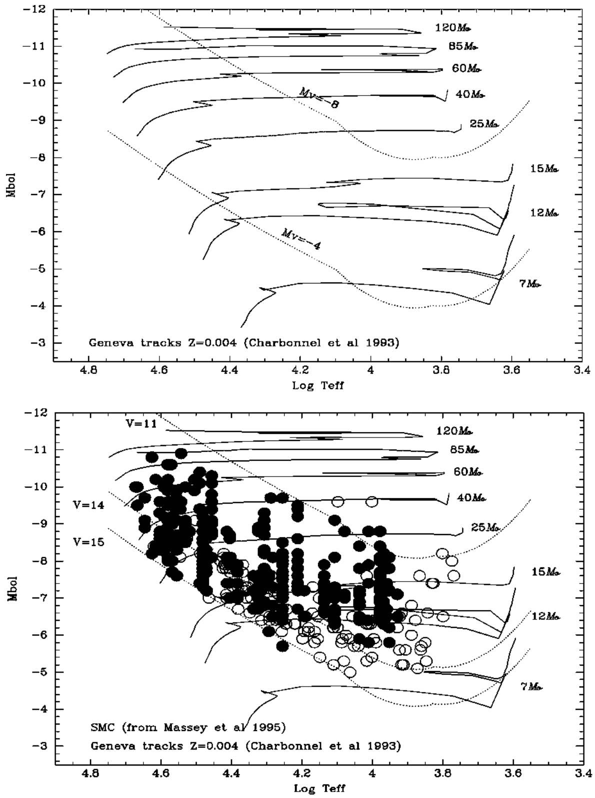

Figure 1 The $Z = 0.004$ evolutionary tracks of Charbonnel et al. (1993) (*solid lines*) are shown with dotted lines of constant absolute visual magnitude in the upper figure. An H-R diagram of the SMC is shown in the lower figure, with the data from Massey et al. (1995c).

a full quantitative analysis of OB stars in nearby galaxies using spectral synthesis: Puls et al. (1996) have studied stars in the Magellanic Clouds, and Muschielok et al. (1999), Monteverde et al. (2000), Smartt et al. (2001), and Bresolin et al. (2002a) have pushed this to a few stars in NGC 6822, M33, M31, and even beyond the Local Group to NGC 300, respectively. Because they are so much brighter visually, a full analysis of A-type supergiants is possible even at at distance of 7 Mpc (Bresolin et al. 2001).

Thus photometry, be it from the ground or from spacecraft, provides a first cut at identifying the luminous, hot star population. But, to go further requires spectroscopy. For this reason, the number of O-type stars is poorly determined even in well-studied regions of nearby systems.

2.2.1. MILKY WAY The closest O-type star is Zeta Ophiuchi, at a distance of 140 pc according to *Hipparcos*. Most O stars have been found via spectroscopy of stars in the Henry Draper (HD) catalog, or as part of studies of nearby clusters, typically at distances of 1–3 kpc. Garmany et al. (1982) compiled a catalog of 765 Galactic O-type stars, which they argued was complete to a distance of 2.5 kpc. Subsequent studies of young clusters, OB associations, and H II regions in the optical (summarized by Massey et al. 1995b, 2001) and the IR (Hanson & Conti 1995) have revealed many additional O-type stars, underlying the difficulty in achieving, and evaluating, completeness.

2.2.2. MAGELLANIC CLOUDS (MCs) Objective prism surveys carried out in the 1960s and 1970s identified many of the less crowded early-type stars in the LMC and SMC. Two catalogs still serve as the primary references: a list of 1822 LMC members by Rousseau et al. (1978) and a list of 524 SMC members of Azzopardi & Vigneau (1982). Neither survey went sufficiently deep to be complete for even lightly reddened O-type stars, and subsequent CCD studies of Magellanic Cloud OB associations revealed many previously unknown early-type stars (compare Massey et al. 1995c, 2000) whenever crowding was an issue—which is, after all, just where we expect to find massive stars. Nevertheless, these two catalogs remain useful for studies of the “field” populations of the MCs. These catalogs were updated with new slit spectral types for several hundred stars by Fitzpatrick & Garmany (1990) and Massey et al. (1995c) in order to study stellar evolution. Most recently, this work has been updated to include *UBVR* CCD measurements by Massey (2002a).

The OB associations of the LMC were identified by Lucke & Hodge (1970); those of the SMC were identified by Hodge (1985). The regions they selected have turned out to contain stars of very similar ages for the most part (see, for example, Massey et al. 2000), suggesting that the boundaries drawn are sensible. Efforts to identify OB associations by various automatic algorithms (e.g., Kontizas et al. 1999) are valuable in that subjectivity is reduced, but the results of such analysis need to be put to similar tests.

The advent of CCDs on telescopes in the mid-1980s ushered in new studies of massive stars in the Magellanic Clouds. Detailed CCD studies (both photometric and spectroscopic) of the stellar content of selected H II regions and OB associations were carried out from the European Southern Observatory (Stasinska et al. 1986, Testor & Lortet 1987, Heydari-Malayeri et al. 1987, Lortet & Testor 1988, Schild & Testor 1992, Testor et al. 1993, Testor & Niemela 1998), CTIO (Massey et al. 1989a,b; Parker et al. 1992; Parker 1993; Parker & Garmany 1993; Garmany et al. 1994; Oey & Massey 1995), and Las Campanas (Oey 1996). These studies provided a wealth of information that have helped further our understandings of massive star evolution, star formation, and the initial mass function as discussed further below.

The availability of very wide field CCD imagers have allowed several new photometric surveys of the Clouds. The Magellanic Clouds Photometric Survey of Zaritsky et al. (1999) goes very deep, and is providing important new data on the star-formation histories of the Clouds, but unfortunately saturates at magnitudes where massive stars dominate, as well as having calibration problems at *U*, which have only been partially addressed (compare Zaritsky et al. 2002, Massey 2002a).

The CCD survey by Massey (2002a) was aimed at rectifying these problems but does not quite cover all of the star-forming regions of the Clouds; its modest scale also causes problems in crowded regions. Parker et al. (1998) used images from *UIT* to identify hot, massive stars. A collaborative spectroscopic survey using all three of these resources is planned.

Because of its small field of view, but extremely high spatial resolution, *HST* imaging and spectroscopy has played a very important but "niche" role in massive star studies in the MCs. Walborn et al. (1999b) used a combination of *HST* imaging and spectroscopy to resolve several bright knots of stars. It has also proven very useful in identifying stars in otherwise unresolved high-excitation compact H II regions (see Heydari-Malayeri et al. 1999 and references therein).

One of the very interesting findings of *HST* has been what we have learned about the R136 cluster in the 30 Doradus region of the LMC. Once thought to be a single, 2000 $\mathcal{M}_{\odot}$ object, we know today that this is actually a very compact, extremely rich cluster of the most massive and luminous stars known (Campbell et al. 1992, Massey & Hunter 1998). The cluster contains more stars of type O3 (the earliest spectral type) than the total known everywhere else.

2.2.3. BEYOND THE MCs Wide-field photographic surveys helped identify many of the brighter stellar members in Local Group galaxies beyond the Magellanic Clouds; some of these remain the definitive reference to this day. Of particular note are the Kayser (1966) study of NGC 6822, the Humphreys & Sandage (1980) study of M33, and the Berkhuijsen et al. (1988) and Ivanov et al. (1993) studies of M31. OB associations were identified by means of multicolor plate material: for M31 by van den Bergh (1964), for M33 by Humphreys & Sandage (1980), for NGC 6822 by Hodge (1977), and for IC 1613 by Hodge (1978). Photoelectric photometry was hampered by crowding, but was carried out for NGC 6822 by van den Bergh & Humphreys (1979), and for a few stars in some of the studies cited above.

Such studies set the stage for the later CCD photometric studies of selected regions by Massey et al. (1986, 1995a), Hodge & Lee (1988), Hodge et al. (1988), Wilson (1992), Regan & Wilson (1993), Gallart et al. (1996), and Massey (1998b), among others. These studies targeted specific regions of these galaxies owing to the small size of the arrays (typically covering a few arcminutes squared) versus the large angular extent of a galaxy. One exception was the Herculean effort by Magnier et al. (1992), who combined 100 small field-of-view CCD frames to obtain *BVRI* photometry of a square degree region of M31. Observations with *UIT* were also used by Hill et al. (1995) and Massey et al. (1996) to identify hot stars in M31 and M33, respectively, from UV colors.

Once photometry was obtained, spectroscopy followed. The difficulty of course is that in the photographic era, only the brightest stars were within reach of even the largest telescopes (4- to 5-m in those days). In a landmark series of papers, Humphreys obtained spectra of the brightest stars in a number of Local Group galaxies (and beyond), and thus was able to place the visually brightest stars on the H-R diagram in a meaningful way, and make comparisons with the Milky Way and Magellanic Clouds (Humphreys 1978, 1979a; Humphreys & Davidson

1979; Humphreys 1979b, 1980a,b,c). These studies determined an upper luminosity boundary for stars in the H-R diagram, now known as the "Humphreys-Davidson" limit. For the hotter stars, there is a limit of decreasing luminosity with decreasing temperature, and an upper limit near $M_{\rm bol} \sim -10$ for stars cooler than 15,000°K. (This limit is now well understood, as described in Section 3.2.)

Nevertheless, the stars studied were primarily A- and B-type supergiants. Spectroscopic studies of the hotter and hence fainter but more luminous stars were carried out once more sensitive detectors became available. The first unambiguous classification spectra of O-type stars beyond the MCs were obtained by Massey et al. (1986), followed by Humphreys et al. (1990), with additional studies of OB stars by Herrero et al. (1994), Massey et al. (1995a), Monteverde et al. (1996), and Massey et al. (1996). These studies identified targets for UV *HST* studies of stellar winds, and found interesting evolved objects (Ofpe/WN9 stars, LBV candidates), but were too incomplete to produce useful H-R diagrams for any specific OB association, much less galaxy-wide. These detectors also allowed better studies of the A- and B-type supergiants (e.g., Herrero et al. 1994; Venn et al. 2000, 2001).

HST has played an important role in the study of massive stars in these nearby systems, primarily through optical and near-UV imaging (a good example is the recent study of the H II regions Hubble V and X in NGC 6822 by Bianchi et al. 2001), and in UV stellar wind studies (Hutchings et al. 1992; Bianchi et al. 1994, 1996; Urbaneja et al. 2002; Bresolin et al. 2002b). Magnier et al. (1997) used *HST* images to study the ages of OB associations in M31, although the results are somewhat contradicted by the results of the spectroscopic studies cited above, presumably because of the intrinsic uncertainties associated with using photometry alone in studies of young massive stars.

We are entering a new era of studies of the unevolved massive stars in Local Group galaxies. Wide-field CCD imagers on 4-m telescopes are allowing accurate photometry galaxy-wide. One such project is the Local Group Galaxies Survey, using the KPNO and CTIO 4-m telescopes to image each of the galaxies listed in Table 1 in *UBVRI*, Hα, [O III], and [S II], to a (broadband) depth of 25 (S/N$\sim$3) and a spatial resolution of $\sim$1 arcsecond. Spectroscopy needs to follow.

2.3. Star Formation, Initial Mass Functions, and Stellar Evolution

What have the above studies taught us? For clusters and associations we gain an understanding of how star formation proceeds as a function of mass, and to what degree it is coeval. By appealing to theoretical evolutionary tracks, we can construct the present day mass function (PDMF), and if the star-formation history of the region is understood, we can deduce the initial mass function (IMF), the distribution of stars as a function of mass at birth. We can see if there is an observed "upper limit" to the mass of a star, and whether this (and the IMF slope) varies with metallicity by comparing results for different regions. And, we can use these studies to test our stellar evolutionary models by comparing the distribution of stars on the H-R diagram with that expected from the models.

2.3.1. CLUSTERS Herbig (1962a,b) was the first to suggest that star formation of low and intermediate mass stars might continue over a prolonged time, followed by the formation of high-mass stars, which halts all star formation caused by (he supposed) the introduction of turbulence via the heating and ionization of the gas. His argument was based on the contradiction in the implied ages owing to the presence of late-type and early-type main-sequence stars in the same clusters (e.g., Orion, NGC 2264, and NGC 6530), given the long contraction times of lower-mass stars compared to the ages of the high-mass stars (see also Iben & Talbot 1966). The Herbst & Miller (1982) study of NGC 3293 supported the view that when stars in a cluster form, the lower- and intermediate-mass stars form first, with the process continuing gradually until the high-mass stars form. The age of a cluster, and the degree of coevality, thus likely depend upon the mass range under discussion.

Recent studies of clusters and associations containing massive stars for the most part support this view, and in particular that the massive stars may form over a very short interval of time, with $\Delta\tau < 1$–2 Myr. The study of NGC 6611 by Hillenbrand et al. (1993) was the first to establish such a tight constraint, but such is the norm rather than the exception (Massey 1998a; Massey & Hunter 1998; Massey et al. 2000, 2001). However, the simple view of an abrupt termination of star formation is not supported in all cases. For NGC 6611, Hillenbrand et al. (1993) state that "... the formation of O stars neither ushered in nor concluded the star-formation process ... ." The O-type stars suggest an age of 2 Myr, but there are also pre-main-sequence stars with ages as young as 0.25 Myr.

Massey (1998a) argues that it is not the actual formation of high-mass stars that stops further star formation, but rather what happens after they have formed. In a cluster whose massive stars formed ~5 Myr ago, any stars above $>50\ \mathcal{M}_\odot$ will have exploded as supernovae, disrupting the surrounding gas. But, a cluster whose massive stars are only 2 Myr old may not have experienced any supernovae. Instead, it is through the actions of their strong stellar winds that massive stars could disrupt further star formation. (This explanation was not available to Herbig, as O-star stellar winds were yet to be discovered.) The winds of a single $100\mathcal{M}_\odot$ contribute $\sim 2 \times 10^{51}$ ergs during the main sequence, and another 5×10^{50} ergs during the WR phase (Abbott 1982b). Compare this to the $\sim 10^{51}$ ergs imparted to the ISM by a Type II supernova. Thus in very rich clusters, star formation may be halted by main-sequence stellar winds after a short period of time (1 Myr?), while in less rich clusters this disruption may not occur until a massive star reaches the WR stage (2.5–3 Myr). In even less rich clusters, nothing may happen until the most massive star explodes (>3 Myr).

For a while, it appeared that the star formation history of the R136 cluster in the LMC challenged this paradigm. The presence of WRs suggested an age of ~4 Myr, and yet there were intermediate-mass stars with ages that ranged from >4 Myr to as young as 1–2 Myr (Hunter et al. 1995). However, de Koter et al. (1997) and Massey & Hunter (1998) independently found that the "Wolf-Rayet stars" were not the usual sort of evolved objects associated with that label, but were instead simply extraordinarily luminous stars that therefore had extreme stellar winds that mimicked some, but not all, of the spectroscopic properties of true WRs. In other words, these were "Of stars on steroids." (Similar stars are also seen in the Galactic cluster NGC 3603; see comparison in Massey & Hunter 1998.) Massey & Hunter (1998) found that R136's brighter stars were predominantly of spectral type O3, the hottest, most luminous, and youngest phase of a massive star; the ages of the massive star population is 1–2 Myr, with the greatest uncertainty in the effective temperatures of the O3 class. Thus a simple story of star formation can now be told for R136, in which the intermediate-mass stars began forming some 6 Myr ago and continued up to the time that the high-mass stars formed 1–2 Myr ago. In a rich, dense cluster like R136, the stellar winds would have a drastic effect on the ISM, and indeed Walborn et al. (1999a) has suggested that the R136 cluster has triggered a second generation of star formation in the greater 30 Doradus region.

The determination of the IMF in a cluster containing massive stars is straightforward in principle, at least to the extent that masses can be inferred from the evolutionary models: the slope of the IMF is identical to that of the PDMF for a coeval region, as long as one realizes that the highest-mass bin has been selectively depopulated by stellar deaths for clusters older than a few million years. In practice, the actual values of the slopes of the IMF depend strongly on how reddening is treated, and what transformations are used from spectral types and photometry to effective temperatures. When we compare the slopes found for clusters analyzed in a consistent manner, we find that there is no difference for the SMC, LMC, and Milky Way, despite the factor of 4 in metallicities (Table 1). This is shown in Figure 2, where the quantity Γ is the slope of the IMF (Scalo 1986), with $\Gamma = -1.35$ corresponding to that found by Salpeter (1955) for stars in the solar neighborhood.

Care should be taken not to overinterpret the error bars in such data. Usually these are no more than the formal 1σ errors of the fits to the mass function. Although typical errors are 0.1–0.2, minor differences in the analysis (such as how reddening is treated) can easily change the derived slope by several times these values. For instance, Hillenbrand et al. (1993) find an IMF slope of $\Gamma = -1.1 \pm 0.1$ for their study of NGC 6611; a reanalysis of the same data by Massey et al. (1995c) yields $\Gamma = -0.7 \pm 0.2$. It is probably differences such as these that account for much, if not all, of the variations in the IMF slope discussed by Scalo (1998). Similarly the significance of $\Gamma = -0.7$ found by Figer et al. (1999) for massive stars in the Arches cluster is unclear, given the lack of spectroscopy and the need to rely purely on a K-band luminosity function for deriving this result.

The IMF slopes shown in Figure 2 are all consistent with the Salpeter (1955) slope found for stars in the solar neighborhood. However, the absolute value of Γ derived from these data cannot be taken purely at face value: there has been, for instance, no correction for binarity. What these findings do show, instead, is that the IMFs are not significantly different in a relative sense.

What then about an upper mass cutoff? The highest-mass stars seen in the OB associations of the SMC, LMC, and Milky Way data are all about the same, 85–120 $\mathcal{M}_\odot$ (Massey et al. 1995c). Here R136 teaches us an important lesson. The masses of the highest-mass stars seen in this cluster are 135–155 $\mathcal{M}_\odot$, depending

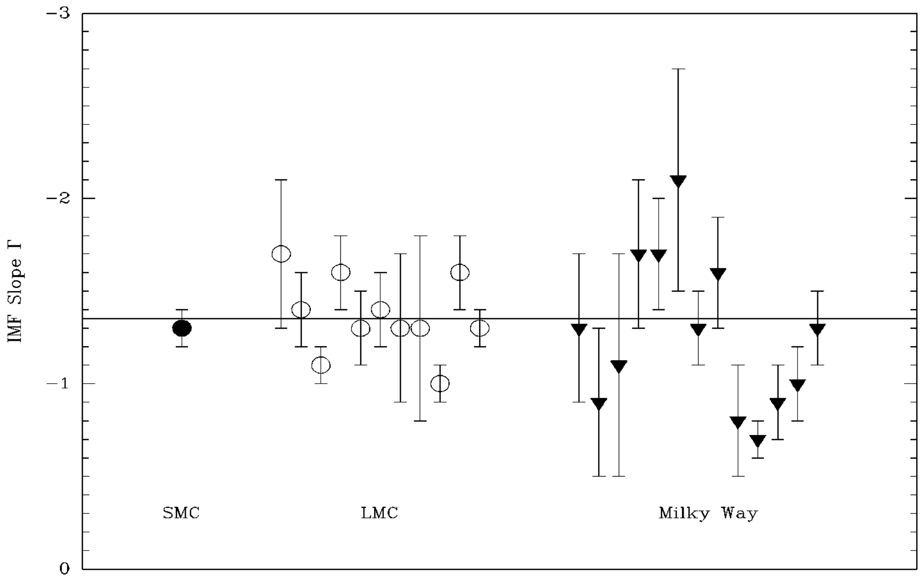

Figure 2 The initial mass function slopes Γ are shown for OB associations and clusters analyzed in a uniform manner; the data are from Table 3 of Massey (1998a) updated to include R136 (Massey & Hunter 1998) and h and χ Persei (Slesnick et al. 2002). The solid line at $\Gamma = -1.35$ indicates a Salpeter (1955) slope.

upon what is assumed for the effective temperature scale for the hottest stars. Yet in fact the number of these very massive stars is just what one would expect from extrapolating the IMF slopes from that of the intermediate-mass stars (Massey & Hunter 1998). The "upper mass limits" observed in these more sparsely populated OB associations are also consistent with the extrapolation of the IMF to higher masses—these limits turn out to have been statistical, rather than physical, and just what happens when the IMF peters down to a single star. Whatever it is that limits the ultimate mass of a star, we have yet to encounter it in nature.[1]

[1]Theory offers us only modest guidance in what the maximum stellar mass allowed by nature is and what the limiting factor may be. An excellent review may be found in Appenzeller (1987), who notes that Eddington (1926) was the first to propose that stars more massive than some amount would be pulsationally unstable, and should blow off their outer layers, thus limiting their mass. Early estimates of this limit were as low as 60 $\mathcal{M}_\odot$ (Schwarzschild & Harm 1959). Modern estimates, however, place this limit as high as 440 $\mathcal{M}_\odot$ (Klapp et al. 1987), although this is still based upon the same classical perturbation linearization methods used by Eddington. Recent "nonlinear" analysis suggests that the mass loss from such instabilities would only be comparable to the mass loss of radiatively driven stellar winds (Appenzeller 1987). Similarly, it was once thought that radiation pressure acting on grains would limit how large a star could form, but we now understand that disks play an important role in the formation of stars, and there may be sufficient shielding by the inner part of the disk to mitigate the effects of radiation pressure.

Clusters that contain evolved massive stars can set limits on the masses of their progenitors, if they are sufficiently coeval. Regions in the Magellanic Clouds and the Milky Way have been studied for this purpose by Massey et al. (2000, 2001). The results are discussed in Section 3. Work is still progressing on extending such studies to the more distant members of the Local Group.

2.3.2. FIELD Although the coeval populations of OB associations and clusters have much to teach us about star formation and the IMF, it is the mixed-age populations that have provided the most critical tests of stellar evolution theory. If a sufficiently large region of a galaxy is observed to have a heterogeneous mix of ages, then the number of stars of each mass can be compared to stellar evolutionary isochrones in order to see if the numbers are consistent with that predicted by theory.

Fitzpatrick & Garmany (1990) used existing photometric and spectroscopic data of the LMC to investigate the distribution of stars in the resulting mixed-age H-R diagram, and compared these with the extant evolutionary tracks. Their data are shown in Figure 3. They recognized that the existing catalogs were grossly incomplete for the younger massive stars (for the reasons described above); instead, the big result of their study was at cooler spectral types, where an abrupt "ledge" was found. Years earlier Meylan & Maeder (1982) had found a serious discrepancy between the evolutionary models and the number of A-type supergiants in clusters,

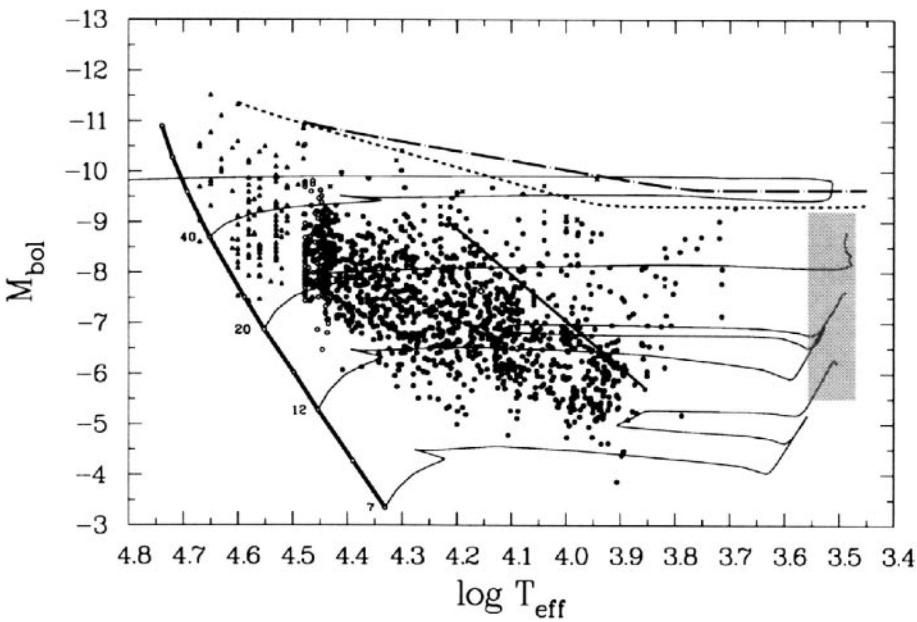

Figure 3 The H-R diagram of the LMC given by Fitzpatrick & Garmany (1990), and used with permission. The absence of young massive stars is a selection effect, corrected for in later work (e.g., Massey et al. 1995c). The "ledge" is shown by a diagonal solid line in mid-figure.

as if the main sequence was much wider than the models could possibly allow. Instead, Fitzpatrick & Garmany (1990) proposed that the large number of stars found between the nominal end of the main sequence and the ledge was in the blue region of the H-R diagram for the second time: that these were stars that had already become RSGs and that their numbers could best be explained by evolutionary models that contain "blue loops"—excursions back to blue after the red supergiant phase. This had followed on the heels of the supernova explosion of the star Sk $-$ 69°202, the otherwise unremarkable B3 I star that became SN1987A, an event which suggested that blue loops were an evolutionary fact (Saio et al. 1988). Recent work by Salasnich et al. (1999) suggests that if a more realistic high mass loss is included during the red supergiant phase, the length of the subsequent blue loop phase is enhanced, and that there is good agreement between the models and the observed distribution in the H-R diagram, although observational data supporting this claim have yet to be shown.

Massey et al. (1995c) investigated a different part of this issue, namely the actual distribution of stars on the nominal main sequence. They had added many more spectral types to the catalog constructed by Fitzpatrick & Garmany (1990) for the LMC, and also obtained new data for the SMC. They attempted to carefully correct for incompleteness for the young massive stars by deep imaging and spectroscopy in a few fields, the results of which were used to provide a statistical correction for the catalogs. What they found was very good agreement between the predictions of the (normal mass loss) evolutionary tracks of the Geneva group (Schaller et al. 1992, Schaerer et al. 1993) and observations.

A curious fact emerged from this study, namely that the IMF appeared to be quite different for massive stars found in the field than in clusters. Rather than a Salpeter $\Gamma \sim -1.3$ slope, the massive stars that are found well away from clusters and association show a much steeper slope, with $\Gamma \sim -4$. Yet, stars as massive as those found within a cluster and association are found in the field. Massey (1998a) shows images of the regions surrounding four O3 stars in the LMC; two are in "normal" environments (OB associations) and two have been selected as being well away from other known massive stars. The O3 stage is very short ($<$1 Myr), and it is hard to see how such a star could have traveled far from its birthplace in such a short time. A radial velocity study (determining one component of the space motion) would both help answer the source of these stars and presumably constrain explanations for the peculiar IMF.

Beyond the MCs sufficient spectroscopy is still lacking, although follow-ups to the photometric survey mentioned above are planned. Instead, studies have intercompared luminosity functions. One of the most ambitious of these is that of Freedman (1985), who attempted to constrain the IMF from observations of luminosity functions in a variety of late-type galaxies in the Local Group and beyond. She was careful to include only the bluest stars, in order to eliminate the effects of foreground contamination. However, as Figure 4 shows, luminosity functions based upon such a restriction are still highly insensitive to the IMF.

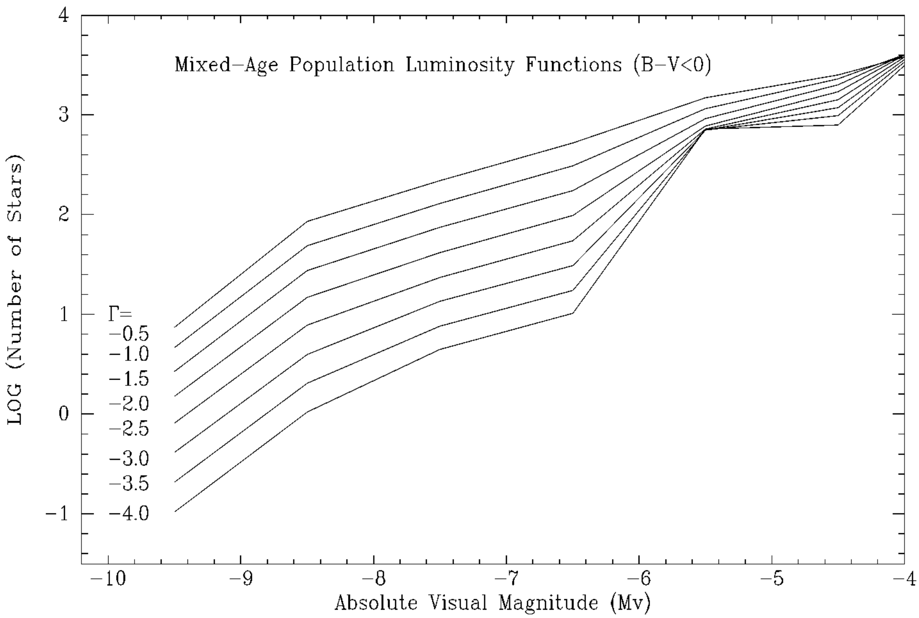

Figure 4 The lack of sensitivity of a V-band luminosity function to the IMF slope Γ is shown; the calculations were based upon the evolutionary models of Schaller et al. (1992) and have been normalized to 10,000 stars with $M_V < -3$. This figure is based upon work in Massey (1998a).

2.4. Mass-Loss Rates

The presence of P Cygni profiles in UV resonance lines of normal OB supergiants demonstrated the existence of stellar winds in these stars (Morton 1967); subsequent studies with the *Copernicus* satellite revealed that such mass-loss was ubiquitous among hot, luminous stars (Snow & Morton 1976). The stellar winds affect even the optical spectra in extreme cases, resulting in Hα emission (Conti & Frost 1977) and velocity differences among the photospheric lines (Hutchings 1968). An excellent early review is given by Conti (1978). Measurement of mass-loss rates are model-dependent, and generally determined from high dispersion UV spectra and/or Hα profile modeling. Sophisticated modeling was used to determine mass-loss rates for a number of Galactic and Magellanic Cloud stars by Puls et al. (1996). Typical mass-loss rates are 0.2–20 $\times$ 10^{-6} $M_\odot$ yr^{-1}. Given the effects of metallicity on the evolution of massive stars, an accurate parameterization of mass loss across the upper-end of the HRD is clearly of interest.

Such mass loss is driven by radiation pressure: the high luminosities of these stars result in momentum transfer through absorption in the resonance lines of highly ionized metal lines. Abbott (1982a) provides one of the clearest discussions of radiatively driven stellar winds. One expects that the mass loss rate $\dot{M}$ will scale with luminosity as

$$\dot{M} \sim L^{1/\alpha} M_{\rm eff}^{(\alpha-1)/\alpha}$$

(Puls et al. 1996, Lamers & Cassinelli 1999), where $M_{\rm eff}$ is the effective mass (the mass modified by radiation pressure), and α is the power-law exponent of the line-strength distribution function of the thousands of lines driving the stellar wind (see also Kudritzki & Puls 2000). However, this deceptively simple expression hides much of the complication, since α depends upon the effective temperature and Z in a manner that can only be determined by detailed modeling (Kudritzki & Puls 2000, Puls et al. 2000). For O-type stars, the value of α is approximately 0.6 (Pauldrach et al. 1986, Kudritzki & Puls 2000), suggesting a luminosity exponent $\sim$1.7, similar to what is observed.

Many efforts have been made to provide an explicit parameterization of mass-loss rates with stellar properties. The empirical relationship of de Jager et al. (1988) has commonly been adopted:

$$\log \dot{M} = -8.158 + 1.769 \log L - 1.676 \log T_{\rm eff}$$

where $T_{\rm eff}$ is the effective temperature in degrees K, and L is the luminosity in solar units. (Note that the opacity sources decrease with higher temperatures, and hence, at a given bolometric luminosity an A-type supergiant will have a higher mass-loss rate than an O-type star; see Section 3.2.) Extension to non-Galactic metallicities has typically been done by scaling by $(Z/Z_\odot)^{0.5}$ following Kudritzki et al. (1989), e.g., Schaller et al. (1992), although others cite a stronger dependence (Lamers & Cassinelli 1996, Vink et al. 2001). An independent mass-loss relationship is usually adopted for Wolf-Rayet stars, as one would expect little dependence of the mass-loss rates on initial metallicity for Wolf-Rayet stars, since these stars have generated their own enhanced chemical abundances (Abbott & Conti 1987), although Crowther et al. (2002b) argue that iron is important in accelerating the stellar winds.

Vink et al. (2001) provides a theoretical prediction for the mass-loss rates as a function of luminosity, mass, terminal velocity, effective temperature, and metallicity. In Figure 5 the mass-loss rates predicted both by their work (filled symbols) and those predicted by the de Jager et al. (1988) relation, scaled by $(Z/Z_0)^{0.5}$ (open symbols), are compared with the observed rates of Puls et al. (1996). We see that both do a reasonably good job of matching the data, although there remains a very significant scatter ($\sim$0.3 dex). Some of this may reflect uncertainty in the observed rates.

Puls et al. (2000), Vink et al. (2001), and Kudritzki (2002) have recently investigated the theoretical dependence of mass-loss rates on metallicity. Puls et al. (2000) argue that Galactic mass-loss rates should scale approximately as $(Z/Z_\odot)^{(1-\alpha)/\alpha}$, with the complication that α will depend upon Z as well as effective temperature. For $\alpha = 0.6$, the scaling factor would be $(Z/Z_\odot)^{0.7}$. Kudritzki (2002) finds that

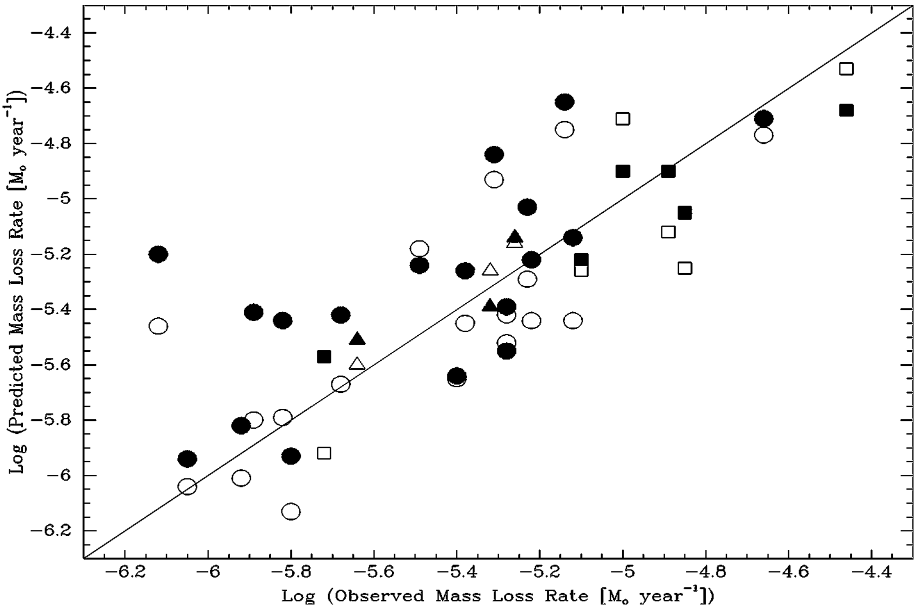

Figure 5 A comparison is shown between the observed mass-loss rates determined by Puls et al. (1996) and that predicted by the empirical fit of de Jager et al. (1988) (*open symbols*) and the theoretical formalism of Vink et al. (2001) (*filled symbols*). Circles denote Galactic stars, squares denote LMC stars, and triangles denote SMC stars.

the situation is more complicated than a simple power-law scaling, with a threshold effect at low Z. However, over the metallicity range usually considered (i.e., SMC to solar neighborhood), a scaling with $(Z/Z_\odot)^{0.5}$ turns out to be a good approximation, at least for O-type stars (see his table 2). However, Vink et al. (2001) concludes that the mass-loss rates scale as $(Z/Z_\odot)^{0.7}$ when one includes the dependence of the terminal velocity with metallicity.

Further observational checks would be useful on mass-loss rates, particularly at higher metallicities. So far, mass-loss rates have been derived in a consistent manner only for Milky Way, LMC, and SMC stars, which cover a range of metallicity of a factor of 3.7 (Table 1). This could be pushed to a factor of ~15 by studies of massive stars throughout the Local Group, and probe regions that are higher in metallicity than the solar neighborhood, e.g., in the Andromeda Galaxy. Fledgling efforts in this direction have been taken by Bianchi et al. (1994, 1996), Smartt et al. (2001), Urbaneja et al. (2002), and even beyond the Local Group by Bresolin et al. (2002a). (See also Prinja & Crowther 1998, who reanalyze much of these data, but stop short of deriving mass-loss rates.) The UV observations needed to measure the terminal velocities are well within the reach of *HST*, and the optical data (needed to determine the other stellar parameters) are obtainable with 8-m class ground-based spectroscopy.

3. EVOLVED MASSIVE STARS

3.1. The Conti Scenario

Conti (1976) was the first to propose that the mass loss that characterizes massive stars could explain the existence of the peculiar Wolf-Rayet stars (WRs), objects whose spectra are marked by strong, broad emission lines of helium and nitrogen (WN type), or of helium, carbon, and oxygen (WC types), with little or no sign of hydrogen in their spectra. In what has become known as the Conti scenario, a massive O-type star loses a significant amount of mass via stellar winds, revealing first the CNO-burning products at its surface, and subsequently the He-burning products. These two stages are spectroscopically identified with the WN and WC phase. Such stars would be over-luminous for their mass (because of mass loss), in accord with the observations of WR stars in binary systems. Their strong, broad emission lines are the product of an expanding, extended atmosphere.

Extremely luminous, unstable stars (near their Eddington limit, where radiation pressure balances gravity) are known: the so-called luminous blue variables (LBVs). S Dor in the LMC and η Car in the Milky Way are perhaps the best known examples. These stars show large visual brightness variations coupled to spectral changes, with episodic bouts of extreme mass loss. In addition, luminous red supergiants (RSGs) are known, corresponding to (initial) masses of $\lesssim 40\,\mathcal{M}_\odot$. So, in one version of this Conti scenario we would have:

$$
\begin{aligned}
m > 85\,\mathcal{M}_\odot&: \mathrm{O} \longrightarrow \mathrm{LBV} \longrightarrow \mathrm{WN} \longrightarrow \mathrm{WC} \longrightarrow \mathrm{SN}\\
40 > m > 85\,\mathcal{M}_\odot&: \mathrm{O} \longrightarrow \mathrm{WN} \longrightarrow \mathrm{WC} \longrightarrow \mathrm{SN}\\
25 > m > 40\,\mathcal{M}_\odot&: \mathrm{O} \longrightarrow \mathrm{RSG} \longrightarrow \mathrm{WN} \longrightarrow \mathrm{WC} \longrightarrow \mathrm{SN}\\
20 > m > 25\,\mathcal{M}_\odot&: \mathrm{O} \longrightarrow \mathrm{RSG} \longrightarrow \mathrm{WN} \longrightarrow \mathrm{SN}\\
10 > m > 20\,\mathcal{M}_\odot&: \mathrm{OB} \longrightarrow \mathrm{RSG} \longrightarrow \mathrm{BSG} \longrightarrow \mathrm{SN}.
\end{aligned}
$$

The mass ranges shown are meant only to be illustrative; even if the overall picture is correct, the mass ranges involved will certainly be a function of metallicity.

Recent work has attempted to determine these mass ranges observationally using the turn-off masses in clusters containing evolved massive stars. This is a classical technique, first used by Sandage (1953) to determine the progenitor masses of RR Lyra stars in globular clusters, and subsequently by Anthony-Twarog (1982) to determine the progenitor masses of white dwarfs in intermediate-age open clusters. To the degree that a cluster is coeval, the mass of the highest-mass stars still on the main sequence should be just slightly less than the mass of the progenitor of any evolved objects in the cluster. However, it is one thing to do this for a cluster with ages of 10 Gyr or even 40–70 Myr, and quite another to do this for a region with an age of only a few Myr. Still, a high degree of coevality has been found for many young, rich regions (Section 2.3.1), and the very data one needs to determine masses also provide a direct measure of coevality. This method was applied to Wolf-Rayet stars by Schild & Maeder (1984) and Humphreys et al.

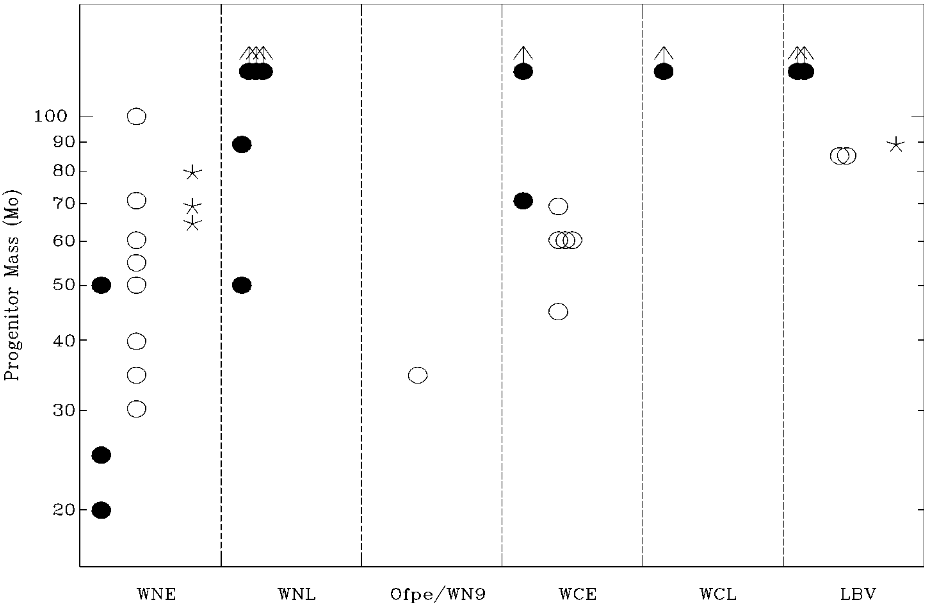

Figure 6 The progenitor masses are shown for various evolved massive stars for the Milky Way (*filled circled*), LMC (*open circles*), and SMC (*asterisks*). These data are from Massey et al. (2001).

(1985) using data on Milky Way clusters from the literature. New observations have been carried out by Massey et al. (2000, 2001), who studied a number of regions in the MCs and Milky Way. The results are shown in Figure 6.

According to these data, Milky Way WN stars come from stars with masses as low as $\sim 20\,\mathcal{M}_\odot$. The higher-mass stars that become WNs in the Milky Way are late-type WNs (WNLs), and the lower-mass ones are early-type WNs (WNEs). In the LMC the minimum mass for becoming a WN is somewhat higher, as we expect, and in the SMC WRs come from only the highest-mass stars. Similarly, the classical LBVs come from the highest-mass stars, consistent with the picture painted above.

3.2. The Eddington Limit and Luminous Blue Variables

Hubble & Sandage (1953) discussed five irregular variables in M31 and M33 that were, at times, among the brightest resolved stellar objects in these galaxies. Photographic plates provided long-term photometry extending back to 1916, which revealed episodic visual brightenings of several magnitudes. The stars were extremely luminous and blue; they showed spectra that were of intermediate F-type (during maxima). A footnote suggested that the LMC star S Doradus might also be the same type of object. The connection to Galactic stars η Carinae and P Cygni came later. Conti (1984) coined the term luminous blue variable, or LBV, to describe these objects.

There has been a great deal written of what constitutes a true LBV. Whereas many luminous blue stars show minor photometric variability (such as B[e] stars), there is a consensus that to be a true LBV, the variability must manifest itself in an outburst, in which the star increases in (visual) brightness by 1–2 mag, coupled to an outburst of mass loss, during which the star also undergoes a significant change in spectral appearance (Bohannan 1997, Conti 1997). These outbursts happen on the timescale of every few decades or, in some cases, centuries, making a complete census a very long-term undertaking.

Lamers & Fitzpatrick (1988) explored the relationship between the LBVs, the observed upper luminosity limit in the H-R diagram (the "Humphreys-Davidson" limit from Humphreys & Davidson 1979, 1984; Garmany et al. 1987), and the Eddington limit, where radiation pressure balances gravity. The observed luminosity limit decreases as a function of effective luminosity for stars with $T_{\rm eff} \gtrsim 10000^\circ$K, after which it is nearly constant at $\log L/L_\odot \sim 5.7$ ($M_{\rm bol} \sim -9.5$). During quiescence, the classical LBVs are found just to the hot side of the diagonal part of the upper luminosity limit, cohabiting a band of width $\Delta \log T_{\rm eff} = 0.3$ with normal blue supergiants, although LBVs display a mass-loss rate 3–10 times higher.

Why does this upper luminosity limit exist? It implies that stars more massive than $\sim$40 $\mathcal{M}_\odot$ do not evolve to the RSG phase but that something stops them dead in their (evolutionary) tracks, forcing the stars once again blue-wards. Enhanced internal mixing could do so (Maeder 1982), but it is highly tempting to interpret the Humphreys-Davidson limit in terms of the Eddington limit: that the ultimate luminosity of a star is set by simple physics (Humphreys & Davidson 1984, Appenzeller 1986, Lamers 1986). However, the classical Eddington limit (in which radiation pressure comes from scattering from free electrons) sets a limit on the stellar luminosity that is nearly 10 times higher than the Humphreys-Davidson line. Lamers & Fitzpatrick (1988) used model atmospheres to explore where radiation pressure and gravity are balanced if the full effects of metal line opacities are included. They found excellent agreement between the observed luminosity limit and the modified (or atmospheric) Eddington limit. Rather than the Eddington limit being at a constant luminosity (as would be expected for electron scattering), the atmospheric Eddington limit is a trough: radiation pressure increases with decreasing temperature, to a minimum at about 10,000°K. The location of this trough changes slightly with metallicity (Lamers & Noordhoek 1993, Lamers 1997). The work was extended to improved models by Pauldrach & Puls (1990) and Lamers (1997), with the results shown in Figure 7.

When a massive star evolves to cooler $T_{\rm eff}$, it does so at nearly constant $M_{\rm bol}$ (see Figure 1). The star's surface gravity decreases as its radius increases, while radiation pressure increases because the photospheric flux-mean opacity increases. For stars with initial masses $\gtrsim$40 $\mathcal{M}_\odot$, the star approaches its (modified) Eddington limit, and the mass-loss rate increases many-fold. However, this does not disrupt the star, as would be expected for stars exceeding the classical Eddington limit, as it is only in the outer layers of the star that line opacity is effective in transferring momentum from the radiation field to matter. This enhanced mass loss sets a natural limit to the upper luminosities of red supergiants.

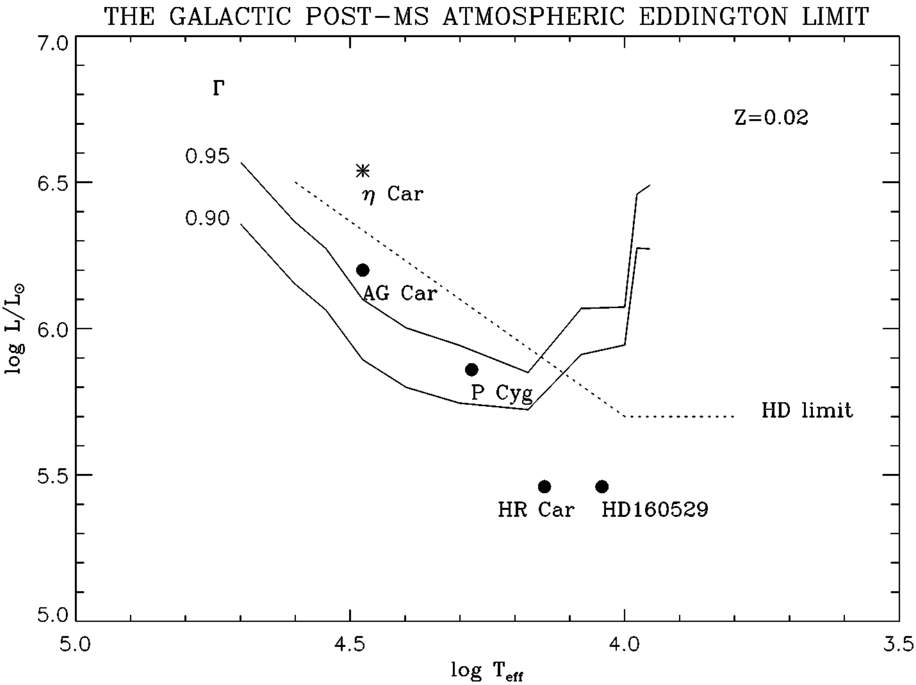

Figure 7 The observed luminosity limit for hot stars (*dotted line*) (Humphreys & Davidson 1979, 1984; Garmany et al. 1987) is well matched by the modified Eddington limit (*solid curve*) determined from model atmosphere calculations by Lamers (1997). Γ represents the fraction of the ratio of the force of radiation pressure to gravity. The location of a few well-known LBVs are shown. This figure is from Lamers (1997) and is used by permission.

We would thus expect LBVs to be located directly to the left of the observed luminosity limit, unstable because they are near their Eddington limits. The longer they remain LBVs the more unstable they become, as their bolometric luminosity remains roughly constant while their surface gravities decrease because of the high mass-loss rates ($10^{-4.5}\mathcal{M}_{\odot}$ year^{-1}). Thus stars of different ages may be in the same part of the H-R diagram, but with drastically different properties, explaining how extreme LBVs and relatively normal blue supergiants can coexist in the same part of the H-R diagram. During the occasional eruptions (triggered by some poorly understood mechanism; see Humphreys & Davidson 1994), their mass-loss rates increase to 10^{-2} or even $10^{-1}\mathcal{M}_{\odot}$ year^{-1}. Their cool, F-type spectral types are a consequence of the stellar wind turning opaque (Davidson 1987), and its interpretation in terms of an effective temperature is misleading.

There are observational challenges to confirming this picture. For one thing, the location of LBVs in the H-R diagram are not that well determined, particularly the bolometric luminosity. We know the $M_{\rm bol}$ relatively well for η Car because it is surrounded by circumstellar material, which has reprocessed most of the far

UV into the IR, where it can be directly measured. Distances for many Galactic LBVs are uncertain if they are not in an obvious group of early-type stars, as is η Car. For instance, the LBV AG Car was for a time considered an example of a low-luminosity LBV ($M_{\rm bol} \sim -9$), well below the observed luminosity limit, but subsequent work placed the distance at greater values and $M_{\rm bol} \sim -10$ to -11 (Humphreys 1989, Hoekzema et al. 1992), at or near its Eddington limit.

A second complication in testing this picture observationally is that LBVs only rarely undergo eruptions—some of the Galactic LBVs are known because of extreme events that took place on historical timescales, such as η Car, which underwent a major eruption in 1830, and P Cyg, whose last major eruption was in 1655. Had these stars been located in the Magellanic Clouds, would we recognize them as LBVs today? These stars would be noteworthy not because of their extreme variability, but because of their extreme luminosities.

Given this, it is difficult to know how to recognize LBVs. Is the star VI Cyg No. 12 an incipient LBV? It is extremely luminous (visually it is the most luminous star in the Milky Way!) and is surrounded by ejecta, as revealed by its high extinction compared to other stars in the Cyg OB2 association. The star shows modest spectral and photometric variability (Massey & Thompson 1991), although it hasn't yet shown the 1–2 mag variations that might occur every few decades or longer. Its position is well above the observed upper luminosity limit, as shown by Massey & Thompson (1991), who estimate its bolometric magnitude at ~ -11.

Candidate LBVs have generally been identified on the basis of single-epoch spectroscopy that shows some similarities to known LBVs, such as Hα stars found in M33 and M31 by Neese et al. (1991) and King et al. (1998), or the Ofpe/WN9 stars found by Massey et al. (1996). As Bohannan (1997) notes, a spectral similarity to known LBVs is not sufficient to consider a star an LBV. However, spectral variability is sometimes confirmed, such as R85 in the LMC (Massey et al. 2000) or UIT 247 (Massey et al. 1996, Monteverde et al. 1996). Humphreys & Davidson (1994), Bohannan (1997), and Parker (1997) all provide lists of confirmed and suspected LBVs in the Milky Way and nearby galaxies. When one recalls the extensive time base needed to identify the original five Hubble-Sandage variables (one of which, Var A, is no longer considered an LBV, as it developed an M-type spectrum in the 1980s, unprecedented for an LBV; see Humphreys 1989), one realizes that our knowledge of LBVs in the Local Group is very incomplete and may require centuries to have firm statistics.

3.3. Red Supergiants

If our current understanding of massive star evolution is correct, stars with initial masses $\lesssim 40\,\mathcal{M}_\odot$ become red supergiants after leaving the main sequence. However, there is a significant problem with all of the current evolutionary models, which is illustrated in Figure 8: none of the models extend to cool enough temperatures at sufficiently high mass. The reason for this is likely related to how the mixing-length is dealt with. (See figure 9 of Maeder & Meynet 1987).

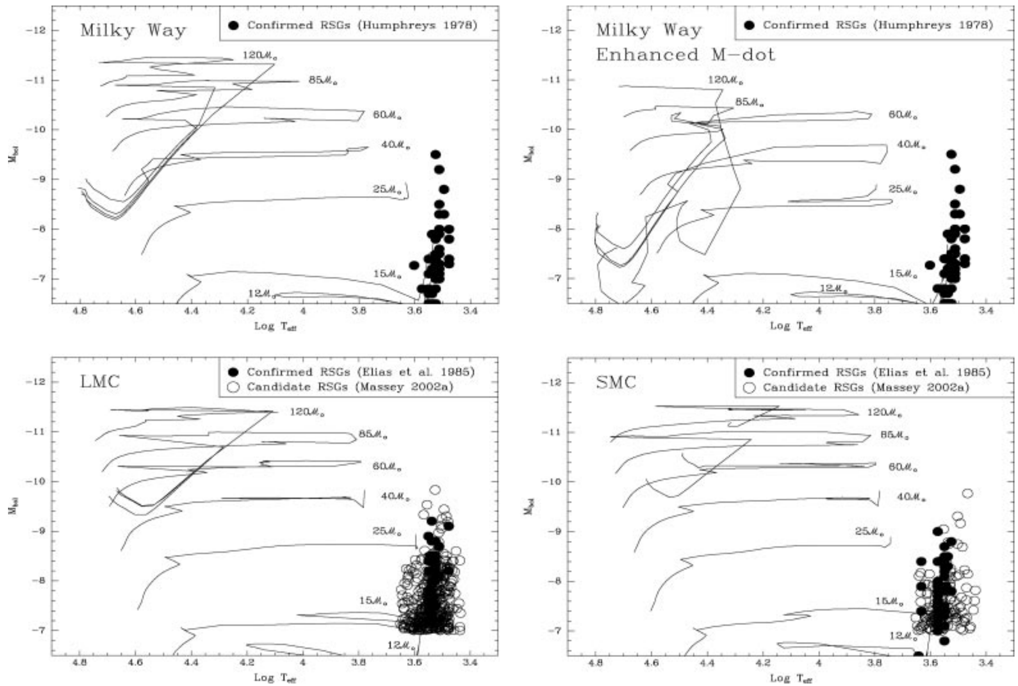

Figure 8 None of the evolutionary tracks extend to sufficiently cool temperatures to produce luminous red supergiants. The problem is true for the Padova evolutionary tracks (not shown), as well as for the four sets of Geneva models shown here. The two at the top are computed for $Z = 0.02$ metallcity appropriate to the Milky Way, using normal mass-loss rates (*upper left*) from Schaller et al. (1992) and enhanced mass-loss rates (*upper right*) from Meynet et al. (1994). The problem also exists at low metallicity, as shown with the normal mass-loss rates for $Z = 0.008$, suitable for the LMC (*lower left*) from Schaerer et al. (1993), and $Z = 0.004$, suitable for the SMC (*lower right*) from Charbonnel et al. (1993). The data for the Galactic RSGs is taken from Humphreys (1978); for the Magellanic Clouds, the data are from Elias et al. (1985) and Massey (2002a).

It was van den Bergh (1973) who first suggested that differences in the relative number of blue and red supergiants (B/R) seen in nearby galaxies were caused by the effects of metallicity on stellar evolution. This was further developed by Brunish et al. (1986). However, statistics on the number of red supergiants in nearby galaxies has been very uncertain because of contamination by foreground Galactic dwarfs. Massey (1998b) showed that about half of the RSGs identified in M33 by Humphreys & Sandage (1980) were in fact foreground Galactic red dwarfs. However, low-gravity (supergiant) and high-gravity (dwarfs) red stars can be distinguished in a $B - V$ versus $V - R$ plot, and follow-up spectroscopy at the Ca II triplet proved effective in confirming the identifications (Massey 1998b). Studies in selective fields of NGC 6822, M33, and M31 showed that as the metallicity increases there are proportionately fewer of the high-luminosity RSGs. However, although there are fewer, there are still some high-luminosity RSGs, even in M31, where the metallicity is the highest. This is consistent with the suggestion by Maeder et al. (1980) that the mass range for becoming an RSG doesn't change with metallicity, but that the relative length of time a massive star spends as an RSG (rather than a WR) decreases with increasing metallicity because of the higher mass-loss rates.

Elias et al. (1985) found that the mean spectral type of RSGs changes with metallicity: the average type in the Milky Way is M2; in the SMC, it is M0. This suggests that we have to be very careful in what we include when we count RSGs in nearby galaxies of the Local Group. Using preliminary data for the LMC and SMC, Massey (2002a) notes that there is a factor of three difference in the relative number of blue and red supergiants between the LMC and SMC if one counts only M-type stars; if one instead includes potential K-type supergiants, the difference between the two galaxies is much lower.

3.4. Wolf-Rayet Stars

In 1867 Wolf and Rayet conducted a survey of stars in Cygnus using a visual spectrometer at the Paris Observatory. They came across three stars (all within a degree of one another) whose spectra were dominated by brilliant, broad emission bands, rather than absorption lines. Within 25 years a total of 55 similar stars had been discovered by Copeland, Fleming, Pickering, and Respighi. These efforts, along with visual work by Vogel in 1885, and the early photographic study of their spectra by Pickering in 1890, are nicely discussed in the contemporaneous review by Scheiner & Frost (1894). These studies succeeded admirably in measuring the wavelengths of the brightest emission features, but it would be 30 years before Beals (1930) and others correctly identified the emitting atoms as ionized helium, carbon, and nitrogen.

What gives WR stars their remarkable spectral appearance? The answer is both their strong stellar winds and highly evolved surface chemical abundances. The WR emission lines are signatures of high mass-loss rates, of order $10^{-4.5}\mathcal{M}_{\odot}$ year^{-1} (Abbott & Conti 1987, Nugis et al. 1998). The chemical compositions of

TABLE 3 Evolution of Wolf-Rayet chemical abundance

	Number ratios				
	H/He	**N/He**	**C/N**	**(C + O)/He**	**References[a]**
Cosmic	11.7	0.001	4.8	0.015	(1)
CNO-burning	0	0.004	0.05	0.0002	(2)
He-burning	0	$0.004 \rightarrow 0$	$0.05 \rightarrow \infty$	$0.0002 \rightarrow \infty$	(2)
WNL	$\lesssim 4$	0.002–0.008	0.01–0.13	—	(3)
WNE	$\lesssim 0.6$	0.003–0.006	0.03–0.05	0.0004	(4)
WN/C	$\ll 0.01$	0.005–0.006	2–3	0.025	(5)
WC	$\ll 0.01$	$\ll 0.001$	$\gg 1000$	0.1–2.7	(6), (7)

[a]References: (1) Maeder (1983), (2) Schaller et al. (1992), (3) Crowther et al. (1995a), (4) Crowther et al. (1995b), (5) Crowther et al. (1995c), (6) Crowther et al. (1995c) and references therein, (7) Crowther et al. (2002b).

Wolf-Rayet stars prove their status as evolved objects, with WN-type WRs showing little or no hydrogen but strong helium and nitrogen; WC-type WRs show no nitrogen but strong carbon and oxygen in addition to helium. Abundance studies have confirmed that the surface composition of WN stars is consistent with the equilibrium products of the CNO cycle H-burning, while the surface composition of WC stars reflect He-burning products (Maeder 1983, Maeder & Conti 1994). A comparison is shown in Table 3.[2] During the CNO process in massive stars, nitrogen is produced at the expense of carbon and oxygen (Maeder 1983). Relative to helium (which is the primary product of the CNO cycle, after all!) the N/He number ratio increases by a factor of four. The carbon to nitrogen ratio decreases by a factor of 100. Late-type WNs have a C/N ratio intermediate between cosmic and the equilibrium ratios, and that of early-type WNs are indistinguishable from the CNO equilibrium ratios. During He-burning, nitrogen is destroyed as part of a highly resonant reaction creating first ^{18}O and then ^{20}Ne. ^{12}C is the main product of the He-burning reaction. Oxygen is subsequently produced from carbon. The C/N ratio thus increases from the CNO equilibrium value of 0.05 to infinity, and the $(C + O)$ mass abundance also increases. The abundance ratios measured in WC stars have the He-burning revealed at their surfaces. The WN/C stars are a relatively rare type of WR that shows abnormally strong carbon lines for a WN star; Massey & Grove (1989) demonstrated that these are not WN + WC binaries as once thought, but that the lines originate in the same star. Their abundances

[2]Note that the standard CNO cycle quoted in most astronomy texts is purely catalytic, in the sense that CNO only facilitates the transformation of hydrogen into helium, with no net gain of other elements. However, modern explorations of the reaction rates have shown that the actual situation is far more complex, with multiple branches occurring, and that the equilibrium abundance ratios are highly dependent upon temperature and pressure. An excellent general review is given by Caughlan (1977), and the application to massive stars is nicely described by Maeder (1983).

were shown by Crowther et al. (1995c) to be intermediate between that of WNs and WCs.

The WN and WC classes are divided up into excitation subclasses: WN3 . . . WN9, and WC4 . . . WC9 (Conti 1988). The late WN types (WNLs) are of lower excitation (N III dominates over N V) and differ from the early WNs (WNEs) by having higher visual luminosities and some hydrogen evident in their spectra. It is not clear if all WNLs evolve to WNEs, or whether they have different mass progenitors. Similarly the WCs are divided into early (WCE) and late (WCL) on the basis of the relative strengths of C III, C IV, and O V. The WC sequence appears to be primarily an abundance sequence with lower excitation (WCLs) types having a lower ratio of (C + O)/He at the surface (Smith & Hummer 1988), although Crowther et al. (2002b) has recently called this into question.

As a massive star evolves it will peel down like an onion because of mass loss, with the progression WNL to WNE to WN/C to WC occurring as deeper layers are reached. If the initial composition was high, then the WCs will be of late-type; if the initial composition was low, then the WCs will be of WCE type (Smith & Maeder 1991). Although this at first seems counterintuitive, it is relatively easy to understand: at low metallicity (smaller mass-loss rate) the onion is not peeled down to the He-burning layer until later in the evolution. (Recently Crowther et al. 2002b has argued that the WC subtypes are caused by differences in mass-loss rates and do not reflect difference in composition.) In practice things are probably a lot more complicated: the WNL phase may occur in the highest-mass stars, while the star is still undergoing core H-burning, with the diluted products of CNO brought to the surface by mixing. This would be consistent with the results of studies of coeval associations in the Milky Way (Massey et al. 2001) described above, and the interpretation of current stellar evolutionary calculations (Maeder & Meynet 2000b). Further, we expect that only the most luminous (massive) stars will achieve sufficient mass loss to peel down to the deeper onion layers to the products of He-burning (WC stars).

A key test of stellar evolutionary theory is provided by the WR content of nearby galaxies: both the numbers of WRs and the relative number of WNs and WCs. We discuss first the detection methods and then what is currently known about the Wolf-Rayet content of the Local Group.

3.4.1. DETECTION METHODS Because of their high luminosity and strong emission lines, detection of Wolf-Rayet stars would appear to be a simple matter, even in the more distant galaxies of the Local Group, almost within reach of amateur telescopes. Whereas this may be only a slight exaggeration for WC-type WRs, the emission lines in WN-type WRs are considerably weaker, and so producing a selection-free sample is a harder undertaking than one might suppose. In the optical, the strongest lines are He II λ4686 and C III λ4650, and in Figure 9 we show the difference in the strengths of these lines. The absolute visual magnitudes are similar on the average (typically $M_V = -4.0$; see Conti & Vacca 1990), although the WNs span a larger range in $M_V = \sim -2.5$ to -7 (van der Hucht 2001). We

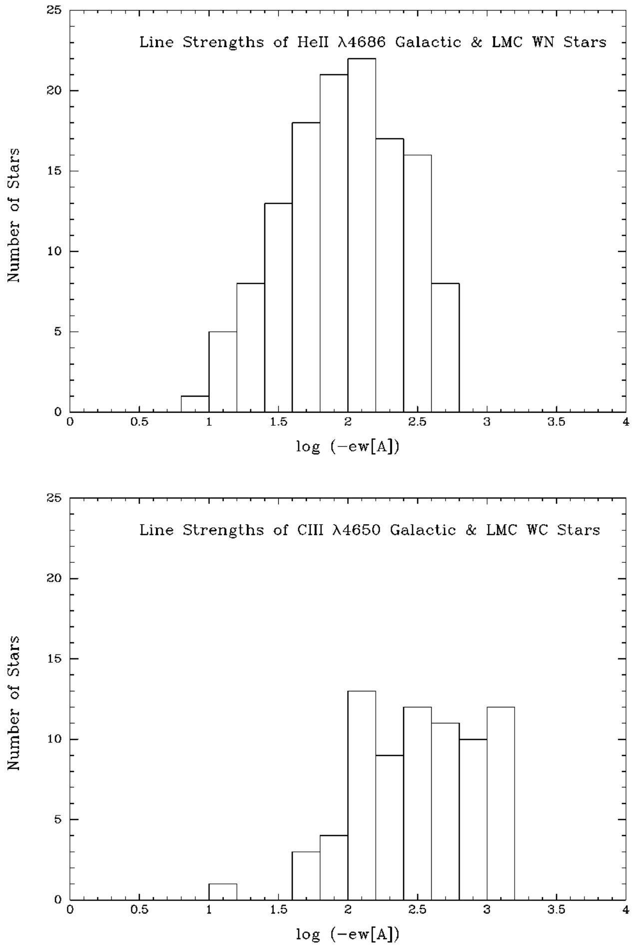

Figure 9 Equivalent widths (ew) of the strongest optical lines in WNs are compared to those for WCs. The data are for He II λ4686 (WNs) and C III λ4650 (WCs). This figure is from Massey & Johnson (1998) and is used with permission.

see that there is a factor of four difference in the average line strengths, with the lines in the weakest-lined WNs more than a 100 times weaker than the lines in the strongest-lined WCs. Thus finding a handful of WCs in a nearby galaxy may be possible from one's backyard, but determining the relative number of WCs and WNs takes more effort.

Various WR detection techniques are critiqued by (Massey 1998c). Most discoveries of WRs have taken place by (often accidental) spectroscopy, via directed objective prism searches, or by interference filter imaging. Armandroff & Massey (1985) describe a filter system optimized for the detection of WRs, and which distinguishes WCs from WNs: the *WC* filter is centered on C III λ4650 (the strongest line in WC stars), the *WN* filter is centered on He II λ4686 (present in both WCs and WNs), and the *CT* filter is centered on a featureless region at λ4750. The filters are ∼50 Å wide and were designed by considering the spectrophotometry of many WRs and normal stars, trying to maximize detectability while minimizing false detections. The filter system is shown in Figure 10. This system has proven very effective in finding WRs in selected regions of the SMC, NGC 6822, M31, and M33, and IC10. Nevertheless, follow-up spectroscopy is needed to weed out false positives, particularly if the detection limits are sensitive enough to find weak-lined WNs. A new survey using this filter system is currently underway with the KPNO 4-m telescope and Mosaic camera, covering most of the Local Group galaxies. Royer et al. (1998, 2001) describe an interesting extension to a five-filter set of narrower filters which purport to also classify WRs into spectral subtypes without the need for followup spectroscopy, but their discovery of WC9 stars in the low-metallicity galaxy IC10 using this system has turned out to be wrong (Massey 2002b, Crowther et al. 2002a), suggesting that spectroscopic followup is still required.

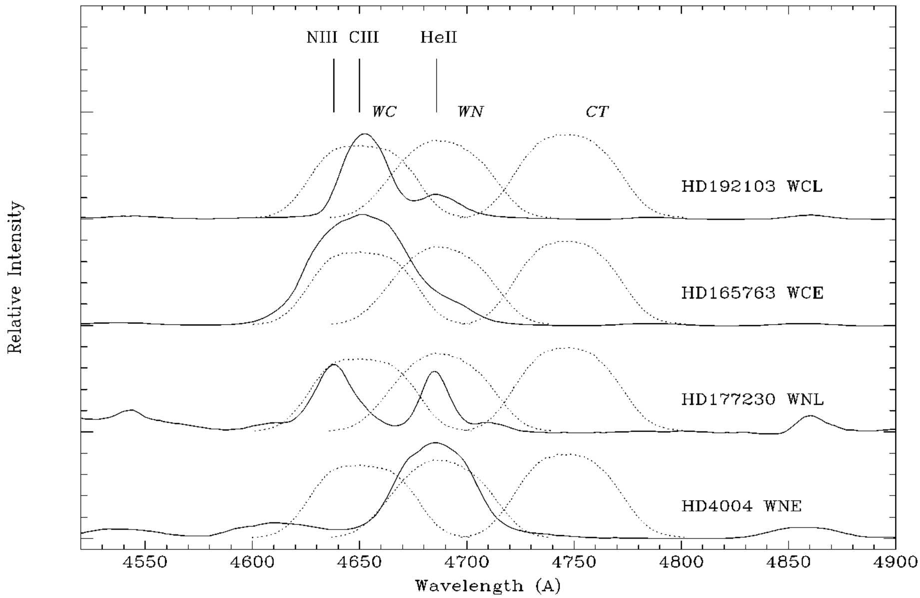

Figure 10 The three filters *WC*, *WN*, *CT* (*dotted lines*) are shown superposed upon normalized spectra of four Wolf-Rayet stars (*solid lines*). The spectrum of HD 177230 has been scaled up by a factor of three to make the features more visible, while the spectra of the two WC stars have been decreased by a factor of two. The locations of He II λ4686 (prominent in both WCs and WNs), C III λ4650 (WCs), and N III λ4634, 42 (WNLs) are indicated.

3.4.2. WRs IN THE MILKY WAY About one third of the 227 WRs known in the Milky Way (van der Hucht 2001) were found by Cannon as part of the HD catalog; the rest were mostly found accidentally by spectroscopy in interesting regions (such as the Galactic Center) rather than as part of directed searches. Of course, this makes the completeness difficult to estimate. Conti & Vacca (1990) have studied the distribution of the known WRs in the Milky Way and emphasized the incompleteness problem. Roughly equal numbers of WNs and WCs are known in the Milky Way. Both early and late WNs and WNs are found, with all of the WCLs inwards of the solar circle, probably the result of the metallicity gradient (Smith & Maeder 1991).

3.4.3. MAGELLANIC CLOUDS Most of the Wolf-Rayet stars in the Small and Large Clouds were found as a result of directed objective prism searches. Breysacher et al. (1999) lists 134 WRs in the LMC, although this number includes some stars classified as "O3If*/WN6-A," which probably represents an extreme Of star rather than a true WR. It also includes the R136 stars, which were also shown to be "Of stars on steroids" (Massey & Hunter 1998). Since the catalog was published one additional WR star, a B0 I + WN3, was found by Massey et al. (2000); it is likely that the occasional additional WR will be found by spectroscopy, particularly in crowded regions where the objective prism method is confused by overlapping images.

For many years only eight WRs were known in the SMC (Azzopardi & Breysacher 1979). A ninth was discovered by Morgan et al. (1991). This is a very low number compared to that of the LMC, given that the galaxy's luminosity is only a factor of 3.6 lower; i.e., we would expect about 35 SMC WRs. The deficiency of WRs could either be caused by a lack of massive star progenitors in the SMC, or it could be telling us that the SMC WRs evolve from higher mass stars than in the LMC, as would be expected given the SMC's low metallicity. Massey & Duffy (2001) argue that the massive star star-formation rate is similar in the two Clouds (per unit luminosity), and that any deficiency of SMC WRs must be because of the latter explanation. They undertook a comprehensive search for WR stars using the optimized filter set and followed up with spectroscopy; they found two more WRs, but more importantly, they were able to rule out the possibility that there is still a considerably large population of WRs waiting to be found in the SMC.

In the LMC, most of the WNs are of early type, with a few WNLs found near 30 Doradus, where very high-mass stars abound. All of the WCs are early type. In the SMC all of the WNs are of early type, and the one WC is also of early type.

3.4.4. BEYOND THE MAGELLANIC CLOUDS The first Wolf-Rayet stars beyond the Magellanic Clouds were found by Wray & Corso (1972) using interference-filter imaging with photographic plates and blink comparison. A summary of subsequent surveys can be found in the appendix of Massey & Johnson (1998), which also contains a complete catalog. At this writing there are 141 spectroscopically confirmed WRs in M33, 48 known in M31, 4 in NGC 6822, 1 in IC 1613, and

26 in IC10, including the 9 newly confirmed WRs by Crowther et al. (2002a) and 2 by Massey & Holmes (2002). Surveys for WRs in NGC 6822 and IC 1613 are probably complete; larger areas of M31 and M33 have now been searched, with many additional candidates that need to be observed spectroscopically. Massey & Holmes (2002) expect another $\sim$70 of their WR candidates in IC10 to be spectroscopically confirmed.

Massey & Johnson (1998) compare the surface density of WRs using only those areas where surveys are complete. NGC 6822, IC 1613, and M31 all have similar surface densities (0.6–0.7 WRs kpc^{-2}). This does suggest that star formation has been less energetic in M31, as only active OB associations have been studied so far in that galaxy, while the others are galaxy-wide samples. In M33, where selected OB associations were imaged, the surface density is $\sim$4 WRs kpc^{-2}. Remarkably the average surface density for IC10 is at least 11 WRs kpc^{-2}, and may be as high as 45 WRs kpc^{-2}, roughly 20 times what is seen for the LMC as a whole, consistent with suggestions that IC10 is undergoing a galaxy-wide starburst (Massey & Holmes 2002 and references therein). Completing galaxy-wide surveys for WRs in M31 and M33 will allow more meaningful comparisons to be made.

The types of Wolf-Rayet stars are consistent with the previous discussions, with few WCs found in low-metallicity environments, and those being WCEs. M33 shows a strong gradient in the number of WC and WN stars with galactocentric distance, consistent with its metallicity gradient (Massey & Conti 1983, Massey & Johnson 1998). Late-type WNs are rare in general, but there are not good statistics on the relative number of WNE/WNL because of the low signal-to-noise ratio of most of the early spectroscopy (e.g., Massey et al. 1987, Armandroff & Massey 1991). Also unknown are statistics on the intermediate WN/Cs.

3.5. Implications for Stellar Evolution

Although not everything is known about the evolved massive star content of Local Group galaxies, the data collected over the years are sufficiently complete to provide a useful challenge to stellar evolutionary models.

First, let us consider the relative number of WC and WN stars as a function of metallicity (Figure 11, upper panel). There is a very smooth progression from low metallicity (relatively few WCs) to high metallicity (more WCs).

Both IC10 and the Milky Way deviate from this trend, in having a higher WC/WN ratio than expected, consistent with the discussion above that the data here are probably incomplete and WNs are harder to find. (The data for the Milky Way are based upon stars within 3 kpc of the sun.) A new survey carried out for IC10 by Massey & Holmes (2002) suggest that the WC/WN ratio may be normal in that galaxy. If not, it could be that the very high star-formation rate in IC10 has somehow affected the statistics, either by producing a top-heavy IMF or by having a burst that is so short in duration ($\Delta\tau \lesssim 200{,}000$ years) that all of the massive stars in IC10 have evolved in lock step. Investigation of the remaining WR candidates in IC10 would resolve this puzzle.

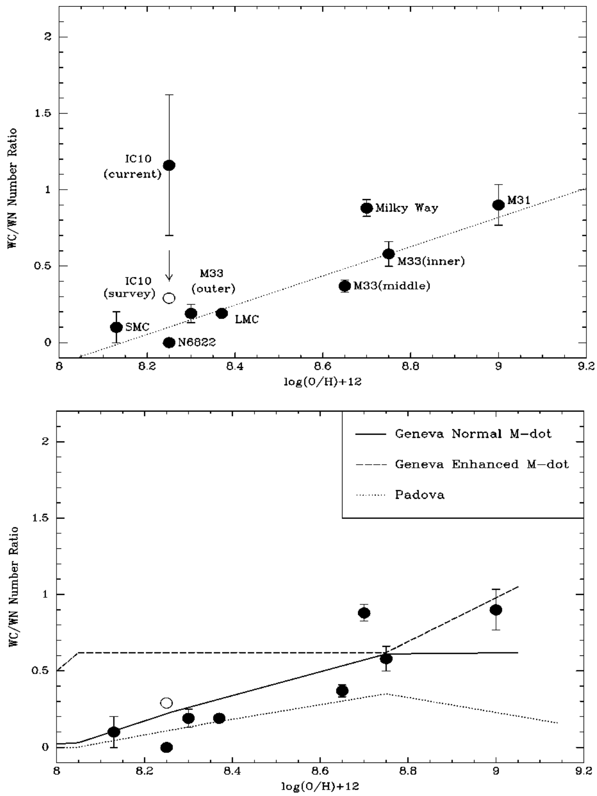

Figure 11 The relative number of WCs and WNs is shown as a function of metallicity. The data come from Massey & Johnson (1998), and have been updated for the SMC (Massey & Duffy 2001) and IC10 (Massey & Holmes 2002). If the anticipated number of IC10 WR candidates are spectroscopically confirmed, the anomalously high WC/WN ratio in that galaxy would become normal, shown by the open circle. The dashed line in the upper panel is a least-squares fit to the data, ignoring both IC10 and the Milky Way points (see text). In the lower panel we compare the predictions of stellar models to these data.

How well do the stellar evolutionary models reproduce this trend? To answer this one must first decide how to recognize a WR star from the models, and in particular how to distinguish WNs from WCs. Various criteria have been used in the literature (Schaerer & Vacca 1998, Leitherer et al. 1999), based upon the surface chemical abundances and effective temperatures. Here I adopt the criteria that the mass fraction of hydrogen be <0.4, that $\log T_{\rm eff} > 3.9$ for a star to be considered a WR, and that the WC stage is reached when the C/N number ratio is ≥ 5, following the precepts of Schaerer & Vacca (1998), except that I've used the C/N ratio rather than the (C + O)/He ratio, as the former is observationally well determined (see

Table 3), and is, after all the primary way that WRs are classified; however, nearly identical results are obtained either way. Following Schaerer & Vacca (1998) the numbers of WRs were determined by integrating the IMF over closely spaced isochrones rather than interpolation over the coarsely spaced evolutionary tracks. (I am grateful to Daniel Schaerer for suggesting this.)

As shown in the lower panel of Figure 11, the normal mass-loss evolutionary tracks of the Geneva group (Schaller et al. 1992, Schaerer et al. 1993, Charbonnel et al. 1993) and of the Padova group (Bressan et al. 1993, Fagotto et al. 1994) both do a reasonable job of reproducing the observed trend with metallicities, although neither predict as many WCs as are actually seen at the highest metallicities. The enhanced mass-loss evolutionary tracks (with $\dot{M}$ twice as large as the de Jager et al. 1988 relation predicts) of the Geneva group (Meynet et al. 1994) do not do very well: These models predict far more WCs at low metallicities than what is actually observed.

The enhanced (2×) mass-loss models were introduced by Meynet et al. (1994), and have (perhaps unfortunately) become a staple of starburst models (e.g., Schaerer & Vacca 1998, Leitherer et al. 1999, Smith et al. 2002). The original motivation for these models was to account for three apparent disagreements between observations and the model predictions that Meynet et al. (1994) suspected might be cured by increasing the mass-loss rates: 1. the existence of WNEs with low luminosities ($\log L/L_\odot \sim 4.5$–5, according to Koesterke et al. 1991), 2. the surface abundances of WC stars compared to the model predictions (Schaerer & Maeder 1992), and 3. the number ratio between the number of blue and red supergiants in clusters of the LMC and SMC.

None of these issues appears to provide much justification for doubling the observed mass-loss rates on the main-sequence, given that there is reasonably good agreement between the standard de Jager et al. (1988) relation and modern mass-loss determinations (Figure 5). In particular, the following counterarguments can be made with the advantage of hindsight. 1. Advances in Wolf-Rayet model atmospheres (including blanketing and stellar wind clumping) have resulted in significantly higher luminosities (0.5–0.7 dex) than that of the older models (Crowther 1999). 2. It is now realized that rotational mixing has a very significant effect on the surface abundances, particularly in the later evolutionary stages (Maeder & Meynet 2000b). However, even in the original description of this problem by Schaerer & Maeder (1992) the suggestion was made that higher mass-loss rates for the post-main-sequence stages (LBVs, RSGs, and WRs) would resolve the disagreement. Given that the mass-loss rates during the LBV and RSG phases are poorly constrained by theory (Sections 3.2 and 3.3), this seems an agreeable solution. 3. The interpretation of the ratio of blue to red supergiants in clusters relies upon assumptions of coevality. A global comparison of B/R as a function of metallicity is still lacking, as described above in Section 3.3.

Next, let us consider how good the agreement is between the minimum mass to become a Wolf-Rayet star in comparison with what we think we know from studies of coeval regions. The evolutionary models predict that at Galactic metallicities, this minimum mass is 25–28 $\mathcal{M}_\odot$ (WN) and 40 $\mathcal{M}_\odot$ (WC); the lowest we observe is 20 and 70 $\mathcal{M}_\odot$. There is no conflict here (there could be WCs with masses less than 70 $\mathcal{M}_\odot$ that we just haven't observed). For the LMC the enhanced mass-loss models predict the minimum masses for becoming a WR are 27 $\mathcal{M}_\odot$ (WN) and 40 $\mathcal{M}_\odot$ (WC); the normal mass loss models predict 42 $\mathcal{M}_\odot$ (WN) and 60 $\mathcal{M}_\odot$ (WC). What we observe is 30 $\mathcal{M}_\odot$ (WN) and 45 $\mathcal{M}_\odot$ (WC). Here we find that the enhanced mass-loss models do match the observations better. For the SMC metallicity the models predict a minimum mass of WNs of 41 $\mathcal{M}_\odot$ (high $\dot{M}$) or 43 $\mathcal{M}_\odot$ (normal $\dot{M}$), and we observe none with masses less than 80 $\mathcal{M}_\odot$, again leading to no conflict between observations and the models. The cluster containing the one SMC WC star did not prove to be coeval, so there are no data constraining the progenitor masses of WCs at SMC-like metallicities.

Finally, let us turn to the relative number of RSGs and WRs. Figure 12 shows an incredibly strong trend with metallicity, a factor of 100 over a range of 0.9 dex in metallicity. This diagram includes only RSGs that lie in the same areas for which surveys are complete for WRs, and counts only RSGs with $M_{\rm bol} < -7.5$ in order to avoid intermediate-mass asymptotic giant branch stars, which would contaminate the same at fainter luminosities (Brunish et al. 1986). Here we see that none of the models does an adequate job of matching the observations.

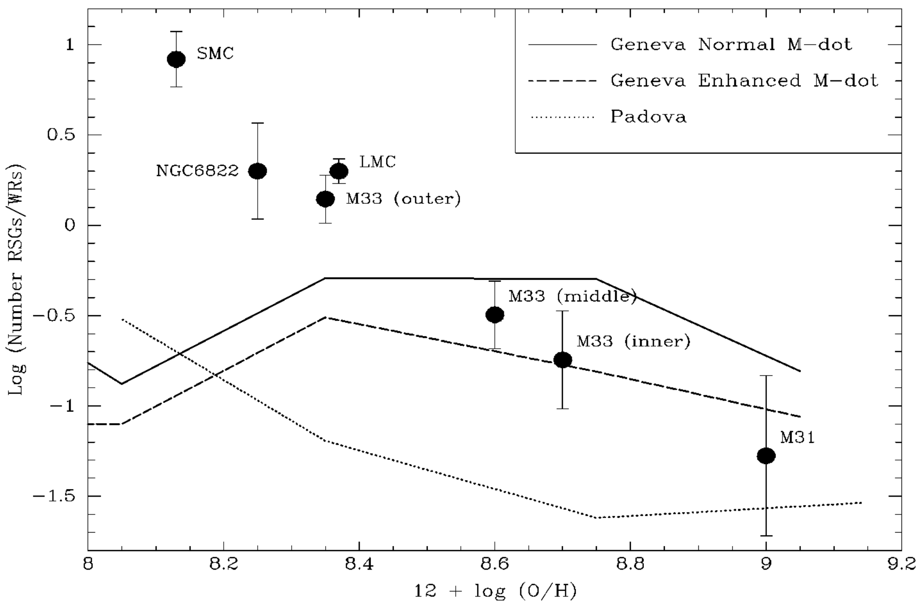

Figure 12 The relative number of RSGs and WRs is shown as a function of metallicity. The data come from Massey (2002a) and references therein. The predictions from stellar models are also shown.

In summary, we find good agreement between the normal mass-loss models and the number ratio of WC/WN stars in nearby galaxies, although the models don't produce quite as many WCs as one observes at the highest metallicities. The enhanced mass-loss models do not match the distribution of WC/WN stars at all, producing far too many WCs at low metallicities. On the other hand, the data on the progenitor masses determined from coeval regions slightly favor the enhanced mass-loss models, at least in terms of what is observed for WC stars in the LMC. Finally, none of the current evolutionary models reproduce the trend of RSG/WR with metallicity. Presumably this is related to the difficulty in producing sufficiently cool high-luminosity RSGs discussed earlier.

Stellar models that include rotation will soon be available (Maeder & Meynet 2000b, Maeder & Meynet 2002) and tested similarly. The observational databases that are being assembled (including galaxy-wide B/R values) will serve as a useful measure of the successes of these and subsequent generations of models.

4. SUMMARY AND REMAINING QUESTIONS

Much is known about the massive star content of the nearest galaxies, and these data show that the initial mass function is independent of metallicity, that the predictions of the Conti scenario (Maeder & Conti 1994) are qualitatively correct, and that the current generation of stellar evolution models do a reasonably good job of matching many of the observational quantities, while still leaving room for improvement. More complete data on the massive star content of the Local Group are becoming available thanks to large-field photometric surveys. These studies must be followed up with corresponding spectroscopy, though, if interesting questions are going to be answered. Here are a few:

(1) Is the IMF of massive stars found in the field the same as that in OB associations?

(2) Something must impose a physical limit on the highest-mass star that can form. How far can we push the observational limits?

(3) How does the B/R supergiant ratio depend upon metallicity?

(4) The ratio of WC to WN stars seems to be very well behaved with metallicity, although the galaxy IC10 poses a problem. Is IC10's high proportion of WCs among its spectroscopically confirmed WRs caused by physics associated with the starburst phenomenon or by incompleteness for WNs? Spectroscopy of the new candidates will answer this.

(5) How does the average spectral type of RSGs depend upon metallicity when pushed to a larger range?

(6) How do mass-loss rates of massive stars scale at higher metallicity?

(7) How do the effective temperatures of O-type stars depend upon metallicity?

(8) How does the relative number of LBVs scale with other evolved massive stars?

(9) How does the relative number of various Wolf-Rayet subtypes vary with metallicity, and in particular what differences are seen in the number of WN/C stars?

(10) How well do the new generation of evolutionary models that include rotation do at matching the observational quantities?

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