Evolutionary models for giant extragalactic HII regions at different metallicities

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Abstract. — We present theoretical evolutionary models for ionizing star clusters and their associated giant HII regions up to an age of 5.4 Myr and metallicities between 1/20 and 2.5 times solar. These young clusters can provide the high energy photons needed to keep the regions ionized. We discuss in some detail the Wolf-Rayet and red supergiant phases, sometimes present in the early stages of the evolution of a star cluster. The spectral features characteristic of these stars, when detected and measured on the spectra of giant extragalactic HII regions (GEHR), can be used to constrain the age of the ionizing stellar population. The emergent ionizing continua of the clusters are used as input for the photoionization code CLOUDY to obtain the corresponding emission line spectra in the optical and infrared ranges. The effect of the metallicity is taken into account in the stellar evolution and atmosphere models as well as in the nebular gas producing a consistent set of models. These models can provide estimates of the mass, age and metallicity of the dominant stellar population ionizing a given GEHR solely from the analysis of its optical emission line spectrum.

Key words: HII regions — stars: early — type; Wolf-Rayet; evolution — galaxies: ISM

1. Introduction

Giant extragalactic HII regions (GEHR) are large and luminous HII regions, observed in external galaxies mainly on the discs of early type spirals and in irregulars. Sometimes they are also found in the circumnuclear region of starburst galaxies. The size of these regions, as measured from their diameters on Hα images, can be as large as 1 Kpc. Recent reviews of the general properties of GEHR can be found in Dinnerstein (1990), Shields (1990), Kennicutt (1991), and Melnick (1992).

The Hα emission luminosities of GEHR are in the range $10^{38} \, - \, 10^{41}$ ergs$^{-1}$, implying values for the number of ionizing photons per second between $10^{50}$ s$^{-1}$ and $10^{53}$ s$^{-1}$, which require the presence of young star clusters as ionization source. These stellar clusters in the cores of GEHR cannot be resolved from the ground and their properties must be inferred from the integrated observations of the ionized nebulae. For example, in 30 Doradus, a GEHR in the Large Magellanic Cloud, a star cluster inside the nebula, NGC 2070, has been observed, and its core (R136) has been resolved into individual stars (WR, O-B stars, blue supergiants..) (Walborn 1991; Melnick 1992). A much more detailed study is offered by Hubble Space Telescope observations (Weigelt et al. 1991; Campbell et al. 1992, Nota et al. 1994). Other GEHR in nearby galaxies have been spatially resolved in part: NGC 5471 in M 101, 3 times more luminous than 30 Dor (Skillman 1985), NGC 604 in M 33 (Díaz et al. 1987) and NGC 2363, in NGC 2366 (González-Delgado et al. 1994). These observations suggest that in most observed GEHR, only the integrated light of the whole complex masked by the gas emission is seen and probably inside them, a 30Dor-like structure is hidden, and only should be seen with better spatial resolution, not possible from ground observations.

In all kinds of emission line objects like galactic HII regions, GEHR, HII galaxies, and even starburst nuclei, the spectrum of the photoionized gas seems to be controlled by the shape of the ionizing continuum, the degree of ionization and the chemical composition. If we assume that the gas is photoionized by the radiation coming from a star cluster it should be possible to derive its mass, age and metallicity from the observation of the emission-line spectrum of the gas, and eventually from suitable stellar features associated with the presence of some particular massive stars in the cluster, like Wolf-Rayet (WR) and red supergiant (RSG) stars.
In this work theoretical evolutionary models for very young star clusters are presented. Their synthesized spectral energy distributions (SED) constitute the input for the photoionization code CLOUDY used to compute the emission line spectrum of the gas. The evolution of the star clusters is followed from 1 to 54 Myr. Five metallicities are considered between 1/20 \( Z=0.001 \) and 2.5 times solar \( Z=0.05 \). The effect of the metal content is taken into account simultaneously in the evolutionary tracks, the atmosphere models, and the surrounding gas, producing a consistent grid of models. The main goals of this work are first to reproduce globally the observations of GEHR of different metallicities and second, to devise an observational method to derive the physical properties — mass, age and metallicity — of the dominant stellar population responsible for the photoionization of GEHR, from their observed integrated optical emission-line spectrum.

The details of the stellar evolution models are discussed in Sect. 2. Section 3 describes the cluster spectral energy distributions computed with our population synthesis code whose main characteristics are also recalled. Section 4 is devoted to the discussion of the photoionization models whose detailed results are presented in García Vargas et al. (1995). The method to determine the physical parameters of the ionizing star clusters in GEHR is discussed in Sect. 5. Section 6 deals with the infrared predictions for high metallicity GEHR. Particular attention is paid to the IRAS excesses by emission line contamination in the filters. The impact of ISO observations to test the predicted infrared emission line spectrum is stressed. Finally, in Sect. 7, we summarize the main conclusions of the present work.

2. Stellar evolution models

2.1. Stellar evolutionary tracks

To obtain the basic properties of the HR diagram (HRD) of a given ionizing cluster we constructed isochrones at several ages by interpolating between the evolutionary sequences calculated by Bressan et al. (1993), and Fagotto et al. (1994a, b). These tracks were computed using the recent radiative opacities of Iglesias et al. (1992) for the initial chemical compositions \( Z=0.001 \), \( Y=0.23 \); \( Z=0.004 \), \( Y=0.24 \); \( Z=0.008 \), \( Y=0.25 \); \( Z=0.02 \), \( Y=0.28 \); and \( Z=0.05 \), \( Y=0.352 \) (Padova models).

The Padova models account for mild overshoot from the convective core, which in the formalism of Bressan et al. (1981) is expressed by taking the mean free path of convective elements across the unstable region defined by the Schwarzschild criterion, to be half the local pressure scale height \( t=\lambda_c \times H_p \). This is equivalent to the amount of over-

shoot considered by the Geneva group (Maeder & Meynet 1991; Schaller et al. 1992), who used a different formalism by defining directly the region of overshoot above the unstable core to be 0.25 \( H_p \). Moreover, the Padova models account for a 0.7 \( H_p \) region of overshoot below the base of the convective envelope (Aloni et al. 1991).

A significant difference in Padova models with respect to the standard mass-loss models by the Geneva group is the adoption of the low rate of the \( 12C(\alpha, \gamma)16O \) reaction in the He-burning phase (Caughlan & Fowler 1988). In the domain of massive stars the effects of a lower \( 12C(\alpha, \gamma)16O \) reaction rate are a slightly shorter He-burning lifetime and a slightly higher central carbon abundance at the end of the He-burning phase.

In the age range of a few Myr considered in the present work, evolutionary effects appear for stars with masses over 20 \( M_\odot \) and therefore mass-loss is the parameter that by far dominates the phenomenology of the HRD at different metallicities. In the evolutionary tracks, mass loss by stellar winds has been accounted for according to the rates given by de Jager et al. (1988) from the main sequence up to the so-called de Jager limit in the HRD. They include the dependence on metallicity given by Kudritzki et al. (1985). Beyond the de Jager limit the most massive stars enter the region where Luminous Blue Variables \( (LBV) \) are observed and, accordingly, the mass-loss rate has been increased to \( 10^{-3} M_\odot \text{ yr}^{-1} \). As the evolution proceeds, the surface hydrogen abundance by mass in the most massive stars eventually falls below the value of 0.3. In this case the model is supposed to become a Wolf-Rayet \( (WR) \) star and the mass-loss rate is derived according to Langer (1989). This critical value of the WR surface hydrogen is slightly lower than the value of 0.4 adopted by Maeder & Meynet (1994), and consequently the WR phase starts later in the models adopted here.

Finally it is worth noticing that for ages lower than about 3 Myr all stars up to the upper limit of the isochrone are still alive because, at the highest masses, the stellar lifetime approaches an asymptotic limit of about 3 Myr.

2.2. Theoretical HR diagram

The ionizing clusters have been assumed to form in a single burst with a standard power-law initial mass function (IMF), \( \phi(m) = m^{-\alpha} \), with \( \alpha = 2.35 \) (Salpeter), considering the masses of formed stars between the lower limit \( m_{\text{low}} = 0.85 M_\odot \) and the upper limit \( m_{\text{up}} = 120 M_\odot \).

There is recent observational evidence that points against the "single instantaneous burst" approximation for large star-forming regions. The HST UV image of the central starburst in He2-10 (Conti 1994; Conti & Vacca
1994) shows nine individual "blobs" of about 10 pc in a region of 125 pc with a strong UV continuum which might correspond to different star associations. Similar concentrations have been observed in the starburst galaxies NGC 3690 (Leitherer 1994), and M 83 by Heap (1994) and in some GEHR spatially resolved in part from ground observations like 30Dor (Mathis et al. 1985), NGC 5471 in M 101 (Skillman 1985), NGC 604 in M 33 (Díaz et al. 1987) and NGC 2363 in NGC 2366 (González-Delgado et al. 1994). The recent HST observations of NGC 604 (Drissen et al. 1993) seem to suggest that star formation might proceed during a few Myr. In these cases a combination of the model star clusters presented here would be a better representation of reality.

Models are computed for total cluster masses, \( M_T = \frac{\int m \phi(m) dm}{m_{\text{low}}} \), between 0.12 \( 10^5 \, M_\odot \) and 0.6 \( 10^5 \, M_\odot \). Clusters in this range of masses can actually provide the necessary ionizing photons, \( Q(H) \), to produce the observed H\(_\alpha\) luminosities (log \( Q(H) \approx 51 \)) of GEHR (see García-Vargas & Díaz 1994). However, the results can be scaled for higher cluster masses as explained in Sect. 5. This is not the case for clusters with masses below 0.1 \( 10^5 \, M_\odot \) since in clusters of such low mass the formation of very massive stars is not possible if an analytical approach for the IMF is considered. This so called richness effect has been analysed by García-Vargas & Díaz (1994) who made a comparison between the populations of star clusters synthesized using an analytical form and Monte-Carlo simulations for the IMF. According to their results, if a Monte-Carlo algorithm is used to produce the mass spectrum, a non zero probability exists of forming massive stars even in the less massive clusters, although in general the probability of forming massive stars decreases with decreasing cluster mass for a given set of IMF parameters. In the range of cluster masses considered in the present work the results are identical if the mass spectrum is computed either analytically or stochastically. The difference between an analytical and Monte-Carlo method has also been studied by Cerviño & Mas-Hesse (1994).

We want to emphasize that the number of massive stars present in a given cluster is definitively important for the total production of ionizing photons and therefore to account for the ionizing power of the cluster. Moreover, some stellar populations, such as WR and RSG, represent evolutionary phases of the most massive stars in the cluster, and can eventually be used as age indicators.

The factors which can alter the number of massive stars are: the mass of the cluster through the richness effect, as mentioned above, and the parameters of the IMF. In fact, the effect of varying the upper limit and the slope of the IMF according to metallicity has been considered in previous works in order to explain the observations of GEHR. Shields & Tinsley (1976) suggested a change in the upper mass limit of the IMF, this being higher at low metallicity. Terlevich & Melnick (1981), on the other hand, suggested an increase in the slope of the IMF with metallicity. Both approaches were tested for solar metallicity GEHR by García-Vargas & Díaz (1994) who suggested possible changes to both upper mass limit and slope of the IMF. However, any definite conclusion about the matter would be strongly affected by stellar evolution uncertainties as is evidenced by the fact that the use of the most recent models makes any changes of the IMF parameters unnecessary, at least for high metallicity regions, as discussed in Sect. 4. On the other hand, for a cluster with a given total mass, the number of stars above 20 \( M_\odot \) is a function of \( m_{\text{low}} \). Indeed the ratio between this number and the total number of stars in a cluster changes from 1.28 \( 10^{-2} \) to 7.1 \( 10^{-4} \) with \( m_{\text{low}} \) decreasing from 0.85 \( M_\odot \) to 0.1 \( M_\odot \).

The change of some properties of synthetic star clusters with the IMF parameters can also be seen in Mas-Hesse & Kunth (1991), Leitherer (1994) and Leitherer & Heckman (1995).

Pre-main sequence evolution has been neglected and all unevolved stars are assumed to be on their zero age main sequence (ZAMS) location. The pre-main sequence lifetime for a star is larger than 1 Myr at about 5 \( M_\odot \) and it reaches 2 \( 10^7 \) yr at 1.5 \( M_\odot \) (see e.g. Iben 1965). However, the models spend only a tiny fraction of their pre-main sequence lifetime far from their ZAMS location and only for the lowest masses this fraction is comparable to the age of the isochrone. Therefore the contribution of these stars to the integrated light of the stellar population can be neglected.

A fine grid of isochrones for very young stellar populations (1 to 5.4 Myr) was computed following the method outlined by Bertelli et al. (1990) and more recently by Bertelli et al. (1994). In particular the latter adopted the same set of evolutionary tracks to compute isochrones in the age range 4 Myr to 20 Gyr. The reason for stopping our calculations at the age of 5.4 Myr is that a single cluster older than this cannot provide the ionizing photons required to reproduce the observations of GEHR without the inclusion of the effects of SN explosions and winds of massive stars, not considered in these first generation of models. There are observations of GEHR in which one sees the effects of winds and SNe, namely NGC 5471 in M 101 (Skillman 1985), GEHR in NGC 2403 (Yang et al. 1994) and NGC 2363 (González-Delgado et al. 1994). Recently, Leitherer (1994) discussed the energy balance in the IS medium by winds and SNe, derived from his evolutionary models, and a detailed description of these models can be seen in Leitherer & Heckman (1995).

Isochrones for different ages are shown in Figs. 1a (2 Myr), 1b (3 Myr), 1c (4 Myr), and 1d (5 Myr) where dif-
Fig. 1. Isochrones at 2 Myr a), 3 Myr b), 4 Myr c) and 5 Myr d). Different symbols correspond to isochrones of different metallicities, as labelled in the plot.

2.3. The Wolf-Rayet and red supergiant phases

Special attention has to be paid to some important evolutionary stages like the Wolf-Rayet (WR) and Red Supergiant (RSG) phases, which can provide important observational clues about the nature of the star clusters embedded in GEHR. To this purpose we have collected in Table 1 the ages and durations of several relevant evolutionary phases of massive stars, at varying metal content. More details can be found in Bressan et al. (1993) and Fagotto et al. (1994a, b). In Table 1, $M$ indicates the initial mass in solar masses, $LBV$ and $\Delta t$ the starting age and the duration of the $LBV$ phase in Myr. The luminosity boundary that defines the beginning of the $LBV$ phase has been approximated by the relation $\log L_{LBV} = \text{MAX}(1.18 \times \log(M_{10}^4) + 5.71, 5.71)$ which roughly represents the observed limit to the bright supergiant stars in the HRD of the Large Magellanic Cloud (Fitzpatrick & Garmany 1990).

The red supergiant phases are indicated by RSG, which is the age at which the star eventually becomes a Red Su-
Table 1. Characteristic times of relevant phases for massive stars

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<td>21.97</td>
<td>20.63</td>
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As far as the WR phase is concerned, WNL, WNE, WC, WO stem for the age at which the models become a WNL, WNE, WC and WO star respectively. The beginning of each different WR subtype is derived from the surface abundance by mass of hydrogen, helium, nitrogen, carbon and oxygen, $X$, $Y$, $N$, $C$ and $O$ respectively. Namely the model is identified as a WNL star if $0 < X^* < 0.3$, a WNE star if $X^* = 0$ and $N^* > C^*$, a WC star if $X^* = 0$ and $N^* < C^*$, and finally a WO star if $X^* = 0$ and the ratio by number of $(C^* + O^*)/Y^*$ is greater than unity.

Finally, the last three columns indicate the final age of the model and the duration of the hydrogen and helium burning phases respectively.

Inspection of Table 1 shows that WRs and RSGs constitute alternative evolutionary paths determined by the age and metal content of a star cluster. Clusters younger than about 5 Myr are rich in WR stars but lack RSG stars. On the contrary, clusters older than this do not possess WR stars but contain RSGs. At the lowest metallicities the age spread between the oldest cluster still showing WR stars and the youngest one containing RSG stars increases, because mass loss is less efficient in producing WR stars ($M \propto \sqrt{Z}$) and, at the same time, helium ignition occurs at hot effective temperatures ($T_{eff} > 4$) preventing...
the star from becoming a red supergiant but for a brief final episode.

In the RSG stage itself the models show a manifold of paths. Toward the lower mass end all stars display the characteristics of the standard evolution at constant mass of intermediate mass stars, namely a more or less extended loop during the central He burning phase. At higher initial masses and depending on the metal content, some models burn helium as RSG stars but do not perform the blue loop (in this case RSG\textsubscript{T} is null), while some others start burning helium as blue supergiants (BSG) and become RSG stars only at the end of the helium burning phase (in this case only RSG\textsubscript{I} is present). The model of 30 $M_{\odot}$ of solar composition is particular in that after becoming a RSG star, due to heavy mass loss it leaves the RSG phase and turns into a WR star. Star clusters of this composition and turnoff mass should contain RSG and WR stars at the same time. Besides this, there are a number of very luminous stars that enter the LBV region and reach for a while effective temperatures as low as log $T_{eff}=3.7$. The precise location of these stars in the red side of HR diagram depends very much on the mass-loss rate adopted during the LBV phase and therefore we have not included these stars as RSG stars in Table 1.

To quantify the above considerations we show in Figs. 2a-e the ratio of the number of different WR subtypes and RSGs to the total number of stars with initial mass greater than 20 $M_{\odot}$, for single stellar populations of different chemical composition as a function of their age. The above number ratios have been obtained by integrating the IMF by number in the mass intervals corresponding to the different evolutionary phases. The smooth line across each figure gives the total number of stars with initial mass greater than 20 $M_{\odot}$, for a cluster of 1 $M_{\odot}$. For a given age, to get the total number of massive stars of a determined type in a cluster with total mass in stars $M_{T}$, it is necessary to multiply the value given in the plot by the value at the same age given in the smooth line and by the factor 0.462 $M_{T}$.

Notice the prominence of the WNL and WC phases in metal rich populations, the former dominating between 2 and 5.2 Myr and the latter appearing at about 3 Myr and lasting up to about 6 Myr. Again these stellar evolutionary phases are distinctly separate in age but for a small time interval of about 1 Myr. The existence of high luminosity WR stars and their relation with the metallicity of the environment is generally well described by the above models. However, as already pointed out by Maeder & Meynet (1994), the predicted number of galactic WNE stars is significantly lower than the observed one; indeed if a standard mass loss law and mixing scheme are adopted, this phase is virtually absent. By contrast, in the recent census by Hamann et al. (1993) WNE stars encompass 50% of all galactic WN stars. Furthermore WN stars with luminosity as low as log $L/L_{\odot}=3.5$ are found, while the standard theory predicts none below log $L/L_{\odot}=5.5$. A better agreement with the overall observed WR number distribution and luminosity function can be obtained either adopting an enhanced mass-loss rate as in Maeder & Meynet (1994) or by assuming a slow diffusion of elements from the convective core to the stellar surface as a result of differential rotation as in Deng et al. (1994). In the former case (Maeder & Meynet 1994) the authors adopt a mass loss rate twice the one obtained by the de Jager parameterization, even if there is a general good agreement between this parameterization and the recent empirical estimates (Lamers & Leitherer 1993). Note however that de Koter et al. (1994), using HST observations, have estimated that the mass loss rate of the O3f-WN star R136a5 in the 30Dor center is 8 times higher than the value predicted by the theory. An enhanced mass-loss rate forces the most massive star to become less luminous and long lived, a fact that, coupled with the resulting different internal chemical profile, is responsible for a better distribution of WR stars among the different subtypes and a better agreement with the WR stars of lower luminosity. In the latter case (Deng et al. 1994) stars of initial mass as low as 9–12 $M_{\odot}$ become WR stars when particular conditions that trigger the diffusive instability are met and the HR diagram of the WR star of low luminosity is quite well reproduced, without the need for an atmospheric correction to the effective temperature. These latter models can reproduce the location in the HR diagram of the majority of the galactic WN stars (Hamann et al. 1993). In summary there is a general difficulty in the interpretation of both the evolutionary status and the number distribution of WR subtypes in the light of the standard models and different alternative paths are being proposed which however are not considered here.

At lower metallicities the mass loss rate is less efficient and, as a consequence, the number of WR stars at a given age is a decreasing function of the metallicity (Figs. 2a-c). The distribution of WR stars among the different subtypes reflects the interplay between the mass loss rate which is metal dependent, and the behaviour of the convective core which determines the inner chemical profile (see also Maeder 1991).

From Table 1 it is seen that the age at which a model of initial mass $M=120 M_{\odot}$ becomes a WR star, increases from a metal rich population ($Z=0.02$, $t_{WNL}=1.91\times10^{6}$yr) to a metal poor population ($Z=0.004$, $t_{WNL}=2.82\times10^{6}$yr). Notice also that in the former case the model first becomes a LBV star and immediately after a WR star, while in the latter case the star first becomes a WR star and only at a later time it enters the LBV region.
Fig. 2. Relative numbers of massive stars in the main relevant evolutionary phases, as a function of the cluster age (in Myr), using the standard IMF ($\alpha = 2.35$, $m_1 = 0.85 \, M_\odot$, $m_u = 120 \, M_\odot$). Different phases of massive stars are plotted with different symbols as labelled in the figure. The spikes of red supergiant stars (RSG) at ages younger than about 5 Myr are due to the rapid excursion towards the red side of the HR diagram of some massive stars models. They are not considered as beginners of the RSG phase in Table 1. The numbers are normalized to the current number of stars with initial masses higher than 20 $M_\odot$. The smooth line with stars represent the current total number of massive stars (initial masses higher than 20 $M_\odot$) for a cluster of 1 $M_\odot$. To get the total number of massive stars of a determined type in a cluster with initial total mass in stars $M_T$, it is necessary to multiply the value given in the figure by the value given in the smooth line at that age, and by the factor 0.462 $M_T$. 

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With regard to the RSG phase, WR stars disappear when the cluster becomes older than about 5 Myr and leave room to the RSG stars. As already anticipated, at low metallicity the appearance of the RSG phase is delayed until the evolution of the less massive stars that ignite helium in this phase. Finally the LBV stars reaching low effective temperatures, which have been marked in Figs. 2a-e with the same symbol as the RSG stars, are clearly visible as spikes at ages younger than about 5 Myr.

About the same conclusions have been reached by Cerviño & Mass-Hesse (1994), who studied the metallicity effects on the WR and RSG phases, using the evolutionary tracks of Schaerer et al. (1993a, b) and Charbonnel et al. (1993a, b). They find a shorter WR phase when metallicity decreases, and a lack of RSGs at very low metallicity. In the present models, the WR phase has a maximum in duration and in number ratios of WR to stars more massive than 20 $M_\odot$ at solar metallicity, although the phase is seen in bursts with metallicities between 0.2 $Z_\odot$ and 2.5 $Z_\odot$. On the other hand, the RSG phase exists at metallicities as low as 1/20 $Z_\odot$, although the phase is shorter at low metallicity.

2.4. Observational signatures of massive stars

Observationally, the presence of Wolf-Rayet stars can be detected through some emission features in the optical spectrum, usually broad due to the velocity of the wind where they are produced. However, in GEHR the detection of WR features is difficult because they are faint and narrow nebular lines are usually superimposed.

WR features have been observed in star-forming complexes (Conti 1994; González-Delgado et al. 1995) and in particular in GEHR (Sargent & Searle 1970; Keel 1982; Díaz et al. 1987; Terlevich et al. 1990; González-Delgado et al. 1994). WR subtypes contribute to different optical emission lines: WN stars present nitrogen lines, like NIII $\lambda\lambda$ 4634–4642 Å, NIV $\lambda$ 4057 Å or NIV $\lambda\lambda$ 4605–4622 Å, and helium lines like He II $\lambda\lambda$ 4339, 4542 Å and the prominent 4686 Å. WC stars also present this line, as well as carbon lines, like CII $\lambda$ 4267 Å and CIII $\lambda\lambda$ 4650, 5696 Å, CIV $\lambda\lambda$ 4658, 5801, 5812 Å and oxygen lines, like OV $\lambda$ 5592 Å and OVI $\lambda\lambda$ 3815, 3835 Å. WN and WC types are sub-divided in WN2-WN9 and WC4-WC9 sequences, depending on the relative strengths of the emission lines (Van der Hucht et al. 1981). These lines often appear as bumps in intermediate resolution spectra. The most common WR bumps detected in GEHR are the bump at 4660 Å ($\Delta\lambda = 70$ Å), and the bump at 5810 Å ($\Delta\lambda = 80$ Å). Empirical calibrations for the luminosity of WR bumps with WR type can be seen in Conti et al. (1983), Conti & Massey (1989) and more recently in Smith (1991), Vacca & Conti (1992) and Conti (1993). From Smith (1991) calibration, one WC4 star produces a luminosity (in logarithm of its value in erg s$^{-1}$) of 36.7 and 36.5 in the 4660 and 5810 bumps respectively, while a WN7 produces luminosities of 36.5 and 35.1 for the same bumps. The detection of some of these spectroscopic features implies the presence of a given kind of WR star, and can constrain the age and the metallicity of star-forming regions.

On the other hand, Red Supergiants affect the $V-K$ colour and, if they are present, strong CaII triplet, CaT, absorption features in the near infrared ($\lambda\lambda$ 8498, 8542, 8662 Å) should be detected, although the CaT shows a biparametrical behaviour with stellar surface gravity and metallicity which has to be taken into account (Díaz et al. 1989; Zhou 1991).

CaT features have been observed in GEHR surrounding the starburst nucleus of NGC 3310 (Terlevich et al. 1990; Pastoriza et al. 1993) and in one of these regions (labelled A and called “Jumbo” in their work) WR features have also been observed. More recently, a systematic search has been made for RSGs in the giant HII region NGC 604 in the spiral galaxy M 33 (GEFE Col. 1994, private communication) where the whole central core of the region was scanned to produce a very high signal-to-noise spectrum. No CaT was detected in this spectrum. However, in a specific region which showed WR features in the blue, CaT was detected in the IR spectrum once the Paschen emission lines were subtracted. This illustrates the difficulties involved in the detection of massive star features in GEHR.

The interpretation of these findings are even more challenging. The evolutionary models predict a small area in the metallicity-age diagram in which the WR and Red Supergiant phases could coexist (see Fig. 2). This fact led Pastoriza et al. (1993) to conclude that the “Jumbo” region probably underwent a prolonged activity of star-formation with ages between 3 and 15 Myr. But this conclusion is more questionable in the case of NGC 604 where the rest of the observations are consistent with a single short duration burst (Drissen et al. 1993).

CaII triplet equivalent widths synthesized for single stellar populations (SSP) of different ages at solar metallicity can be seen in García-Vargas et al. (1992, 1993).

3. Spectral energy distributions of ionizing star clusters

We have synthesized the emergent spectrum of an evolving star cluster by calculating the number of stars in each point of the isochrone and assigning to it the most ad-
equate stellar atmosphere model, i.e. the closest one in effective temperature and surface gravity. The stellar spectrum has then been scaled to the luminosity of the corresponding theoretical star in the HRD.

To build our stellar spectral library we assembled different stellar atmosphere models. We adopted those of Clegg & Middlemass (1987) for stars with $T_{\text{eff}} \geq 50000$ K (corresponding to the last evolutionary stages of massive stars) and those of Kurucz (1992) for stars with $35000 \leq T_{\text{eff}} < 50000$ K. We have chosen these models for two reasons: first, because there is a good agreement between Kurucz’s and Clegg’s & Middlemass’ models of a same $T_{\text{eff}}$, which provides a consistent effective temperature scale, and second, because Kurucz’s grid includes models in a wide range of metallicities, between $10^{-5} Z_{\odot}$ and 10 $Z_{\odot}$, allowing us to build a consistent grid of models, in which the effect of the metallicity is taken into account not only in the evolutionary tracks but also in the atmosphere models. However, these models do not include blanketing effects which could be important for the highest metallicity stars.

Figures 3a-d show the synthesized spectra of a cluster of $0.6 \times 10^3 M_{\odot}$ with the “standard” IMF ($\alpha=2.35$, $m_{\text{low}}=0.85 M_{\odot}$, $m_{\text{up}}=120 M_{\odot}$) at 2, 3, 4 and 5 Myr respectively. Different lines correspond to SEDs for clusters at different metallicities ($Z=0.001, 0.004, 0.008, 0.02$ and 0.05), as labelled in the figure.

The most remarkable characteristic of the spectra is the ionizing continuum beyond 1 Ryd which becomes harder when the extreme WR stars appear, a fact which depends on both the age and the metallicity of the cluster, as discussed in the previous section.

The number of Lyman ionizing photons per solar mass of the cluster, $Q_m(\text{H})$, is controlled by massive O-B main sequence stars and, as shown in Fig. 4, does not change much either with metallicity or during the evolution of a given cluster. Only for the highest metallicity case this change is larger than an order of magnitude. This fact implies that the intensities of recombination emission lines of the gas like those of the Balmer series, will be mainly dominated by the mass of the cluster, but not by its age or metallicity.

However, since $Q(\text{He})$ is dominated by the extreme WR stars, whose appearance and lifetime are governed by the metallicity of the cluster, the ratio of the number of helium ionizing photons, $Q(\text{He})$ ($\nu \geq 1.8$ Ryd), to total ionizing photons, $Q(\text{H})$ ($\nu \geq 1$ Ryd), i.e. the “hardness” of the spectrum, changes appreciably, both with the cluster metallicity and age.

In the present models, the atmosphere models used to represent WR stars are those of Clegg & Middlemass (1987) which do not include the existence of a wind. Leitherer & Heckman (1995) have built a grid of synthetic clusters, using the theoretical SEDs for WR atmospheres given by Schmutz et al. (1992). These models include a thick wind, characterized by its velocity law and the mass loss rate. We have compared our results for the numbers of $Q(\text{H})$ and $Q(\text{He})$ respectively, with those obtained by these authors and the agreement is better than 20% for the metallicities in common (0.008, 0.02, 0.05), in spite of using different codes, and different (Geneva group) stellar evolutionary tracks.

In previous works (Mas-Hesse & Kunth 1991; García-Vargas & Díaz 1994) an equivalent ionizing temperature, $T_{\text{eq}}$, for each cluster has been defined as the $T_{\text{eff}}$ of a star whose spectrum has the same value of $Q(\text{He})/Q(\text{H})$ as that of the whole cluster. This definition is of course very model-dependent. In particular it depends, not only on the physical properties of the single stars (gravity and metallicity), but also on the chosen atmosphere models, and both things should be specified when assigning effective temperatures to GEHR.

For unvelevated, zero age clusters, the equivalent temperature, $T_{\text{eq}}$, probably denotes the temperature of the most massive, and therefore the hottest, stars in the clusters. The temperature of the hottest stars in the main sequence increases as $Z$ decreases. This provides a natural explanation to the low values of equivalent temperatures inferred for high metallicity HII regions (see McGaugh 1991). However this approach is valid only during the period when the ionization is controlled by the O-B main sequence stars, i.e. during the 2.5–3 Myr after the burst of star formation. This is no longer the case for older clusters in which evolved stars can actually be hotter and more luminous than the most massive stars still on the main sequence. They are certainly less numerous than the hot main sequence stars, but their high luminosities can compensate this handicap letting them control the shape of the ionizing radiation and therefore dominate the ionization of the nebula.

Figure 5 shows the evolution of $T_{\text{eq}}$ with the age of the cluster at different metallicities. For ages less than 3 Myr a trend is observed in the sense of $T_{\text{eq}}$ being lower at higher metallicities as expected. For ages greater than 3 Myr, when WR stars form part of the cluster, this trend disappears and any interpretation of the shape of the ionizing radiation in terms of the metallicity alone becomes impossible.
4. Photoionization models

The cluster ionizing spectra computed as described in the previous section have been used as input for the photoionization code CLOUDY (Ferland 1991). The resulting emission-line spectra of the ionized regions have been calculated, under certain hypotheses about the geometry and the physical and chemical conditions of the gas. A spherical distribution for the gas has been assumed. In this approach the gas is considered to be placed in a thin shell at a certain distance, called the inner radius, from the ionizing source. In these calculations, given the large value of the inner radius used, this geometry is equivalent to an effectively plane-parallel geometry. In these models, the regions are considered to be ionization bounded.

The ionization parameter, $u$, is defined as the ratio between the density of the ionizing photons and the density of particles, and it is equal to $u = Q(\text{H})/4 \pi c n_H r^2$. The density of the ionizing photons is calculated from the total number of Lyman continuum photons ($\nu \geq 1$ Ryd), $Q(\text{H})$, which reach the region of radius $r$ at a velocity $c$ and $n_H$ is the density of ionized hydrogen, taken to be equal to the electronic density, $n_e$, if the region is completely ionized. Contributions to $n_e$ from ionized helium amount to only 10% in the most excited regions. Constant density throughout the nebula has been assumed for simplicity. The assumed value of 10 cm$^{-3}$ is in fact a lower limit to the density. Some emission lines are density sensitive. However this effect is only significant for values of $n_e \geq 500$ cm$^{-3}$. Therefore, these models should not be used for star forming systems in which the measured densities are higher than this value (e.g. most starburst nuclei).

Values for the average radius of the regions between 76 and 229 pc have been considered to cover a wide range in ionization parameter $-4.1 \leq \log u \leq -1.6$. These sizes are characteristic of extragalactic HII regions (e.g. Van der Hulst et al. 1988; Díaz et al. 1991; González-Delgado et al. 1994, 1995).
Fig. 4. Number of ionizing photons per unit mass of the cluster versus the age (in Myr). Different symbols correspond to different metal abundances.

Fig. 5. Evolution of $T_{eq}$ of the ionizing cluster versus the age at different metallicities. Symbols are labelled as in Fig. 4.
Some amount of dust is expected to be associated with the star cluster, but its effect has not been considered in this set of models, except for a depletion of the heavy elements. The most important effect would be the absorption of part of the Lyman continuum photons by this dust, which would be re-emitted in the infrared.

The values of solar abundances assumed in our set of models are the same as in Stašińska (1990) (including Ca, Mg, Si and Fe, which were depleted by a factor 10, to take into account the effect of dust), except the sulphur abundance, for which the ratio S/O given by Grevesse & Anders (1989) has been taken.

Table 2 summarizes the input parameters for the grid of models whose results are presented in great detail in García-Vargas et al. (1994). Some results can be inspected by means of the diagnostic diagrams of Figs. 6 to 10, in which the ratios of the most common optical emission line intensities, relative to Hβ have been plotted versus the value of the ionization parameter, u. Results for different metallicities (Figs. 6 (Z=0.001), 7 (Z=0.004), 8 (Z=0.008), 9 (Z=0.02) and 10 (Z=0.05)) and ages: 2 Myr (a), 3 Myr (b), 4 Myr (c), and 5.2 Myr (d) are shown.

Table 2. Model input parameters

<table>
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<th>IMF parameters:</th>
<th>( \alpha = 2.35 ), ( m_{\text{low}} = 0.85 ) ( M_\odot ),</th>
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<tr>
<td>( m_{\text{up}} = 120 ) ( M_\odot )</td>
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<tr>
<td>( Z = 0.001, 0.004, 0.008, 0.02 ) (solar), 0.05</td>
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<tr>
<td>( M_T / M_\odot = 0.12, 0.2, 0.4, 0.6 \times 10^5 )</td>
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To test the adequacy of the present models to reproduce actual high metallicity GEHR we have plotted in Fig. 11 the ratio log \( ([\text{O III}]4959+5007/H_\beta) \) versus log \( ([\text{O II}]3726+3729)/[\text{O III}]4959+5007 \) for models and observations. Different types of open symbols represent the evolution of star clusters of different mass, as labelled in the figure. These points, representative of the clusters at 1.5, 2, 2.5, 3, 3.5, 4, 4.5, 5, 5.1, 5.2, 5.3 and 5.4 Myr are joined with a solid line or a dashed line for solar and twice solar metallicity respectively. Observational data are plotted as filled symbols and constitute a compilation from several works published in the last ten years (Phillips et al. 1984; Díaz 1985; McCaill et al. 1985; Vilchez et al. 1988; Díaz et al. 1990, 1991; Henry et al. 1992; Oey & Kennicutt 1993 and Zaritski et al. 1994). Their associated errors, not shown in the plot, can be quite large in some cases, since high metallicity GEHR usually present very low surface brightness.

In our previous work (García-Vargas & Díaz 1994) this excitation diagram showed a disagreement between models and observations in the sense of a lack of observed high excitation regions which were theoretically predicted. This disagreement could, in principle, be explained by changes in the IMF parameters. As can be seen from Fig. 11, our new set of models, computed with a standard IMF, seems to represent the observations of GEHR rather well.

The filling factor, \( \epsilon \), corresponding to each model can be derived from the ratio \( \epsilon = \left( \frac{r}{r_1} \right)^3 \), where \( r \) is the average radius of the ionized region, and \( r_1 \) is the Strömgren radius, defined as \( r_1 = \left( \frac{3Q(H)}{4\pi\alpha_B n_H^2} \right)^{1/3} \) (Osterbrock 1989). The values of Case B recombination coefficients for hydrogen, \( \alpha_B(n_H, T) \) at 100, 4.54, 2.59, and 2.52 (in units of \( 10^{-13} \) cm\(^3\)s\(^{-1}\)) for electron temperatures of 5000, 10000, and 20000 K respectively (Osterbrock 1989). As an illustrative example let us consider the models for a cluster of 0.6 \( M_\odot \) at \( Z = 0.02 \) (solar). The average value of the gas electron temperature over the evolution of the cluster between 1.5 and 5.4 Myr is 7000 K, which implies a value of \( 3.76 \times 10^{-13} \) cm\(^3\)s\(^{-1}\) for \( \alpha_B \). The filling factors calculated with the two extreme values of ionizing photons, (log \( Q(H) = 50.4, 51.6 \)) are 0.003 and 0.05 respectively. These values are similar to those usually found in GEHR (e.g. see Díaz et al. (1991) for the HII regions in M 51).

The mass of the ionized gas can be calculated as the sum of the ionized hydrogen \( M(\text{H}^+) \) and single ionized helium \( M(\text{He}^+) \). Following Osterbrock (1989) the total mass of ionized gas under the assumed geometrical hypotheses is:

\[
M(\text{H}^+) + M(\text{He}^+) = \frac{1 + 4y^+}{1 + y^+} \frac{m_p \epsilon Q(H)}{n_e \alpha_B(H,T)}
\]

where \( y^+ \) is \( \text{He}^+/\text{H}^+ \) by number, \( m_p \) is the proton mass, \( \epsilon \) the filling factor, \( Q(H) \) the total number of ionizing photons, \( n_e \) the electron density and \( \alpha_B(H, T) \) the total hydrogen recombination coefficient.

As an example, the masses of ionized gas for the models with \( Z = 0.02 \) and \( M_T = 6 \times 10^4 \) \( M_\odot \) are between 0.2 \( 10^4 \) \( M_\odot \) and 60 \( 10^4 \) \( M_\odot \) which are 0.03 and 10 times the mass of the corresponding star cluster.

Massive stars release large amount of gas to the surrounding interstellar medium in the course of their lifetime, through stellar winds and supernovae. In the present model range of ages (1–5.4 Myr) most of injected gas comes from winds. Figure 12 shows the ratio between the total injected mass by winds and the total mass of the star cluster in the case of solar composition. This value increases with age, reaching 0.15 at 4–5 Myr. This means that the star cluster can depose up to 15% of its initial
mass in the interstellar medium, only by stellar winds, which will mix with the ionized gas. Part of this material will come from WR winds and could therefore be enriched in He, C, N and O. However, this gas constitutes only a little fraction of the total and will be eventually diluted in the IS medium producing little global enrichment (Leitherer 1994). This result agrees with spatially resolved observations of some GEHR regions in which no substantial enrichment is derived from the analysis of the emission line spectrum of the gas, even at the positions coincident with WR features (e.g González-Delgado et al. 1994, 1995).

### 5. Physical parameters for giant extragalactic HII regions

In a previous work (García-Vargas & Díaz 1994) we defined the physical parameters of GEHR as the mass, age and metallicity of the single star cluster assumed as ionization source. In this section we give a simple method to get these cluster parameters by using the ionized gas emission line spectrum as observational tool.

The number of photons ionizing a given HII region, \( Q(H) \), can be derived from the extinction corrected \( H_\alpha \) or \( H_\beta \) intensity if the distance to the parent galaxy is known (Osterbrock 1989):

\[
Q(H)(s^{-1}) = 7.31 \times 10^{11} I_{H_\alpha} (\text{ergs}^{-1})
\]

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If the region is assumed to be ionization bounded and no dust is present, this number coincides with the number of ionizing photons emitted by the central source. Since the number of ionizing photons per solar mass does not change much with the age of a given cluster, its total mass can be estimated if its metallicity is known (see Fig. 4 and Sect. 3).

The metallicity of a given HII region can be derived quite accurately for the cases in which the electron temperature can be calculated. This requires the observation of weak temperature-sensitive lines, being [OIII] λ 4363 Å the most commonly used one. There are observational difficulties for the measurement of this line, like its proximity to the Hβ, Balmer line and the HgI night sky line at λ 4359 Å which demand good quality and relatively high spectral resolution observations. Once the electron temperature is known, ionic abundances can be obtained, in the low density limit, for the intensities of collisionally excited lines (e.g. Pagel et al. 1992). A final correction has to be made for unseen ionization states of a given element.

The cases in which the λ 4363 Å is not measured require the use of empirical calibrations of strong line intensities with metallicity. The one relating the intensities of the [OIII] λ 4959, 5007 Å and [OII] λ 3727, 3729 Å lines to the oxygen content (Pagel et al. 1979) is widely used. The calibration has been made empirically for low metallicity regions (Edmunds & Pagel 1984; Skillman 1989) and with the use of photo-ionization models for high metallicity ones (McCall et al. 1985; McGaugh 1991). These calibrations can be used with some caution to estimate metallicities up to the solar value.

Higher metallicity regions require a more careful study as explained in Díaz et al. (1991) and Díaz (1994) since conventional methods do not allow a sufficiently accurate abundance determination (Oey & Kennicutt 1993). However, for our purposes, the finding that a given HII region possesses a metallicity higher than solar is enough to allow the selection of the appropriate set of models.

Following the method described above, the total mass of the ionizing cluster can be estimated. Since in our models
the possible presence of dust is not taken into account this estimate constitutes a lower limit to the actual mass.

The most difficult parameter to derive is, without doubt, the age of the star cluster. Some observables can be used as age tracers. Obviously, the detection of WR bumps or RSG features can help to constraint the age of a cluster, as discussed in Sect. 2. Another important parameter could be the Hβ equivalent width, which decreases with the age of the burst (Mas-Hesse & Kunth 1991; García-Vargas & Díaz 1994; García-Vargas et al. 1995). However, in the presence of a mixture of populations, as would be the case if a previous burst had taken place, the older population would dominate the Hβ continuum, and the Hβ equivalent width would not be representative of the very young cluster ionizing the gas. As an alternative approach, we suggest the use of the ionization parameter, $u$, as an indicator of the ionizing power of the star cluster. Once the mass and metal content of the cluster are known, the ionizing power is a function of the age alone.

The ionization parameter, $u$, can be derived from our models by using suitable emission line ratios. At low metallicity logu can be calibrated against the sum of the S$^+$ lines, SII]+6716+6731, once the value of $Z$ is known (Fig. 13a). A least-square polynomial fitting gives:

$$\log u = -0.82r_1^2 - 4.14r_1 - 6.56; \quad Z = 0.001 \quad (3)$$

$$\log u = -0.80r_1^2 - 3.30r_1 - 4.86; \quad Z = 0.004 \quad (4)$$

where $r_1 = \log([\text{SII}]6716 + 6731/H_{\alpha})$.

At high metallicity (Fig. 13b), the ratio ([SII]+6716+6731)/([SIII]9069+9532) provides a metallicity independent calibration of $u$ at least for $\log u \leq -2.5$ (Díaz et al. 1991; Díaz 1994). A linear least-square fit to
the data gives:

$$\log u = (-1.18 \pm 0.02)r_2 + (-3.08 \pm 0.01); \ Z \geq 0.008$$

(5)

where $r_2 = \log([[\text{SII}]6716 + 6731]/[\text{SIII}]9069 + 9532)$.

In summary the physical parameters of the ionizing star clusters in GEHR can be obtained in the following way: 1) the metallicity can be deduced from the following line intensities of the different elements by deriving the electron temperature of the gas when possible, e.g. in low metallicity regions, or using empirical calibrations. In the case of regions of high metallicity the diagnostic diagram showed in Fig. 11 can also be used to discriminate if the abundance is solar or higher than solar. 2) The mass of the star cluster can be derived directly from the $H_\alpha$ luminosity, through the calibration of the number of ionizing photons, in absence of dust. 3) The age can be estimated from the ionization parameter through the $[\text{SII}] / H_\alpha$ or $[\text{SII}] / [\text{SIII}]$ ratios, for low and high metallicities respectively. Other observational features, like WR bumps, CaII triplet lines or $H_\beta$ equivalent width can provide additional information about the age of the dominant ionizing population in GEHR.

Finally, the integrated spectrum of a cluster with the derived values of mass, age and composition can be used as input for a photoionization code in order to reproduce the whole emission line spectrum of the gas, for a more detailed analysis.

6. The IR emission-line spectrum for GEHR and the implications for the IRAS colors

Usually, IRAS excesses observed in GEHR and HII galaxies have been claimed to be produced by dust. In particular 25 $\mu$m excesses in star-forming regions have been explained as emission from very hot or shock heated dust (Dultzin-Hayyan et al. 1988), and 12 $\mu$m excesses as

However, some contribution to the IR flux is expected by the infrared emission lines which, when the metallicity increases, begin to control the cooling of the nebula taking over the optical ones. This effect can be seen in our grid of models for $Z \geq 0.008$ (García-Vargas et al. 1995). Some of these IR emission lines are in the wavelength range of IRAS bands and their contribution has been suggested as an important pollution agent of IRAS colours in the case of planetary nebulae (Preite-Martínez & Pottasch 1987). We have computed the corresponding contributions for GEHR using the results of our models.

We have calculated the total luminosity of the main emission-lines in each IRAS band, relative to H$_\alpha$, by taking the intensities of the IR lines from our models and weighting them by the corresponding response curves. The values of the system spectral response for IRAS bands have been taken from Neugebauer et al. (1984).

Figures 14(a-d) show the results for each of the IRAS bands. The points in the plot represent the values for clusters of 0.12, 0.2, 0.4 and 0.6 $10^5$ $M_\odot$ and the whole range of considered ages. Different symbols correspond to different values of the metallicity. The contribution of IR lines to the 25 $\mu$m band is found to be negligible, but this is not the case for the other bands where the luminosity in emission lines can reach values as high as the H$_\alpha$ luminosity. If dust is present, its contribution to the FIR (60 and 100 $\mu$m) can render the emission-line pollution negligible. Some theoretical predictions about FIR luminosity in star-forming regions can be seen in Mas-Hesse & Kunth (1991).

We would like to stress the importance of infrared spectroscopical observations for the analysis of GEHR, especially for the high metallicity ones, in which the optical
Fig. 11. Diagnostic diagram. Comparison between our evolutionary models for high metallicity GEHR and the observations. Observational data are plotted as filled symbols and they are a compilation from several works published in the last ten years (see references in the text). Different types of open symbols represent the evolution of star clusters of different mass, as labelled in the figure. These points, representative of the clusters at 1.5, 2, 2.5, 3, 3.5, 4, 4.5, 5, 5.1, 5.2, 5.3 and 5.4 Myr are joined with a solid line or a dashed line for solar and twice solar metallicity respectively.

Fig. 12. Mass lost by stellar winds, respect to the initial cluster mass in stars as a function of the age of the cluster (in Myr)
diagnostic diagrams are not the most adequate to study the properties of the gas, and consequently the physical parameters of the star clusters responsible for the ionization. Unfortunately, the ISO spatial resolution will not be able to resolve most individual GEHR and only for galactic nuclei and GEHR in nearby galaxies relevant information will be obtained. However the integrated ISO observations will be very important to test the IR contribution of star-forming regions.

As a first suggestion we propose the use of [CII]158 μm line as tracer of star-formation in high metallicity environments. Figure 15 shows the emission-line intensity ratio of [CII]158 μm to H_α versus the age of the cluster, for clusters of different metallicity. For the highest metallicity models this ratio is as high as 3. This fact implies that IR images in [CII]158 μm could be used as high metallicity star-forming tracers in galaxies, instead of the traditional H_α images. We suggest narrow band imaging with ISOPHOT as a test of this result.

7. Summary and conclusions

We have followed the evolution of very young star clusters (1−5.4 Myr) in a wide range of metallicities (1/20 − 2.5 solar) using the stellar evolution models computed by the Padova group. These clusters are assumed to be the ionization source of GEHR, since clusters older than these ages cannot provide the required ionizing photons. For the range of ages considered, evolutionary effects appear for stars with masses higher than 20 M_☉ with metallicity dependent mass loss being an important parameter which controls the properties of WR stars. These stars are responsible for the hardening of the cluster SED. The present models reproduce the observed integrated nebular spectra of GEHR without changes in the standard IMF parameters.

The main goal of the present work is to offer a simple way to derive the physical parameters — mass, age and metallicity — of the stellar clusters responsible for the ionization of GEHR, using the observed emission line spectrum of the photoionized gas.

- The mass in stars is derived from H_α luminosity, almost independently of the metallicity. This mass constitutes a lower limit if dust is present.
- The metallicity can be derived by standard methods except in the cases of higher than solar abundances which the present models can help to identify.
- Finally, for a given metallicity, the age of the cluster can be estimated by the ionization parameter, u, obtained from [SII]/H_α or from [SII]/[SIII] ratios for low and high metallicities respectively. The presence of stellar features from different types of WR and RSG can actually help to restrict the possible ranges both in age and metallicity.

The importance of the infrared GEHR analysis is stressed since at high metallicity (Z > 0.008) the cooling of the gas through IR emission lines becomes increasingly dominant. Therefore IR diagnostic diagrams would be more useful to analyze the properties of high metallicity gas. Finally, [CII]158 μm line narrow imaging is suggested as star-formation tracer in high metallicity scenarios, instead of the traditional H_α imaging.

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Fig. 14. Luminosity in each of the IRAS band due to contamination from the IR emission line spectrum, at 12 μm a), 25 μm b), 60 μm c), and 100 μm d). Different symbols correspond to different values of the metallicity as labelled in the plot.

Fig. 15. [CII]158 μm /Hα as a function of the age of the cluster for high metallicity models
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