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MASSIVE STAR POPULATIONS IN NEARBY GALAXIES

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1. INTRODUCTION

Massive stars are among the main drivers of the evolution of galaxies. These O type stars, along with their highly evolved descendants, the even more energetic Wolf-Rayet objects, are major contributors to the UV radiation and power the far-infrared luminosities through the heating of dust. Their stellar winds are important sources of mechanical power. As progenitors of supernovae, massive stars are agents of nucleosynthesis and may be intimately involved in the initiation of new star formation processes. Hence, massive star evolution is a key study in the exploration of the nearby and distant Universe.

The laws of physics are, so far as we know, the same throughout the Universe. Why should we study massive stars in other galaxies, which is certainly more difficult than studying these objects nearby? We do so because the initial compositions of those stars, in particular their modes of star formation and environments, may well differ from place to place. This leads to different evolutionary histories with a number of observable consequences. For example, it has been known for quite some time that the number ratio of blue to red supergiants shows a gradient in the Milky Way and seems to be different from the ratios found in the Magellanic Clouds. Also, the relative frequency of Wolf-Rayet (W-R) stars to their O-type progenitors appears to be much larger in inner Galactic regions compared to some low-metallicity galaxies. Similarly, the ratio of W-R stars of subtypes WN to those of type WC also changes by a factor of about 20 or more between metal-rich and metal-poor environments. Furthermore, the studies of starburst galaxies containing recently born massive stars show the existence of conspicuous differences in their massive star population statistics. Finally, spectroscopic abundance determinations in AGN and QSOs give us evidence of a very different chemical history among their constituent gaseous and stellar content. These few striking examples illustrate that targe differences may exist in massive star populations among galaxies. It is thus essential to present a good description of such differences and to have a proper understanding of them.

Until about 20 years ago, it was generally thought that the evolution of massive stars was fully understood. With an internal physics governed by electron scattering opacities and a simple equation of state, the stars were supposed to gently leave the main sequence (MS) and finally explode as red supergiants, giving rise to SN II. More recent years have demonstrated the major role of mass loss and initial metallicity, hi addition to the initial mass function (IMF), and the star formation rate (SFR) for shaping massive star evolution and population statistics. The color-magnitude diagrams of young clusters, the stellar abundances of He and CNO elements, and the studies about SN 1987A and its precursor have led to numerous additional investigations on the role of convection and mixing in massive star evolution.

In Section 2 we present some of the statistical properties of massive stars that can be studied individually, and consider what differences have been found between those in three relatively well-known galaxies (the Milky Way and the Magellanic Clouds). We examine the evolution models of OB stars and supergiants in Section 3, and compare them with the observations. We consider the properties of W-R stars, those highly evolved descendents of the most massive stars, in confrontation with the predictions of stellar evolution models in Section 4. Observations and models of even more distant galaxies containing starburst phenomena are considered in Section 5. In these cases, we are usually dealing only with integrated properties of stars in galaxies. We intimate some directions for the future in Section 6.

2. DISTRIBUTION OF INDIVIDUAL OB STARS

2.1. Overview

Massive stars are, for the most part, located in stellar associations born in giant molecular clouds; a powerful enough birth event would be called a "starburst." Initially these stars are surrounded by the dense molecular gas cloud and surrounded by the commonly associated dust. These ensembles will radiate strongly in the IR and radio regions due to the heating of dust and gas excitation, but might be completely hidden optically (e.g. W51 in our Galaxy). After some time, the molecular clouds are dissociated and the dust is dissipated by the radiation and stellar winds from the O stars within, and the region becomes visible as an optical H II or giant H II (GH II) region (e.g. 30 Doradus). The appearance of the spiral arms of galaxies in the visible is primarily determined by the distributions of the H II and GH II regions within them. For the nearer galaxies, the individual stars may be investigated, but for more distant ones, only the integrated properties of the association as it affects the excited gas (and dust) can be studied.

Actual counts of massive stars in associations can be used directly to give estimates of the slope of the IMF, along with the related upper and lower mass limits $M_{\rm upper}$ and $M_{\rm lower}$. We consider these parameters for a group of associations in our Galaxy and the Magellanic Clouds in Section 2.2. Tn more distant GH II regions and starburst galaxies (Section 5), we turn to indirect methods used to confront the questions of "How many?" and "What kinds of massive stars are present?". In such distant galaxies, we make use of the integrated spectra and of global properties, such as their far-infrared (FIR) luminosities, and their optical and UV imaging. For indirect methods, we examine the use of 30 Doradus as a fundamental calibrator.

2.2 Direct Star Counts/Census

Pioneering efforts to elucidate the numbers and types of massive stars in various environments have been made primarily by Massey and associates, using both photometry and spectroscopy. As Massey (1985) has shown, even the unreddened U BV colors for the hottest stars are degenerate, i.e. one cannot distinguish between the hottest and coolest O type stars on the basis of their photometry alone, as is commonly done with luminosity functions, Massey and his associates' homogeneous approach to the determination of the IMF for various associations of the Galaxy and Magellanic Clouds assures us that their comparisons among various stellar groupings ought to be consistent with each other.

2.2.1 PROCEDURE One first acquires deep CCD U BV frames of the relevant stellar associations. Accurate photometry (to 0.02 mag) must be accomplished, and color-color plots used to estimate the extinction and identify the bluest stars. The U BV colors are used to determine the brightness of the stars but spectra suitable for classification are needed to determine the effective temperature, $T_{\rm eff}$, of all stars earlier in type than BlV or so. Obtaining spectra is a time-consuming effort requiring large telescopes; the photometry can be done on modest ones.

Distances of the associations by classic spectroscopic parallax methods are obtained for the Galactic clusters; for the Magellanic Clouds, the standard distances are used. $M_{\rm V}$ and spectral type (or unreddened color) are converted to $M_{\rm bol}$ and $T_{\rm eff}$ using a calibration procedure. These "observed" parameters for the association stars are plotted on a "theoretical" HR diagram with evolutionary tracks. Finally, one counts the numbers of stars in each mass interval along the track, and plots the values as a function of mass. The slope of this relationship is referred to as Γ, defined in the equation

$$f(M) = AM^{\Gamma-1} \qquad (1)$$

where *f(M)* is the fractional number of stars per unit mass interval M, A is a scaling constant, and the Salpeter value for Γ is -1.35. By the above procedure, one is essentially measuring the slope of the Present Day Mass Function (PDMF). An important assumption is that this number is identical to the slope of the IMF; in other words, stellar deaths can be ignored. This is reasonable for the very youngest O associations, and one can, if necessary, account for the already highly evolved W-R stars that are present in some of the regions studied. A further typical assumption is that the spread in formation time is of the order of, or less than, the evolution time, which seems reasonable for O associations (but see Section 2.2.2). Finally, one ignores the binary membership. Unless the binary fraction, typically considered to be 40% for hot stars (Garmany et al 1980), is different from place to place, this assumption is also not unreasonable in the context of seeking similarities and differences in the numbers and types of massive stars in various environments.

2.2.2 MASSIVE STAR CENSUS IN O-ASSOCATIONS In Table 1 (adapted from Conti 1994), we summarize the statistics for ten associations in the Galaxy and Magellanic Clouds, along with those for the solar vicinity. They have been grouped by galaxy abundance, thus sampling the composition of the environment out of which these stars formed.

Table 1. Massive star statistics for various regions

Association	-Γ	# > 60 $M\odot$	Galaxy	References
Field < $R\odot$	1.3		Milky Way	Garmany et al 1982
Field > $R\odot$	2.1			Garmany et al 1982
Cyg OB2	1.0	7		Massey & Thompson 1991
Car OB1	1.3	7		Massey & Johnson 1993
Ser OB1	1.1	1		Hillenbrand et al 1993
LH9	1.6	0	LMC	Parker et al 1992
LH10	1.1	4		Parker et al 1992
LH58	1.7	0		Garmany et al 1993
LH117	1.8	2		Massey et al 1989a
LH118	1.8	0		Massey et al 1989a
30 Dor	1.4	21		Parker 1991
NGC 346	1.8	3	SMC	Massey et al 1989b

The two entries for the solar vicinity were obtained by a method similar to the others but the models used were older and the numbers not quite comparable; the values are just noted for completeness. The authors of the various papers cited suggest that Γ has been determined to an accuracy of ± 0.2 in

each association. If there is a dependence on metal abundance, Z, which, given the uncertainties, is by no means clear, then Γ gets *shallower* as the metal abundance increases. This is in opposition to the prevailing view (Shields & Tinsley 1976) that Γ becomes steeper with increasing abundance.

The entries in Table 1 also give a quantitative indication of $M_{\rm upper}$ for the most massive objects by listing the actual numbers of stars with masses larger than 60 $M\odot$ as inferred from the $M_{\rm bol}$ and $T_{\rm eff}$. We obtained this by inspection of the HR diagrams plotted in the various papers cited. The values, which are more or less proportional to the total numbers of stars in each region, range farther upward in mass to somewhere between 80 to 100 $M\odot$ this is a reasonable estimate of $M_{\rm upper}$ for starburst modeling purposes (see also Section 3.4). There is no dependence of either the numbers of massive stars or the $M_{\rm upper}$ on the galaxy environment, or on metallicity Z.

It has been suggested from indirect arguments that the $M_{\rm lower}$, limit in some starburst regions might be significantly larger than the canonical 0.1 $M\odot$ found near the Sun, and more like a few $M\odot$ (e.g. M 82 - Rieke 1991, McLeod et al 1993). Is there any direct evidence for this? Among well-studied energetic GH II regions, 30 Dor would be the place to look. However, Parker's (1991) survey was complete only to an apparent magnitude corresponding to a few $M\odot$, i.e. just where it begins to get interesting. Despite the (crowding) difficulties, a deeper CCD photometric survey of 30 for needs to be made to investigate its $M_{\rm lower}$ limit and this issue in general.

In their study of NGC 6611 = Ser OB1, Hillenbrand et al (1993) have gone sufficiently deep in their CCD survey to be able to say something about the stellar population at 3-7 $M\odot$. Remarkably, they find that these stars are above the main sequence, in the pre-main sequence phase. Furthermore, the ages of these still contracting stars are a few $\times 10^5$ years, appreciably less than the turn-off time of the upper main sequence, which is a few $\times 10^6$ years! Thus in this association, the presence of (at least) 13 O stars has not inhibited further star formation of lower mass stars. Whether ten or one hundred times that many O stars would inhibit subsequent lower mass star formation remains problematic.

Hillenbrand et al (1993) also call attention to several luminous stars that sit well to the right of the main body of massive stars in NGC 6611. They argue that these stars are indeed cluster members, which must have formed before most of the rest. Eye examination of the rest of the associations referenced above also invariably reveals a few stars in similar, advanced, evolutionary stages. Hillenbrand et al suggest star formation might proceed much like "popcorn" when it is heated; a few kernels "pop" before the main body, and a few more lag behind.

2.2.3 LUMINOSITY FUNCTIONS Massey (1985) noted that using only photometry, it is difficult to distinguish among the most massive stars employing luminosity functions alone. In particular. Massey et al (1989b) show that had they taken only their U BV photometry for the analysis of NGC 346, they would have found Γ to be -2.5, instead of the -1.8 value listed in Table 1. Hill et al (1994) have derived Γ for 14 OB associations in the Magellanic Clouds using CCD photometry. They too find no difference in this parameter between the LMC and SMC, thus no dependence on Z. While their mean values of Γ are somewhat larger than those listed in Table 1 (and probably for the reason mentioned above), this distinction is probably not significant. Their uncertainties in Γ are also larger than those found by Massey and associates with their techniques.

Massey et al (1986) have done CCD photometry of several associations in M31. Using plots and evolution tracks similar to those discussed above, they obtained the curious result (shown in their Figures 31 and 32) that there are no stars in the luminous associations OB78 and OB48 more massive than 40 $M\odot$, although both have W-R stars present. This inference is probably faulty as it is based on photometry only. Thus the use of luminosity functions for massive stars must be treated with caution,

as both the derived Γ and the $M_{\rm upper}$ might be suspect. For luminous stars of M31 and other Local Group galaxies, spectra are difficult but not impossible to obtain on the largest telescopes, especially - with multiple slit instrumentation.

3. STELLAR MODELS AND OBSERVATIONS

3.1 Recent Progress in the Input Physics

Only in the Milky Way and a few galaxies of the Local Group are individual observations of massive stars feasible. Before interpreting the integrated spectra of more distant galaxies, we must first proceed to careful tests of the current models by observing stars in nearby galaxies and checking whether these models are realistic.

Over recent years many grids of stellar models of massive stars at different metallicities *Z* have been produced with various physical assumptions (Brunish & Truran 1982 a,b; Chin & Stothers 1990; Maeder 1990; Arnett 1991; Baraffe & El Eid 1991; Schaller et al 1992; Schaerer et al 1993 a,b; Charbonnel et al 1993: Woosley et al 1993; Alongi et al 1993; Bressan et al 1993; de Loore & Vanbeveren 1994; Meynet et at 1994). The general input physics for stellar models has been extensively discussed (Iben 1974, Iben & Renzini 1983, Chiosi & Maeder 1986. Maeder & Meynet 1989, Chiosi et al 1992a, Schaller et al 1992). We shall limit our review to those points most critical for massive stars. Evaporation by stellar winds is a dominant feature of massive star evolution and all model predictions are influenced. Stellar wind models have been developed by several groups (Abbott 1982, Pauldrach et al 1986, Owocki et al 1988, Kudritzki et al 1987, Schaerer & Schmutz 1994). However, the observed mass loss rates, M, and the wind momentum in O-type stars are generally still larger than predicted (Lamers & Leitherer 1993). Thus, taking theoretical $\dot{M}$ is probably not yet a comfortable choice and it seems preferable to base the models on the empirical values. Those compiled by de Jager et al (1988) have commonly been adopted. These mainly depend on luminosity and to a smaller extent on $T_{\rm eff}$; the role of rotation is still uncertain (Nieuwenhuijzen & de Jager 1988, 1990; Howarth & Prinja 1989). These average $\dot{M}$ might be too low (Schaerer & Maeder 1992). Considerable uncertainties remain in the adopted $\dot{M}$, particularly for the red supergiants which lose mass at very high values (de Jager et al 1988, Stencel et al 1989, Jura & Kleinmann 1990). Presently there is no complete theory for the winds of red supergiants. For these, evidence of dust ejection is provided by IRAS observations, which show that some of them possess extended circumstellar shells (Stencel et al 1989) potentially leading to OH/IR sources (Cohen 1992). Evidence for strong winds in the previous red supergiant phase of SN 1987A has also been presented by Fransson et al (1989), and for the more recent SN 1993J by Hoflich et al (1993).

Different treatments of convection and mixing in stellar interiors have been advocated, giving a major uncertainty in massive star models. We identify the following different assumptions regarding convection and mixing in massive star models:

- Schwarzschild's criterion,
- Schwarzschild's criterion and core overshooting,
- Overshooting below the convective envelope,
- Ledoux criterion,

- Semiconvection or semiconvective diffusion,
- Turbulent diffusion or other forms of rotational mixing.

All of these models are claimed by their authors to fit the observations and the debate has been lively in recent years (Chiosi & Maeder 1986; Maeder & Meynet 1989; Brocato et al 1989; Lattanzio et al 1991; Stothers 1991 a,b; Stothers & Chin 1990, 1991, 1992 a,b). There is at present no definite theoretical or observational proof in favor of any model. However, a few useful indications on the limits and possibilities of the various models must be mentioned.

Although claims have been made in favor of substantial overshooting from convective cores with respect to what is predicted by Schwarzschild's criterion, it now seems clear that the overshooting distance is limited to about (0.2-0.4) $H_{\rm p}$ (Maeder & Meynet 1989, Stothers & Chin 1991, Napiwotzki et al 1993, Meynet et al 1993). The main effect is to increase the main sequence width and lifetimes, while the helium burning lifetimes are reduced (due to higher L); the blue loops are also shorter. In contrast to overshooting, which extends convective zones, the Ledoux criterion tends to prevent convective mixing in zones with variable mean molecular weights (Kippenhahn & Weigert 1990). Some recent comparisons with observations seem to favor Ledoux's rather than Schwarzschild's criterion for convection (Stothers & Chin 1992 a,b); however, the result may depend on the adopted $\dot{M}$. Other recent theoretical work (Grossman et al 1993) shows that the Ledoux criterion has no bearing at all in stratified stellar layers. Thus, both at the theoretical and observational levels, the convective criteria remain uncertain.

Semiconvection occurs in zones that are convectively unstable according to Schwarzschild's criterion, but not according to Ledoux's. Semiconvection may thus produce some mixing in zones with a gradient of the mean molecular weight. Various treatments of the problem have been made (Chiosi & Maeder 1986, Langer at al 1989, Arnett 1991, Chiosi et al 1992a, Langer 1992, Alongi et al 1993). In a semiconvective zone, the nonadiabatic effects (radiative losses) produce a progressive increase of the amplitudes of oscillations at the Brunt-Vaisala frequency around a stability level (Kippenhahn & Weigert 1990). The growth of amplitudes is generally rapid compared to the evolutionary timescale, so that a situation equivalent to Schwarzschild's criterion is established. However, this might not be true in massive stars, as shown by Langer et al (1985), who discussed the timescales involved in semiconvective mixing. They propose a diffusion treatment that is equivalent to Schwarzschild's criterion when the mixing timescale is short compared to the evolutionary timescale, and to the Ledoux criterion in the opposite case. Models with such diffusion are of special relevance to the discussion about the blue progenitor of SN 1987A (Langer et al 1989; Langer 1991 a,c) as well as about the evolutionary status of blue supergiants (cf Section 3.5).

The effects of rotationally induced mixing may be important for massive stars. The radiative viscosity is so large that dissipative processes may have a timescale comparable to the evolutionary timescalcs of massive stars (Maeder 1987b). Mixing could produce chemically homogeneous or nearly homogeneous evolution on the main sequence and thus lead directly as a result of nuclear burning to the formation of He stars, which would be observed as W-R stars. Models of massive stars losing mass and angular momentum have been calculated by Sreenivasan & Wilson (1985). As Langer (1992b) emphasized, models with semiconvective and rotational mixing may solve several problems: the existence of the WN+WC stars (Langer 1991b; cf Section 4.2), the origin of nitrogen enhancement in OB supergiants, the alleged mass discrepancy for OB main sequence stars (cf Section 3.3), and the nature of the blue progenitor of SN 1987A. Curiously enough, the claims in favor of semiconvection mean less mixing in the convective zone with varying mean molecular weight, while the claims in favor of rotational mixing mean more mixing from inner material into the outer radiative zone (Langer 1993). The situation is still uncertain, but we think it likely that mass and metallicity are not sufficient

to describe massive star evolution and that rotational velocity will be an unavoidable additional parameter, as well as (for some) membership in close binary systems.

3.2 Metallicity Effects in Massive Stars

Metallicity, like other effects such as nonconstant star formation rates and peculiar initial mass functions (Section 5), is a key factor influencing massive star populations in galaxies. Metallicity effects can enter evolution through at least four possible doors:

1. Nuclear production. Metallicity Z may influence the nuclear rates; a good example occurs for the CNO cycle. A very slight contraction or expansion to a new equilibrium state may compensate for a change in nuclear rates (Schwarzschild 1958). In massive stars, a lower Z also produces a more active H-burning shell in the post-main sequence evolution and this favors a blue location in a part or the whole of the He-burning phase (Brunish & Truran 1982 a,b; Schaller et al 1992). This was one of the initial explanations proposed for the blue precursor of SN 1987A (Truran & Weiss 1987).

2. Opacity effects. In the interiors of massive stars, electron scattering, which is independent of Z, is the main opacity source. Thus, in contrast to the case of low and intermediate mass stars, metallicity has no great direct effect on the inner structure of massive stars.

3. Stellar winds. In the very external layers, Z may strongly influence the opacity and thus the atmospheres and winds. Wind models for O stars by Abbott (1982) suggested a Z-dependence of the mass loss rates $\dot{M}$ of the form $\dot{M} \propto Z^{\alpha}$, with α = 1.0. Other models gave a value of α between 0.5 and 0.7 (Kudritzki et al 1987, 1991; Leitherer & Langer 1991; Kudritzki 1994). It is likely that this is the main effect by which Z may influence massive star evolution (Maeder 1991a). For yellow and red supergiants, there are no models (Lafon & Berruyer 1991) nor observatidas (Jura & Kleinmann 1990) giving reliable $\dot{M}$ vs Z information; thus a major uncertainty in post-MS evolution remains.

4. Helium content. A ratio $\Delta Y / \Delta Z$ greater than 3 between the relative enrichments in helium and heavy elements has been established from low-Z H II regions (Peimbert 1986, Pagel et al 1992). Thus, changes in Z imply large changes in Y, which have a direct effect on the models.

3.3 Main Sequence Evolution

3.3.1 HR DIAGRAM, LIFETIMES, MASSES Let us examine a few of the main properties of the models of massive stars at various metallicities. At low metallicity Z, the zero-age main sequence (ZAMS) is shifted to the blue due to the lower opacity in the external layers. Between the sequences at Z 0.001 and 0.04. the shift in log $T_{\rm eff}$ amounts to +0.06 dex at 20 $M\odot$ (Schaller et al 1992, Schaerer et al 1993b) and a lowering in luminosity by 0.10 dex. The reason is that at low Z the hydrogen content is higher, thus the electron scattering opacity is larger. The width of the MS band is predicted to change considerably according to metallicity. The main feature is a prominent "paunch," which is displaced to lower luminosities for lower mass loss rates and metallicities. Two physical effects are responsible for this paunch (Maeder 1980). First, the large mass fraction of the He core, resulting from the removal of the outer layers, favors the redward extension of the tracks, Second, when the surface hydrogen content becomes lower than $X_{\rm s}$ = 0.3 or 0.4 as a result of mass loss (Chiosi & Maeder 1986), the lowering of the surface opacity moves the star back to the blue. Thus, the paunch appears in the range of masses where mass loss is sufficient to increase the core mass fraction, but not high enough to lower $X_{\rm s}$ below the critical limit. An increase of overshooting or opacity may enhance the paunch. Models with enhanced opacities may have a MS band covering all the HR diagram (Stothers

& Chin 1977, Nasi & Forieri 1990).

The lifetimes in the various nuclear phases change with Z. For the H-burning phase, the lifetimes t(H) are typically longer by 35% for a 20 $M\odot$ model at Z = 0.001 compared to Z = 0.040. The reason rests on the lower luminosity and the larger reservoir of hydrogen. The lifetimes t(He) in the He-burning phase are generally longer in models with higher Z, due to the higher mass loss rates which lead to a drastic decrease of the luminosities in this phase. For models of 15 to 120 $M\odot$, the t(He) / t(H) is typically 9 to 10%; these ratios are between 11 and 19% at Z = 0.04 and they may amount to 50% if the mass loss rates are increased by a factor of 2 (Meynet et al 1994). These large factors show how our ignorance of the exact mass loss rates at various Z may affect massive star models.

There is an apparent lack of O stars close to the theoretical zero-age sequence (Garmany et al 1982). This is also quite clear in recent gravity and $T_{\rm eff}$ determinations by Herrero et al (1992). We notice that for massive stars, the accretion timescale of the protostellar cloud is longer than the Kelvin-Helmholtz timescale (Yorke 1986). The consequence is that no massive pre-MS star should be visible (Palla et al 1993) - a fact that could contribute to obscured stars close to the ZAMS, Wood & Churchwell (1989) and Chiosi et al (1992a) suggest that 10% to 20% of the O stars are still embedded in their parent molecular clouds. An alternative explanation is that there is no true ZAMS corresponding to a chemically homogeneous stage for O stars, because nuclear reactions ignite early during the contraction phase (Appenzeller 1980) and may thus make stars inhomogeneous before the end of the contraction phase.

Another potential problem is the so-called mass discrepancy for O stars. Spectroscopic masses derived from gravity and terminal velocity determinations were claimed to be smaller than predicted by stellar models (Bohannan et al 1990, Groenewegen et al 1989, Herrero et al 1992, Kudritzki et al 1992). In other words, spectroscopy suggests that O stars are overluminous for their masses and the discrepancy amounts up to about 50%. Langer (1992) interprets the overluminosity of O stars as a sign of rotational or tidal mixing enlarging the helium core. Apart from the fact that the force multiplier may not be correctly predicted by non-LTE wind models, the reality of the mass discrepancy has been questioned recently by Lamers & Leitherer (1993). They show that large discrepancies exist between theoretical and observed mass loss rates, as is true for the terminal velocities; they also argue that the discrepancies cannot be solved by adopting smaller masses for O stars. According to Schaerer & Schmutz (1994), the use of plane-parallel models for O stars may lead to significant errors for spectroscopic gravities, masses, and helium abundances. It is thus possible that the mass discrepancy is due to the inadequate modeling of stellar atmospheres. This view seems confirmed by the most recent work of Pauldrach et al (1994) and Kudritzki (1994), who do not find the mass discrepancy once additional wind opacity due to iron transitions is taken into account.

3.3.2 ABUNDANCES ON THE MS The surface abundances in He and CNO elements offer a powerful test of stellar evolution. Evidence of CN processing is provided by He and N enhancements together with C depletion, while O depletion only occurs for advanced stages of processing. The abundances may cover a range from solar values (C/N = 4, O/N 10) to CNO equilibrium values in the extreme case which is reached in WN stars (C/N = 0.02, O/N = 0.1; Maeder 1983, 1987a). Models with mass loss but no extra-mixing predict He and N enrichment in MS stars only for initial masses larger than about 50 $M\odot$ depending on the mass loss rates. Models with rotational mixing may lead to a precocious appearance of the products of the CNO cycle (cf Maeder 19875, Langer 1992). The observations of 25 OB stars by Herrero et al (1992) show that most MS stars have normal He and N abundances. The same is true for MS B-type stars (Gies & Lambert 1992). For example, even the most massive object, Melnick 42 (O3f), appears to show normal abundance ratios (Pauldrach et al 1994; but see Heap et al 1991). However, there are also exceptions for O and B stars. For example, the O4f star

ζ Pup presents evidence of an atmosphere with CNO burned material (Bohannan et al 1986, Pauldrach et al 1994). Fast rotators are also an exception and they generally show He and N enhancements (Herrero et al 1992). Another ease is the group of ON stars, i.e. O stars with N-enrichments (Walborn 1976, 1988; Howarth & Prinja 1989); this group contains at least 50% short-period binaries (Bolton & Rogers 1978).

An analysis of the association Per OB1 (Maeder 1987b) suggests that there is a bifurcation in stellar evolution: While most stars follow the tracks of inhomogeneous evolution, a fraction of about 15%, mainly composed of fast rotators and binaries, may evolve homogeneously and become ON blue stragglers.

3.4 The Eddington Limit and LBV Stars

The value of the mass of the most massive stars in galaxies has been a much debated subject. Recent photometric and spectroscopic studies suggest stellar masses up to about 100 $M\odot$ (Section 2.2.2; Table 1; Divan & Burnichon-Prevot 1988, Kudritzki 1988, Heydari-Malayeri & Hutsemekers 1991, Massey & Johnson 1993). Recently Pauldrach et al (1994) have suggested that the most massive star known is Melnick 42 in the LMC, which may have a mass of up to 150 $M\odot$.

There is an upper luminosity limit to the distribution of stars in the HR diagram. It runs from log $L / L\odot$ = 6.8 at $T_{\rm eff}$ = 40,000 K to log $L / L\odot$ = 5.8 at 15,000 K and it stays constant at lower $T_{\rm eff}$ (Humphreys & Davidson 1979: Humphreys 1989, 1992). The theoretical location of the Eddington limit has been examined by Lamers & Fitzpatrick (1988) on the basis of model atmospheres including metal line opacities. The limit was shown to agree with the observed limit in the Milky Way and in the LMC. Subsequent investigations indicate that the Eddington limit rises again at low $T_{\rm eff}$ since the opacities decrease considerably there (Lamers & Noordhoek 1993). Thus, the lowest part of the limit was called the "Eddington trough"; its location in the HR diagram is, of course, higher for stars of lower Z since they have lower opacities in the external layers.

The Eddington limit or "trough" may prevent the redward evolution of very massive stars in the HR diagram (Maeder 1983, Lamers & Noordhoek 1993). The region inside the trough will be empty except for unstable stars during their outbursts. The upper luminosity limit is determined by stars that can just pass under the Eddington trough. Thus, because the location of the trough depends on Z, the upper luminosity of red supergiants may not be an ideal standard candle, contrary to expectations (Humphreys 1983b).

The Luminous Blue Variables (LBVs, Conti 1984), also called hypergiants, S Dor, or Hubble-Sandage Variables, are optically the brightest blue supergiants. They show irregular and violent outbursts, with average mass loss rates up to about 10^{-3} $M\odot$ yr^{-1} (Davidson 1989, Lamers 1989). The group continues toward lower $T_{\rm eff}$ as the so-called cool hypergiants (Humphreys 1992), which are the most luminous F, G, K, and M stars. These also show evidence of variability, of high mass loss, and extensive circumstellar dust. The He and CNO abundances in LBVs (Davidson et al 1986) are in agreement with products of the CNO cycle at equilibrium, which confirms that LBVs are post-MS supergiants (Maeder 1983). About 30 LBVs have been identified by various authors in nearby galaxies (see list by Humphreys 1989) including the LMC, M31, M33, NGC 2403, M81, and M101. Among hypergiants, the OH/IR supergiants, revealed by radio and IR observations, are the most extreme M supergiants, likely having optically thick dust shells. About two dozen cool hypergiants are currently known in the Milky Way (Humphreys 1991).

The bolometric luminosities of LBVs are constant (Appenzeller & Wolf 1981) during an outburst. However, the matter ejection, particularly during the outbursts, modifies the photospheric radius and $T_{\rm eff}$, and as a consequence also the bolometric correction and visual luminosity. During their outbursts, LBVs essentially move back and forth horizontally along the HR diagram. Obscuration by gas and dust may also affect the emitted light (Davidson 1987). The circumstellar environment of these stars is peculiar and may affect the distance estimates (Viotti et al 1993). The evolutionary changes of P Cyg over the past two centuries have been recently discussed by Lamers & de Groot (1992), de Groot & Lamers (1992), and El Eid & Hartman (1993) and have been shown to correspond to recent theoretical estimates. At its minimum visual light, the star is hotter (T = 20,000-25,000 K) than at its maximum light (where T = 9000 K).

In the past, LBVs have been assigned to all possible evolutionary stages, but they are currently interpreted as a short stage in the evolution of massive stars with initial $M > 40\ M\odot$. A likely scenario is

O star -> Of / WN -> LBV -> Of / WN -> LBV ... -> WN -> WC.

After central H-exhaustion, the star undergoes redward evolution in the HR diagram and is likely to reach the Eddington limit or the "trough". Strong mass loss occurs with shell ejection (LBV). As a result, stability and bluer location in the HR diagram are restored (Of/WN). Internal evolution again brings the star to the red in a few centuries - a time that may depend on the stellar mass and amount of ejected mass (cf Maeder 1989, 1992b). The star again moves toward the Eddington limit, and the cycle of evolution between the Of/WN and LBV stages continues until, as a result of mass loss, the surface hydrogen content is low enough ($X_{\rm s} \leq 0.3$) so that the star definitely settles in the Wolf-Rayet stage. The overall duration of the LBV phase is fixed by the amount of mass ΔM to be lost between the end of the MS phase and the entry in the W-R phase. For a typical $\Delta M = 10\ M\odot$ and an average $\dot{M}$ of 10^{-3} to $10^{-4}\ M\odot\ {\rm yr}^{-1}$, the typical duration would be $\approx 10^4$ to 10^5 yr. This general scenario is consistent with several properties of LBVs: their location in the Hertzsprung-Russell diagram (Humphreys 1989, Massey & Johnson 1993), their $\dot{M}$ rates (Lamers 1989), their high N/C and N/O abundance ratios (Davidson et al 1986), and the existence of transition objects, as discussed below.

Many observational studies have been made of these transition objects, which are often of spectral type Of/WN and present spectral variability. Examples are S Dor (Appenzeller & Wolf 1981, Wolf et al 1988), R 71 (Appenzeller & Wolf 1981, Wolf et al 1981), AG Car (Caputo & Viotti 1970, Viotti et al 1993), R 127 (Stahl et al 1983; Stahl 1986, 1987; Wolf 1989), R 84 (Schmutz et al 1991), and He 3-519 (Davidson et al 1993). It is also possible that after the LBV phase, some stars go to the stage of OH/IR object and then become W-R stars. This different, but not contradictory scenario, could happen to stars with masses low enough to enable them to go below the "trough". The special cases of Var A in M33 (Humphreys 1989) and IRC+104020 - an extreme galactic F-supergiant with a very large IR excess from circumstellar dust (Jones et al 1993) - might correspond to such a scenario.

The physical origin of the outbursts in LBVs and hypergiants is still a matter of controversy and several models have been considered (e.g. Stothers & Chin 1983; Doom et al 1986; Appenzeller 1989; de Jager 1992; Maeder 1989, 1992b). The most striking property of these models is the strong density inversion occurring in the outer layers, where a thin gaseous layer floats upon a radiatively supported zone. This zone results from the opacity peak which leads to supra-Eddington luminosities in some layers. The idea of a density inversion has a 40 year history (Underhill 1949, Mihalas 1969, Osmer 1972, BisnovatyiKogan & Nadyozhin 1972, Stothers & Chin 1983). A review of the literature shows that essentially three different kinds of conclusions were drawn: 1. A Rayleigh-Taylor instability occurs as a result of the density inversion, which is therefore washed out by the instability. 2. The

supra-Eddington luminosity drives an outward acceleration and mass loss without a density inversion. 3. Strong convection and turbulence develop and the inversion is maintained.

A difficulty with most models is that they look for a hydrostatic solution to the problem. However, the resolution likely lies in the context of hydrodynamical models. Although the second of the above conclusions seems preferable, it is still unclear whether or not the density inversion is maintained. Another noticeable peculiarity in the physics of LBVs is that the thermal timescale in the outer layers is shorter than the dynamical timescale. During an outburst, which is at the dynamical timescale, the ionization front is able to substantially migrate inward (Maeder 1992b), so that some layers of matter may participate in the ejection and produce the observed shells (Hutsemekers 1994).

3.5 Blue and Red Supergiants

Conti (1991b) recently reviewed the observations of hot massive stars in galaxies and a complete list of the observations of red supergiants in galaxies has been given by Humphreys (1991). Humphreys (1983b, 1991) has also reviewed the potential role of red supergiants as distance indicators. Amazingly, many problems and controversies remain about supergiants, for which evolution is even more uncertain than for W-R stars! The reason is that W-R stars are dominated by the overwhelming effect of mass loss, which washes out most effects related to uncertainties in convection and mixing. Supergiants are often close to a neutral state between a blue and a red location in the HR diagram (Tuchmann & Wheeler 1989, 1990); even minor changes in convection and mixing processes may greatly affect their evolution.

3.5.1 CHEMICAL ABUNDANCES Walborn (1976, 1988) proposed that ordinary OB supergiants have an atmospheric composition enriched in helium and nitrogen and depleted in carbon, as a result of CNO processing. According to Walborn, it may just be the small group of the so-called OBC supergiants that have normal cosmic abundances (Howarth & Prinja 1989). Herrero et al (1992) showed that most OB supergiants and Of stars show helium enhancements. As for all rules, there are exceptions: A few B-supergiants do not show He and N excesses (Dufton & Lennon 1989). Herrero et al also show that fast rotators of all luminosities present evidence of CNO processing. Enhancements of nitrogen and helium abundances have also been found for post-MS B type stars by Gies & Lambert (1992), and by Voels et al (1989) in the 09.5 Ia star α Cam. As expected, the so-called OBN stars show evidence of He and N excesses with C depletion (Walborn 1988, Schonberner et al 1988).

Abundance determinations have also been made for B supergiants in the LMC and SMC, particularly interesting in relation to the progenitor of SN 1987A. These supergiants generally show He and N enhancements (Reitermann et al 1990, Kudritzki et al 1990, Lennon et al 1991). A recent high-dispersion study of LMC B-supergiants also confirms such enrichments (Fitzpatrick & Bohannan 1993). Among 62 stars of types B0.7 to B3, only 7 are OBC stars (Fitzpatrick 1991). These authors conclude, in agreement with the Walborn hypothesis, that the "typical" supergiants show contaminated surfaces, and only the rare nitrogen weak stars (OBC) have retained their original main sequence composition. The progenitor of SN 1987A, which was a B-supergiant, had N/C and N/O ratios larger than solar values by 37 and 12, respectively (Fransson et al 1989). From all these results, it is clear that most B-type supergiants in the Galaxy, the LMC, and the SMC generally show evidence of CNO processing on their surfaces.

The above observations place severe constraints on stellar models, which do not usually predict He and N enrichments in blue supergiants at solar composition. At solar Z, blue loops with the associated He and N enrichments (as a result of dredge-up in red supergiants) only occur for $M \leq 15\ M$. This is the case for the models with Schwarzschild's criterion and overshooting (Schaller et al 1992), and with the Ledoux criterion (Stothers & Chin 1992 a,b; Brocato & Castellani 1993). Models with

semiconvection (Arnett 1991) have the same difficulty: At solar composition, the evolution goes straight to the red supergiant phase and there are no enriched blue supergiants. At lower metallicity, the blue loops are generally more developed and thus blue supergiants are predicted with He and N enrichments. However, even in this case it seems necessary (Langer 1992) to advocate some, rotational mixing to account for the observed abundances.

The study of CNO abundances in three A-type supergiants (Venn 1993) reveals N-enrichments larger than predicted by the first convective dredge-up, if these stars have first gone to the red supergiant stage. This supports the idea of additional mixing. Analyses of four F-type supergiants by Luck & Lambert (1985) show material processed by the CN cycle at a level that may be higher than predicted. Analyses of some F and K supergiants in the SMC by Barbuy et al (1991) indicate solar N/Fe and C/Fe ratios, and thus no evidence of CNO processing. A further study by Barbuy et al (1992) of 14 Galactic F-supergiants shows an absence of CNO processed material in stars with low rotational velocities. For F supergiants with high rotation, the derivation of CNO abundances is unfortunately masked by the line broadening. Yellow supergiants also show sodium overabundance by a factor of 3 to 4 (Boyarchuk et al 1988: cf also Lambert 1992). Boyarchuk et al have suggested an increase of this overabundance with initial stellar masses. An interpretation put forward by Denissenkov & Ivanov (1987) and Denissenkov (1988, 1989) rests on proton capture by the isotope Ne^{22}, supposed to be overabundant. However, it is not clear why Ne^{22} should be overabundant, whether it is present initially, or whether it results from N-burning in the helium core.

Red supergiants of type G-K Ib exhibit some sodium overabundances, but less pronounced than in F supergiants (Lambert 1992). For red supergiants, the presence of CNO processed elements, as a result of dredge-up in the deep convective envelopes, is both expected and observed (Lambert et al 1984, Harris & Lambert 1984). Comparisons show a general agreement (Maeder 1987a), with possible indications that some extra mixing may be needed.

3.5.2 THE BLUE HERTZSPRUNG GAP There are many more stars outside of the MS band than predicted (Meylan & Maeder 1982). The problem is particularly serious in the SMC and LMC. In the Milky Way, excesses of A-type supergiants have also been suggested (Stothers & Chin 1977, Chiosi et al 1978). The observed and theoretical numbers can be brought into agreement if the MS phase would also include the B- and A-type supergiant stages. This discrepancy is related to the problem of the so-called blue Hertzsprung gap (BHG), which is predicted by most stellar models to occur at the end of the MS and is not observed. Instead, the true star distribution appears continuous from the MS to the A-type supergiants (Nasi & Forieri 1990, Fitzpatrick & Garmany 1990, Chiosi et al 1992b).

Various explanations have been proposed for the lack of a BHG. Opacity effects may produce a "paunch" on the MS as discussed above (Section 3.3), but with present opacities (Iglesias et al 1992) and mass loss rates, the paunch occurs at luminosities too high to account for the observations (Schaller et al 1992). Extended atmospheres and the role of binaries (Tuchmann & Wheeler 1989, 1990) have also been advocated as explanations. Mixing reduces, but does not suppress, the gap (Langer 1991c). The temperature scale may also be a problem given the photometric nature of most of the observations of stars. A gap between $T_{\rm eff}$ = 35,000 K and 20,000 K corresponds only to a difference of 0.04 in (B - V) color, which is quite small and may be blurred by other effects. The adjustment of individual isochrones on star clusters (Meynet et al 1993), together with a mapping of He and CNO abundances and the use of (U - B) colors, may eventually inform us as to the reality of the gap problem and the exact status of blue supergiants.

3.5.3 THE SUPERGIANT DISTRIBUTION AND THE SN 1957A PROGENITOR A drop-off in the distribution of LMC supergiants in the HR diagram to the right of an oblique line between log $T_{\rm eff}$ = 4.2 and 3.9 was noted by Fitzpatrick & Garmany (1990), and called a "ledge": A further study of

5050 LMC stars with new calibrations and reddening corrections (Gochermann 1994) shows that the "ledge" might be less significant. In data from the Galaxy it appears marginally (Blaha & Humphreys 1989). Two kinds of models are able to produce high numbers of blue supergiants and to produce a "ledge" or at least a marked decrease in the star distribution in the HR diagram: *(a)* models with low mass loss, *(b)* models with blue loops.

Models with low mass loss (Brunish & Truran 1982 a,b: Schaller et al 1992) predict that most of the He core burning phase is spent in the blue supergiant phase directly after the MS. This may give a ledge; however such models do not provide red supergiants, in disagreement with observations in the LMC and SMC. The blue location of models with low mass loss is due to the large intermediate convective zone which homogenizes a part of the star (Stothers & Chin 1979, Maeder 1981). Mass loss, even if small, reduces this zone and favors the redward motion in the HR diagram, and thus the star becomes a red supergiant early during the He phase. However, the uncertainties about mass loss are critical. As an example, 25 $M\odot$ models at $Z = 0.008$ with typical $\dot{M}$ (Schaerer et al 1993a) spend most of their He-burning phase in the blue with log $T_{\rm eff}$ between 4.3 and 3.9. An enhancement of $\dot{M}$ by a factor of 2 leads to a red location of the whole He-burning phase (Meynet et al 1994). Thus, as long as the $\dot{M}$ rates are imprecise, it may be difficult to derive conclusions about semiconvection, diffusion, and rotational mixing from the distribution of supergiants.

Models with blue loops also enhance the number of blue supergiants, and are simultaneously able to account for some He and N enrichments in blue supergiants, but often not as much as required by the observations (Section 3.5.1). Models with Schwarzschild's criterion and overshooting at $Z \leq 0.008$ have well-developed blue phases at all masses (Schaerer et al 1993a). This is also the case for models with the Ledoux criterion for $Z < 0.004$ (Brocato & Castellani 1993) and for models with semiconvection by Arnett (1991) at $Z < 0.007$, which well reproduce the numbers of blue and red supergiants in the LMC and lead to a blue location of the supernova progenitor at 20 $M\odot$, as must be the case for SN 1987A (Arnett 1991, Langer 1991c). The physical connection leading to the blue progenitor is most interesting (Langer 1991 a,c) and illustrates a general stellar property. The mild mixing reduces the He content in the He burning shell, which is therefore less efficient. At the end of the He core burning phase, when the CO core contracts, the He shell acts as a weak minor and produces only a moderate expansion in the intershell region between the He and H shells. Thus, the H shell is not extinct and it keeps very active. It acts as a strong mirror responding to the moderate intershell expansion by a strong contraction of the external envelope; therefore a blue final location results.

In conclusion, uncertainties in mass loss rates may allow many models to fit some features of the supergiant distributions. However, the situation is not fully satisfactory regarding the He and N enhancements, the BHG, and the "ledge." These features are usually not predicted at solar *Z*. At lower *Z*, most models predict blue loops and better fit the observations, but even there the agreement does not seem complete for either the BHG, or the He or N abundances. It is essential that the models reproduce the observations at all metallicities and this is not yet achieved.

3.5.4 THE RATIO OF BLUE TO RED SUPERGIANTS (B/R) The B/R of supergiants was among the first stellar properties to be shown to vary through galaxies (van den Bergh 1968, Humphreys & Davidson 1979, Humphreys 1983a, Meylan & Maeder 1983, Humphreys & McElroy 1984, Brunish et al 1986). Studies have been made in the Milky Way, the LMC, the SMC, and also in M33 (Humphreys & Sandage 1980, Freedman 1985). The B/R depends on the range of luminosities considered, being found to be slightly larger at higher luminosities. The main trend is that B/R increases steeply with *Z*: for $M_{\rm bol}$ between -7.5 and -8.5. B/R is up to 40 or more in inner Galactic regions and only about 4 in the SMC (Humphreys & McElroy 1984). A difference in B/R by an order

of magnitude between the Galaxy and the SMC was also found on the basis of well-selected clusters (Meylan & Maeder 1982). Further studies of the young SMC cluster NGC 330 confirm the high number of red supergiants (Carney et al 1985).

Many star models are able to account for the occurrence of blue supergiants, with predicted B/R more or less in agreement with the observations (Brunish et al 1986). As already mentioned, one reason for this is the flexibility offered by the uncertain mass loss rates. Indeed, B/R may change from infinity in the case of no mass loss to about 0 for high mass loss rates. Thus, we emphasize that the real difficulty is not to account for some average observed B/R in the LMC or the Galaxy, but to account also for its change with metallicity. The models with Schwarzschild's criterion and overshooting (Schaller et al 1992, Alongi et al 1993), the models with the Ledoux criterion (Brocato & Castellani 1993), and the models with semiconvection (Arnett 1991), even if they are able to fit some average B/R, all appear to predict higher B/R at lower Z, in contradiction to the observations. This is a major problem, which is not solved by any of the published models.

We also point out an interesting change is the $T_{\rm eff}$ of red supergiants according to the metallicity of the parent galaxy. Red supergiants are hotter at lower Z, the difference amounting to about 800 K between models at Z = 0.001 and at Z = 0.040 (Schaller et al 1992). This is consistent with observations (Humphreys 1979, Elias et al 1985) which show that red supergiants in the Milky Way have spectral types between MO and M5, while in the SMC they are between types K3 and M2. This difference in the range of the spectral types of red supergiants is mainly the result of increased opacities at higher metallicities. We recall that a significant star formation rate is also a necessary condition for the presence of red supergiants in a galaxy. As an example, the paucity of red supergiants in M31 has been assigned to the low star formation rate rather than to the metallicity (Humphreys et al 1988). However, when B/R ratios are considered in galaxies, the effects of possible differences in the SFR and IMF are quite small and mainly result from effects of Z. These various tests show just how essential the studies of galaxies with different metallicities are for stellar evolution.

4. W-R STARS: OBSERVATIONS AND PREDICTIONS

4.1 Overview

Recent general reviews on Wolf-Rayet stars have been made by Abbott & Conti (1981), Willis (1987, 1991), Conti & Underhill (1988), Smith (1991a), van der Hucht (1991, 1992), Maeder (1991c), and Massey & Armandroff (1991). W-R stars are nowadays considered as "bare cores" resulting mainly from stellar winds peeling off of single stars initially more massive than about 25 to 40 $M\odot$. Close binaries might also lose their outer layers from Roche lobe overflow (RLOF). The main evidence for the bare core model as reviewed by Lamers et al (1991) are the following:

1. H/He ratios in W-R stars are low or zero.

2. The CNO ratios are typical of nuclear equilibrium (Section 4.2.1) in WN stars.

3. The continuity of the abundances in the sequence of types O, Of, WNL, WNE, WCL, WCE, and WO corresponds nicely to a progression in peeling off the outer material from evolving massive stars.

4. The observed $\dot{M}$ in progenitor O stars and in supergiants are high enough to remove the stellar envelopes within the stellar lifetimes. Also, the average $\dot{M}$ in W-R stars (Conti 1988) are able to accomplish further significant mass loss.

5. W-R stars have low average masses (between 5 and 10 $M\odot$; Abbott & Conti 1987); moreover, they fit well the mass-luminosity relation for He stars (Smith & Maeder 1989).

6. W-R stars are present in young clusters and associations with ages smaller than 6 Myr (Humphreys & McElroy 1984, Schild & Maeder 1984).

7. Transition objects Of/WN between Of and W-R stars and between LBV and WN stars exist (Section 3.3).

8. He- and N-rich shells are present around some W-R stars (Esteban & Vilchez 1991).

9. The W-R/O and WN/WC number ratios are consistent with theoretical expectations in galaxies with different Z (Section 4.5).

With their bright emission lines and their high luminosities, W-R stars are observable at large distances and are thus the stars for which we have the best sampling in other galaxies. Their emission lines also can become visible in the integrated spectrum of galaxies with active star formation, which enables us to extend the studies of young massive stars even farther out in the Universe. The discovery of the differences in W-R populations in galaxies has a long history starting with Roberts (1962) in the Milky Way and later Smith (1968 a,b,c) for the Magellanic Clouds. Further studies in the LMC and SMC (Azzopardi & Breysacher 1979, 1985; Breysacher 1981) confirmed the variations of W-R populations. These were also noticed (e.g. Kunth & Sargent 1981) in a sample of blue compact galaxies, which are dwarf galaxies with very active star formation, and in the so-called H II or W-R galaxies (Conti 1991a). The observed variations concern mostly the statistics, and in particular, ratios such as W-R/O or WC/WN.

The existence and origin of the variations of W-R populations in galaxies has been extensively debated over the past decade (see Section 4.6). Indeed, properties and statistics of W-R stars depend on many parameters: metallicity Z, star formation rate (SFR), initial mass function (IMF), age and duration of the bursts, binary frequency, etc. It is essential to distinguish between *(a)* regions in galaxies where the assumption of an average constant star formation rate over, say, the past 20 Myr is valid, and *(b)* regions or galaxies where a strong recent burst has recently occurred so that the assumption of a constant SFR does not apply. The first case concerns selected volumes in the Milky Way and in other galaxies (where individual W-R stars may be counted), where a stationary situation for star formation can be assumed. Within these regions, the effects of metallicity intrinsic to stellar evolution can be assumed to be the dominant factor responsible for the differences in W-R populations. This case is examined here first, as its proper understanding is a prerequisite for the studies of the second case (Section 5), which concerns distant GH II regions. W-R galaxies, and other starbursts.

4.2 Subtypes and Chemical Abundances

The basic physical parameters of W-R stars, i.e. their masses M, luminosities L, mass loss rates $\dot{M}$, radii R, and temperatures $T_{\rm eff}$, have been discussed by Conti (1988), Abbott & Conti (1987), and van der Hucht (1992). The M are in the range of 5 to 50 $M\odot$ with an average of about 10 $M\odot$: the L are between $10^{4.5}$ and 10^{6} $L\odot$; the $\dot{M}$ between 10^{-5} and 10^{-4} $M\odot$ / yr with an average of 4×10^{-5} $M\odot$ / yr; and the observed $T_{\rm eff}$ are between 30,000 K and 90,000 K, There are two main groups of Wolf-Rayet stars: subtypes WN and WC; a small additional subset is labeled WO (Barlow & Hummer 1982).

4.2.1 WN STARS The WN types are nitrogen-rich helium stars (Smith 1973) showing the "equilibrium" products of the CNO hydrogen burning cycle. The late WN stars of types WN6ÄWN9, abbreviated WNL (Vanbeveren & Conti 1980, Conti & Massey 1989), generally still contain some hydrogen with H/He ratios between 5 and 1 by number (Conti et al 1983a; Willis 1991; Hamann et al 1991, 1993; Crowther et al 1991). The early WN stars of types WN2-WN6, or WNE, generally show no evidence of hydrogen in their spectra. (There are a few exceptions to this general abundance relation with spectral subtype.) In a smaller sample of objects, Hamann et al (1993) found a correlation of the hydrogen content with the $T_{\rm eff}$ of WN stars, rather than with their subtypes; the coolest WN stars showed hydrogen and the hottest ones had none. The presence or absence of hydrogen certainly should be a determining factor influencing the opacity, temperature, and the structure of the outer layers.

The observed abundance ratios in mass fraction are C/He = (0.21-8) $\times$ 10^{-2}; N/He = (0.035-1.4) $\times$ 10^{-2}; and C/N (0.6-6.0) $\times$ 10^{-2} (cf Willis 1991, Nugis 1991). The corresponding solar ratios are respectively 1.1 $\times$ 10^{-2}, 3.3 $\times$ 10^{-3}, and 3.25 (Grevesse 1991 and references therein). The well-studied WN5 star HD 50896 gives similar results (cf Hillier 1987 a,b, 1988). Not much is known concerning the oxygen content of WN stars. Studies of ring nebulae around some WN stars also show strong overabundances of N and He with respect to the Sun (Parker 1978; Esteban & Vilchez 1991, 1992; Esteban et a1 1992, 1993). These authors suggest that the ring nebulae have been ejected at the end of the red supergiant phase. In our opinion, it is more likely that a large part of the shell ejection, or even most of it, occurs during the first thousand years after the entry in the WNE and WC stages, which are marked by extreme mass loss, as predicted by the M vs M relation for W-R stars (cf Langer 1989b). Among WN stars, eight have inordinately strong CIV lines (Conti & Massey 1989). Labeled WN/WC stars, these are suggested to be transition objects between WN and WC stars (Section 4.5).

The observed abundance ratios span the range of equilibrium values of the CNO cycle (Maeder 1983, 1991b), with C/N and O/N ratios two orders of magnitude smaller than solar. Interestingly enough, such values are essentially independent of the various model assumptions and mainly reflect the nuclear cross sections and initial composition. The good agreement between the observed and predicted values of CNO equilibrium indicates the general correctness of our understanding of the CNO cycle and of the relevant nuclear data.

The initial CNO content, which depends of the initial *Z*, determines the amount of nitrogen in WN stars (Maeder 1990, Schaller et al 1992). The N abundance is thus lower for lower initial *Z*, but equilibrium ratios such as C/N are predicted to be independent of the initial Z. In this connection, one can understand the result by Smith (1991a) who noticed that the ratio of the λ 4686 He II line to the λ 4640 N III line is stronger for WNL stars in the LMC compared to the Galaxy.

4.2.2 WC STARS WC stars contain no hydrogen and, as a result of mass loss, are mainly He, C, and O cores as a result of mass loss (Smith & Hummer 1988, Torres 1988, de Freitas-Pacheco & Machado 1988, Hillier 1989, Willis 1991, Nugis 1991, Eenens & Williams 1992, de Freitas Pacheco et al 1993). These stars represent objects in which we see at the surface the result of triple-α and other helium burning reactions. A most interesting finding is the one by Smith & Hummer, who showed that the C/He ratio is increasing for earlier WC subtypes. Smith & Maeder (1991) emphasize that a measured (C+0) / He ratio is to be preferred to the C/He and C/O ratios which go up and down during helium processing. They propose the following calibration in (C+O) / He number ratios: WC9, 0.03-0.06; WC8, 0.1; WC7, 0.2; WC6, 0.3; WC5, 0.55; WC4, 0.7-1.0; WO, > 1.

The sequence of types WC9 to WO appears as a progression in the exposure of the products of He burning. The rare WO stars (Barlow & Hummer 1982, Kingsburgh & Barlow 1991, Polcaro et al 1992, Kingsburgh et al 1994) simply appear to be the most extreme type in this sequence.

Comparisons of observations and model predictions show a generally good agreement (Willis 1991, 1994): however, closer comparisons (Schaerer & Maeder 1992) suggest that $\dot{M}$ in previous evolutionary phases could be higher by a factor of 2 with respect to current values by de Jager et al (1988). The above connection between WC sub-types and the (C+O) / He ratios is the key to understanding the Z-dependence of the distribution of WC stars in galaxies of different metallicities (Section 4.6).

A present uncertainty concerns neon in WC stars. Models predict a substantial abundance of neon (larger than 0.03 in mass fraction) - essentially Ne^{22} at high initial Z and Ne^{20} at low Z (Maeder 1991a). However, the only available data, which come from *IRAS* observation of Ne II at 15.5 microns, indicate an abundance of 0.005 in the WC8 star γ Vel (Barlow et al 1988). Some of the nuclear cross sections in the chain leading to Ne^{22} are still very uncertain and the ashes of N^{14} could possibly be stocked in the form of O^{18} (indistinguishable from O^{16} in W-R stars) rather than in the form of Ne^{22}. The problem is of importance for explaining the role of WC stars as a possible site for s-elements (Prantzos et al 1990), since these elements should be formed by Ne^{22} (α , n)Mg^{25}. The role of W-R stars as producers of radioactive Al^{26}, detectable through y ray observations, does not seem important according to Signore & Dupraz (1990), but according to Meynet (1994a) their role could be as important as that of supernovae.

4.3 Physical Properties

In addition to the usual model ingredients, the W-R star models specifically require special attention on a number of points. Concerning microphysics, the W-R models demand Rosseland opacities for the appropriate He-C-O mixtures (Iglesias & Rogers 1993) and detailed calculations of the ionization balance for heavy elements (Langer et al 1986, Schaller et al 1992). The Ms in stages previous to the W-R phases and their dependence on Z are very critical. Also, the adopted definitions (based on surface abundances) for the transitions from LBV to WNL, WNL to WNE, and WNE to WC are of importance for the comparison of models and observations.

For $\dot{M}$ in the W-R stage, the average observed rates (Abbott et al 1986, Conti 1988) have often been used. However, these rates have led to masses and luminosities that are too high with respect to the observations (Schmutz et al 1989). A number of convergent suggestions have been provided recently in favor of a mass dependence of $\dot{M}$ in WNE and WC stars. In particular, models by Langer(1989b) suggest a relation of the form: $\dot{M}$ (W-R) = (0.6-1.0) $\times$ 10^{-7} $(M / M\odot)^{2.5}$ $M\odot$ / yr, where the first coefficient applies to WNE and the second to WC stars. Similar mass-dependent $\dot{M}$ s have been provided from binaries (Abbott et al 1986, St Louis et al 1988), and from modeling the wind properties (Turolla et al 1988, Bandiera & Turolla 1990, Schaerer & Maeder 1992). The $\dot{M}$ vs M relation generally leads to an enormous mass loss at the entry in the WNE stage, which results in very low final W-R masses. This has a considerable impact on the chemical yields of massive stars (Maeder 1992a), also resulting in an increase of the W-R lifetimes. A question remains as to whether the WN luminosities predicted from models with standard $\dot{M}$ (de Jager et al 1988) are not too high with respect to the observed ones (Howarth & Schmutz 1992). Indeed, the relatively low observed luminosities of some WN stars support larger $\dot{M}$ in previous stages. There are at present no indications of a mass dependence of $\dot{M}$ for WNL stars, although such a relation would not be too surprising.

The maximum mass for the vibrational stability of a He star is about 16 $M\odot$ (Noels & Maserel 1982, Noels & Magain 1984). Thus, if a star enters the helium configuration with a mass larger than critical (which occurs in current models), it may be expected to be vibrationally unstable with high mass loss as a consequence (Maeder 1985). The fact that regular pulsations have been recently observed in the WN8 star WR 40 (Blecha et al 1992) might give some support to this claim. However, the nature of

these pulsations is still under discussion (Kirbiyik 1987) and different pulsation modes have been proposed by Glatzel et al (1993) and Kiriakidis et al (1993). Attempts are also being made to explain the strong W-R winds by multi-scattering and purely radiative processes (Pauldrach et al 1988, Cassinelli 1991), by radiation and turbulence (Blomme et al 1991), or by radiation and Alfven waves (Dos Santos et al 1993). The main difficulty, which is not satisfactorily resolved, is to explain why the wind momentum of W-R stats may be up to 30 times the photon momentum (Barlow et al 1981, Cassinelli 1991; but see Lucy & Abbott 1993).

The problems of the atmospheres of hot stars have been reviewed by Kudritzki & Hummer (1990) and the different definitions of the radii and $T_{\rm eff}$ in extended atmospheres by Bascheck et al (1991). Values of hit have been given recently by Conti (1988), Schmutz et al (1989, 1993), and Koesterke et al (1992): They range between about 3×10^4 and 10^5 K. In order to compare the observed $T_{\rm eff}$ with data from interior models, a simple correction scheme has been proposed to roughly account for the optically thick winds of W-R stars (de Loore et al 1982, Langer 1989a). More refined procedures have been established by Kato & Iben (1992), by Schaller et al (1992), and in particular by Schaerer & Schmutz (1994). The net result is that from a surface temperature of 1-1.5 $\times 10^5$ K (without the wind), the W-R stars are shifted down to $T_{\rm eff}$ 3-10 $\times 10^4$ K according to their $\dot{M}$ rates, and thus also according to their masses and luminosities since there is a M-L-$\dot{M}$ relation.

Since W-R stars of types WNE and WC are He-C-O cores, they have a rather simple internal structure with little compositional difference between center and surface. W-R properties and the relations between the subtypes are mostly independent of their formation. Evolutionary models predict M-L-$\dot{M}$ -$T_{\rm eff}$ relations (Schaerer & Maeder 1992) for W-R stars without hydrogen. The mass-luminosity M-L relation is (Maeder 1983, Langer 1989a, Beech and Mitalas 1992, Schaerer & Maeder 1992):

$$\log L/L_\odot = 3.03 + 2.695(\log M/M_\odot) - 0.461(\log M/M_\odot)^2. \qquad (2)$$

For $M > 10\ M\odot$, a linear relation may be appropriate. On the observational side, the M-L relation has been confirmed (Smith & Maeder 1989, Smith et al 1994). The $\dot{M}$ vs M relation is supported by binary observations as discussed above. The M-L-$\dot{M}$ -$T_{\rm eff}$ relations also indicate that WNE stars and WC stars should follow well-defined tracks in the HR. diagrams. Such alignments seem to be present in the data of Hamann et al (1991, 1993; but see also Maeder & Meynet 1994).

For WNL stars, the luminosities are generally higher than for WNE stars (Conti 1988). Models indi1cate that WNL luminosities are mainly related to the initial masses. The reason is that the luminosity depends on the size of the He cores, which is determined mainly by the initial mass rather than by the actual mass, as long as the He cores are not themselves peeled off. Models also suggest, in agreement with observations (cf Hamann et al 1993), that the WNL $T_{\rm eff}$ are mainly determined by the remaining hydrogen content.

4.4 Initial Masses, Lifetimes, Formation

Observationally, most W-R stars appear to originate from stars initially more massive than about 40 $M\odot$ (Conti et al 1983b, Conti 1984, Humphreys et al 1985, Tutukov & Yungelson 1985). From the presence of W-R stars in clusters down to type BO, it is clear that a few W-R stars may originate from initial masses down to 20-25 $M\odot$ (Firmani 1982, The et al 1982, Schild & Maeder 1984). The modeling of W-R ring nebulae (Esteban et al 1992) also supports the above values of initial masses.

The minimum mass for forming WC stars does not seem significantly higher than that for WN stars.

Maeder & Meynet (1994) have obtained the lifetimes in the W-R stage for two different cases: 1. the standard case with mass loss rates by de Jager et al (1988) in pre-W-R stages and the scaling with $Z^{0.5}$ at other metallicities; and 2. the case with M arbitrarily twice as large as in pre-W-R stages. Indeed, several observations, in particular the chemical abundances in WC stars (Section 4.2), the W-R luminosities (Section 4.3), and the number ratios of W-R stats (Section 4.5) clearly support the case of enhanced mass loss, for which the W-R lifetimes are shown in Figure 1. From this figure we note that:

- The minimum mass for forming W-R stars ranges from about 20 to 25 $M\odot$ at Z = 0.04 to about 80 $M\odot$ at Z = 0.001. The lower mass value is in agreement with the observations in the Milky Way. The change of the minimum initial mass with Z is a key effect to explaining the W-R statistics in galaxies.

- The lifetimes in the W-R stage go up with M and Z, which is just the opposite of the behavior on the main sequence. The average W-R lifetime weighted by the IMF is about 0.6 Myr at Z = 0.02 (for $\dot{M}$ rates twice the standard ones). Figure 1 shows lifetimes up to about 2 Myr in extreme cases.

- A detailed inspection of the model results indicates that the largest initial stellar masses spend most of the W-R phase in the WNL stage (Langer 1987, Maeder 1991a). If the WNL phase is defined from the H abundance at the surface (Section 4.2.1), then the WNL stage can even be entered during the *main sequence phase of the most massive stars*.

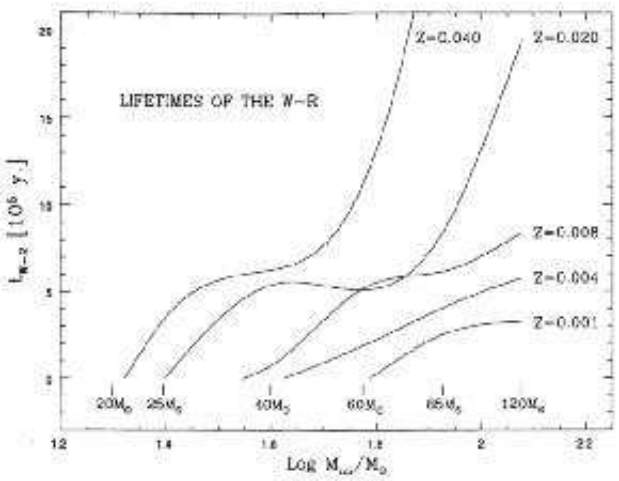

Figure 1. Durations of the W-R phase as a function of initial stellar masses for different Z in the case of enhanced mass loss rates (from Maeder & Meynet 1994).

- The lower initial masses have very short WNL phases, and spend much more time in the WNE and WC phases.

The formation of W-R stars is largely dominated by the overwhelming effects of mass loss, as first proposed in the "Conti" scenario (Conti 1976). Mixing processes due to rotation or tidal distortion in binaries may favor in some cases the formation of W-R stars and increase their lifetimes (Maeder 1981, 1987b). Semiconvection or some mild mixing at the edge of the He core seems necessary to account for the existence of intermediate WN/WC stars (Langer 1991b). These stars represent about 4% of the W-R stars (Conti & Massey 1989), while models without extra mixing only predict 1% or

less of WN/WC stars. These stars cannot be explained by binary evolution (Vrancken et al 1991). Some additional mixing is necessary, but at the same time the small observed fraction of WN/WC stars implies that the part of the stellar mass that is actually mixed is quite small, and this puts a limit on the role of mixing at the edge of the He core.

Detailed investigations of W-R binaries have been carried out in the Galaxy by Massey (1981) and in the Magellanic Clouds by Moffat 1988 and Moffat et al 1990. Binary mass transfer by RLOF, which is an extreme case of tidal interaction, may contribute to the formation of WR+O binaries (de Loore 1982; De Greve et al 1988; Vanbeveren 1988, 1991, 1994; Schulte-Ladbeck 1989; De Greve 1991, de Loore & Vanbeveren 1994). The fraction of all stars (single + binaries) undergoing RLOF is estimated to be between 20 and 40% (Podsiadlowski et al 1992). We may note that this percentage also includes binaries that could be mixed by tidal interactions and would thus evolve homogeneously, without large increase of their radius and thus without RLOF, Indeed, the importance of RLOF in W-R+O binaries is still unclear. From the similarity of the relatively large orbital eccentricities in W-R+O and O+O binaries, Massey (1981) concluded that mass transfer probably did not play a *major* role in the formation of W-R+O binaries. We may conjecture that several effects contribute to the formation of W-R stars; it is likely that the relative importance of these effects changes with Z as discussed below.

4.5 W-R Statistics

4.5.1 BASIC DATA AND ITS INTERPRETATION W-R stars are observed in several galaxies of the Local Group, and provide statistical data on their relative frequencies at various metallicities. In the Milky Way, the catalogs by van der Hucht et al (1988), and by Conti & Vacca (1990) provide rather complete samples up to about 2.5 kpc. These data show that the number density of W-R stars projected onto the Galactic plane is strongly increasing with decreasing galactocentric distance. Deep surveys are extending the sampling (Shara et al 1991). The LMC and SMC catalogs are cornerstones for data at other Z (Azzopardi & Breysacher 1979, 1985; Breysacher 1981, 1986). A few additional W-R stars have also been identified (Morgan & Good 1985, Testor & Schild 1990, Schild et al 1991, Morgan et al 1991). Most (75%) of the W-R stars in the center of the 30 Dor Nebula are WNL stars of types WN6-WN7 (Moffat et al 1987), while in the surroundings the proportion is much lower. The excess of WNL stars in giant H II regions is a common feature, as evidenced by those in M33 (Drissen et al 1990, 1991). The subtype distribution of W-R stars in the Magellanic Clouds has been considered by Smith (1991b). Data on W-R stars in M31 have been obtained by Moffat & Shara (1983, 1987), Massey et al (1986, 1987a), Armandroff & Massey (1991), and Willis et al (1992). For M33, studies have been made by Wray & Corso (1972), Conti & Massey (1981), Massey & Conti (1983), Massey et al (1987 a, b), Schild et al (1990), and Armandroff & Massey (1991). In both M31 and M33, the samples are still incomplete. In the two small galaxies NGC 6822 and IC 1613 of the Local Group, many W-R stars were proposed by Armandroff & Massey (1985) and Massey et al (1987a), but further analyses by Azzopardi et al (1988; see also Smith 1988) confirmed only four W-R stars in NGC 6822 and one, which was already found by Davidson & Kinman (1982), in IC 1613. The study of W-R stars in other galaxies is continuing. The galaxy IC 10 at 1.5 Mpc exhibits a very high density of W-R stars with a WC/WN number ratio of about 0.5 (Massey et al 1992). Ten individual W-R stars have been detected in the galaxy NGC 300 at a distance of 1.5 Mpc (Schild & Testor 1991, 1992).

Analyses of W-R star statistics in nearby galaxies have been made by Azzopardi et al (1988), Smith (1988), Massey & Armandroff (1991), Maeder (1991a), and Maeder & Meynet (1994). Table 2 gives the available data on the W-R/O, WC/W-R, and WC/WN ratios for galaxies of the Local Group, when an indication of the metallicity Z is available. The table is adapted from Maeder (1991a) with revisions according to recent data from Conti & Vacca (1990) for the Milky Way; for M31 the W-R/O is from Cananzi (1992); for the SMC the new W-R star found by Morgan et al (1991) is included; the WN/WC ratios are in agreement with those found by Armandroff & Massey (1991). Due to small

number statistics, the ratios for NGC 6822 and IC 1613 are not significant.

Such number ratios are to be preferred to surface densities, which would depend not only on stellar evolution but also on the current SFR. For the Galaxy, the statistics for O stars is based on the survey by Garmany & Conti (1982). In other galaxies, the numbers of O stars were estimated by Azzopardi et al (1988) on the basis of UV data from Geneva and Marseille balloon experiments and on the basis of Lequeux's (1986) luminosity function. These indirect estimates lead to larger uncertainties in the numbers of O stars than in those for W-R stars.

The origin of the observed variations of the relative number of W-R stars in different environments was attributed to metallicity by Smith (1973) and Maeder et al (1980), who suggested that high Z favors mass loss, which in turn favors the formation of W-R stars. The variations of W-R subclasses in M31 were also attributed to a metallicity effect by Moffat & Shara (1983). The total dependence on Z was criticized by several authors, who attributed the differences in the W-R populations mainly to changes in the IMf and SFR (Bertelli & Chiosi 1981, 1982; Garmany et al 1982; Armandroff & Massey 1985; Massey 1985; Massey et al 1986; Massey & Armandroff 1991). As suggested by these authors, it is possible that prominent departures from the assumption of constant star formation, such as is the case in 30 Dor in the LMC (Moffat et al 1987) or in giant HII regions of M33 (Conti & Massey 1981, Drissen et at 1990) where recent bursts of SFR have occurred, may produce peculiar W-R number ratios (Section 5). However, we note that no systematic difference in the IMF slope has been found between the Galaxy, the LMC, and SMC (Humphreys & McElroy, 1984, Mateo 1988, Massey et at 1989a, Parker et al 1992, Section 2). Also, the Galactic gradient of the surface density of W-R stars is much steeper than that of their precursor O stars (Meylan & Maeder 1982, van der Hucht et al 1988) - a fact that is reflected by the changes of the W-R/O in Table 2. Thus, it is likely that the basic effect in stellar evolutionary models is Z, and that effects connected to the SFR and perhaps to the IMF are population parameters that may also influence the relative frequencies of W-R stars in G HII regions and bursts.

Table 2. Observed W-R/O, WC/W-R, and WC/WN in galaxies of various metallicities

GALAXY	Z	W-R/O	WC/W-R	WC/WN
M31	0.035	0.24	0.44	0.79
MILKY WAY				
ring 6-7.5 kpc	0.029	0.205	0.55	1.22
ring 7.5-9 kpc	0.020	0.104	0.48	0.92
ring 9.5-11 kpc	0.013	0.033	0.33	0.49
M33	0.013	0.06	0.52	1.08
LMC	0.006	0.04	0.20	0.26
NGC 6822	0.005	0.02	-	-
SMC	0.002	0.017	0.11	0.13
IC 1613	0.002	0.02	-	-

4.5.2 PREDICTED W-R/O AND WC/WN VALUES As illustrated in Table 2, the W-R/O ratio increases with the metallicity of the parent galaxy (Maeder et al 1980; Azzopardi et al 1988; Smith 1988, 1991a). This general trend is also confirmed by studies of the integrated properties of H II or W-R galaxies (Arnault et al 1989; Conti 1991 a,b; Smith 1991a; Vacca & Conti 1992; Mas-Hesse & Kunth 1991 a,b; Mas-Hesse 1992). In stellar models a growth of the W-R/O with Z is predicted (Maeder 1991 a, Maeder & Meynet 1994), resulting from the lowering of the minimum initial mass for forming W-R stars and from the increase of the lifetimes with increasing Z (and mass loss). Figure 2 compares the observations of Table 2 with theoretical values for models with enhanced $\dot{M}$ as defined in Section 4.4 and for stars with a Salpeter IMF. The general agreement is quite good, confirming that metallicity is a key factor in the variations of the W-R/O ratio. Although some scatter appears, it might be due to local departures from the simple assumption of a past constant SER and to the averaging over some range of Z in large galaxies (Smith 1991a). No satisfactory agreement between observed and predicted W-R/O can be achieved for models with standard $\dot{M}$ (de Jager et al 1988). The differences would be especially large at high Z, while at $Z = 0.002$, the W-R/O values are in both cases of mass loss equal to about 0.005. Such negligible W-R/O values at low Z are in agreement with the low fraction of W-R stars observed in metal deficient galaxies. Also, the study of the integrated spectrum and He II 4686 Å feature in dwarf galaxies shows a general absence of W-R contribution for galaxies with very low oxygen content, corresponding to about $Z = 0.002$ (Arnault et al 1989, Smith 1991a).

The observed WC/WN and WC/W-R numbers in Table 2 show a general growth with increasing Z, but it is not monotonic and shows an appreciable scatter. This was noted by Armandroff & Massey (1991) and Massey & Armandroff (1991) as an argument supporting the fact that metallicity is not the only determining factor for explaining the W-R statistics. From models, the change in WC/WN results from the higher $\dot{M}$ which leads to an earlier visibility of the products of He burning (Maeder 1991a, Maeder & Meynet 1994). Interestingly enough, for larger M the predicted WC/WN ratios, instead of further increasing as expected, go down again as Z becomes greater than 0.01. This occurs because the WN phase of the most massive stars has already been entered during the main sequence phase and is therefore much longer. This model result accounts for the nonmonotonic behavior of WC/WN found by Armandroff & Massey (1991).

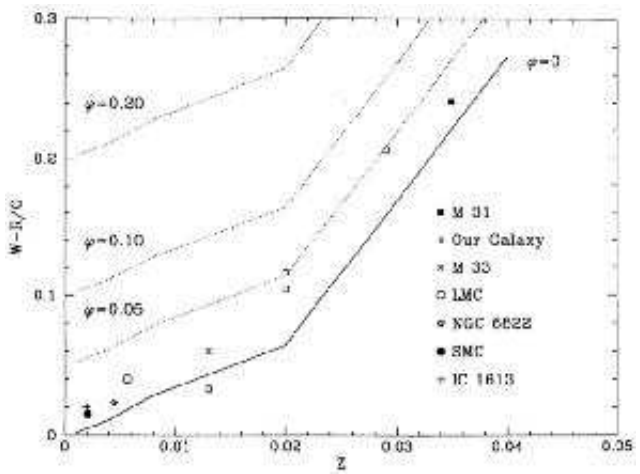

Figure 2. W-R/O as a function of Z in nearby galaxies compared to model predictions by Maeder & Meynet (1994). The solid line represents the predictions of single star models with enhanced mass loss rates (Section 4.4) and Salpeter's IMF. The dotted lines show the same for different values φ, the fraction of O stars undergoing mass transfer in binaries.

The comparison between observed and theoretical WC/WN or WC/W-R shows, as for the W-R/O, that a better fit is obtained for models with enhanced $\dot{M}$ in previous phases. Nevertheless, the scatter is still there, and reflects departures from the assumption of a constant SFR. Such departures are prominent in some giant H II regions, such as 30 Dor in the LMC, or NGC 592, 595, and 604 in M33, which show evidence of intense star formation (Conti & Massey 1981; Drissen et al 1990, 1991). Armandroff & Massey (1985) and Massey & Armandroff (1991) noticed that regions with metallicity similar to that of the LMC and SMC have different W-R numbers. Smith & Maeder (1991) suggested that in large spirals the W-R populations will he heavily weighted toward properties of high *Z* values. Thus both the effects of bursts of star formation (Section 5) and of the averaging over *Z* may contribute to increases in the W-R/O and WC/WN ratios, a situation which may apply particularly to M33.

Close comparisons between models and observations must also account for the various channels of W-R formation and in particular the Roche lobe overflow (RLOF) in close binaries (Vanbeveren 1991, 1994; Vanbeveren & de Loore 1993). The fraction of O stars becoming W-R stars as a result of RLOF was estimated by the latter authors to be about 35% (see also Podsiadlowski et al 1992). A new analysis quoted above (Maeder & Meynet 1994) suggests that the fraction of W-R stars owing their existence to RLOF is highly variable with the metallicity of the parent galaxy; it is nearly 100% at low *Z*, like in the SMC (Smith 1991b), and lower than 10% in the inner regions of the Milky Way.

The above results are compatible as shown in Figure 2 with a relatively low fraction Φ , at most 10%, of the ensemble of the O stars that become W-R as a result of binary mass transfer. Such a low fraction is consistent with by Massey's (1981) result on the distribution of the eccentricities of WR+O binaries. A close investigation of the young cluster Tr 14 (Penny et al 1993) reveals a general absence of close binaries among the brightest O stars. In further support of a low fraction of O stars undergoing RLOF, we remark that at the low *Z* in the SMC, the W-R/O ratio is about 0.017, while it is 0.21 in inner Galactic regions. Thus, if the fraction of O stars becoming W-R as a result of RLOF is the same in both areas, this fraction would be at most 0.017/0.21 = 8%, assuming that all W-R stars in the SMC are formed by RLOF. This fixes the upper limit of the fraction of O stars undergoing RLOF under the mentioned assumption.

Stellar winds produce fewer W-R stars at low *Z*, and only the binary channel seems efficient in forming WR stars. As an example, in the SMC, eight of the nine observed W-R stars may be binaries; five are confirmed (Moffat 1988, Smith 1991a). These binaries were likely formed by mass transfer (De Greve et al 1988). They are of type WNE, which suggests that the binary channel may mainly lead to WNE stars. Consistent with the SMC observations, the proportion of W-R binaries has been found to be larger toward the anticenter in the Milky Way (van der Hucht et al 1988). The hypothesis that a small fraction of the W-R stars is formed by RLOF also gives a better fit to the observed behavior of the WNE/W-R and WNL/W-R ratios with *Z* (Maeder & Meynet 1994). We may also note that the average ages of WNL and WNE stars are probably not the same; Moffat et al (1991) suggest that the former are relatively younger than the latter; WNL stars are also more luminous than WNE stars (Conti 1988). These facts are consistent with the result (Section 4.5) that WNL stars mainly originate from the most massive stars.

4.5.3 WC SUBTYPE STATISTICS The distributions of WC and WO stars in galaxies exhibit a number of distinct properties.

1. It is well known that WC stars are relatively more numerous in inner galactic regions (cf van der Hucht et al 1988, Conti & Vacca 1990). More specifically, the later WC subtypes - WC9 and WC8 - are only found in inner galactic regions with higher *Z*, while outer regions with lower *Z* contain WC stars of earlier subtypes, mainly WC6-WC4. This is also true in the LMC, where

only WC6-WC4 subtypes are found, with very uniform properties (Breysacher 1986, Smith et al 1990, Smith 1991b). In more extreme low-Z dwarf galaxies, the rare WC stars only belong to subtypes WC4 and WO. In M31 as in the Milky Way, late WC stars (WCL) are found in inner galactic regions and early ones (WCE) in outer regions (Moffat & Shara 1987). However, the metal-rich galaxy M31 contains no WC9 and WC8 stars (Massey et al 1987a).

2. The luminosity of earlier WC subtypes is lower than that for later WC subtypes (Lundstrom & Stenholm 1984, van der Hucht et al 1988, Conti 1988).

3. Stars of a given WC subtype seem brighter in a galaxy with lower *Z* as suggested by Smith & Maeder (1991), who point out that LMC WC4 stars are brighter than Galactic WC5-WC6 stars.

These various facts can be easily understood on the basis of recent models and of the relation between WC subtypes and the (C+O)/He ratios as shown by Smith & Maeder (1991). The entry points and lifetimes in the WC9 to WO sequence are very dependent on M and *Z*. At high *Z* and high M, due to high mass loss the WC stage is entered very early during the He burning phase, so that the surface (C+O)/He ratio is very low, which implies a type WC9 or WC8. As evolution continues, mass and luminosity decline and (C+O)/He decreases and thus the sequence of types WC9 -> WO is described. The entry point in that sequence occurs at lower L and earlier WC types for lower masses. At lower *Z*, as in the LMC, the entry in the WC phase occurs at a later stage of central He-burning, i.e. with higher O/C ratios (Smith et al 1990) and also with higher (C+O)/He ratios at the surface, which means earlier WC types, typically WC5-WC4 (Smith & Maeder 1991). Then, the further evolutionary sequence described is short. This behavior also explains the two above-mentioned points concerning the luminosity of WC stars at different *Z*. That the luminosity for a WC subtype depends on the initial *Z* may have consequences for the interpretation of WC lines in integrated spectra of galaxies.

From W-R stars in clusters and in galaxies, some relationships between subtypes can be understood (van der Hucht et al 1988; see also Schild & Maeder 1984, Moffat et al 1986, Moffat 1988, Schild et al 1991): For galactocentric radius $R < 8.5$ kpc, we have WNL -> WCL; while for $R > 6.5$ kpc, we find WNL -> WCE -> WO and WNE -> no WC stars.

These connections are supported by the model results (Maeder 1991a). The first connection is typical of high mass and high *Z*: A long WNL phase, followed by a negligible WNE phase, emerges on the late and luminous part of the WC sequence, typically at WC9 or WC8. The second connection is typical of large masses with solar or lower metallicity, while the third one corresponds to the lowest part of the mass range able to form W-R stars.

We conclude this section by underlining that the observations and understanding of W-R stars in nearby galaxies has brought the clarification of many problems, which also have far-reaching consequences for the injection of mass, momentum, and energy into the interstellar medium (Leitherer et al 1992a).

5. MASSIVE STARS IN STARBURSTS

5.1 Integrated Spectra of Galaxies

The spectra of galaxies are primarily those of the underlying stellar population (absorption lines) with the addition of nebular emission lines from the GH II and young starburst regions which are present if recent star formation has occurred. We first consider how one might infer the distribution of massive stars from the nebular line analyses, and then discuss spectral synthesis of the stellar features.

5.1.1 NEBULAR LINE ANALYSES The number of exciting stars of ionized hydrogen regions (NO*) can readily be estimated from analysis of the emission line spectra (e.g. Kennicutt 1984, Shields 1990). One observes the spatially integrated Hα flux (or the Hβ or radio recombination measure), accounts for the extinction (if necessary), and from simple recombination theory infers the total number of Lyman continuum photons (*N* Lyc) being emitted. This last step contains the nominal nebular analysis assumptions of "Case B" - no dust, and ionization bounded. The *N* Lyc is a direct measure of the number of exciting stars. The total number released is a product of the numbers of hot stars, the slope of the IMF, and the *N* Lyc at each spectral subtype.

Vacca (1991, 1994) has thoroughly reviewed and quantified this procedure. He introduces the parameter η_0 as defined in the equation

$$NO^{\star} = NO7V/\eta_0 \qquad (3)$$

where *N O7V* is the number of "equivalent" O7V type stars, and *NO** is the total number of O stars. The *N O7V* is simply the observed *N* Lyc divided by the number of Lyman continuum photons emitted by an O7V star (1×10^{49} s^{-1}). The η_0 can be calculated as a function of Γ, $T_{\rm eff}$, and log *g* using the Kurucz stellar atmosphere models. Its value, tabulated by Vacca, depends also on $M_{\rm upper}$, and the $M_{\rm lower}$ for Lyman photon production - roughly the OB star boundary. This latter parameter depends on the metal abundance *Z* via the stellar structure models. This procedure makes the assumption that all O stars in H II and GH II regions are main sequence; one can separately allow for massive star evolution.

How accurately does this method work? It has been calibrated using 30 Dor in the LMC. Parker (1991, 1993) and Parker & Garmany (1993) have made a detailed census of the hot star population in an 7' × 7' area centered on R136. They find about 400 O stars. Vacca (1991, 1994) has analyzed the nebular spectrum of a trailed, 15-minute spectrophotometric exposure of an 8' × 8' area scan centered on R136 (taken by M. Phillips). Using the Z value for the LMC and Parker's value for Γ (- 1.4), he finds η_0 to be 0.44 and estimates that about 330 O stars are present, within 30% of the actual count (allowing for the slight difference in areas)! This gives us confidence in the procedure which has been applied to W-R and other emission line galaxies by Vacca & Conti (1992) as we show below (Section 5.1.4).

5.1.2 EMISSION LINE GALAXIES - STARBURSTS These galaxies with emission line spectra like those of H II regions were noted by Sargent & Searle (1970). Given the often substantial numbers of O-type stars found within these galaxies, they can be understood to be examples of very young "starbursts." Using slit spectroscopy, one may obtain the numbers of exciting stars by a procedure similar to that outlined above for 30 Dor. For galaxies, however, we might have additional complications if the slit width does not include the entire region of interest, if the nebulosity is density bounded, or if dust is present in sufficient quantities to absorb a considerable fraction of the Lyman photons.

5.1.3 STARBURST MODELS Several models for starburst populations have been produced to simulate various properties of the galaxy spectra such as: the overall stellar and nebular spectrum (Leitherer 1990, 1991; Leitherer et at 1992b; Bernlohr 1992, 1993), the strengths of the Si IV and CIV UV wind features emitted by O stars (Leitherer & Lamers 1991; Mas-Hesse & Kunth 1991 a,b), the widths of these UV lines (Robert et al 1993), the far-infrared and radio emission by the dust (Mas-Hesse 1992, Desert 1993), the emission line ratios such as λ4686 He II / Hβ or the so-called W-R "bump" λ4650 / Hβ (Arnault et al 1989, Meyner 1994b), or other line ratios such as λ4686He

II / λ4650CIII, which is sensitive to the WN/WC ratio (Kruger et al 1992).

The basic physical parameters for a starburst model are the star formation rate - in particular the intensity and duration of the burst, the age after its beginning, the IMF, and Z. The hope of starburst models is to disentangle these various parameters. Let us consider the didactical, but not unrealistic, case of an instantaneous "burst" (formation over a time of up to 1 Myr - small with respect to the massive star evolution time). In this case, an evolved W-R population with its prominent emission features results from only a part of the mass range of the potential W-R progenitors, while in the case of a constant SFR the W-R population results from an equilibrium mixture of the whole potential mass range. We may schematically distinguish four different epochs in the evolution of a burst according to recent models of massive stars and W-R stars (Maeder & Meynet 1994):

1. O-phase: For an age $t \leq 2$ Myr, massive stars are in their O-type phase, giving rise to H II regions without W-R features.

2. WNL phase: From $t = 2$ Myr to about 3 Myr, a large number of W-R stars are present, nearly all of them are of WNL subtype.

3. WC+WNL+WNE phase: After $t = 3$ Myr, the three subtypes may coexist with fractions depending on the mass loss and/or *Z*. At solar *Z* and standard M, WC stars dominate, followed by WNL in the first part of the period and by WNE stars in the second part. The higher the $\dot{M}$ and *Z*, the more numerous are the WC stars. At lower *Z* and $\dot{M}$, single star evolution only leads to WNL stars and very few WC stars, but we may also expect (Section 4.5.2) a fraction of W-R stars of WNE type from RLOF binary evolution.

4. O-phase (again): For $t \geq 7$ Myr, the W-R stars have disappeared but at up to 10 Myr there are still O-type stars to produce H II region nebular emission line features. One way to distinguish between this phase and the first O-phase might be to examine the equivalent width of Hβ nebula emission; according to Copetti et al (1986) this parameter steadily decreases in value with age in model H II regions (while the number of Lyc photons remains more or less steady, the starlight at λ4860 steadily increases).

During these phases of a burst, the various ratios of W-R subtypes to O stars are much larger than in the case of a constant SFR. As an example, at solar *Z* the average WNL/O ratio is up to six times larger, and the enhancement is even greater at lower *Z*. The reason is that the duration of the W-R-rich phase is shorter at lower *Z*, meaning that stars of only a narrow mass range become W-R stars, thus the contrast between the cases of a burst and of constant SFR is much larger. For bursts longer than 1 Myr, the situation is intermediate between the instantaneous burst and constant SFR. We also notice that if an observed H II region consists of a burst plus a region of lower but constant SFR, we have for the 2 Myr after the burst a much lower W-R/O ratio than for constant SFR, since at this time the W-R stars from the burst have not yet appeared. On the whole, then, one must he cautious before making quantitative inferences from number ratios alone.

5.1.4 WOLF-RAYET GALAXIES These are a subset of emission line galaxies in which, in addition to the nebular line spectrum, one observes broad emission at λ4686 Å due to the presence of Wolf-Rayet *stars* (Conti 1991 a,b, and references therein). The starburst phenomena illustrated by W-R galaxies represent an extreme burst of star formation, in which hundreds to thousands (or more) of massive stars have been born. There are currently about 50 examples of such systems, most of which have been discovered serendipitously. Many are Markarian or Zwicky galaxies and exhibit disturbed morphologies which may be the result of interactions or mergers. Examples of W-R galaxies may be found among Blue Compact Dwarf Galaxies (Sargent 1970), isolated extragalactic H II

regions (Sargent & Searle 1970), dwarf irregulars (Dinerstein & Shields 1986), "amorphous" galaxies (Walsh & Roy 1987), spiral galaxies containing knots or GH II regions (Keel 1982), recent galaxy mergers (Rubin et al 1990), or powerful *IRAS* galaxies (Armus et al 1988).

The broad emission features are seen in contrast to the galaxy stellar population continua. The dilution is such that the few hundred Å equivalent width of an emission line of typical single W-R stars is usually only a few Å in W-R galaxies. Examples of the spectra are given in Vacca & Conti (1992) and Conti (1993). Many of these galaxies are metal-weak but this may be a selection effect, given that starburst galaxies with more normal composition are likely to have a brighter underlying stellar population, which could "drown out" the W-R stars even if present.

Vacca & Conti (1992) have recently analyzed optical spectra of ten Wolf-Rayet and four other emission line galaxies. The nebular line ratios indicate that the excitation is caused by stars. The strength of the λ 4686 Å may be used to infer the numbers of W-R stars present; this uses a calibration of the line flux for single W-R stars in the LMC, dividing that number into the measured line flux in the galaxy. Typically, tens to hundreds of W-R stars are inferred to be present in the starbursts. This procedure has been quantitatively checked by using the area spectrophotometry of 30 Dor in which a broad λ 4686 He II line is measured; the inferred numbers of W-R stars (20) are similar to the census of Moffat et al (1987). The contribution of the W-R stars to the observed *N* Lyc is subtracted from the observed value for the galaxy, and the remainder is treated, as above, to determine the number of O-type stars present. As it is not possible to derive the slope of an IMF from this procedure, a Γ equivalent to that of 30 Dor was adopted for each galaxy; similarly the was assumed to be the same (100 $M\odot$). With the usual assumptions that low mass stars have formed along with the high mass ones, typical star formation rates range from 2×10^{-2} to 3 $M\odot$ per year (the former is the value for 30 Dor).

Following the nebular line analysis taken for 30 Dor, spectrophotometric studies have provided quantitative values of WNL/O for W-R galaxies (Vacca & Conti 1992, Conti 1993). Figure 3 shows the comparison of these observations with continuous star formation and newly constructed burst models (Meynet 1994b; see also Maeder & Meynet 1994). The observed WNL/O ratios in W-R galaxies are much higher than the predictions of models with constant SFR, in agreement with previous estimates (Arnault et al 1989). In Figure 3, burst models with two different IMF slopes are shown which bracket the observations nicely. Notice also the importance of *Z* to the predictions. The observed values are, strictly speaking, the number of WN stars derived from the strength of λ 4686 He II, but in most of these W-R galaxies, there is little or no evidence of WC stars. The largest source of uncertainty in the observed values is a possible mismatch between the slit width (1.5") and the starburst (typically 2 to 3"). This underestimates the number of O stars. Arbitrarily correcting for this would *lower* the ratio by a factor 2 (see a more thorough discussion in Conti 1993).

W-R binaries might enhance the production of W-R stars, but this effect is believed to be small except at the lowest *Z*. While Masegosa et al (1991) suggest that SN contamination could affect the starburst properties, particularly the "blue bump" at λ 4650, this may be unimportant as no other effects (e.g. shocks) are seen in the nebular spectra of W-R galaxies (Vacca & Conti 1992). Could there be different values for Γ in various starburst regions? This possibility has been raised in recent years (Scalo 1989; Mas-Hesse & Kunth 199l a,b; Rieke 1991; Joseph 1991; Bernlohr 1993). We have already noted (Section 2) that in regions where the O star populations can be counted *directly*, Γ does appear to differ from region to region by a small, but significant amount. Thus differences in Γ from starburst region to starburst region are possible, but final conclusions are still uncertain.

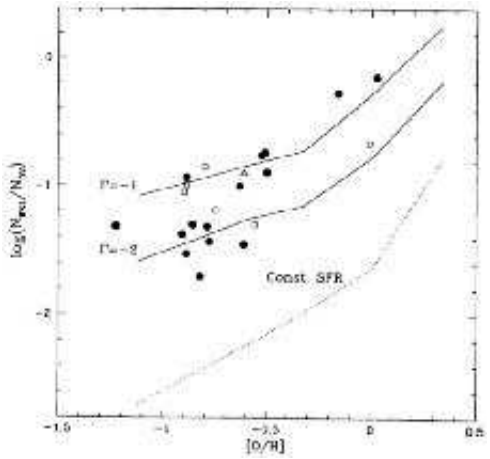

Figure 3. WNL/OV ratio in W-R galaxies, as a function of the oxygen abundance [O/H]; data adapted from Conti (1993); stellar models from Maeder & Meynet (1994). (bullet) Galaxies of Vacca & Conti (1992); (square) galaxies of W.D. Vacca (private communication); (triangle) 30 Dor (from Vacca 1991). The dotted line denotes the predictions for "continuous" star formation; the solid lines are for values estimated for a "burst" (see text). Γ is the slope of the IMF as defined in Section 2.2.

From Figure 3 we conclude that for W-R galaxies, the W-R/O ratios are well above the predictions for "continuous" star formation, and nicely match the predictions of "burst" models. We infer that the starbursts observed in W-R galaxies are typically going on for only a relatively "brief" interval, typically $\lesssim 10^6$ yr. The energetics are similar to 30 Doradus at the faint end to more than 100× larger. Are these nearby starbursts paradigms for the first phases of massive star evolution in very young galaxies?

Phillips & Conti (1992) discovered evidence for WC9 stars in a GH II region at the edge of the bar in the metal-rich (+0.5 dex) galaxy NGC 1365. As their spectrum only covered the yellow region, there is currently no information on the numbers of WN stars in their strong starburst region. However, the presence of late type WC stars in a strong *Z* environment nicely follows the models predictions (Section 4.5.3).

5.1.5 SPECTRAL SYNTHESIS Optical spectroscopy of most galaxies shows absorption line spectra from an old evolved population of late type giant stars. In the W-R galaxies studied by Vacca & Conti (1992) the optical absorption spectra are those of late B and A type stars (narrow K line, upper Balmer series in absorption, no other features). In these objects, even though there are large numbers of very hot stars, the optical spectra are dominated by other stars. However, in the UV the O (and W-R) stars become the dominant contributors to the continua. In emission line galaxies containing OB stars, P Cygni spectral features at λ 1400 Si IV and λ 1550 C IV are seen; if W-R stars are present, an emission line at λ 1640 He II is found.

In Figure 4 we show IUE spectra of a 3’ × 3’ area of the GH II region 30 Dor in the LMC and NGC 1741, a W-R galaxy. UV spectra such as these can be used directly to estimate the numbers of hot stars present by spectral synthesis techniques (Leitherer et al 1992 a,b; Robert et al 1993). One can also ascertain the age of the starburst and tell whether or not the formation of the massive stars has been "continuous" or a "burst": In work in progress, Vacca et at (1994) have found a reasonable fit for the Si IV, C IV, and He II profiles for ages between 1.5 and 2 Myr for the UV spectrum of 30 Dor illustrated in Figure 4. For younger ages, insufficient W-R stars have been produced to match the λ 1640 emission line; for older ages, the Si IV and C IV profiles begin to diverge from the observations.

It is also possible to estimate the number of O stars from the extinction-corrected IUE continuum of 30 Dor, as the models predict this quantity. The agreement with the census from Parker (1991, 1993) and Parker & Garmany (1993) is good. Unfortunately, even the brightest starburst galaxies are at the limit of IUE sensitivity. The 10" × 20" slit is typically larger than the starburst, so dilution of the spectral signatures can be a problem. For the number counts, but not for the age estimate, the UV extinction is critical and remains a serious problem for more general application of this method.

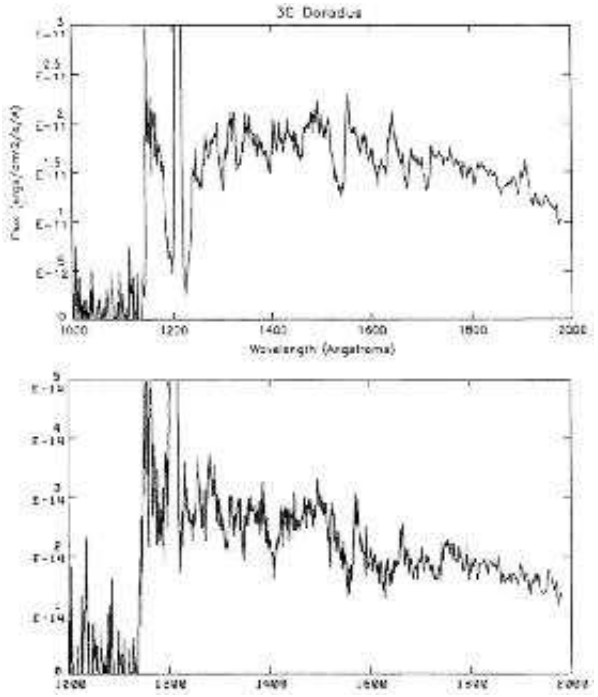

Figure 4. Observed *IUE* SWP spectra of 30 Dor (*top*) and NGC 1741 (*bottom*), a W-R galaxy. Stellar line features at λ 1400 Si IV, λ 1550 C IV, λ 1640 He II, and λ 1718 N IV are seen in both objects. The lines shortward of Si IV are primarily interstellar.

5.2 Global Properties

5.2.1 FAR-INFRARED LUMINOSITIES Radiation from dusty H II and GH II regions and galaxies in the far-infrared appears to come from heating of the dust by their stellar populations. Devereux & Young (1991) have argued that the far-infrared (FIR) luminosities relate directly to the content of *hot stars*. They have derived the FIR luminosity of 124 spiral galaxies from the *IRAS* archives, which also have Hα measurements. From the individual *IRAS* colors, they find that this dust is at a temperature of 30 to 40 K, corresponding to heating by *hot* stars: they suggest that ambient heating by stars of all types would be more like 15-20 K. In their Figure 5 they show a correlation between the FIR and the Hα emission extending over two orders of magnitude. Part of the dispersion in their relation could come from the differences in $T_{\rm eff}$ of the exciting stars. Although Devereux & Young claim that one can infer the actual $T_{\rm eff}$ values, we don't believe that the modeling used is adequate.

With considerable caution, one could perhaps use the FIR luminosities of star forming regions to infer the total numbers of hot stars present, but, absent other information, one cannot infer an IMF slope (Γ) or an $M_{\rm upper}$. The FIR luminosities for normal spiral galaxies range from 10^9 to 10^{11} $L\odot$ and, according to Devereux & Young (1991), show no dependence on the Hubble type. This seems at odds with the appearance of stellar images on direct plates. It would be nice to confirm that the relationship

between FIR luminosity and numbers of hot exciting stars is as orderly as claimed. This might be accomplished with identifications and counts of O type stars in otherwise obscured GH II regions of our Galaxy, such as W51, through classification K band spectroscopy (Conti et al 1993).

5.2.2 IMAGING OF EMISSION LINE AND W-R GALAXIES Optical CCD photometry of galaxies is a growth industry: Spectral syntheses of the continua can give us valuable information on stellar populations of all but the hottest and most massive stars. As we have already noted, even in W-R galaxies, demonstrably the youngest examples of starburst phenomena, the optical light comes primarily from stars of type A. Thus in this wavelength region we are sampling stars with lifetimes (not necessarily ages) of 50 million years or more.

Conti & Vacca (1994) have used the *HST* with FOC camera and the λ 2200 Å filter to obtain a UV image of the W-R galaxy He 2-10; the spatial resolution with image restoration is $\approx$ 0.1". At this wavelength one is primarily sampling the OB stars, which have lifetimes of a few $\times 10^6$ to 10^7 yr. Given the presence of W-R stars in this galaxy, the age is closer to the lower of those numbers. This W-R galaxy has recently been found by Corbin et al (1993), using deep CCD photometry, to be at the center of a faint elliptical galaxy (3 kpc diameter)! There are three starburst regions, the strongest (containing W-R stars) at the center. Figure 5 is a reconstructed HST UV image of the central starburst in He 2-10. One can see ten individual knots of activity, several just at the limit of resolution (10 pc in diameter), with separations of about two to three times this number.

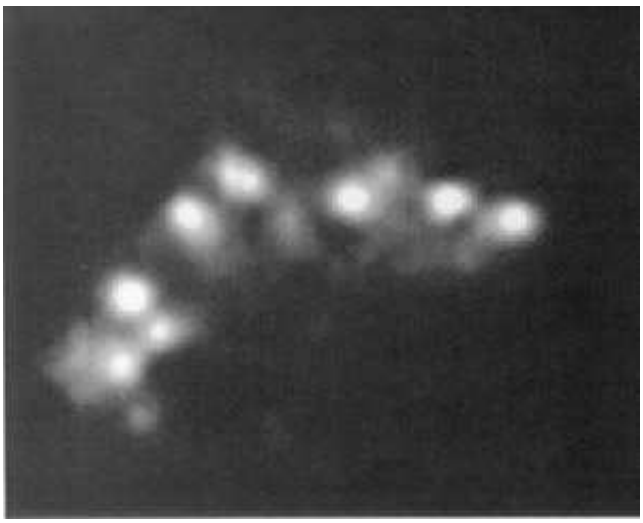

Figure 5. Reconstructed UV image of the central starburst in He2-10 (from Conti & Vacca 1994). Nine individual starburst knots are readily seen. The spatial scale across the figure is 3". corresponding to 125pc. The parent elliptical galaxy (Corbin et al 1993) is about 80" in diameter, which is much larger than the scale of this starburst.

Conti & Vacca (1994) estimate the average λ 2200 luminosities of the individual starburst knots in He 2-10 to be $\approx 10^{38}$ erg s^{-1} $Å^{-1}$. This corresponds to M_V of -13.9, and mean total masses (assuming a normal IMF) of a few $\times 10^5$ $M\odot$. These masses are similar to those of globular clusters, but in objects with ages of less than 10 Myr! Whitmore et al (1993) have recently identified about 40 blue objects in the galaxy merger remnant NGC 7252 as being globular clusters of age about 500 Myr, on the basis of their optical colors and absolute magnitudes. Their average M_V are about one magnitude fainter than those of He 2-10, which could be an evolution effect since the starburst in NGC 7252 is older.

Conti & Vacca (1994) have noted that at least a dozen other W-R galaxies for which they have HST UV imaging also show multiple knots of star forming activity. The luminosities of the knots in several of these galaxies are substantially larger than those of He 2-10. Many of these other galaxies are clear examples of merging or interacting galaxies (the status of He 2-10 is not quite clear on this point). They thus suggest that galaxy merger events may lead to star formation episodes which produce knots of activity and eventual production of what we now recognize as globular clusters. Of course, it is not certain that where large numbers of massive stars have been produced, as in these starburst regions, low mass stars are also born. In nearby galactic associations we do consistently see a lower main sequence, but in highly energetic star formation episodes, this might not always be the case.

6. CONCLUSIONS

Theoretical considerations and extensive modeling suggest that substantial distinctions in the evolutionary history of massive stars will arise because of the importance of mass loss and its strong dependence on metallicity. We have reviewed and discussed the properties, chemical abundances, and populations of individual OB stars, supergiants, and Wolf-Rayet stars in various galaxy environments and find order of magnitude differences from place to place. These are understood to depend on *Z*, along with the IMF and SFR, but not all these parameters are completely sorted out at the present time. From detailed star counts in stellar associations (Section 2), it does appear that there is no simple dependence of the IMF slope, Γ , on *Z*, but there may be differences from place to place. There is no evidence that the $M_{\rm upper}$ limit depends on *Z*; data concerning potential differences in the $M_{\rm lower}$ limit with location is currently lacking. A major unresolved question is that concerning the SFR: What does this parameter depend upon? Clearly one needs sufficient molecular gas; future studies should address the relationship between this parameter and the actual numbers of massive stars in very quantitative terms.

We have lightly touched upon the connection between studies of nearby massive star formation regions where stellar statistics may be accomplished, and measurement and analyses of their integrated properties. These objects, such as 30 Dor in the LMC and NGC 604 in M33, can be used as "stepping stones" in our understanding of similar phenomena in more distant galaxies. It will be important to improve the "calibration" of these types of massive star groupings in various galaxy environments to better understand those even more energetic and less common starbursts found at larger distances.

In starbursts we are dealing nearly exclusively with integrated spectral properties. For those containing massive stars of O and W-R type we have an advantage that the lifetimes are less than 10 Myr, and the formation time scales appear to be only 1 Myr. The modeling can be thus simplified to that of a "burst." Probably SN will not have played much of a role, as yet, in the energetics of the phenomena we are observing; the excitation of the gas will be primarily due to stars. Multiwavelength studies of many galaxies have already been made but there has been little quantitative integration of all these data for the same galaxies and the same starbursts. In addition to the inferences concerning hot stars, one would like to know the cool star population, and the quantity of neutral and molecular gas. This requires observations at IR wavelengths and in the sub-mm and cm regimes. One certain caution is that of aperture; it is critically important that these are matched, independent of the wavelengths, so that the same volume elements are being examined in each case. Even more significant is the probability that as one goes to shorter and shorter wavelengths, differential internal galactic extinction might necessarily shield from our view the "back side" of a starburst episode.

With our new-found knowledge of the properties and evolution of massive stars, we now can begin to study and understand the appearances of ever more distant galaxies, and one would hope, to delineate their past history. "In the beginning" there were undoubtedly many massive stars in newly forming galaxies. Our improved comprehension of newly born local massive stars can help clarify our knowledge of some of the earliest stages in the evolution of our Universe.

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