Rev. Mexicana Astron. Astrof., 5, 109-124 (1981).

CHEMICAL EVOLUTION OF GALAXIES. I. CONSTRAINTS IMPOSED BY THE $\Delta Y/\Delta Z$ RATIO

A. Serrano and M. Peimbert

Instituto de Astronomía
Universidad Nacional Autónoma de México
Received 1980 September 3

RESUMEN

Hemos calculado modelos de evolución química de galaxias considerando: a) pérdida de masa y producción de helio en estrellas de masa intermedia basándonos en modelos estelares de evolución en la rama gigante asintótica; b) varias tasas de pérdida de masa en estrellas masivas; c) varias funciones iniciales de masa, como la de Serrano (1978), y d) la evolución con y sin reciclaje instantáneo.

De estos modelos hemos derivado la razón de abundancias $\Delta Y/\Delta Z$ y concluido que la razón observada de $\Delta Y/\Delta Z$ nos permite reducir de una manera considerable el número de modelos viables de evolución estelar y galáctica.

En particular, se discuten en el texto restricciones a la evolución estelar en la rama gigante asintótica, a la función inicial de masa, a la escala de distancias de las nebulosas planetarias, a la pérdida de masa en estrellas masivas y a la teoría estándar de la gran explosión.

ABSTRACT

We have computed chemical evolution models considering: a) mass loss and He production in intermediate mass stars based on models of asymptotic giant branch evolution; b) several rates of mass loss for massive stars; c) various initial mass functions, including that by Serrano (1978), and d) evolution with and without instantaneous recycling.

From these models we have derived the $\Delta Y/\Delta Z$ abundance ratio and concluded that the observed $\Delta Y/\Delta Z$ ratio provides strong restrictions for galactic and stellar evolution models.

In particular, we discuss constraints for asymptotic giant branch stellar evolution models, for the initial mass function, for the distance scale to planetary nebulae, for mass loss in massive stars and for the standard big-bang theory.

Key words: COSMOLOGY – GALAXIES-EVOLUTION – GALAXIES-STELLAR CONTENT – INTERSTELLAR-ABUNDANCES – STARS-MASS LOSS.

I. INTRODUCTION

Hacyan *et al.* (1976) produced chemical evolution models of the solar neighborhood without being able to fit the observed helium to heavy element enrichment ratio by mass, $\Delta Y/\Delta Z$, and the yield of primary elements, p, at the same time. For p values compatible with observations, the theoretical $\Delta Y/\Delta Z$ ratios turned out to be significantly smaller than the observed one; while for $\Delta Y/\Delta Z$ ratios compatible with observations, the p values turned out to be significantly smaller than the observed one.

In the meantime there have been six new developments, not considered by Hacyan *et al.* (1976), that increase the predicted $\Delta Y/\Delta Z$ ratio: i) helium production by planetary nebulae (Peimbert and Serrano 1980*a*); ii) helium enrichment due to the mixing of helium into the envelopes of double-shell source stars undergoing helium shell flashes (Gingold 1977); iii) hot-bottom burning in intermediate mass stars (Renzini and Voli 1980); iv) a new initial mass function which gives higher weight to the 2 to 10 $M_{\odot}$ objects (Serrano 1978); v) evolution with mass loss in massive stars (Chiosi *et al.* 1978; Chiosi and Caimmi 1979); and vi) binary evolution (Vanbeveren and Olson 1979). Moreover, recent p determinations indicate lower values than those generally accepted a few years ago.

Due to these new developments we decided to compute chemical evolution models to try to reach agreement between observed and computed $\Delta Y/\Delta Z$ values. A discussion on the heavy element yield is given elsewhere (Peimbert and Serrano 1980*b*, hereinafter Paper II).

In §II we summarize the observed pregalactic helium abundance, Y_p, and the $\Delta Y/\Delta Z$ values; in § III we describe the models of chemical evolution that will be used in this and subsequent papers of this series; in §IV we discuss the results from the calculations and some cosmological implications, and finally in §V we present the conclusions.

II. OBSERVED Y_p AND $\Delta Y/\Delta Z$ VALUES

Table 1 presents what we considered to be the best determinations of Y_p and the $\Delta Y/\Delta Z$ abundance ratio. Yang *et al.* (1979) have also presented a discussion of the helium abundances based on stellar data. In what follows we will discuss some of the entries in Table 1.

We decided to present the results for gaseous nebulae for two values of the mean-square temperature variation along the line of sight. The case of uniform temperature along the line of sight, $t^2 = 0.0$, provides us with absolute lower limits to the heavy element abundances derived from the forbidden lines; while $t^2 = 0.035$, the value derived for the Orion Nebula (Peimbert and Torres-Peimbert 1977), might be close to the t^2 upper limit for gaseous nebulae, considering that the Orion Nebula shows extreme spatial density fluctuations and is probably composed of two disconnected H II regions ionized by stars of different effective temperatures (Peimbert and Torres-Peimbert 1981*a*). Some of the extragalactic H II regions are actually H II region complexes where the lines of high degree of ionization come from regions ionized by early type O stars and the lines of low degree of ionization come from regions ionized by late type O stars (e. g. Alloin *et al.* 1979). For these objects t^2 is relatively large, while for constant density objects photoionized by one or several stars inside the same gaseous nebula t^2 should be closer to 0.0 than to 0.035.

From independent methods, newer values of t^2 have been obtained. By comparing the $C^{++}\lambda\lambda 1907 + 1909$ intensity to that of $C^{+}\lambda 4267$ it has been found that $t^2 = 0.016$ for the Orion Nebula and $t^2 = 0.02$ for the planetary nebula IC 418 (Torres-Peimbert, Peimbert and Daltabuit 1980), while from $\lambda\lambda 1663$, 4363 and 5007 of [OIII] it has been found that $t^2 = 0.00$ for the planetary nebula NGC 3918 (Torres-Peimbert, Peña and Daltabuit 1980). From radio observations and optical data of galactic H II regions it has been found that $t^2 = 0.015$ (Shaver 1980). From this discussion it follows that for H II regions, $t^2 \cong 0.02$; while for planetary nebulae of high degree of ionization, $t^2 \cong 0.00$.

The entries in Table 1 for extragalactic H II regions, planetary nebulae and binary stars correspond to linear regression solutions and their standard deviations.

Based on the data for extragalactic H II regions in

TABLE 1

OBSERVED VALUES OF Y_p AND $\Delta Y/\Delta Z$

Object	Y_p	$\Delta Y/\Delta Z$	t^2 (10^{-2})	Source
Extragalactic H II regions	0.233 ± 0.005	1.7 ± 0.9	3.5	1
"	0.230 ± 0.005	3.2 ± 1.4	0.0	1
"	0.216	2.8 ± 2.7	0.0	2
"	0.216	3.2	0.0	3
"	0.237 ± 0.005	2.2 ± 1.2	0.0	4
LMC and SMC H II regions	0.231	2.1	3.5	5
"	0.229	3.6	0.0	5
Planetary Nebulae	0.254 ± 0.009	2.2 ± 0.5	3.5	6
"	0.249 ± 0.010	3.6 ± 0.8	0.0	6
Main sequence FGK stars	....	5 ± 3	....	7
"	0.2	3.5:	....	8
Supergiant stars	....	4.5:	....	9
Binary stars	0.235 ± 0.010	3.8 ± 1.5	....	10
Globular clusters	0.23 ± 0.02	....	....	11
Dwarf spheroidal galaxies	0.24 ± 0.02	....	....	12

(1) Lequeux *et al.* 1979, (2) Talent 1980, (3) French 1980, (4) Kunth and Sargent 1980, (5) Peimbert and Torres-Peimbert 1974, 1976, (6) Peimbert and Serrano 1980a, (7) Perrin *et al.* 1977, (8) Faulkner 1967, (9) Chiosi and Nasi 1974, (10) Saio 1978, (11) Caputo *et al.* 1978, (12) Hirshfeld 1978.

Lequeux *et al.* (1979) we computed the values for $t^2 = 0.0$ giving equal weight to each galaxy. The values for the Magellanic Clouds were obtained from the observations by Peimbert and Torres-Peimbert (1974, 1976) using more recent atomic data and the H II regions average value for each galaxy (see Lequeux *et al.* 1979). To derive the Y_p and $\Delta Y/\Delta Z$ values from binary stars we used the results by Saio (1978) without considering ADS 520 which shows the largest uncertainties in his sample.

The pregalactic helium abundance is almost independent of t^2. The values for H II regions are of higher accuracy than the planetary nebulae ones due to the better sample of metal poor objects in H II regions and to the possibility of He and O enrichment in the PN envelopes produced by their own evolution. In what follows, we will adopt $Y_p = 0.23$ which is in agreement with the entries in Table 1.

The $\Delta Y/\Delta Z$ values derived from gaseous nebulae are of higher accuracy than those derived from stellar data. In what follows, a value of $\Delta Y/\Delta Z = 3 \pm 1$ will be used as representative of the values in Table 1.

III. THE MODEL OF CHEMICAL EVOLUTION

Let us define the stellar birthrate B(m,t) so that

$$dN = B\, dm\, dt \qquad (1)$$

is the number of stars born with masses between m and m + dm at times between t and t + dt. Then,

$$\dot{N}(t) = \int_0^\infty B\, dm \qquad (2)$$

represents the rate of star formation, by number, at time t. We can then define the initial mass function (IMF) by

$$\phi(m, t) = B(m, t)/\dot{N}(t)\ , \qquad (3)$$

so that

$$\int_0^\infty \phi(m)\, dm = 1 \qquad (4)$$

and ϕdm can be interpreted as the probability of forming a star of mass between m and m + dm.

Notice that the rate at which mass is being converted into stars (the SFR) is

$$\dot{S}(t) = \dot{N} \int_0^\infty m\, \phi\, dm\ ; \qquad (5)$$

i.e. $\dot{S}(t) = \langle m \rangle \dot{N}$,

where $\langle m \rangle$ is an average stellar mass.

To follow the chemical evolution of this spatially homogeneous model means that we obtain the fraction by mass of gas in species i in the interstellar medium (ISM), $X_i(t)$, as a function of time.

Equivalently, one may follow the evolution of the mass of gas in form of species i in the ISM, G_i, as proposed by Talbot and Arnett (1971). The total mass of gas G is

$$G = \sum_i G_i \qquad (6)$$

and the fractional abundances can be expressed by

$$X_i = G_i/G \qquad (7)$$

The solution to the problem of chemical evolution can be obtained by the following scheme:

i) Start off with a mass of gas composed of primordial H and He.

ii) Let stars form according to a certain IMF and a poorly known star formation rate.

iii) Let the stars burn and fertilize the ISM with eventual ejection of part of the ashes.

iv) Allow for exchange of material (gas and/or stars) in and out of the system, altering then the existing chemical composition.

This simple procedure becomes more complicated in the case of the light elements (see e.g., Audouze and Tinsley 1974; Reeves *et al.* 1973) and in the case of elements heavier than Fe. In these cases, other processes of formation and destruction of the nuclei can be important. Nevertheless, for the purposes of this paper in which we will be interested in He and in elements between C and Fe, the model described above will be adequate.

Thus, once a certain nucleus forms, it can only be destroyed either by leaving the system, or by transmutation into a heavier nucleus in the interior of a star.

From the point of view of the gas, we have that

$$\dot{G}_i = -\dot{S} X_i + \dot{O}_i + \dot{E}_i \qquad (8)$$

where, assuming that stars form with the composition of the gas, $\dot{S}X_i$ is the rate at which mass in the form of species i is transformed into stars; $\dot{O}_i$ is due to mass motions in and out of the system; and $\dot{E}_i$ is the rate at which stars return material of species i back to the ISM.

We will largely ignore $\dot{O}_i$, which corresponds to the physics external to our system, and we will concentrate here on the rate of ejection $\dot{E}_i$.

It can be shown that if the time it takes a star of initial mass m to eject a mass $m_{ej}(m)_{(i)}$ of species i is short compared with the timescale at which the SFR is changing, $S/\dot{S}$, then we can express $\dot{E}_i$ by

$$\dot{E}_i(t) = \int_{m(t)}^{m_U} \dot{N}(t-\tau(m))\, m_{ej}(m)_{(i)}\, \phi\, dm. \quad (9)$$

Where $\tau(m)$ is the age of a star of mass m at the time of ejection (it was then born at the time $t-\tau(m)$), and stars of masses less than m(t) have not had time to evolve and eject material of species i yet.

Since we expect stars having more fuel to produce more ashes, we use

$$m_{ej}(m)_{(i)} = \sum_j m_{ij} X_j \quad (10)$$

(Talbot and Arnett 1973), where

m_{ij} = mass of the region of a star of initial mass m in which species j has been converted into species i and ejected to the ISM. (11)

m_{ij} must not be confused with the actual mass ejected as i with a parent j which is $m_{ij}\, X_j$.

We only need to specify:
i) Initial conditions, $X_i(0)$, and stellar lifetimes, $\tau(m)$, ii) IMF, iii) SFR law and mass exchange law, O, iv) $m_{ij}(m)$, and solve the set of integro-differential equations (8) by a method similar to that proposed by Talbot and Arnett (1971).

a) *Initial conditions and stellar lifetimes.*

According to the discussion in §II we will adopt the initial abundances $(X, Y, Z)_0 = (0.77, 0.23, 0.00)$. Also, for the stellar lifetimes

$$\tau(m) = \tau_\infty + \tau_\odot\, m^{-3} \quad (12)$$

with $\tau_\infty = 0.0015\ (10^9$ yrs) and $\tau_\odot = 11\ (10^9$ yrs).

Note that heavy stars would have $\tau(m) \propto m^{-2}$, but we are mainly interested in the non-instantaneous recycling case which depends more on the lighter stars in which (12) is valid. We will also assume (for ease of computation) that $\tau(m)$ is independent of chemical composition.

b) *Initial mass function (IMF).*

We have assumed a power-law IMF independent of time and of chemical composition (see however Paper II for the dependence of very low mass stars with composition). This IMF has two different slopes for light and for heavy stars.

The IMF for massive stars has been derived independently by Serrano (1978) and Lequeux (1979). It is much steeper than that due to Salpeter (1955) and has been obtained in the case of Serrano from the CGO Catalogue (Cruz-González *et al.* 1974), solar neighborhood data (McCuskey 1966) and the luminosity function for open clusters by Taff (1974).

Determinations of the slope of the IMF in this mass range in other galaxies (Sandage and Katem 1973; Lequeux *et al.* 1980) seem to agree with that of the solar neighborhood.

The low mass part of the IMF has been derived by Serrano (1978) from a detailed discussion of the local stellar luminosity function.

As in Lequeux *et al.* (1979) we have arbitrarily assumed continuity of the two portions of the IMF at $1.8 m_\odot$. Setting the high mass and low mass limits of stellar formation at 110 $m_\odot$ and 0.007 $m_\odot$ respectively, the IMF is given by

$$\phi(m) = 0.0714\, m^{-3}, \text{ for } m > 1.8\, m_\odot$$
$$= 0.0313\, m^{-1.6}, \text{ for } m < 1.8\, m_\odot \quad (13)$$

Note that since $\dot{S}$ appears directly in equations (8), it is more convenient to rewrite eq. (9) as

$$\dot{E}_i = \int_{m(t)}^{m_U} \dot{S}(t-\tau(m))$$
$$\times \sum_j \frac{m_{ij}(m)}{m} X_j(t-\tau(m))\, \Phi(m)\, dm, \quad (14)$$

where

$$\Phi(m) = \frac{m\,\phi(m)}{\langle m \rangle} \quad (15)$$

is the IMF by mass, rather than by number, and also obeys

$$\int_0^\infty \Phi(m)\, dm = 1. \quad (16)$$

In our case

$$\Phi(m) = 0.56\, m^{-2}, \text{ for } m > 1.8\, m_\odot;$$
$$= 0.25\, m^{-0.6}, \text{ for } m < 1.8\, m_\odot. \quad (17)$$

As mentioned in Lequeux *et al.* (1979) the steeper IMF for massive stars enhances the role of the intermediate mass stars in the chemical evolution of a galaxy. In particular, it enhances their role as He producers.

c) *The star formation rate (SFR).*

For a long time it was thought that the SFR would decay approximately exponentially with time (e.g. Schmidt 1959; Talbot and Arnett 1971; Tinsley 1972; Thuan *et al.* 1975; Pagel and Patchett 1975; and many others). This behavior of the SFR is present, for example, under the simplest assumption that the SFR is proportional to the gas density.

However, the paucity of G dwarfs with low metallicities in the solar neighborhood, led Lynden-Bell (1975) to suggest a SFR not monotonic with time: increasing first up to a maximum and then decreasing for later times.

Under the assumption that the IMF is continuous, Serrano (1978) concluded that the SFR in the solar neighborhood has been approximately constant or even non-monotonic. A similar conclusion was obtained by Miller and Scalo (1979) when studying the IMF, and by Twarog (1980) from the age-metallicity relation for the disk in the neighborhood of the sun.

It is interesting to notice that all models that successfully reproduce the properties of the solar neighborhood (e.g. Thuan *et al.* 1975; Larson 1976; Chiosi and Matteucci 1980) share this behavior of the SFR; obviously, because accretion in one form or another was predominant when the solar neighborhood started forming stars (see also Paper II for a discussion of this point).

Since, on the other hand, the chemical evolution is only weakly dependent on the SFR (as far as ratios between abundances are concerned), we have adopted a constant SFR in most cases. Whenever we want to relate abundances with time or to count stars, the SFR becomes important. This is illustrated by some models that were calculated using a non-monotonic SFR.

Notice that with a constant SFR a typical spiral galaxy with 10% of its mass as gas and a life of 10^{10} yrs. would have SFR $\cong 0.01$ $M_{gal}/10^9$ yrs. However, Serrano (1978) estimated from star counts that in the solar neighborhood SFR $\cong 0.07$, a result similar to that found by Miller and Scalo (1979) from the same data. From the age-metallicity relation, Twarog (1980) found SFR $\cong 0.1$ $M_{gal}/10^9$ yrs.. This discrepancy arises because of the material returned by the stars (see § IVc below).

d) *The chemical evolution matrix.*

Our knowledge of the chemical evolution matrix

$$f_{ij} = m_{ij}(m)/m \qquad (18)$$

comes from stellar evolution models.

In the case of He only two coefficients are relevant:

m_{HeHe} = mass ejected without having burned He,

and

m_{HeH} =

mass of the stellar region where H was burned but He was not, and which was eventually ejected. (19)

A significant improvement in our understanding of f^{ej}_{He} has come from: i) incorporation of mass loss in evolutionary studies of massive stars (e.g., Chiosi and Nasi 1974; Dearborn and Eggleton 1977; Chiosi *et al.* 1978, 1979; de Loore *et al.* 1977, 1978) and ii) studies of advanced stages of stellar evolution, i.e. asymptotic branch evolution, in intermediate mass stars (e.g. Iben 1977 and references therein; Becker and Iben 1979, 1980; Iben and Truran 1978, hereinafter IT78; Renzini and Voli 1980, hereinafter RV80).

We are adopting the simplifying assumption that f_{ij} do not vary with chemical composition. This approximation, adopted by Talbot and Arnett (1971) and ever since, is not necessarily the best. However, to change it for a better description one would need a complete grid of stellar evolution models for a variety of initial chemical compositions. At the moment, only a few of this kind of calculations have been performed (e.g., Alcock and Paczynski 1978; Becker and Iben 1980; RV80), and they are not enough yet to determine the variation of f_{ij} with composition.

(i) The massive stars ($m > 10\ m_\odot$)

As it is usual in the field, we have adopted the results by Arnett (1978) on the evolution of He stars as the basic elements to calculate the chemical coefficients f_{ij} (m).

But, as pointed out by Chiosi and Caimmi (1979), Arnett's results give f_{ij} as a function of the mass, m_α of the He core at the point at which it decouples from the envelope (for more about this decoupling see § IV.*b*). That is, we obtain f_{ij} (m_α).

To proceed further, one must adopt a relationship between the mass, m_α of this core and the initial mass of the star. It has been shown by Chiosi and Caimmi (1979) that the main effect of taking into account mass loss in stars with $m > 20\ m_\odot$, is to change this m_α(m) relationship. A massive star loses matter from the He rich envelope, without affecting directly the chemical composition of the ISM; but a star losing mass will have a smaller He core than in the case of constant mass evolution.

The evolution of the core itself is always assumed to be as in Arnett (1978). We have approximated his He production by

$$
\begin{aligned}
m_{HeH} &= -1.044 + 1.108\, m_\alpha - 0.084\, m_\alpha^2, && \text{for } m_\alpha < 6\, m_\odot; \\
&= 0.897 + 0.378\, m_\alpha - 0.0166\, m_\alpha^2 + 0.00034\, m_\alpha^3, && \text{for } 6\, m_\odot < m_\alpha < 24\, m_\odot; \\
&= 0.192 + 0.201\, m_\alpha, && \text{for } 24\, m_\odot < m_\alpha < 32\, m_\odot; \\
&= 3.552 + 0.096\, m_\alpha, && \text{for } 32\, m_\odot < m_\alpha < 48\, m_\odot;
\end{aligned}
\tag{20}
$$

with similar expansions for heavier species.

On the whole, we have adopted Chiosi and Caimmi's (1979) m_α (m) relation, although slightly modified in the range between 10 and 20 $m_\odot$ in order to make a smoother extrapolation to less massive stars, where this kind of mass loss becomes less important.

(ii) Intermediate mass stars (IMS) ($1\ m_\odot < m < 8\ m_\odot$)

Following RV80 we shall call IMS all those stars developing a degenerate C-O core and experiencing He shell flashes during their asymptotic giant branch evolution.

Stars with $m < 0.8\ m_\odot$ do not experience He shell flashes as they lose their H rich envelope before they have the chance to have flashes (Renzini 1977). Stars more massive than $\cong 8\ m_\odot$ ignite C in a non-degenerate core. This upper limit of the IMS range depends on the star's initial chemical composition as Becker and Iben (1979) have shown.

Contrary to what happens during main sequence mass loss in massive stars, in IMS there are changes in the surface chemical composition caused by convective dredge-up of processed material from the interior to the surface and by nuclear burning of the deepest layers of the convective envelope itself. These changes are reflected in the interstellar medium because stars lose mass through a stellar wind, and in some cases, through a rapid ejection of the stellar envelope when terminating their asymptotic giant branch phase (PN) and in other cases, by a possible carbon detonation event if C is ignited explosively under degenerate conditions (SN I ?).

Three dredge-up episodes can be distinguished (IT78): a first dredge-up when the convective envelope penetrates inward as stars reach the tip of the red giant branch for the first time, a second dredge-up which follows the ignition of the He burning shell as the convective envelope penetrates into the core, and a set of third dredge-up episodes when, following each He shell flash during the double shell evolution, the convective envelope penetrates the H-He discontinuity.

Peimbert and Serrano (1980*a*) have identified the stars giving rise to planetary nebulae of Types I and II (PNI and PNII) as defined by Peimbert (1978), as those having $m > 2.4\ m_\odot$ and $m < 2.4\ m_\odot$ respectively. This has been confirmed theoretically by RV80 by showing that the first dredge-up dominates in less massive stars giving rise to relatively small He and N surface enhancements while the third dredge-up dominates for stars in the range $2.5\ m_\odot < m < 5\ m_\odot$ resulting in much higher composition enhancements at the surface. Although IT78 evolution is in some respects similar to that by RV80, they did not get PNII because they adopted a relation between the mass of a PN and the mass of the core starting at $\cong 2\ m_\odot$. It was shown by RV80 that this in turn comes from Wood and Cahn's (1977) determination of the luminosity at which a star of a given mass gives rise to a PN. Wood and Cahn assumed a linear relationship between log L_{PN} and m and this resulted in stars with $m < 2\ m_\odot$, having cores less than 0.7 $m_\odot$, not producing PN. Modifying the $m_{PN}(m)$ relationship, RV80 were able to account for PNII.

In order to see the effects of the uncertainties in $m_{PN}(m)$ RV80 treated two cases: case A in which $m_{PN}(m_{core})$ comes essentially from Wood and Cahn (1977) plus the PN K648 in the globular cluster M15, and case B in which they arbitrarily halved m_{PN} corresponding to a given core mass. A discussion of the implications of this on the PN distance scale is given in § IV *a*.

In stars more massive than $\cong 5\ m_\odot$, C ignites under degenerate conditions and, although the subsequent evolution is uncertain, it probably results in a supernova event (see Renzini 1978; also the discussion of SN I by Wheeler 1978).

To calculate the chemical evolution coefficients for IMS we will adopt the results by RV80. Although Becker and Iben (1979) made similar calculations, all the data necessary to calculate the f_{ij} are not presented there; in any case, no major changes are expected with respect to the calculations by RV80.

We have approximated RV80 results by polynomials and adopted for the mass of the PN

$$
m_{PN} = -0.2382 + 0.2692\, m + 0.0516\, m^2, \quad \text{for } 0.8\, m_\odot < m < 2.5\, m_\odot;
$$

$$= -0.7847 + 0.8002\, m - 0.0706\, m^2, \quad \text{for } 2.5\, m_{\odot} < m < 4.8\, m_{\odot};$$

$$= -9.123 + 2.703\, m - 0.104\, m^2, \quad \text{for } 4.8\, m_{\odot} < m < 8\, m_{\odot}. \tag{21}$$

For the overabundance of He in PN we have adopted

$$\Delta Y(PN) = 0.04303 - 0.02648\, m + 0.007093\, m^2, \quad \text{for } 0.8\, m_{\odot} < m < 2.5\, m_{\odot};$$

$$= 0.04137 - 0.02618\, m + 0.00742\, m^2, \quad \text{for } 2.5\, m_{\odot} < m < 4.8 m_{\odot};$$

$$= 0.04947 + 0.00873\, m, \quad \text{for } 4.8\, m_{\odot} < m < 8\, m_{\odot}. \tag{22}$$

Finally, for the ratio of the mean overabundance of the wind with respect to that of the PN we have adopted

$$\frac{\Delta Y(\text{wind})}{\Delta Y(PN)} = 1.6765 - 0.8345\, m + 0.1237\, m^2, \quad \text{for } 0.8\, m_{\odot} < m < 4.8\, m_{\odot}$$

$$= -1.245 + 0.5127\, m - 0.0293\, m^2, \quad \text{for } 4.8 m_{\odot} < m < 8\, m_{\odot} \tag{23}$$

With overabundances (22) and (23) we calculated the equivalent mass fractions f^{ej}_{He} and used these for the model calculations.

e) *The equations for a normal galaxy under a constant SFR (or halfway to instantaneous recycling).*

The success of the instantaneous recycling approximation in describing chemical evolution (Talbot and Arnett 1971; Searle and Sargent 1972) led us to consider the case of a constant SFR.

From instantaneous recycling it is obtained that the chemical evolution is independent of the SFR. Thus, adopting a constant SFR simplifies the equations, is probably consistent with observations (see §III. *c*) and allows us to gain more insight of the relevant processes in the chemical evolution.

Using equations (6), (7) and (8), it can be shown that for any primary species (including He)

$$G\dot{X}_i = \dot{E}_i - X_i\dot{E} \tag{24}$$

That is, the evolution of the chemical composition is determined by the difference between the mass being ejected by stars in form of species i and the mass of species i that the ejecta would have with the present composition of the gas. In other words, $\dot{X}_i$ depends on the difference between the mean abundances of species i in the ejecta. $\langle X^{ej} \rangle = \dot{E}_i/\dot{E}$, and in the gas, X_i.

In the case of a constant SFR

$$\dot{E} = \dot{S} \int_{m(t)}^{m_U} m_{ej}(m)\, \phi\, (m)/\langle m \rangle\, dm,$$

i.e.

$$\dot{E} = \dot{S} \langle f_{ej}(t) \rangle, \tag{25}$$

so that

$$\dot{G} = -\dot{S} + \dot{E} = -\dot{S}\,(1 - \langle f_{ej}(t) \rangle) \tag{26}$$

Then equations (14) and (24) become

$$\frac{-dY}{d \ln G} = -Y\langle f_{ej}(t) \rangle + \int_{m(t)}^{m_U} [X\, m_{HeH} + Y\,(m_{ej} - m_{ZH})]\, \phi\, (m)/\langle m \rangle dm$$

and

$$\frac{-dZ}{d \ln G} = -Z\langle f_{ej}(t) \rangle + \int_{m(t)}^{m_U} [(X+Y)\, m_{ZH} + Z\, m_{ej}]\, \phi\, (m)/\langle m \rangle\, dm, \tag{27}$$

where X, Y, Z are evaluated at the time of birth, $t - \tau(m)$. The first term on the right hand sides of equations (27) gives the newly produced helium and heavy elements while the rest relates to the balance of the already present mass of the element under consideration.

These are the equations actually solved in the models, but to get an idea of the dominant processes, let us rewrite the mean ejected fraction

$$B(t) = \dot{E}/\dot{S} = \langle f_{ej}(t) \rangle,$$

and the locked up fraction

$$A(t) = 1 - B(t); \qquad (28)$$

in terms of which equations (27) become

$$-A(t)\frac{dY}{d\ln G} = <(1-Z) f^{ej}_{He}> - \Delta Y <f_{ej}> - Y_o (<f^{ej}_{He}> + <f^{ej}_{Z}>) + <\Delta Y (f_{ej} - f^{ej}_{He} - f^{ej}_{Z})>$$

and

$$-A(t)\frac{dZ}{d\ln G} = <(1-Z) f^{ej}_{Z}> + <Z f_{ej}> - Z <f_{ej}>. \qquad (29)$$

For any normal galaxy, $Z << 1$, and

$$-A(t)\frac{dY}{d\ln G} = X_o <f^{ej}_{He}> - Y_o <f^{ej}_{Z}> + \delta ,$$

and

$$-A(t)\frac{dZ}{d\ln G} = <f^{ej}_{Z}>, \qquad (30)$$

where

$$\delta = <\Delta Y f^{ej}_{H}> - \Delta Y <f_{ej}> \qquad (31)$$

is a correction due to the difference in ΔY between the mean hydrogen rich envelopes of stars, and the ISM. Since most of this mean hydrogen rich envelope comes from low mass stars, time lag effects will be important when these stars eject significantly, and then $\delta \neq 0$.

The second term, $- Y_o <f^{ej}_{z}>$, takes into account the He destroyed in forming Z. This term is of the same order but smaller than the first one since Y_o is relatively large.

We must stress that equations (30) are always correct to first order in Z. However, notice that A, $<f^{ej}_{He}>$, $<f^{ej}_{Z}>$, and $<f^{ej}_{H}>$ are time dependent functions through m(t) in equations (25) and (27). In this formulation true instantaneous recycling means taking $\delta = 0$ and all other quantities independent of time, so that

$$\frac{-dY}{d\ln G} = p_Y - Y_o p_z , \qquad (32)$$

$$\frac{-dZ}{d\ln G} = p_z .$$

Here the yields of heavy element and helium are

$$p_Z = <f^{ej}_{Z}>/A,$$

$$p_Y = X_o <f^{ej}_{He}>/A, \qquad (33)$$

and contain information both on stellar nucleosynthesis (through the f's) and on the mean fraction locked up in long lived objects (planets, very light stars, white dwarfs, neutron stars and black holes).

IV. DISCUSSION

Before discussing the results of the models (in §IV.*c* and §IV.*d*), and in view of the importance that having good models for IMS has on chemical evolution, we will draw some conclusions on the distance scale of PN and the relation $m_{PN}/(m)$ (§ IV.*a*), and on the importance of the stellar mass loss in the He enrichment of the interstellar medium (§ IV.*b*).

a) *The average mass of PN and their distance scale.*

Unfortunately, direct estimates of the mass of the emitting region in a PN are very difficult to obtain since we need to know the distance to the nebula.

Based on Shklovski's (1956) hypthesis that the mass of the nebula is the same for all objects, a statistical distance scale can be set up for optically thin PN. Basically, the Hβ flux and the angular radius (and to a lesser extent the chemical composition and the temperature of the nebula), determine the product m_{PN}(rms) $d^{-5/2}$, where m_{PN}(rms) is the root mean-square mass and d is the distance to the object.

Two distance scales commonly used are those due to Seaton (1968) and Webster (1969) (revised by Cahn and Kaler 1971), and to Cudworth (1974). Cudworth's scale (C) is 1.5 times larger than Cahn and Kaler's (CK), giving rise to a larger average mass: 0.5 $m_\odot$ and 0.18 $m_\odot$, respectively.

The electron densities derived from forbidden line ratios, N_e (FL), are in most objects greater than N_e (rms). This implies the presence of density fluctuations and indicates that m_{PN} (rms) is an upper limit to the true mass.

Another estimate of the mass can be obtained assuming that a fraction of the volume (the filling factor) has high density N_e(FL), and that the rest is empty, i.e.,

$$\epsilon = N_e^2 \text{ (rms)} / N_e^2 \text{ (FL)}, \qquad (34)$$

and

$$m_{PN}\,(FL) = \epsilon^{1/2}\, m_{PN}\,(rms). \qquad (35)$$

Torres-Peimbert and Peimbert (1977) found that for optically thin PN $\epsilon \cong 0.05$ with the C distance scale and $\epsilon \cong 0.075$ with the CK distance scale which imply mean masses of 0.11 $m_\odot$, and 0.05 $m_\odot$, respectively. It can be shown that m_{PN} (FL) is a lower limit to the true mass, m_{PN}, since the density fluctuations are not as extreme as assumed in equation (34), therefore $m_{PN}\,(FL) < m_{PN} < m_{PN}\,(rms)$.

The models of intermediate mass stars by RV80 discussed in §III*d* and the IMF given by equation (13) yield: $<m_{PN}> = 0.6 m_\odot$ (RV80, case A) and $<m_{PN}> = 0.33\, m_\odot$ (RV80, case B). These results, together with the observational $<m_{PN}\,(FL)>$ and $<m_{PN}(rms)>$ values, indicate that the C distance scale and case B of RV80 should be preferred over the CK distance scale and case A of RV80.

There are other arguments in favour of the C distance scale: a) the total number of PN in the galaxy and M31 (Alloin *et al.* 1976; Jacoby 1980), b) multiple-shelled PN (Trimble and Sackman 1978), and c) PN and white dwarf birth rates (Weidemann 1977; Peimbert 1981).

Let us finally stress that ultimately case A of RV80 (as well as the $m_{PN}(m)$ relationship by IT78) comes from the semiempirical determination by Wood and Cahn (1977) of the luminosity L_{PN} (m) at which a star of a given mass would give rise to a PN. This estimate is based in turn on the observed period distribution of Mira variables plus some assumptions about the SFR. RV80 adopted case B (arbitrarily halved m_{PN} for a given *core* mass) just to study the effects of the uncertainties in m_{PN}. From the discussion above, it appears that the determination of L_{PN} (m) needs to be revised.

b) *Effects of stellar mass loss on the fraction of newly synthesized He.*

Let us adopt as our "standard model" one with mass loss in heavy stars corresponding to $\alpha = 0.9$ of Chiosi *et al.* (1978) and with mass loss in IMS as in RV80.

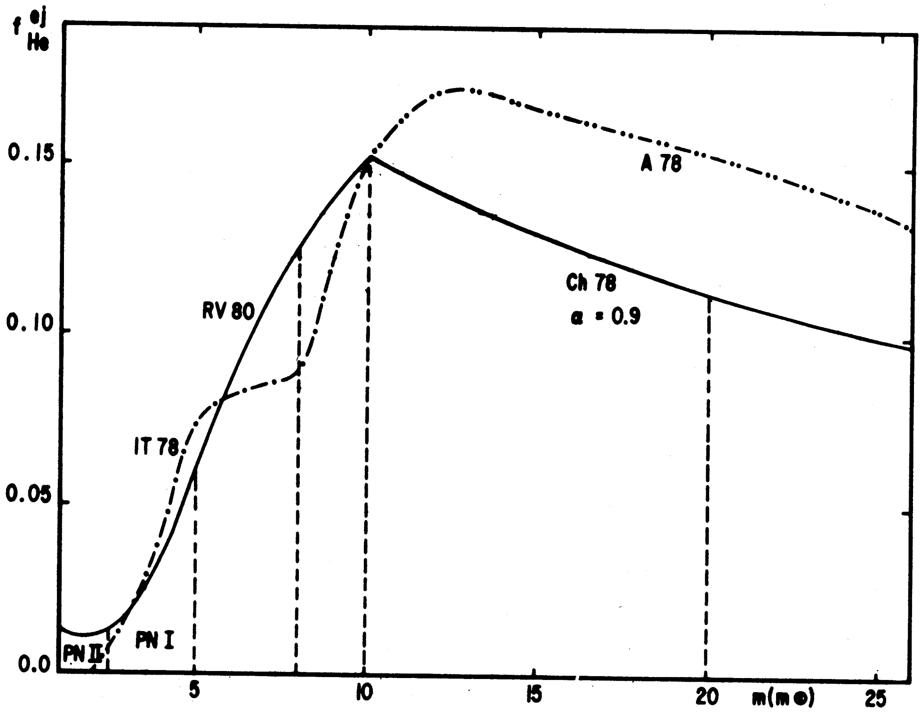

Fig. 1. Fraction of the mass of the star contributing to the He production, as a function of the initial stellar mass. Continuous curve is the standard model (RV80 for m $\lesssim$ 8 $m_\odot$ and $\alpha = 0.9$ of Chiosi *et al.* 1978 for m $\gtrsim$ 10 $m_\odot$). Regions where planetary nebulae of Types I and II contribute are indicated. IT78 = Iben and Truran 1978, A78 = Arnett 1978, Ch78 = Chiosi *et al.* 1978, RV80 = Renzini and Voli 1980.

In Figure 1 we present the fraction f^{ej}_{He} of the mass of a star where H has been burned but He has not and which is ejected, as a function of stellar mass. The most conspicuous overall feature of this diagram is that f^{ej}_{He} has a maximum of about 15% of the stellar mass in the region 10-15 $m_\odot$. There is a steep decline in f^{ej}_{He} towards lower masses as well as a more gentle decline for heavier stars.

As mentioned in § III.*d*, the right hand side of Figure 1 comes essentially from Arnett's classical calculations on the evolution of He stars. The implicit assumption is that the He core evolution is independent of the envelope. In fact, the coupling between the He core and the envelope is a decreasing function of mass. More massive stars have higher internal temperatures and so have higher fractions of the total flux carried by neutrinos, leaving less energy to be carried by radiation and convection. Lamb *et al.* (1977) found that for $m > 20\, m_\odot$ the "bare core" assumption was indeed valid, but that for a 15 $m_\odot$ star it was not, because the He core mass continued increasing significantly due to the action of the H burning shell. Thus it is possible that more detailed stellar models in the range 8 $m_\odot < m < 15$ $m_\odot$ will change this region of Figure 1.

Even if we assume a "bare core" evolution, we still need to assign a given He core mass to an initial stellar mass. In Figure 1 we used the relation given by Chiosi and Caimmi (1979) for $\alpha = 0.9$. For mass loss rates smaller than model $\alpha = 0.9$ of Chiosi, the fraction f^{ej}_{He} increases for a given initial stellar mass. This comes because a smaller mass loss rate increases the mass of the He core (and so of m_{HeH}) of the star. As can be seen in Figure 1, the extreme case of no mass loss at all increases f^{ej}_{He} by $\cong 4\%$ of the stellar mass, which is a significant fraction of f^{ej}_{He}.

On the left hand side of Figure 1 ($m < 8$ $m_\odot$), the effects of the different dredging-up mechanisms are present (IT78, RV80). In PNII and their associated previous winds, with 1 $m_\odot < m < 2.5$ $m_\odot$, the first

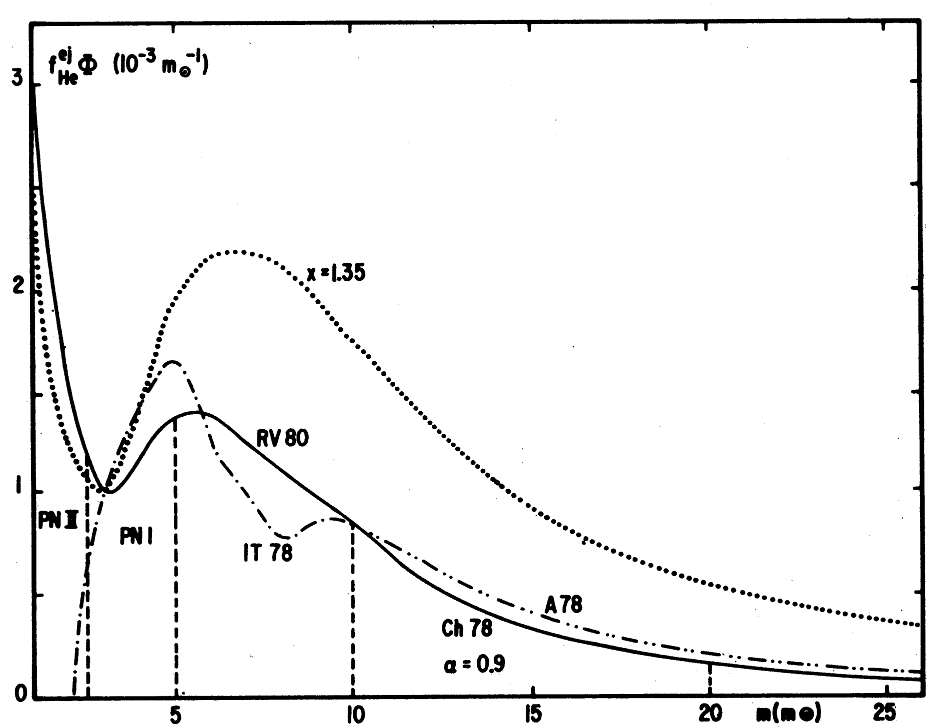

Fig. 2. Weighted fraction of the mass of the star contributing to the He production, per unit stellar mass. The area under the curve gives the fraction itself. Curves have the same meaning as in Figure1, but a curve for Salpeter's IMF has been added.

dredge-up dominates and f^{ej}_{He} is equivalent to about 1% of the stellar mass. In PNI (2.5 $m_\odot < m < 5\ m_\odot$) the third dredge-up contributes a major fraction of the increased He abundance. Finally, in the most massive IMS (5 $m_\odot < m < 8\ m_\odot$) we can see the effect of the second dredge-up episode and of hot-bottom burning as discussed by RV80 (as these authors point out, IT78 underestimate this effect). The composition assumed in this mass range is that of the core when it reaches 1.4 $m_\odot$ and any further processing has been neglected. As has been mentioned in §III.*d* these stars ignite C under degenerate conditions and their further evolution is very uncertain (Iben 1978*a, b*), but if they indeed end up as SN I (Renzini 1978) then it will mean that these objects can contribute significantly to the He enrichment. Notice finally that if we use IT78 results, PNII do not contribute to the He enrichment at all.

From the previous discussion it would appear unnecessary to discuss IMS because the larger f^{ej}_{He} are for stars with $m > 10\ m_\odot$. However, light stars are much more numerous than heavier stars so that the relevant quantity for the chemical evolution is not f^{ej}_{He} but rather this fraction weighted by the IMF by mass. In Figure 2 we show this weighted fraction. The area under the curves in this figure gives directly the fraction of the average stellar mass $<m>$ that burns H but does not burn He and is ejected. From this figure several important points can be made:

i) The maximum of the curve has moved down to $\cong 5\ m_\odot$ compared with the 10 $m_\odot$ maximum in Figure 1.

ii) On the contrary of what happens using IT78 models, PNII contribute significantly to the He enrichment. Although f^{ej}_{He} was fairly small and nearly constant with mass, stars in this mass range are the most numerous.

iii) There is also a significant contribution coming from PNI and from IMS in the range 5 $m_\odot < m < 8\ m_\odot$ (SNI?). IMS on the whole make up $\cong 2/3$ of the newly ejected He per generation of stars.

iv) As far as massive stars are concerned, the range 10 $m_\odot < m < 20\ m_\odot$, where stellar mass loss is less important for the structure, contributes most of the He given by massive stars. The assumptions made to assign stellar masses to He cores in this mass range can artificially increase f^{ej}_{He} (m). For example, Chiosi and Caimmi (1979) and Chiosi (1979) assume that all stars between 10 $m_\odot$ and 20 $m_\odot$ have a 3 $m_\odot$ He core, the most efficient in producing He. However, when one takes into account IMS, this effect loses importance.

Finally, notice that for a flatter IMF more weight is given to massive stars with respect to IMS as He producers.

c) *Effects of stellar mass loss and of the IMF on* $\Delta Y/\Delta Z$.

What interests us is not the total mass of newly produced He but rather its ratio to the total mass of newly produced heavy elements. From equations (30) we have that

$$\frac{dY}{dZ} = \frac{X_o < f^{ej}_{He} > - Y_o < f^{ej}_Z > + \delta}{< f^{ej}_Z >} . \quad (36)$$

To a first approximation in intermediately evolved systems ($\delta \cong 0$), $\Delta Y/\Delta Z$ depends only on the ratio $< f^{ej}_{He} > / < f^{ej}_Z >$. This is a fortunate circumstance because different mass ranges predominate in each mass fraction and because stars with $m < 1\ m_\odot$ do not affect $\Delta Y/\Delta Z$. Observational constraints then can be set on the models. If the mass range for both averages were the same, no such constraints would arise. More precisely, the heavy element production is significant only for high mass stars with $m > 10\ m_\odot$ (although in IMS there is some production of C that will be ignored here but treated in Paper III of this series).

As in the case of He discussed in § IV.*b*, the change in stellar structure brought about by mass loss in these heavy stars tends to decrease the production of heavy

TABLE 2

$\Delta Y/\Delta Z$ FOR DIFFERENT ASSUMPTIONS ABOUT MASS LOSS

Models adopted for mass loss		Properties at G = 0.1		Maximum $\Delta Y/\Delta Z$
Intermediate mass stars	Massive stars (α)	ΔY	t(10^9 y)	
RV80 (Std.)	0.9	0.032	12.19	3.06
"	0.8	0.032	"	2.28
"	0.0	0.031	"	1.01
IT78	0.9	0.026	12.22	2.47
no	0.9	0.013	"	1.23

RV80 = Renzini and Voli 1980; IT78 = Iben and Truran 1978; α of Chiosi *et al.* 1978.

elements for a given initial stellar mass. Thus, $<f_Z^{ej}>$ is a decreasing function of stellar mass loss.

We can see then how processes of mass loss, both in the IMS and in the heavy stars, cooperate to increase ΔY/ΔZ over the value of 0.4 obtained in the first attempts to explain the observed ΔY/ΔZ = 3. Mass loss and dredge-up in the wind and PN tend to increase $<f_{He}^{ej}>$ while stellar mass loss in heavy stars decreases $<f_Z^{ej}>$. Consequently, ΔY/ΔZ increases.

In Table 2 we present the values of ΔY/ΔZ calculated from the models (no instantaneous recycling) under a SFR = 0.1 $M_{gal}/10^9$ y and with the standard IMF given in equation (17). Decreasing α from 0.9 (extreme mass loss) to 0.0 (constant mass evolution) causes ΔY/ΔZ to decrease from 3 to 1, emphasizing the role of mass loss in massive stars, essentially by decreasing the heavy element production.

On the other hand, we can see from Table 2 that the model with no helium production in IMS has ΔY/ΔZ = 1.23, an indication of the importance of the dredging-up episodes and of mass loss in IMS for the He production. Notice also that the IT78 models indicate that PNII contribute about about 1/6 of the new He and that the effective SFR is slightly decreased.

In Figure 3 we show the effect of mass loss in massive stars on ΔY/ΔZ. An interesting feature of this figure is that Y depends only on time and not on α. This is a consequence of the predominance of IMS in the He production which implies that the effective SFR does not vary with α.

Adopting a flatter IMF (smaller x) tends to increase the relative importance of massive stars as He producers, and to decrease ΔY/ΔZ. In Table 3 it can be noted that Salpeter's (1955) IMF with x = 1.35 is not compatible with observations and that x cannot be larger than ≅ 2.0 since ΔY/ΔZ starts increasing very rapidly with x.

In Tables 2 and 3 we also present the helium enrichment, Δ Y, and the time of evolution when the galaxy has 10% of its mass as gas.

Notice that, due to astration, the time required to reach a given fraction of gas decreases as x increases. For example, in the "standard model" at t = 12.56 × 10^9y, the total mass that has gone into stars is $\Delta S = \dot{S}t = 1.26\ M_{gal}$. But there is still 0.07 of M_{gal} as gas, meaning that the total mass that has been ejected is ΔE = 0.33 M_{gal}; which in turn implies an average ejected fraction of 26.3% of the matter

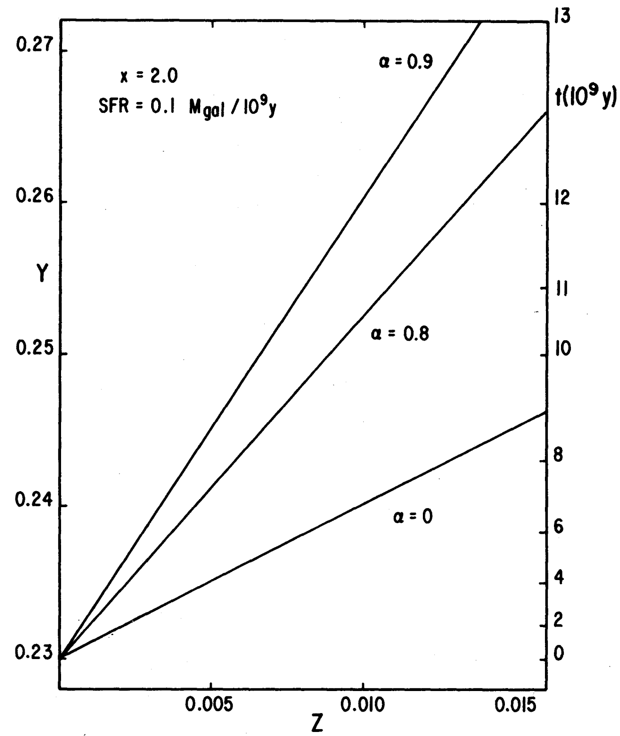

Fig. 3. Helium versus heavy element abundance for models of galactic evolution having different mass loss rates in massive stars, as parametrized by α of Chiosi *et al.* 1978. In extreme mass loss, α = 0.9, while in models with constant mass evolution α = 0.0. On the right hand side the ordinate is the time elapsed (since Y(t) is independent of α).

going into stars. If we look now at a model with Salpeter's law, 42% of the mass of gas going into stars is returned back on this occasion. Thus, the total mass ejected is now 0.27 M_{gal}. We see then that a change by a factor of less than 1.6 in $\dot{E}/\dot{S}$ causes the fraction of gas at t = 12.56 × 10^9y to change by almost a factor of 4.

It must be said that although in Tables 2 and 3 we have presented the maximum value of ΔY/ΔZ, once the model has evolved for ≅ 10^9 y, ΔY/ΔZ does not change by more than ≅ 10%. This can also be seen in Table 4 and Figures 3 and 4.

d) *The effect of the SFR.*

Even though the SFR is constant, the model does not behave exactly as in instantaneous recycling. In the "standard model", for example, the "pseudo-yields"

$$p_z' = -Z/\ln G$$

and (37)

$$p'_y = -Y/\ln G + Y_o\, p'_z$$

are not constant as in instantaneous recycling (see Table 4). Up to about 10^9y the change in time is mainly due to less massive stars contributing more and more to the ejection, hence decreasing the locked-up fraction from 1 to ≅ 0.8. From then on, this effect becomes less important and the effect of dilution ($\delta \neq 0$ in equation 36) grows in importance, being dominant at 13×10^9y. This does not mean that Y starts decreasing with Z but only that there is a slight deviation from a straight line.

Models with higher SFR will have a similar development. The main change is in the gas fraction remaining at a given time, which decreases strongly as the SFR increases.

In Figure 4 we show the behaviour of models with varying SFR (0.1,1,10) $M_{gal}/10^9$ y. The most striking feature of this diagram is that as the SFR increases, the $\Delta Y/\Delta Z$ ratio decreases, particularly at low Z. This is so because IMS, which are responsible for most of the He, have not had time to eject most of their material back to the interstellar medium.

In a certain sense all the models are homologous: at a given time the pseudo-yields (equations 37), $\Delta Y/\Delta Z$ and the ejected fraction ($B = \dot{E}/\dot{S}$) are very nearly the same, independently of the SFR. In other words, if $p'_z = -Z/\ln G$ is the same at a given time in all models, then the model with the higher SFR has a lower G, and so a larger Z.

Foı a high SFR the Y versus Z curve deviates from a straight line. This effect could explain the very low Y_p value (≅ 0.216) found by Talent (1980) and French (1980). This value, if true, is incompatible with the standard big-bang theory.

A SFR = 10 $M_{gal}/10^9$y corresponds to a timescale of gas consumption of ≅ 10^8y. Colors in blue compact galaxies indicate bursts with timescales of 10^7y to a few 10^8y (Lequeux *et al.* 1979), and independent evidence for such a burst in II Zw70 has been presented by Bergeron (1976) and O'Connell *et al.* (1978). As pointed out by Searle *et al.* (1973) this burst may hide an older stellar component formed at a slow rate.

Lequeux *et al.* (1979) favoured an old origin of blue compact galaxies since they shared the same region as normal irregulars on the Y versus Z and N/O versus O diagrams.

Let us assume however, that we have a mixture of old and young galaxies in a sample. On the average old galaxies have large values of Z and fall on the upper half of the upper curve in Figure 4. Young galaxies, which are the least numerous, are located on the lower curve of Figure 4. Clearly, their effect on the sample is more important in the low Z region. If we were to measure the abundances in these galaxies and tried to fit a linear regression to Y versus Z, then we would underestimate Y_p (Y at Z = 0). Taking the most extreme possible case, this underestimate of Y_p can amount to ≅ 0.01.

TABLE 3

$\Delta Y/\Delta Z$ FOR DIFFERENT SLOPES OF THE INITIAL MASS FUNCTION

Slope x	Properties at G = 0.1: ΔY	Properties at G = 0.1: t(10^9y)	Maximum ΔY/ΔZ
1.3	⩾ 0.057	> 15.	0.89
1.4	⩾ 0.057	> 15.	1.06
1.5	⩾ 0.057	> 15.	1.26
1.6	0.055	13.83	1.49
1.7	0.048	13.29	1.78
1.8	0.042	12.85	2.13
1.9	0.037	12.48	2.55
2.0	0.032	12.19	3.06
2.1	0.028	11.93	3.67
2.2	0.025	11.72	4.42

TABLE 4

TIME EVOLUTION OF THE STANDARD MODEL

t (10^9y)	G	p'_Z	p'_Y	ΔY/ΔZ	$B = \dot{E}/\dot{S}$
0.1	0.99	0.0039	0.007	1.74	.07
0.5	0.95	0.0041	0.011	2.43	.12
1.0	0.91	0.0043	0.012	2.66	.14
5.0	0.58	0.0047	0.015	2.99	.22
10.0	0.21	0.0047	0.016	3.05	.25
13.0	0.04	0.0043	0.014	3.03	.26

e) *Cosmological implications.*

It has been shown that stellar nucleosyntesis and mass loss contribute to the He enrichment of the gas in a stellar system at the rate $\Delta Y = 3\ \Delta Z$ as required by observations. Most systems have $Z \ll 1$ and the He produced is small compared with the values of ≅ 0.2 – 0.3 found in astronomical objects. Hence the proposition that before stellar formation took place, already existed a Y_p of ≅ 0.23. Notice, that in very metal rich galaxies, like M83 studied by Dufour *et al.* (1980), ΔY could be comparable to Y_p (see Paper II for further discussion).

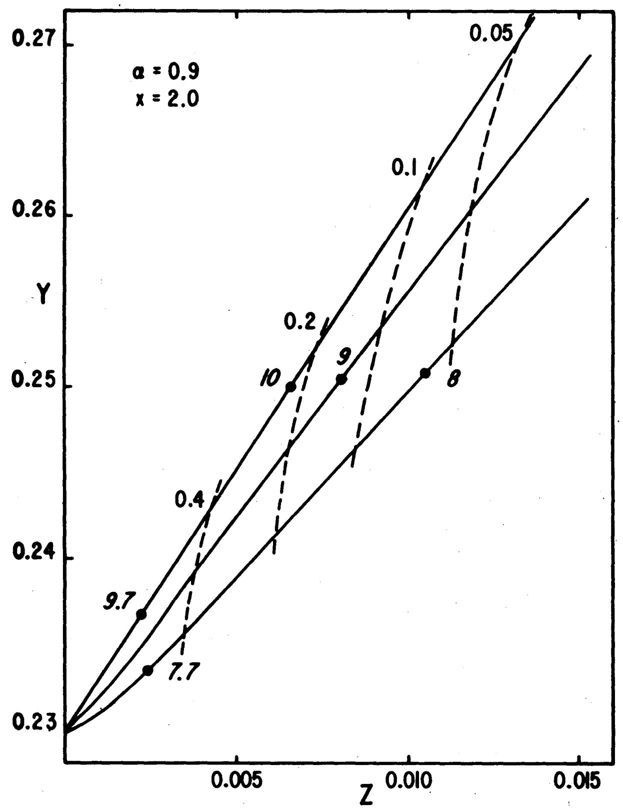

Fig. 4. Helium versus heavy element abundance for models having different values of SFR. Upper curve corresponds to SFR = 0.1 ("standard model") middle curve to SRF = 1, and lower curve to SRF = 10 (in units of $M_{gal}/10^9$ y). Dashed lines are labeled by constant fractions of the gas remaining in the system and filled circles indicate logarithms of elapsed times.

Standard big-bang models (Dicke *et al.* 1965), take into account e^+, e^-, ν_e, $\bar{\nu}_e$, ν_μ, $\bar{\nu}_\mu$ and γ, and predict that if $Y_p = 0.23$ then the nucleon density in the Universe must be $\rho_N \cong 1.5 \times 10^{-31}$ g cm^{-3} (Wagoner *et al.* 1967; Wagoner 1973; Yang *et al.* 1979).

Measuring this density in terms of $\Omega = \rho/\rho_{crit} = 8\pi G\rho/3 H^2$, this means that $\Omega_N h_o = 0.08$ (where h_o is the Hubble constant in units of 100 km s^{-1} Mpc^{-1}) i.e. an open Universe. Some estimates of the nucleon mass in the Universe gave $\Omega_N h_0^2 \approx 0.01$ (e.g. Oort 1958; Peebles 1971; Gott *et al.* 1974) which seemed to end the case in a consistent manner.

The situation for the standard big-bang has become uneasy since. Perl *et al.* (1975, 1976) discovered another lepton more massive than the electron and the muon; the τ lepton (see review by Tayler 1980 for a general discussion of this point). If the neutrino associated with the τ lepton were "light" ($m_\nu < 1$ Mev) the expansion rate of the Universe would be longer, and more He would be predicted than in the standard big-bang model (Steigman *et al.* 1977; Yang *et al.* 1979).

Yang *et al.* (1979) found that if $Y_p \cong 0.25$ then only three lepton pairs could exist. However, we have argued in § II that the stellar data that Yang *et al.* use are of lower accuracy than the He abundances derived from H II regions. Under this assumption and adopting $Y_p = 0.23$ and Yang *et al.* models, it follows that the only possibility for the ν_τ to be "light" is for $\Omega_N h_0^2 = 0.0035$, i.e., $\rho_N \cong 7 \times 10^{-32}$ g cm^{-3} which seems rather low compared with the value of $\Omega_N h_0^2 \gtrsim 0.01$ obtained by Gott *et al.* and certainly with Peeble's 1979 value of 0.4 ± 0.2. Thus, ν_τ is probably heavy.

The very same constraints $Y_P = 0.23$ and $\Omega_N h_0^2 \gtrsim 0.01$ imply that the Universe is not nucleon dominated (Yang *et al.* 1979) and suggest that heavy stable leptons ($m > 2$ GeV) might dominate the mass. The case for an open Universe is weakened, but it is still preferred if the heavy leptons cluster with the nucleons (Gunn *et al.* 1978).

Furthermore, recent determinations of Y_p from extragalactic H II regions (Talent 1980; French 1980) shown in Table 1 give $Y_p = 0.216$. If these observations are correct, standard big-bang theory would need $\Omega_N h_o^2 = 0.003$ which seems extremely low. Other non conventional processes must be invoked to reduce Y_p to these values: a Universe dominated by stable heavy leptons (Lee and Weinberg 1977; Gunn *et al.* 1978), or having ν_e degenerate (e.g., Beaudet and Goret 1976; Yahil and Beaudet 1976; Beaudet and Yahil 1977), or an anisotropic Universe (Gisler *et al.* 1974; Barrow 1976; Barrow and Matzner 1977).

We have seen however in § IV.d that in a mixed sample of old galaxies with moderately high Z and young galaxies with low Z (see Figure 4), we tend to underestimate Y_p by as much as ≈ 0.01. Thus, assuming no errors in the observations, the standard big-bang can be saved if there are very young galaxies in the samples of Talent (1980) and French (1980).

V. CONCLUSIONS

a) We have compiled recent determinations of the pregalactic helium abundance, Y_p, and of the abundance enrichment ratio $\Delta Y/\Delta Z$. They indicate that $Y_p \approx 0.23$ (although some determinations give $Y_p = 0.22$) and $\Delta Y/\Delta Z = 3$.

b) We have been able to produce a "standard model" of chemical evolution which gives $\Delta Y/\Delta Z = 3$. In this model we have adopted Serrano's (1978) IMF with a slope $x = 2$ in the high mass range. Also, we have used a constant SRF≅ 0.1 $M_{gal}/10^9$ yrs similar to the value in the solar neighborhood. In massive stars, we used stellar evolution models by Chiosi *et al.* (1978). In intermediate mass stars, on the other hand, we have adopted models of asymptotic giant branch evolution by Renzini and Voli (1980); these models are consistent with helium abundance determinations in planetary nebulae. The "standard model" has the following implications:

i) Stars with $m < 0.8$ $m_\odot$ do not affect $\Delta Y/\Delta Z$.

ii) Intermediate mass stars contribute with 2/3 of the newly formed He ejected per generation of stars, $<f^{ej}_{He}>$. (1) Stars with 0.8 $m_\odot < m < 2.5$ $m_\odot$ in which the first dredge-up dominates contribute with 1/6 to $<f^{ej}_{He}>$. (2) Another 1/6 is contributed by stars with 2.5 $m_\odot < m < 5$ $m_\odot$. Here the third dredge-up dominates and PNI are produced. (3) The first dredge-up and hot bottom burning in stars with 5 $m_\odot < m < 8$ $m_\odot$ help to produce 1/3 of $<f^{ej}_{He}>$.

iii) Massive stars contribute with only 1/3 of $<f^{ej}_{He}>$ mostly from the 10-20 $m_\odot$ range where mass loss is still not dominating the structure. Although stars with $m > 20$ $m_\odot$ eject small amounts of He, a large mass loss is needed to reduce the Z production and give rise to a high $\Delta Y/\Delta Z$ value.

c) Once it is accepted that stars with $m < 0.8$ $m_\odot$ produce neither Y nor Z, any change in the stellar evolution input (for any given IMF and SFR) tends to decrease $\Delta Y/\Delta Z$. If, for example, He shell flashes and mixing in stars with $m < 2$ $m_\odot$ is not taken into account (as in Iben and Truran 1978) then $\Delta Y/\Delta Z$ is reduced. A similar reduction is obtained if we ignore hot-bottom burning or neglect mass loss in massive stars.

d) If one accepts as reasonable the stellar evolution with mass loss, both in IMS and in massive stars, then a slope of the IMF flatter than x ≅ 2 (as e.g., Salpeter's) gives rise to smaller values of $\Delta Y/\Delta Z$, while a steeper slope gives higher values of $\Delta Y/\Delta Z$. The point is that $\Delta Y/\Delta Z$ is sensitive to the slope because Y and Z are produced by stars in different mass ranges.

e) With the adopted stellar evolution and IMF, a comparison with observed PN masses tends to favor Cudworth's distance scale and case B of Renzini and Voli (1980).

f) Models calculated with low SFR (0.1 $M_{gal}/10^9$ y) produce, within 10%, a linear relation between Y and Z after 10^9 y. It must be stressed that the instantaneous recycling approximation has not been used here.

g) The Y versus Z curve deviates from a straight line only for high SFR, similar to those found in blue compact galaxies. Thus, galaxies with low Y and relatively high Z can be understood as very young galaxies. An empirical determination of Y_p from a sample of old metal rich and very young metal poor galaxies could yield a spurious Y_p value lower (by as much as 0.01) than the true value (see Figure 4). In this way an empirical $Y_p = 0.216$ could be consistent with a 'true' $Y_p = 0.23$ if such very young galaxies exist. Notice however, that the value $Y_p = 0.23$ obtained from the Magellanic Cluds is not affected by the previous discussion since the SFR in these objects are relatively low.

REFERENCES

Alcock, B. and Paczynski, B. 1978, *Ap. J.*, **223**, 244.
Alloin, D., Collin-Souffrin, S., Joly, M., and Vigroux, L. 1979, *Astr. and Ap.*, **78**, 200.
Alloin, D., Cruz-González, C., and Peimbert, M. 1976, *Ap. J.*, **205**, 74.
Arnett, W.D., 1978, *Ap. J.*, **219**, 1008.
Audouze, J., and Tinsley, B.M. 1974, *Ap. J.*, **192**, 487.
Barrow, J.D. 1976, *M.N.R.A.S.*, **175**, 359.
Barrow, J.D. and Matzner, R.A. 1977, *M.N.R.A.S.*, **181**, 719.
Beaudet, G. and Goret, P. 1976, *Astr. and Ap.*, **49**, 415.
Beaudet, G. and Yahil, A. 1977, *Ap. J.*, **218**, 253.
Becker, S.A. and Iben, I. Jr. 1979, *Ap. J.*, **232**, 831.
Becker, S.A. and Iben, I. Jr. 1980, *Ap. J.*, **237**, 111.
Bergeron, J. 1976, *Ap. J*, **210**, 287.
Cahn, J.H. and Kaler, J.B. 1971, *Ap. J. Suppl.*, **22**, 319.
Caputo, F., Castellani, V., and Wood, P.R. 1978, *M.N.R.A.S.*, **184**, 377.
Chiosi, C. 1979, *Astr. and Ap.*, **80**, 252.
Chiosi, C. and Caimmi, R. 1979, *Astr. and Ap.*, **80**, 234.
Chiosi, C. and Matteucci, F. 1980, preprint.
Chiosi, C. and Nasi, E. 1974, *Astr. and Ap.*, **35**, 81.
Chiosi, C., Nasi, E., and Sreenivasan, S.R. 1978, *Astr. and Ap.*, **63**, 103.
Chiosi, C., Nasi, E., and Bertelli, G. 1979, *Astr. and Ap.*, **74**, 62.
Cruz-González, C., Recillas-Cruz, E., Costero, R., Peimbert, M., and Torres-Peimbert, S. 1974, *Rev. Mexicana Astron. Astrof.*, **1**, 211.
Cudworth, K.M. 1974, *A. J.*, 79, 1384.
Dearborn, D.S.P. and Eggleton, P.P. 1977, *Ap. J.*, **213**, 448.
de Loore, C., de Greve, J.P., Lamers, H.J.G.L.M. 1977, *Astr. and Ap.*, **61**, 251.
de Loore, C., de Greve, J.P., and Vanbeveren, D. 1978, *Astr. and Ap.*, **67**, 373.

Dicke, R.H., Peebles, P.J.E., Roll, P.G., and Wilkinson, D.T. 1965, *Ap. J.*, **142**, 414.

Dufour, R.J., Talbot, R.J., Jensen, E.B., and Shields, G.A. 1980, *Ap. J.*, **236**, 119.

Faulkner, J. 1967, *Ap. J.*, **147**, 617.

French, H.B. 1980, *Ap. J.*, **240**, 41.

Gingold, R.A. 1977, *M.N.R.A.S.*, **178**, 569.

Gisler, G.R., Harrison, E.R., and Rees, M.J. 1974, *M.N.R.A.S.*, **166**, 663.

Gott, J.R., III, Gunn, J.E., Schramm, D.N., and Tinsley, B.M. 1974, *Ap. J.*, **194**, 543.

Gunn, J.E., Lee, B.W., Lerche, I., Schramm, D.N., and Steigman, G., 1978, *Ap. J.*, **223**, 1015.

Hacyan, S., Dultzin-Hacyan, D., Torres-Peimbert, S., and Peimbert, M. 1976, *Rev. Mexicana Astron. Astrof.*, **1**, 355.

Hirshfeld, A. 1978, Ph. D. Thesis, Yale University.

Iben, I., Jr. 1977, in *Advanced Stages in Stellar Evolution*, ed. P. Bouvier and A. Maeder (Geneve).

Iben, I., Jr. 1978*a*, *Ap. J.*, **219**, 213.

Iben, I., Jr. 1978*b*, *Ap. J.*, **226**, 996.

Iben, I., Jr. and Truran, J.W. 1978, *Ap. J.*, **220**, 980.

Jacoby, G.H. 1980, *Ap. J. Suppl.*, **42**, 1.

Kunth, D. and Sargent, W.L.W. 1980, private communication.

Lamb, S.A., Iben, I., Jr., and Howard, W.M. 1976, *Ap. J.*, **207**, 209.

Larson, R.B. 1976, *M.N.R.A.S.*, **176**, 31.

Lee, B.W. and Weinberg, S. 1977, *Phys. Rev. Lett.*, **38**, 1237.

Lequeux, J. 1979, *Astr. and Ap.*, **80**, 35.

Lequeux, J., Martin, N., Prevot, L., Prevot-Burnichon, M.L., Rebeirot, E., and Rousseau, J. 1980, *Astr. and Ap.*, **85**, 305.

Lequeux, J., Peimbert, M., Rayo, J.F., Serrano, A, and Torres-Peimbert, S. 1979, *Astr. and Ap.*, **80**, 155.

Lynden-Bell, D. 1975, *Vistas in Astronomy*, **19**, 299.

Mc Cuskey, S.W. 1966, *Vistas in Astronomy*, **7**, 141.

Miller, G.E. and Scalo, J.M. 1979, *Ap. J. Suppl.*, **41**, 513.

O'Connell, R.W., Thuan, T.X., and Goldstein, S.J. 1978, *Ap. J.*, **226**, L11.

Oort, J.H. 1958, in Eleventh Solvay Conf. on *The Structure and Evolution of the Universe*, p. 163.

Pagel, B.E.J. and Patchett, B.E. 1975, *M.N.R.A.S.*, **172**, 13.

Peebles, P.J.E. 1971, *Physical Cosmology*, (Princeton U.P.: Princeton, N.J.).

Peebles, P.J.E. 1979, *A. J.*, **84**, 730.

Peimbert, M. 1978, in *Planetary Nebulae: Observations and Theory*, IAU *Symp.* 76, ed. Y. Terzian (Dordrecht: D. Reidel), p. 215.

Peimbert, M. 1981, in *Physical Processes in Red Giants*, ed. I. Iben Jr. and A. Renzini (Dordrecht: D. Reidel), in press.

Peimbert, M. and Serrano, A. 1980 , *Rev. Mexicana Astron. Astrof.*, **5**, 9.

Peimbert, M. and Serrano, A. 1980*b*, *M.N.R.A.S.*, submitted, (Paper II).

Peimbert, M. and Torres-Peimbert, S. 1974, *Ap. J.*, **193**, 327.

Peimbert, M. and Torres-Peimbert, S. 1976, *Ap. J.*, **203**, 581.

Peimbert, M. and Torres-Peimbert, S. 1977, *M.N.R.A.S.*, **179**, 217.

Peimbert, M. and Torres-Peimbert S. 1981, in preparation.

Perl, M.L., *et al.* 1975, *Phys. Rev. Letters*, **35**, 1489.

Perl, M.L., *et al.* 1976, *Phys. Lett.*, **63B**, 466.

Perrin, M.-N., Hejlsen, P.M., Cayrel de Strobel, G., and Cayrel, R. 1977, *Astr. and Ap.*, **54**, 779.

Reeves, H., Audouze, J., Fowler, W.A., and Schramm, D.N. 1973, *Ap. J.*, **179**, 909.

Renzini, A. 1977, in *Advanced Stages in Stellar Evolution*, ed. P. Bouvier and A. Maeder, (Geneve), p. 149.

Renzini, A. 1978, Proceedings of the Workshop *Supernovae and Supernovae Remnants, Mem. Soc. Astron. Italiana*, **49**, 389.

Renzini, A. and Voli, M. 1980, preprint (RV80).

Saio, H. 1978, *Publ. Astron. Soc. Japan*, **30**, 139.

Salpeter, E.E. 1955, *Ap. J.*, **121**, 161.

Sandage, A.R. and Katem, V. 1976 *A.J.*, **81**, 743.

Schmidt, M. 1959, *Ap. J.*, **129**, 243.

Searle, L. and Sargent, W.L.W. 1972, *Ap. J.*, **173**, 25.

Searle, L., Sargent, W.L.W., and Bagnuolo, W.G. 1973, *Ap. J.*, **179**, 427.

Seaton, M.J. 1968, *Ap. (Letters)*, **2**, 55.

Serrano, A. 1978. Ph. D. Thesis, University of Sussex.

Shaver, P. 1980, private comunication.

Shklovski, I.S. 1956, *Soviet Astr.-AJ*, **33**, 315.

Steigman, G., Schramm, D.N., and Gunn, J.E. 1977, *Phys. Lett.*, **66B**, 202.

Taff, L.G. 1974, *A. J.*, **79**, 1280.

Talbot, R.J. and Arnett, W.D. 1971, *Ap. J.*, **170**, 409.

Talbot, R.J. and Arnett, W D. 1973, *Ap. J.*, **186**, 51.

Talent, D.L. 1980, preprint.

Tayler, R.J. 1980, *Rep. on Prog. in Phys.*, **43**, 253.

Thuan, T.X., Hart, M.H., and Ostriker, J.P. 1975, *Ap. J.*, **201**, 756.

Torres-Peimbert, S. and Peimbert, M. 1977, *Rev. Mexicana Astron. Astrof.*, **2**, 181.

Torres-Peimbert, S., Peimbert, M., and Daltabuit, E. 1980, *Ap. J.*, **238**, 133.

Torres-Peimbert, S., Peña, M., and Daltabuit, E. 1980, in *The Universe in Ultraviolet Wavelenghts: The First Two Years of IUE*, NASA *Symp.*, ed. R.D. Chapman, in press.

Trimble, V. and Sackman, I.-J. 1978, *M.N.R.A.S.*, **182**, 97.

Twarog, B.A. 1980, preprint.

Vanbeveren, D. and Olson, G.L. 1980, *Astr. and Ap.*, **81**, 228.

Wagoner, R.V. 1973, *Ap. J.*, **179**, 343.

Wagoner, R.V., Fowler, W.A., and Hoyle, F. 1967, *Ap. J.*, **148**, 3.

Weidemann, V. 1977, *Astr. and Ap.*, **61**, L27.

Webster, B.L. 1969, *M.N.R.A.S.*, **143**, 79.

Wheeler, J.C. 1978, *Mem. Soc. Astron. Italiana*, **49**, 349.

Wood, P.R. and Cahn, J.H. 1977 *Ap. J*, **211**, 499.

Yahil, A. and Beaudet, G. 1976, *Ap. J.*, **206**, 26.

Yang, J., Schramm, D.N., Steigman, G., and Rood, R.T. 1979, *Ap. J.*, **227**, 697.

Manuel Peimbert and Alfonso Serrano: Instituto de Astronomía, UNAM, Apdo. Postal 70-264, México 20, D.F., México,