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GALACTIC EVOLUTION OF D AND ^{3}He

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ABSTRACT

The evolution of D and ^{3}He is considered in the framework of a consistent model for the chemical evolution of the Galaxy. The destruction of D in the course of galactic evolution is found to be modest, of the order of a factor ~ 2; the good agreement found with the observed abundance of D in the presolar material and in the interstellar medium makes possible to derive narrow constraints on the primordial D, or, equivalently, on the allowed range of the baryon-to-photon ratio η according to standard big-bang nucleosynthesis. Conversely, the evolution of ^{3}He remains one of the major problems in the field of chemical galactic evolution: results based on the predictions of updated stellar models are shown to lead to an overproduction of ^{3}He by a factor $\sim$5–7 at the time of formation of the Sun, and by a factor $\sim$5–20 with respect to measured abundances in galactic H II regions. Several possibilities to reduce this discrepancy are presented and discussed quantitatively; none of them is found to produce the desired result, with the exception of a hypothetical low-energy resonance in the cross section of the ^{3}He(^{3}He, $2p$)^{4}He reaction, advocated also as a possible solution to the solar neutrino puzzle. Otherwise, the observed underabundance of ^{3}He in the presolar material and in H II regions must be considered as evidencing local inhomogeneities in the interstellar medium, in which ^{3}He is strongly depleted by some unknown process.

Subject headings: elementary particles — galaxies: evolution — nuclear reactions, nucleosynthesis, abundances

1. INTRODUCTION

The standard hot big bang model of nucleosynthesis (SBBN) is the one characterized by the lowest number of adjustable parameters. Given the number of relativistic neutrino species N_ν and the neutron lifetime τ_n, the abundances of D, ^{3,4}He, ^{6,7}Li, ^{9}Be, and ^{10,11}B in the primordial gas depend on one cosmological parameter only, the baryon-to-photon ratio $\eta \equiv n_b/n_\gamma$. Taking the cosmic background temperature $T_0 = 2.726$ K (Mather et al. 1993), the fraction of the critical density contributed by baryons, Ω_b, is related to η through

$$\Omega_b = 1.46 \times 10^{-2} \eta_{10} h_{50}^{-2} , \qquad (1)$$

where $\eta_{10} = 10^{10}\eta$ and h_{50} is the Hubble constant in units of 50 km s^{-1} Mpc^{-1}. Recent experimental data have provided stringent bounds on N_ν and τ_n: $N_\nu = 3.10 \pm 0.09$ (see e.g., Walker et al. 1991) and $\tau_n = 889.1 \pm 2.1$ (see, e.g., Smith, Kawano, & Malaney 1993). Thus, only the value of η_{10} has to be determined experimentally. In practice, this is done by comparing the predicted abundances of D, ^{3,4}He, and ^{7}Li from a grid of SBBN models characterized by different η_{10} with the corresponding values inferred from their abundances in the present-day interstellar medium (ISM) and in the Sun (Yang et al. 1984; Krauss & Romanelli 1990; Walker et al. 1991; Smith et al. 1993). SBBN has been very successful in accounting for the observed abundances of D, ^{3}He, ^{4}He and ^{7}Li: the window of consistency for η_{10} is bounded from below by the upper limit on [D + ^{3}He]$_P$ (hereafter, primordial abundances will be indicated with a subscript P) and from above by the upper limit to ^{4}He$_P$ (Smith et al. 1993). The value of η_{10} may also be bounded from above by the constraint set by the Li abundance in metal-poor stars (Steigman 1991; Walker et al. 1991).

The value of ^{4}He$_P$ is usually obtained by extrapolating to zero metallicity the observed correlation of the abundance of ^{4}He with the abundance of O or N in H II regions. The importance of an accurate determination of ^{4}He$_P$ cannot be overemphasized. A large amount of papers has been devoted to this problem and the results have been summarized many times (see, e.g., Yang et al. 1984; Pagel et al. 1992). On the other hand, the abundance of D and ^{3}He is known only at the time of formation of the Sun ($t = t_\odot$) and at the present epoch ($t = t_0$); therefore the determination of D$_P$ and ^{3}He$_P$ requires some degree of theoretical modeling to evaluate the amount of D destruction during galactic evolution ("astration") and the amount of ^{3}He destruction/production by stellar processes. Yang et al. (1984) derived an upper limit to (D + ^{3}He)$_P$ using a simple "one-cycle approximation" (OCA), subsequently adopted by most workers in the field:

$$\left(\frac{\mathrm{D} + {}^3\mathrm{He}}{\mathrm{H}}\right)_P \leq \left(\frac{\mathrm{D}}{\mathrm{H}}\right)_t + \frac{1}{g_3}\left(\frac{{}^3\mathrm{He}}{\mathrm{H}}\right)_t , \qquad (2)$$

where t is any epoch (usually the time of formation of the Sun), and g_3 is the mass fraction of ^{3}He that survives processing through a single generation of stars. Olive et al. (1990) and Walker et al. (1991) went a step forward replacing the "one-cycle" approximation with the "instantaneous recycling approximation" (IRA), which accounts for continuous stellar processing. Further progress can be achieved only with the use of detailed models of chemical galactic evolution (CGE), as recently shown by Steigman & Tosi (1992) and Vangioni-Flam, Olive, & Prantzos (1994). In Table 1 we compare the constraints on η_{10} resulting from the abundances of D and ^{3}He at the time of formation of the Sun and at the present epoch,

TABLE 1

CONSTRAINTS ON η_{10}

Tracer	Bounds	Method	Reference
$(D/H)_\odot$	$\eta_{10} \leq 7$	OCA	1
	$\eta_{10} \leq 8.52$	OCA	2
	$\eta_{10} \leq 6.8$	IRA	3
	$3.4 \leq \eta_{10} \leq 5.6$	CGE	4
	$3.8 \leq \eta_{10} \leq 4.8$	CGE	5
$(D/H)_0$	$4.2 \leq \eta_{10} \leq 4.9$	CGE	5
$(D + {}^3He)/H_\odot$	$\eta_{10} \geq 4$	OCA	1
	$\eta_{10} \geq 2.86$	OCA	2
	$\eta_{10} \geq 2.8$	IRA	3
	$3.7 \leq \eta_{10} \leq 5.7$	CGE	4

REFERENCES.—(1) Yang et al. 1984; (2) Smith et al. 1993; (3) Walker et al. 1991; (4) Steigman & Tosi 1992; (5) this work.

obtained with these different approaches. The results obtained in this paper are also given in Table 1. As it is clearly shown, *the most restrictive bounds on η_{10} follow from the results of CGE models.*

Recently, two groups (Songaila et al. 1994; Carswell et al. 1994) have reported the detection of a possible D feature in the spectrum of a Lyα absorption system at $z = 3.32$ in the line of view of the quasar Q0014 + 813, indicating $(D/H)_P \simeq 2 \times 10^{-4}$. These important observations represent the first attempt to determine *directly* the primordial abundance of this element. If confirmed, these measurements imply a primordial D abundance ~ 10 times larger than the one presently observed in the solar system and the ISM, thus requiring a severe depletion of this isotope during galactic evolution (see discussion in Vangioni-Flam & Cassé 1995). It is important to stress that such a high value of $(D/H)_P$ poses a challenge to galactic evolution models, as shown in Palla, Galli, & Silk (1995), *not* to big bang nucleosynthesis; actually, if $(D/H)_P \simeq 2 \times 10^{-4}$ the consistency between the primordial abundance of ^{4}He inferred from observations of very metal poor H II regions and the predictions of SBBN would improve significantly. However, according to both observing groups, the possibility that an errant hydrogen cloud with the appropriate velocity shift might be responsible for the observed absorption feature cannot be ruled out (see also Steigman 1994). Clearly, observations along different lines of views are needed to resolve this fundamental ambiguity.

Our goal in this paper is to reexamine the problem afresh on the basis of the most updated set of available observations and with the help of a reliable model of galactic evolution. It is worth noticing that in previous studies the evolution of ^{3}He has been treated always in an incomplete way, neglecting any stellar production of this isotope. We will see in § 5 that when this assumption is relaxed, as indicated by current models of stellar evolution, serious discrepancies arise between model predictions and observations. The questions we would like to answer here are (1) Is the production of ^{3}He in low- and intermediate-mass stars, as predicted by stellar evolutionary models, consistent with observations? (2) Are the observed abundances of D, ^{3}He, and ^{4}He consistent with SBBN predictions? and, if yes, (3) Which is the allowed range of values for η_{10}?

We will start by briefly reviewing the successes and failures of models of galactic chemical evolution applied to the study of the evolution of the light elements (§ 2). We will then discuss the observed abundances of D and ^{3}He in the ISM and in the Sun (§ 3). The model of chemical galactic evolution is described in § 4 and the results presented in § 5.

2. PREVIOUS WORK ON THE EVOLUTION OF D AND ^{3}He

Some of the problems raised in the pioneering works of Truran & Cameron (1971), Talbot & Arnett (1973), Reeves et al. (1973), Audouze & Tinsley (1974), Tinsley (1974) are today still subjects of debate: for example, (1) Is ^{3}He destroyed or produced by a single stellar generation? (2) How much astration of D is allowed by chemical evolution models without significant overproduction of ^{3}He?

The general conclusion reached on point (1) on the basis of simple "closed box" models with IRA was that the primordial ^{3}He plus that resulting from D burning in stars can account for the presolar ^{3}He, leaving little room for any stellar contribution. The conflict between the prediction of stellar evolution models and the observed abundance of ^{3}He at the time of formation of the solar system was made clear by Rood, Steigman, & Tinsley (1976), who found an overwhelming contribution to the enrichment of ^{3}He in the ISM by low-mass stars in the red giant phase. Using simple estimates, Rood et al. (1976) showed that the abundance of ^{3}He in the ISM increases dramatically in the course of galactic evolution, reaching the presolar value already after a few billion years. As possible solutions to this problem, they suggested the presence of inhomogeneities in the ISM, or uncertainties in the observed presolar abundance of ^{3}He (see also Tinsley 1977).

The problem of the D astration is more direct to analyze: since D can only be destroyed in stars, one simply needs to integrate its astration during galactic evolution. In practice, results obtained by different authors have favored either the case of a *low* astration ($D_P/D_{now} \sim 2$–3), or *large* astration ($D_P/D_{now} \sim 5$–10 or larger). Vidal-Madjar & Gry (1984) obtained an astration factor for D of the order of ~ 6, and derived a primordial D mass fraction $X_{2P} \leq 8.6 \times 10^{-5}$ (corresponding to $\eta_{10} \geq 3.5$). Delbourgo-Salvador et al. (1985, 1986, 1987) raised the primordial D to $X_{2P} \leq 3 \times 10^{-4}$ (corresponding to $\eta_{10} \geq 1.2$) by allowing a large astration factor (~ 15). In contrast, Clayton (1985) estimated a D astration factor of ~ 2–3 from simple models with IRA. Steigman & Tosi (1992) from numerical models found a similar result, determining $X_{2P} \leq 9 \times 10^{-5}$ (corresponding to $\eta_{10} \leq 3.7$).

The recent observations of a possible high $(D/H)_P$ have provoked a revival of models of chemical galactic evolution favoring a high D astration (see, e.g., Cassé & Vangioni-Flam 1994). It appears clear from the above discussion that some of the problems raised in the earlier investigations of the evolution of D and ^{3}He are still unsolved today and that the situation is rather confused. But before discussing our approach in detail, it is convenient to summarize the available observational constraints on the abundances of D and ^{3}He.

3. OBSERVED ABUNDANCES

The observational data on the abundance of the light elements in the ISM today and at the time of formation of the Sun have been reviewed in many places (see especially Geiss & Reeves 1972, 1981; Boesgaard & Steigman 1985; Geiss 1993). In this section we extend and update their data set, including the most recent observations, in order to provide the best constraints to the theoretical models.

It is often necessary to convert the observed number abundances with respect to H, indicated as (N_i/N_H), to mass fractions X_i which are commonly used in models of galactic evolution. The conversion formula is

$$X_i = A_i\left(\frac{N_i}{N_H}\right)X , \qquad (3)$$

where A_i is the mass number for the species i (as customary, we identify with $X \equiv X_1$ and $Y \equiv X_4$ the mass fraction of H and He, respectively). The mass fractions of the light elements produced by SBBN according to Walker et al. (1991) are given in Table 2. The abundance of ^{4}He given in Table 2 corresponds to a neutron lifetime $\tau_n = 889$ s.

3.1. *Abundances of* D *and* ^{3}He *at the Time of the Formation of the Sun*

The abundance of D and ^{3}He at the time of the formation of the Sun are reviewed in Appendix A. The values adopted throughout this paper, resulting from the discussion in Appendix A, are

$$\left(\frac{D}{H}\right)_\odot = (2.52 \pm 0.49) \times 10^{-5} , \qquad (4)$$

$$\left(\frac{^3He}{H}\right)_\odot = (1.48 \pm 0.13) \times 10^{-5} , \qquad (5)$$

$$\left[\frac{(D + {}^3He)}{H}\right]_\odot = (4.08 \pm 0.36) \times 10^{-5} . \qquad (6)$$

In terms of mass fractions, assuming $X_\odot = 0.707 \pm 0.025$ (see Appendix A), we obtain

$$X_{2\odot} = (3.56 \pm 0.78) \times 10^{-5} , \qquad (7)$$

$$X_{3\odot} = (3.14 \pm 0.35) \times 10^{-5} . \qquad (8)$$

3.2. *Abundances of* D *and* ^{3}He *in the ISM*

The abundance of the light elements in the ISM has been reviewed recently by, e.g., Wilson & Matteucci (1992) and Wilson & Rood (1994). A recent reanalysis (McCullough 1991) of the D/H ratio obtained from satellite UV spectroscopy has shown that the range

$$\left(\frac{D}{H}\right)_0 = (1.5 \pm 0.2) \times 10^{-5} , \qquad (9)$$

is consistent with all the data. This value is in agreement with the (D/H) ratio measured along the line of sight toward Capella, (D/H) = $1.65^{+0.07}_{-0.18} \times 10^{-5}$ (Linsky et al. 1993). Taken at face value, these results support the belief in a single (D/H) value in the local ISM and provide a strong constraint on the galactic evolution of D. We will assume in this paper that the McCullough (1991) value is representative of the (D/H) ratio at $t = t_0$.

Less clear is the current status of ^{3}He abundance determinations. Singly ionized ^{3}He has been detected in several H II regions via its hyperfine spin-flip emission at 3.46 cm (Rood, Steigman, & Tinsley 1976; Rood, Bania, & Wilson 1984; Bania, Rood, & Wilson 1987, Balser et al. 1994). The ^{3}He abundance can only be derived after modeling the properties and geometry of the source. The most accurate values of $(^3He/H)_{ISM}$ are presented by Balser et al. (1994), who found abundances in the range ^{3}He/H $\simeq 1$–4×10^{-5} with considerable source to source variation, and no clear dependence on galactocentric distance. Future work will probably shed more light on this rather unsettled situation. For instance, Lemoine et al. (1993) have proposed a profile fitting method for determining the ^{3}He/^{4}He isotopic ratio in the local ISM from measurements of ^{4}He absorption lines in the far UV. Even though the expected isotopic ratio may be very small ($\sim 10^{-4}$) the method requires instrumental performances within the capabilities of the planned Lyman-FUSE mission.

As a reference value at $t = t_0$, we will take in this paper the average abundance in galactic H II regions determined by Balser et al. (1994),

$$\left(\frac{^3He}{H}\right)_0 = (2.42 \pm 1.45) \times 10^{-5} . \qquad (10)$$

TABLE 2

PRIMORDIAL ABUNDANCES ACCORDING TO SBBN[a]

η_{10}	X	Y	X_2	X_3
1	0.780	0.220	7.3(−4)	1.7(−4)
2	0.767	0.234	2.1(−4)	4.4(−5)
2.2	0.765	0.236	1.8(−4)	4.1(−5)
2.4	0.764	0.236	1.6(−4)	3.9(−5)
2.6	0.763	0.237	1.4(−4)	3.7(−5)
2.8	0.762	0.238	1.2(−4)	3.6(−5)
3	0.762	0.238	1.1(−4)	3.4(−5)
3.2	0.761	0.239	9.8(−5)	3.3(−5)
3.4	0.760	0.240	8.7(−5)	3.2(−5)
3.6	0.759	0.240	7.9(−5)	3.1(−5)
3.8	0.759	0.241	7.1(−5)	3.1(−5)
4	0.758	0.242	6.5(−5)	3.0(−5)
4.2	0.758	0.242	6.0(−5)	2.9(−5)
4.4	0.758	0.242	5.6(−5)	2.9(−5)
4.6	0.757	0.243	5.2(−5)	2.8(−5)
4.8	0.757	0.243	4.8(−5)	2.8(−5)
5	0.756	0.244	4.5(−5)	2.7(−5)
6	0.754	0.246	3.3(−5)	2.5(−5)
7	0.753	0.247	2.6(−5)	2.2(−5)
8	0.752	0.248	2.0(−5)	2.2(−5)
9	0.750	0.249	1.6(−5)	2.1(−5)
10	0.750	0.250	1.3(−5)	2.0(−5)

[a] From Walker et al. 1991.

4. A MODEL FOR THE EVOLUTION OF THE LIGHT ELEMENTS

4.1. *Definitions and Equations*

The numerical code we employ is described in detail in Ferrini et al. (1992, hereafter FMPP). The equations for the evolution of the abundance of the chemical species i are

$$\frac{d(X_{i,h}\,g_h)}{dt} = -\psi_h X_{i,h} - fg_h X_{i,h} + W_{i,h} , \qquad (11)$$

$$\frac{d[X_{i,d}(g_d + c_d)]}{dt} = -\psi_d X_{i,d} + fg_h X_{i,h} + W_{i,d} , \qquad (12)$$

where the subscripts h and d refer to the halo and disk regions, respectively. Here, ψ is the star formation rate, f the infall parameter, g and c are the mass fractions in the gas and cloud phases, respectively. The restitution rate $W_i(t)$ includes the contribution of single stars at the end of normal main-sequence evolution (MS), Type I supernovae (SN I) and type II supernovae (SN II):

$$W_i(t) = W_i^{MS}(t) + W_i^{SN\,I}(t) + W_i^{SN\,II}(t) , \qquad (13)$$

for both disk and halo. These quantities are computed in the framework of the matrix formalism introduced by Talbot & Arnett (1976; henceforth TA). Let us briefly recall the essential aspects of this approach.

First, each star is assumed to be chemically homogeneous *at birth*, having the same (average) composition of the ISM at the instant of formation. At the end of the evolution, each star is subdivided into one or more concentric mass shells, where different nucleosynthesis processes have occurred. Particularly relevant are the mass fractions associated to the quantities $q_3(m)$, $q_4(m)$, $q_C(m)$, corresponding, respectively, to the zones within which ^{3}He has been burnt into ^{4}He and ^{7}Be, H into ^{4}He, and ^{4}He into C, O, and heavier species. Finally, the quantity $d(m)$ defines the mass fraction which remains as a stellar remnant upon the death of the star [i.e., all the matter external to $d(m)$ is ejected in the ISM]. Clearly, the ordering is $0 < d(m) \leq q_C(m) \leq q_4(m) \leq q_3(m) \leq 1$.

The mass of the element i ejected by a star of mass m after a time $\tau(m)$ from birth is assumed to be a linear combination of the masses of elements j initially present in the star:

$$m_i^{\rm ej} = \sum_j Q_{ij} m_j^{\rm in} , \tag{14}$$

where the matrix elements Q_{ij} depend only on the mass of the star (the dependence upon the initial metallicity of the star may also be treated with this formalism; see Sandrelli 1993). The restitution rate from MS stars results then

$$W_i^{\rm MS}(t) = \int_{m(t)}^{m_{\rm upp}} \sum_j Q_{ij}(m') X_j[t - \tau(m')] \psi[t - \tau(m')] \phi(m') dm' . \tag{15}$$

This formalism has the advantage that the enrichment of the ISM produced by stars of different masses can be expressed in a compact way simply as a difference between mass fractions. For example, the fraction (in mass) of the original H that is transformed into ^{4}He and ejected is $q_4 - q_C$, as matter inside q_C undergoes ^{4}He burning whereas matter outside q_4 does not burn H; this automatically defines $Q_{41} = q_4 - q_C$. The other matrix elements Q_{ij} can be determined in the same way.

We will consider in detail only the specifications given by TA on the matrix elements relative to D and ^{3}He, based mostly on the numerical models of Bodenheimer (1966) and Iben (1965, 1966a, b, 1967a, b). We have partially reformulated the production matrix Q_{ij} for H, D, and ^{3}He in the light of the recent theoretical results on the early phases of stellar evolution (Palla & Stahler 1993; D'Antona & Mazzitelli 1994) and on the basis of updated calculation of the survival of ^{3}He in the post–main-sequence phase of evolution (Dearborn, Schramm, & Steigman 1986). We maintain the nucleosynthesis prescriptions adopted by FMPP on q_4, q_C, and on all the quantities necessary to evaluate the evolution of elements more massive than ^{4}He (for simplicity referred to as a single element Z). Our detailed nucleosynthesis prescriptions are discussed here below.

4.2. *Stellar Nucleosynthesis of Hydrogen and Helium*

Following TA, we assume that a fraction $q_4(m)$ of the initial mass of H is processed through hydrogen burning, a fraction w_3 is transformed into ^{3}He, and a fraction $1 - q_4 - w_3$ survives stellar processing and is ejected. Since all H inside q_C is transformed into C, O and heavier species, the fraction of H transformed into ^{4}He and ejected is equal to $q_4 - q_C$. Hence, in terms of Q_{ij} matrix elements, we have $Q_{11} = 1 - q_4 - w_3$, $Q_{21} = 0$, $Q_{31} = w_3$, $Q_{41} = q_4 - q_C$, $Q_{Z1} = q_C - d$.

Also, we assume that all the initial ^{4}He external to q_C is not affected by stellar nucleosynthesis, and is ejected at the end of the life of a star, whereas the initial ^{4}He internal to q_C is entirely converted into C, N, O, and more massive nuclei. Therefore, $Q_{14} = Q_{24} = Q_{34} = 0$, $Q_{44} = 1 - q_C$, $Q_{Z4} = q_C - d$.

The values of the quantities $q_4(m)$, $q_C(m)$, amd $d(m)$ are the same adopted by FMPP. Notice that, with the above definition, the sum of matrix elements along each column is equal to $1 - d$, to comply with mass conservation.

4.3. *Stellar Nucleosynthesis of Deuterium*

Deuterium can be destroyed only in stars. The formation of some D in supernova shocks (Colgate 1973), if at all possible, is in any case irrelevant in the galactic context (Epstein, Schramm, & Arnett 1974). However, the amount of actual destruction of D in stars has been rather controversial. TA assumed that in stars less massive than 2 $M_\odot$ all D is converted into heavier species whereas in stars more massive than 4 $M_\odot$ all D survives outside the mass fraction $q_3(m)$, and is completely burnt inside $q_3(m)$. Recently, however, Palla & Stahler (1991, 1993) have shown that this prescription is incorrect, because D is completely destroyed during the protostellar and pre–main-sequence (PMS) phases in stars of all masses. More precisely, in the mass range 2.4 $M_\odot \leq m \leq 8$ $M_\odot$ for a rate of accretion 10^{-5} $M_\odot$ yr^{-1}, or in the mass range 4.3 $M_\odot \leq m \leq$ 15 $M_\odot$ for a rate of accretion 10^{-4} $M_\odot$ yr^{-1}, PMS stars possess very thin outer radiative layers containing unburnt D. If the initial D is completely burnt to ^{3}He, a mass equal to 3/2 of the initial mass of D is added to the initial mass of ^{3}He. Therefore, the initial (on the MS) mass of ^{3}He is equal to $[m_3^{\rm in} + (3/2)m_2^{\rm in}]$. In terms of the matrix elements Q_{ij}, this means $Q_{22} = 0$ and $Q_{32} = (3/2)(1 - q_3)$. The burning of a mass $m_2^{\rm in}$ of D to ^{3}He requires the destruction of a mass equal to $\frac{1}{2}m_2^{\rm in}$ of H, a fraction $1 - d(m)$ of which is ejected. Hence $Q_{12} = -\frac{1}{2}(1 - d)m_2^{\rm in}$.

4.4. *Stellar Nucleosynthesis of* ^{3}He: *production*

This isotope is both produced and destroyed in stars. In TA's formalism, the newly formed (and ejected) ^{3}He is a fraction $w_3(m)$ of the original H mass, its value being estimated on the basis of Iben's numerical results. The destruction of this species occurs inside the mass fraction $q_3(m)$, so that the fraction (in mass) of the original ^{3}He ejected back in the ISM is given by $1 - q_3$. Since the accuracy of the prescriptions adopted by TA has been largely improved, and given the potential importance of this element as a test of cosmological models, we examine here the processes of production of ^{3}He in some detail.

First, ^{3}He is produced in a star by the burning of the initial D during the PMS phase. Second, it is also produced via the *p-p* chain in the MS phase, during which new ^{3}He is created in any sufficiently cool H burning zone independently of the initial D and ^{3}He abundances. In fact, ^{3}He is produced in stars via D$(p, \gamma)^3$He, which occurs at temperatures larger than $\sim 10^6$ K, and destroyed via ^{3}He$(^3$He$, 2p)^4$He or ^{3}He$(^4$He$, \gamma)^7$Be in layers with $T \geq 7 \times 10^6$ K. Therefore, significant abundance of ^{3}He in a stellar interior can be found only in a shell where the temperature is high enough for D burning but low enough for ^{3}He survival.

Indeed, the abundance profile of ^{3}He in MS stars of a few solar mass exhibits a characteristic peaked shape (the peak corresponding to a temperature of 7×10^6 K), with the maximum located at a mass fraction of ~ 0.6 and half-width of ~ 0.2. The peak ^{3}He abundance is of the order of few 10^{-4}–10^{-3} by mass, depending on stellar mass and composition.

TABLE 3

MASS FRACTION OF ^{3}He IN THE EJECTA OF STARS OF VARIOUS MASSES

$m/M_\odot$ (1)	Iben[a] $Z = 0.02$ (2)	RST[b] $Z = 0.02$ (3)	VW[c] $Z = 0.016$ (4)	Straniero[d] $Z = 0.02$ (5)	Straniero[e] $Z = 0.02$ (6)	Straniero[f] $Z = 0.02$ (7)
1	1.62(−3)	9.86(−4)	6.71(−4)	1.38(−3)	1.30(−4)	5.43(−6)
1.25	9.07(−4)	7.15(−4)	...	...	...	...
1.5	6.07(−4)	6.30(−4)	5.33(−4)	...	...	...
2	...	...	2.96(−4)	...	...	...
2.25	2.47(−4)	...	...	...	...	...
2.5	...	...	1.58(−4)	...	...	...
3	1.30(−4)	...	...	1.58(−4)	3.66(−5)	6.94(−6)
3.5	...	...	5.92(−5)	...	...	...
5	5.63(−5)	...	...	8.46(−5)	3.01(−5)	9.34(−6)
9	1.74(−5)	...	...	...	...	...

[a] From Iben 1965, 1966a, b; 1967a, b.
[b] From Rood et al. 1976.
[c] From Vassiliadis & Wood 1993.
[d] From Straniero 1994, private communication; nonresonant case.
[e] From Straniero 1994, private communication; resonant case with $E_r = 10$ keV.
[f] From Straniero 1994, private communication, resonant case with $E_r = 3$ keV.

When the star rises the red giant branch (RGB) for the first time, this freshly produced ^{3}He is diluted over the region covered by the convective envelope at its maximum inward extent. The resulting surface ^{3}He abundance in Iben's "final" (after the first dredge-up) models is given in Table 3. For stars of $m \geq 3\ M_\odot$, the "final" ^{3}He surface abundance has been computed by dividing the "final" ^{3}He mass shown in Table 4 of Iben (1967a) by the mass of the star covered by the convective envelope at its maximum extent. Even though the cross section assumed by Iben for the ^{3}He(^{3}He, 2p)^{4}He reaction (the primary reaction responsible for ^{3}He destruction) was smaller by a factor ~ 5 than the one currently adopted, the corresponding underestimate on the amount of ^{3}He production is only a factor ~ 1.5. In fact, a global scaling of the cross section affects more the position of the peak in the abundance profile (which moves outward as the cross section increases) than the amount of ^{3}He production (Schatzman 1987).

Later on, Rood (1972) and Rood et al. (1976) explored the dependence of the "final" ^{3}He surface abundance in low-mass stars of differing initial compositions, finding that this quantity increases with the time the star takes to get to the base of the RGB, which is a function of the star's initial composition. The values of $X_3^{\rm surf}$ computed by Rood et al. (1976) for $m < 1.5\ M_\odot$ and $Z = 0.02$ are shown in column (3) of Table 3. Clearly, Iben's final abundances computed with the updated cross section for the ^{3}He(^{3}He, 2p)^{4}He are in excess of Rood's only by a factor ~ 1.6 at worst, as predicted by Schatzman (1987). Indeed, the peak abundance of ^{3}He in the closest of Iben's models to the present-day Sun is 4.20×10^{-3} (Iben 1967), lower only by a factor ~ 1.5 than more refined computations for the standard solar models (3.01×10^{-3} in Turck-Chièze et al. 1986; 3.17×10^{-3} in Lebreton & Dappen 1988; 3.22×10^{-3} in Bahcall & Ulrich 1988; 2.86×10^{-3} in Cox, Guzik, & Kidman 1989).

The fate of ^{3}He during the latest stages of stellar evolution has been studied less extensively. According to Rood et al. (1976), "^{3}He survives to the He-flash and presumably after," and the material ejected during RGB winds or in planetary nebulae "should have at least half the ^{3}He abundance as the material from the first excursion up to the RGB." Sackmann, Smith, & Despain (1974) and Scalo, Despain, & Ulrich (1975) have followed in detail the evolution of ^{3}He in red giants of 1, 2, and 5 $M_\odot$ during the phase of deep mixing, starting with the value determined by Rood (1972) at the base of the RGB. Even though their results show that X_3 eventually declines after 10^3–10^4 yr from the start of deep mixing, it is quite unlikely that the star remains in this phase for long times. In particular, at the time when H-burning releases an amount of energy equal to the binding energy of the envelope, and, presumably, a sudden, possibly violent, mass-loss episode occurs, the ^{3}He abundance is still very close to the value on the base of RGB. An important clue is provided by the recent results of Vassiliadis & Wood (1993), in which the fate of ^{3}He is followed through the first and second dredge-up and the thermal pulses phase. The abundance of ^{3}He in the envelope after the thermal pulses phase is listed in Table 3 for the case of solar metallicity. The striking conclusion of Vassiliadis & Wood (1993) is the confirmation of Rood et al.'s (1976) prediction, namely that the abundance of ^{3}He established in the outer envelope of stars in the mass range $0.8 \leq m \leq 5$ remains virtually unchanged even after the thermal pulses phase. Thus, it appears that the ^{3}He enrichment of the ISM from stars of low and intermediate mass is an unavoidable consequence of stellar evolution.

In this context, a particular relevance assumes the first detection of ^{3}He in the ejecta of the planetary nebula NGC 3242, recently reported by Rood, Bania, & Wilson (1992) and Bania, Rood, & Wilson (1993), in which ^{3}He/H $\simeq 1.4 \times 10^{-3}$ corresponding to $X_3^{\rm surf} \simeq 3 \times 10^{-3}$. The mass of the progenitor of NGC 3242 being $1.0 \pm 0.2\ M_\odot$ (Stanghellini, private communication), this value is in agreement with the abundance predicted by Iben's models for a 1 $M_\odot$ star. Figure 1 shows the predicted surface abundance of ^{3}He according to the different stellar models discussed before as function of the stellar mass in the interval $0.5 \leq m \leq 10$. Also shown is the ^{3}He abundance observed in NGC 3242 according to Bania et al. (1993).

The net production of ^{3}He in our model is expressed by the function $w_3(m)$. If $X_3^{\rm surf}(m)$ is the abundance of ^{3}He brought to surface, then $w_3(m)$ results in

$$w_3(m) = \frac{[X_3^{\rm surf}(m) - X_3^{\rm in}(m) - (3/2)X_2^{\rm in}(m)]m^{\rm ej}}{X^{\rm in}m}$$
$$= \frac{X_3^{\rm surf}(m) - X_3^{\rm in}(m) - (3/2)X_2^{\rm in}(m)}{X^{\rm in}}[1 - d(m)] \ . \quad (16)$$

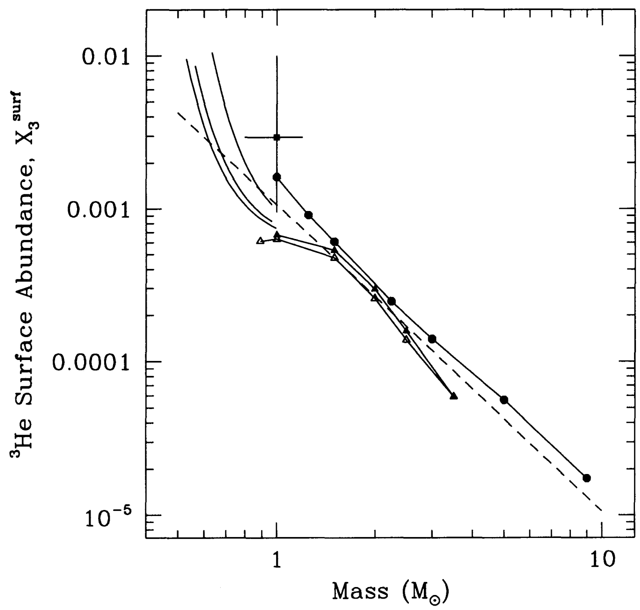

FIG. 1.—Final surface abundance of ^{3}He as function of the stellar mass according to different authors. *Circles*: Iben (1965, 1966a, b, 1967a, b); *thin solid curves*: Rood et al. (1976) for different metallicities ($Z = 2 \times 10^{-2}$, *upper*; 2×10^{-3}, *middle*; 2×10^{-4}, *lower*); *dashed curve*: analytical expression given by Schatzman (1987); *triangles*: Vassiliadis & Wood (1993) for $Z = 0.016$ (*filled symbols*) and $Z = 0.004$ (*open symbols*). Also shown as the *filled square* is the observed ^{3}He abundance in the planetary nebula NGC 3242 (Bania et al. 1993), with the uncertainty on $X_3^{\rm surf}$ from Bania et al. (1993) and on the mass from Stanghellini (1993, private communication).

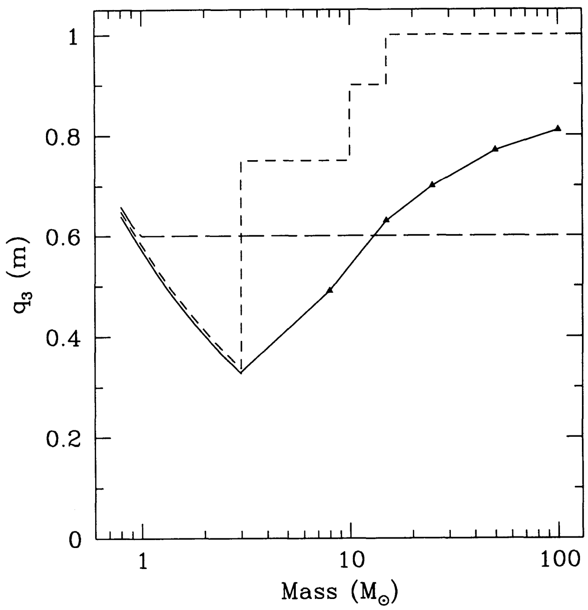

FIG. 2.—The fraction q_3 of stellar mass inside which ^{3}He has been burnt to ^{4}He or ^{7}Be, according to different authors. *Short-dashed curve*: Truran & Cameron (1971); *long-dashed curve*: Talbot & Arnett (1973); *triangles*: Dearborn et al. (1986); *solid line*: adopted in this work.

4.5. *Stellar Nucleosynthesis of* ^{3}He: *Destruction*

The situation we have so far considered is valid for low-mass stars (m smaller than $\sim 2\ M_\odot$). Massive stars ($m \geq 8\ M_\odot$) destroy rather than produce ^{3}He (Dearborn et al. 1986). Let $q_3(m)$ be the mass fraction of a star processed at temperatures larger than $\sim 7 \times 10^6$ K and therefore depleted in ^{3}He. Since, as we have already seen, in Iben's MS stellar models the ^{3}He distribution is peaked around a mass fraction equal to ~ 0.6 almost independently on stellar mass, TA adopted $q_3(m) = \max\,[0.6, d(m)]$ over the entire range of stellar masses (the maximum is to comply with the ordering discussed above). Dearborn et al. (1986) have run stellar models at selected masses (8, 15, 25, 50, and 100 $M_\odot$) specifically to compute the fraction $g_3(m)$ of the stellar mass in which temperatures larger than $\sim 7 \times 10^6$ K are never achieved; their quantity $g_3(m)$ is thus equivalent to our quantity $1 - q_3(m)$. For stars of $m < 8\ M_\odot$, lacking more detailed modeling, a good estimate of $q_3(m)$ is the fraction of the initial mass which is outside the H burning shell. Our complete prescription on $g_3(m)$ is then $q_3(m) = q_4(m)$ if $m < 3\ M_\odot$, and $q_3(m) = 1 - g_3(m)$ at the masses considered by Dearborn et al. (1986), and a linear interpolation in between. The resulting curve is shown in Figure 2 where we compare it with the prescriptions adopted by TA and Truran & Cameron (1971). The complete matrix Q_{ij} is shown in Table 4.

5. RESULTS

5.1. *The Evolution of* D *and* ^{3}He *in the Standard Model*

Starting from the predictions of SBBN, we are now able to follow accurately the evolution of D and ^{3}He during the galactic history. The calculations are based on a particular model selected from our previously published studies (Ferrini & Galli 1988; Galli & Ferrini 1989; FMPP) that gives the best-fit to the observational constraints for our Galaxy (age-metallicity relation, G-dwarfs distribution, etc.). The adopted age of the Galaxy is $t_0 = 13$ Gyr and the epoch of formation of the Sun is $t_\odot = 8.5$ Gyr. For comparative purposes, we also present in Appendix B a simplified, analytical model of the time evolution of D and ^{3}He.

The time evolution of (D/H) for various values of η_{10} is shown in Figure 3; the number abundance of D with respect to H at $t = t_\odot$ and $t = t_0$ is listed in Table 5. Results for ^{3}He are presented in Figure 4 and Table 6 for two cases. In one, stellar production of ^{3}He is computed on the basis of the analytical expression for $X_3^{\rm surf}(m)$ given by Schatzman (1987),

$$X_3^{\rm surf}(m) = 1.063 \times 10^{-3}\left(\frac{m}{M_\odot}\right)^{-2} . \tag{17}$$

which is ultimately based on the solar model described in Schatzman & Maeder (1981). This simple expression represents a good fit to Iben and Vassiliadis & Wood (1993) results for intermediate masses, and a reasonable average of the two at lower masses. This case will be referred to as the

TABLE 4

MATRIX Q_{ij}

	H	D	^{3}He	^{4}He	Z
H	$1 - q_4 - w_3$	$-\frac{1}{2}(1 - d)$	0	0	0
D	0	0	0	0	0
^{3}He......	w_3	$\frac{3}{2}(1 - q_3)$	$1 - q_3$	0	0
^{4}He......	$q_4 - q_C$	$\frac{3}{2}(q_3 - q_C)$	$q_3 - q_C$	$1 - q_C$	0
Z	$q_C - d$	$\frac{3}{2}(q_C - d)$	$q_C - d$	$q_C - d$	$1 - d$

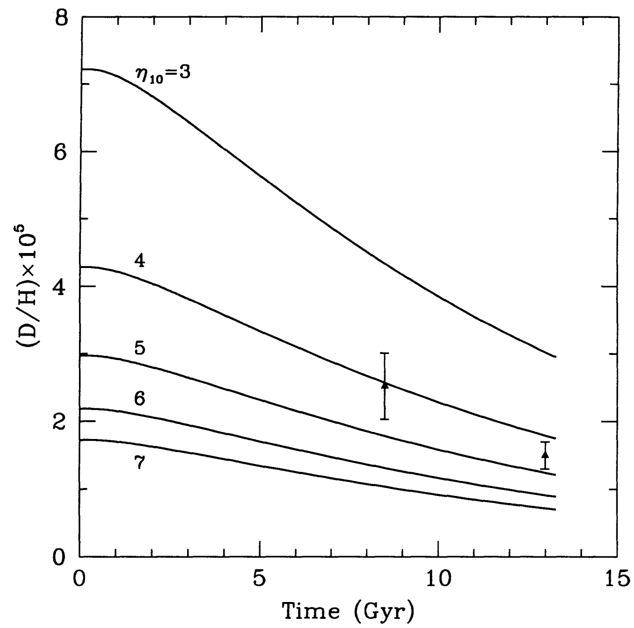

FIG. 3.—Evolution of the deuterium number abundance vs. time for different values of η_{10}. The filled triangles give the observed presolar (t = 8.5 Gyr) and ISM (t = 13 Gyr) abundances, together with their 1 σ uncertainties (data sources are discussed in Appendix A).

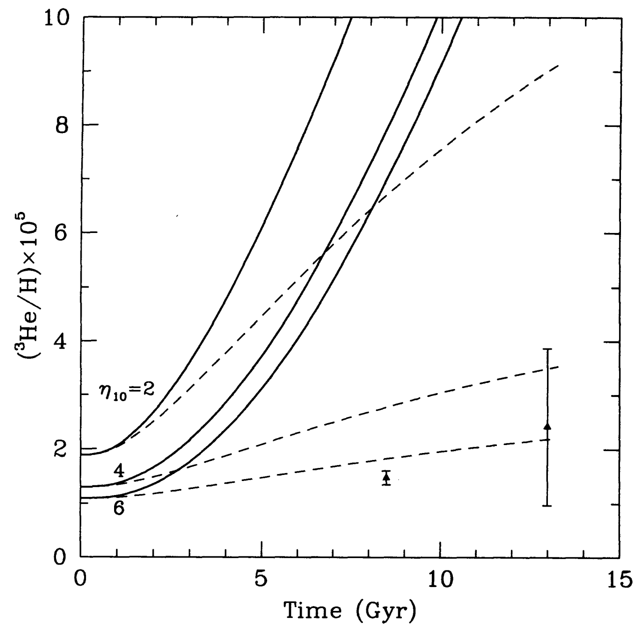

FIG. 4.—Evolution of the ^{3}He number abundance vs. time for selected values of η_{10}. The *solid* curves show the evolution in the standard model that assumes production of ^{3}He in stars, according to the prescription of Schatzman (1987). The *dashed curves* are for the case with no production of ^{3}He $w_3 = 0$. The *filled triangles* give the observed presolar (t = 8.5 Gyr) and ISM (t = 13 Gyr) abundances, together with their 1 σ uncertainty (data sources are discussed in Appendix A).

"standard model." In the other case, we assume no stellar production of ^{3}He and we simply take $w_3(m) = 0$.

From the curves of Figure 3, we see that the astration of D is in all cases relatively modest: the surviving D is less than the primordial value by a factor ~1.5 at $t = t_\odot$ and ~2.5 at $t = t_0$. Our results are thus in agreement with Audouze & Tinsley (1974), Vangioni-Flam & Audouze (1988) (their canonical case), and Steigman & Tosi (1992) but deviate significantly from those of Delbourgo-Salvador et al. (1985, 1986, 1987), and Vangioni-Flam et al. (1994).

The abundance of D in the presolar material and in the ISM sets stringent bounds on the value of η_{10}. If the values adopted in this paper are taken at face value, then the bounds on η_{10} from $(D/H)_\odot$ and from $(D/H)_0$ are $3.8 \le \eta_{10} \le 4.8$ and $4.2 \le \eta_{10} \le 4.9$, respectively. The 1 σ uncertainty on the SBBN abundances of D and ^{3}He is of the order of 30%, according to the analysis of Smith et al. (1993). Since in this range of the baryon-to-photon ratio the relative error on η_{10} is $\Delta\eta_{10}/\eta_{10} \simeq 0.5\,\Delta X_{2P}/X_{2P}$, these bounds should be fairly precise *if* other sources of error are absent.

The run of the ^{3}He abundance versus time of Figure 4 clearly shows the serious conflict with observations, unless stellar production of this isotope is zero, i.e., $w_3 = 0$. Even in this favorable circumstance, however, it is not possible to reproduce the solar value unless $\eta_{10} \simeq 7$, too high a value to provide acceptable results for D (and ^{4}He). If stellar production is taken into account, the overproduction of ^{3}He is so large that the initial abundance becomes unimportant. Thus, the possibility that η_{10} could be of the order of 20 or larger, as advocated by Schatzman (1987) on this basis, would not resolve the discrepancy either.

Finally, in Figure 5 we show the time evolution of the sum (D + ^{3}He). Once again, consistency with observations can only be obtained if there is no net stellar production of ^{3}He. In this case, the slight overproduction of ^{3}He at $t = t_\odot$ shown in Figure 4 is compensated by the effective decrease of D (cf. Fig. 3) such that their sum remains almost constant. Otherwise, the enrichment of ^{3}He is unavoidable. We are thus left with the fundamental problem of finding an effective mechanism capable of eliminating, or at least reducing, the inconsistency. In the following sections, we will consider several effects that are likely to influence the galactic evolution of ^{3}He.

TABLE 5

VALUES OF (D/H) AT $t = t_\odot$ AND $t = t_0$

η_{10}	$t_\odot$	t_0
2......	8.68(−5)	5.73(−5)
3......	4.58(−5)	3.02(−5)
4......	2.72(−5)	1.79(−5)
5......	1.88(−5)	1.24(−5)
6......	1.38(−5)	9.12(−6)
7......	1.09(−5)	7.24(−6)

TABLE 6

VALUES OF (^{3}He/H) AT $t = t_\odot$ AND $t = t_0$

	$w_3 = 0$		STANDARD MODEL	
η_{10}	$t_\odot$	t_0	$t_\odot$	t_0
2......	6.35(−5)	9.12(−5)	1.23(−4)	2.56(−4)
3......	3.85(−5)	5.33(−5)	9.79(−5)	2.17(−4)
4......	2.71(−5)	3.58(−5)	8.66(−5)	2.00(−4)
5......	2.14(−5)	2.74(−5)	8.10(−5)	1.92(−4)
6......	1.80(−5)	2.23(−5)	7.76(−5)	1.87(−4)
7......	1.52(−5)	1.86(−5)	7.49(−5)	1.83(−4)

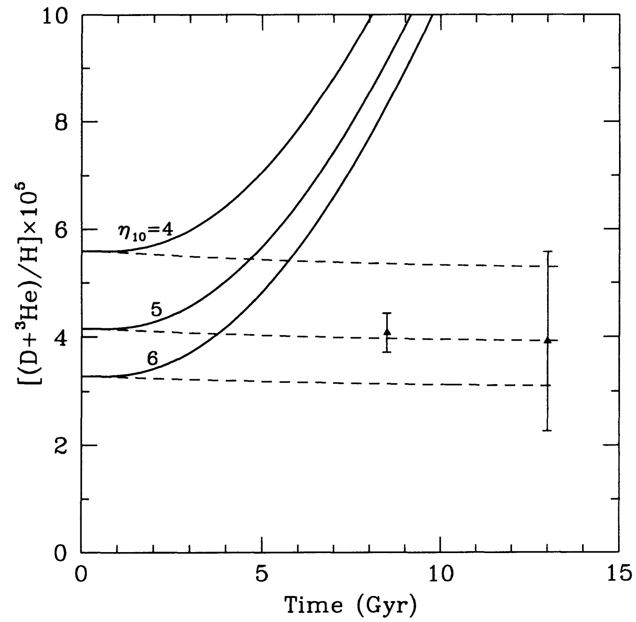

FIG. 5.—Evolution of the (D + ^{3}He) number abundance vs. time for selected values of η_{10}. The *solid curves* show the evolution in the standard model with stellar production of ^{3}He. The *dashed curves* are for the case with no production of ^{3}He, $w_3 = 0$. The *filled triangles* give the observed presolar ($t = 8.5$ Gyr) and ISM ($t = 13$ Gyr) abundances, together with their 1 σ uncertainty (data sources are discussed in Appendix A).

5.2. *Infall*

A key feature of the FMPP model of galactic chemical evolution is the role played by "infall," interpreted as the physical coupling between the halo and the disk zones of the solar cylinder. Indeed, the disk is treated here as a purely secondary structure, formed by the accumulation of gas released from (and partially processed by) the halo. Since different authors have very different views on the nature and relevance (if any) of infall in models of chemical galactic evolution, it is necessary to understand its effects on the evolution of the light elements. To this goal, we have run a modified standard model in which material infalling on the disk is taken as ^{3}He-free, keeping everything else unchanged. Clearly, this choice is an extreme one, aimed at determining the *maximum* dilution of ^{3}He that can be expected to arise from mixing with extragalactic primordial material.

In the resulting evolution, we found that the abundance of ^{3}He at $t = t_\odot$ is reduced by 42% with respect to the standard model. Thus, the effect of unenriched infall cannot account for the reduction of about a factor ≥ 5 required to bring the ^{3}He abundance in agreement with the observed value. In addition, the dilution due to the unenriched material becomes increasingly less efficient with time, whereas stellar production, being larger for less massive stars, has the opposite behavior: thus, at $t = t_0$, the reduction in the ^{3}He abundance is only 35% with respect to the standard model.

5.3. *An Initial Starburst?*

In this section we explore the possibility that the Galaxy had a "starburst" phase at very early times, characterized by the rapid evolution of a large number of massive stars. We test this hypothesis by considering its consequences on the evolution of the light elements and on the metallicity distribution. Following the approach of Charlot et al. (1993), we simulate the effect of a starburst with a modified IMF with respect to the one adopted in the standard model. Leaving everything else unchanged, we assume that during the starburst phase, lasting from $t = 0$ to $t = 2$ Gyr, no stars less massive than 3 $M_\odot$ were formed ("truncated" IMF); from $t = 2$ Gyr to t_0, the evolution is followed as in the standard model (a similar scenario has been considered by Vangioni-Flam & Audouze 1988). According to Charlot et al. (1993), this picture, although very crude, may be relevant for the interpretation of the red-spectrum galaxy population recently discovered at low- and intermediate-redshifts (Aragón-Salamanca et al. 1993, and references therein). However, it is questionable whether such an extreme model could be applied to our Galaxy.

The rationale behind this test is the expected reduction in the ^{3}He abundance that should result in this case (low- and intermediate-mass stars are the most prominent ^{3}He producers). In reality, since in the starburst phase the gas is processed by short-lived massive stars, a "truncated" IMF leads unavoidably to a larger astration rate, which results in a high destruction rate of the initial D (reduced by a factor ~ 17 at $t = t_0$). In turn, this excess of destroyed D is recovered in the form of ^{3}He, and more than compensates for the reduced stellar production of this element. At $t = t_\odot$, the D and ^{3}He mass fractions are $X_2(t_\odot) = 9.04 \times 10^{-6}$ and $X_3(t_\odot) = 4.36 \times 10^{-4}$, respectively, for the standard case $\eta_{10} = 4$. The full time evolution of D and ^{3}He is shown in Figure 6. As we can see, not only the discrepancy with the observed value of presolar ^{3}He is made even worse, but D is now deficient by a large amount, a factor of ~ 4. Also, the resulting age-metallicity relation and the metallicity distribution of disk G-dwarfs are in complete disagreement with observations: solar metallicity is reached already at $t \sim 0.5$ Gyr, and the peak of the G-dwarf distribution is shifted toward [Fe/H] $\simeq 0.5$.

We conclude from this crude test that the possibility of a "truncated" IMF during the early evolution of the Galaxy is

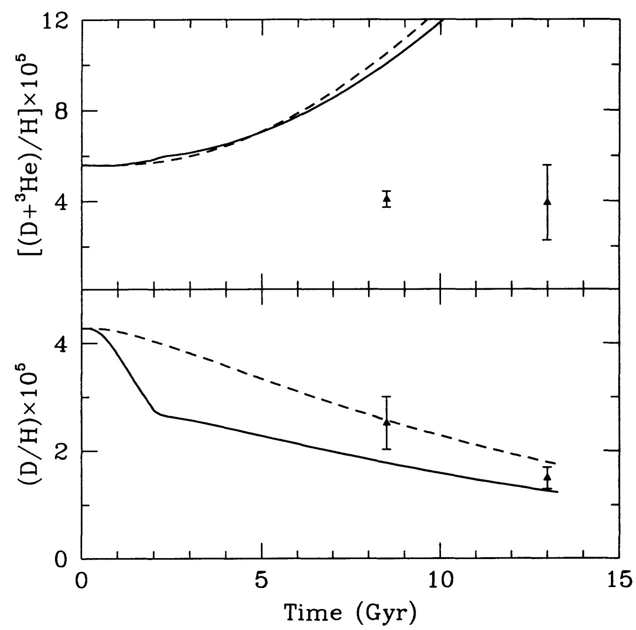

FIG. 6.—Effect of a truncated IMF on the D (*bottom*) and (D + ^{3}He) (*top*) evolution. The evolution in the starburst case is given by the *solid* curve, while the *dashed* curve is for the standard model. In either case, $\eta_{10} = 4$.

strongly contradicted by observations, and, in any case, its effect on the evolution of the light elements would not alleviate the observed discrepancies, but make them worse.

5.4. *Effect of Varying Stellar Lifetimes*

The MS stellar lifetimes adopted in the standard model are obtained from the power-law approximations given by Burkert & Hensler (1987). However, many effects may modify the lifetime of stars in the relevant mass range for production of ^{3}He, making necessary to test the sensitivity of our results on this parameter. To evaluate quantitatively this effect, we have run models based on the main-sequence lifetimes tabulated by VandenBerg (1985) ($Y = 0.25$, $Z = 0.02$), Maeder & Meynet (1989) ($Y = 0.28$, $Z = 0.02$) and Schaller et al. (1992) ($Y = 0.30$, $Z = 0.02$). In all cases the tabulated values have been fitted with a cubic polynomial in log τ–log m.

VandenBerg stellar evolutionary models are calculated without overshooting with updated (at 1985) opacities and nuclear reaction rates. Maeder & Meynet lifetimes are based on stellar models including overshooting ($d/H_p = 0.25$) which tends to increase stellar lifetimes. Indeed, the values tabulated by Maeder & Meynet are higher by a factor from 1.5 to 2.7 than VandenBerg's. It is important to notice, however, that isochrones calculated with the new OPAL opacities and little or no overshooting (see, e.g., Schaller et al. 1992) are in agreement with the results of VandenBerg.

In Table 7 we present the results of our standard model ($\eta_{10} = 4$) for different assumptions on the stellar lifetimes. From the comparison between the results obtained in our "standard" case and those obtained with VandenBerg lifetimes, we see that the expected variations due to different fitting approximations, opacities, and assumptions on the mixing length parameter amount to $\sim$20% at worst. Much larger is the decrease in the predicted abundances of ^{3}He obtained with Maeder & Meynet (1989) lifetimes, owing to the increased delay in the restitution rate of metals to the ISM. The effect can be seen in Figure 7, which shows the time evolution of ^{3}He. The net ^{3}He reduction is a factor 1.3 at $t = t_\odot$ and 1.8 at $t = t_0$, but the disagreement with the observed presolar value is still a factor of 4. Finally, note that the value of the metallicity at the time of the formation of the Sun is almost insensitive to the exact choice of $\tau[m]$, as indicated in column (6) of Table 7.

5.5. *Effect of Varying* w_3

As shown in Figure 1, the abundance of ^{3}He in stellar ejecta is not a very well determined quantity from stellar evolutionary models. For instance, the flattening and the possible turnover found by Vassiliadis & Wood (1993) at low masses appears to be at variance with the steep behavior obtained by Rood et al. (1976) for $m \leq 1\ M_\odot$. Nevertheless, the rough agreement of the predictions based on widely differing methods and assumptions (e.g., opacities) is reassuring on the overall consistency of independent stellar nucleosynthesis calculations.

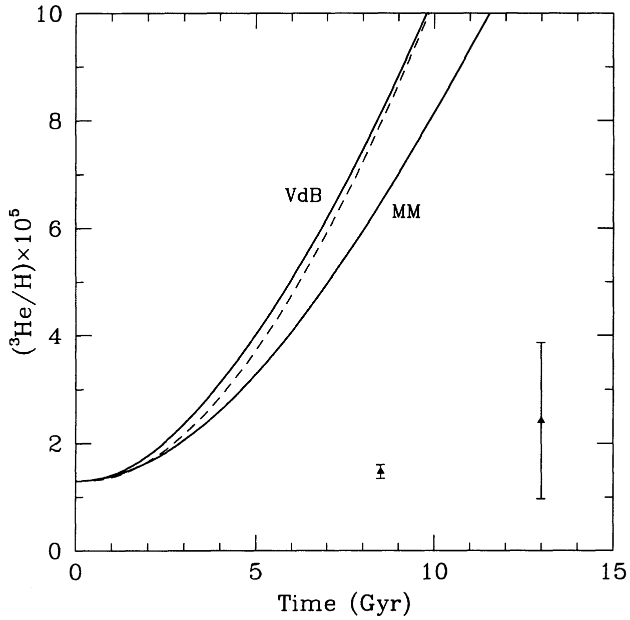

FIG. 7.—Effect of stellar lifetimes on the ^{3}He evolution. The *solid curves* are from the stellar evolution models of VandenBerg (1985, VdB) and Maeder & Meynet (1989, MM), respectively. The *dashed curve* is for the standard model ($\eta_{10} = 4$).

To test the sensitivity of our results to the ^{3}He yield we have run two models with the numerical value of $w_3(m)$ given by the expression (17) divided by a constant factor of 3 and 10, respectively. The results for the ^{3}He abundance at the time of the formation of the Sun and at the present time are given in Table 8. Even in the case of the largest reduction, the discrepancy is still high, more than a factor of 2.

On the basis of the results shown in Table 8, we can also gauge the effect of metallicity on the ^{3}He yield. Rood et al. (1976) and Vassiliadis & Wood (1993) have calculated stellar evolution models of decreasing metallicity and determined the reduction of the surface ^{3}He abundance, resulting from the shorter time necessary to reach the base of the RGB. Rood et al. (1976) found a reduction of only $\sim$30% in a $m = 0.8\ M_\odot$ star, when passing from $Z = 10^{-2}$ to $Z = 10^{-4}$. Correspondingly, Vassiliadis & Wood (1993) determined a $\sim$20% decrease of $X_3^{\rm surf}$ for a $m = 1$ star when passing from $Z = 0.016$ to $Z = 10^{-3}$. Such a weak dependence on Z has a negligible effect on the overall galactic evolution of ^{3}He. Finally, it is important to recall that after $\sim$1–2 Gyr the metallicity of the ISM in the galactic disk has already reached a fraction $\frac{1}{4} - \frac{1}{3}$ of the solar value (Twarog 1980; FMPP).

5.6. *Effect of Nova Outbursts on* ^{3}He *Evolution*

In this section we present an estimate of the fate of ^{3}He in close binary stars that ultimately lead to a nova system. Theo-

TABLE 7

EFFECT OF VARYING $\tau(m)$

Reference (1)	$X_2(t_\odot)$ (2)	$X_2(t_0)$ (3)	$X_3(t_\odot)$ (4)	$X_3(t_0)$ (5)	$Z(t_\odot)$ (6)
Standard $\tau(m)$	3.96(−5)	2.57(−5)	1.91(−4)	4.34(−4)	0.0156
VandenBerg 1985	3.70(−5)	2.05(−5)	1.80(−4)	3.42(−4)	0.0192
Maeder & Maynet 1989	3.90(−5)	2.13(−5)	1.42(−4)	2.62(−4)	0.0188

TABLE 8

EFFECT OF VARYING $w_3(m)$

Model	$X_3(t_\odot)$	$X_3(t_0)$
Standard $w_3(m)$......	1.91(−4)	4.34(−4)
$w_3/3$	9.97(−5)	2.08(−4)
$w_3/10$	7.86(−5)	1.31(−4)
$w_3 = 0$................	5.92(−5)	7.68(−5)

retical models and observations suggest that in these objects, matter transformed from the red giant secondary to the white dwarf primary accumulates in a hydrogen-rich, degenerate envelope that suffers thermonuclear runaway giving rise to the nova outburst. According to several studies, any ^{3}He present in the accreted matter is ultimately transformed into ^{7}Li and ejected back to the ISM (Arnould & Norgaard 1975; Starrfield et al. 1978; Shara 1980; Iben & Tutukov 1984; Matteucci & D'Antona 1991). For example, the calculations by Shara (1980) show that on a timescale of a few tens of years the accreted ^{3}He is completely destroyed at the base of the envelope. According to this picture, in the framework of chemical galactic evolution, novae represent a sink for ^{3}He and a source for ^{7}Li. It is necessary, therefore, to consider the reduction of the ^{3}He yield accounted for by nova systems.

As a simple estimate, let us adopt the conservative assumption that red giants belonging to a nova system do not contribute at all to ^{3}He enrichment of the ISM. Then, we should correct the ^{3}He yield of a star of mass m by the ratio $p(m)$ of the number of nova systems (with secondary of mass m) to the number of single stars of the same mass. For simplicity, let us define a "nova system" as a binary made by a secondary of mass $m_2 = m$ and a primary of mass $m_1 \geq m_2$ evolving ultimately to a white dwarf, with period of revolution ranging between P_1 and P_2. The fraction of novae is thus given by

$$p(m) = \alpha(m) \int_{\mu_1}^{\mu_2} f_\mu(\mu) d\mu \int_{P_1}^{P_2} f_P(P) dP \,, \tag{18}$$

where $\alpha(m)$ is the fraction of stars of mass m that have a companion, $f_\mu(\mu)$ is the (normalized) distribution of mass ratios $[\mu = m_2/(m_1 + m_2)]$, and $f_P(P)$ the (normalized) distribution of periods. For $f_\mu(\mu)$ we take the expression given by Greggio & Renzini (1983),

$$f_\mu(\mu) = 2^{1+\gamma}(1+\gamma)\mu^\gamma \,, \tag{19}$$

with $\gamma = 2$. The progenitors of white dwarfs are assumed to be stars of mass $0.8 \leq m_1 \leq 8$. The limiting mass ratios for a nova system are then $\mu_1 = m/(8 + m)$ and $\mu_2 = \frac{1}{2}$. For the distribution of periods, we take the expression given by Duquennoy & Mayor (1991) for MS stars, neglecting any possible subsequent change during the post-MS evolution,

$$f_P(P) = \frac{1}{\sqrt{2\pi}\sigma_P} \exp\left[-\frac{(\log P - \overline{\log P})^2}{2\sigma_P^2}\right], \tag{20}$$

where P is in days, $\overline{\log P} = 4.8$, and $\sigma_P = 2.3$. Periods of nova systems are in the range between $P_1 = 2$ hr and $P_2 = 12$ hr. As for $\alpha(m)$, following Abt (1983), we assume $\alpha = 0.5$ independently of mass.

With these values, we obtain the final expression

$$p(m) \simeq 0.5 \times \left[1 - \left(\frac{2m}{m+8}\right)^3\right] \times 0.008 \tag{21}$$

(here m in units of $M_\odot$) for $m \leq 8\ M_\odot$, and $p(m) = 0$ for $m \geq 8$ $M_\odot$. Thus, the resulting correction on the ^{3}He yield, equal to $1 - p(m)$ within our assumptions, is negligible for all masses.

5.7. *Hypothetical Resonance in the* ^{3}He(^{3}He, 2p)^{4}He *Cross Section*

Finally, we will explore the possibility that the solution to the problem of the galactic evolution of ^{3}He might come from nuclear physics. In particular, we consider the effects that a resonance at low energies ($E_r \lesssim 100$ keV) in the ^{3}He(^{3}He, 2p)^{4}He reaction might have on the stellar nucleosynthesis of ^{3}He. The existence of such a resonance was postulated independently by Fowler (1972) and Fetysov & Kopysov (1972) as a desperate solution to the solar neutrino puzzle. In fact, they argued that the resonance could significantly reduce the production of ^{7}Be and ^{8}B, and the discrepancy between the predicted and the observed neutrino fluxes might be accounted for, or at least reduced. Fetysov & Kopysov (1972, 1975) set very stringent restrictions on this hypothetical resonance: it must lie less than 20 keV above the ^{3}He + ^{3}He threshold ($\sim$11.6 MeV) and must have a total width less than $\sim$2 keV. Laboratory measurements of the ^{3}He(^{3}He, 2p)^{4}He cross section by Dwarakanath (1974) extended down to $E_r \simeq 30$ keV gave negative results. Equally unsuccessful was the search for an excited level of ^{6}Be near the ^{3}He + ^{3}He threshold (McDonald et al. 1977). In the most recent experimental study of this reaction, Krauss et al. (1987) found no evidence for a resonance down to $E_r \simeq 18$ keV.

However, the possibility of a low-energy resonance cannot be completely dismissed (Rolfs & Rodney 1988). In a recent reanalysis of the solar neutrino puzzle, Castellani, Degl'Innocenti, & Fiorentini (1993) find that a sufficient reduction of ^{7}Be neutrinos can result if the resonance lies well below $\sim$20 keV, an energy region almost unexplored experimentally. Experiments currently in progress in underground laboratories should be able to extend cross section measurements down to $E_r \simeq 10$ keV with the required sensitivity (Arpesella et al. 1991).

Clearly, the existence of a resonance in the ^{3}He(^{3}He, 2p)^{4}He cross section at energies of a few keV would have a dramatic consequence on the stellar yield of this isotope. To evaluate quantitatively this effect, we used the results obtained from a set of stellar models computed with the FRANEC code (O. Straniero, private communication), in which the rate for the ^{3}He(^{3}He, 2p)^{4}He reaction was modified to include the effect of a resonance of strength $(\omega\gamma)_r = 1.2 \times 10^{-19}$ MeV at energy $E_r = 10$ keV, and $(\omega\gamma)_r = 5.0 \times 10^{-22}$ MeV at $E_r = 3.0$ keV. The stellar models were evolved from the Hayashi track to the first dredge-up. The resulting ^{3}He abundances in the outer envelope of stars of different masses after the first dredge-up are listed in the columns (6) and (7) of Table 3, along with the corresponding values obtained in the nonresonant case (col. [5]). The time evolution of the ^{3}He abundance obtained with the stellar yields evaluated in the resonant case is shown in Figure 8 for $\eta_{10} = 4$. The results are startling: the ^{3}He abundance remains almost constant during galactic evolution and the presolar value may be easily accounted for by either values of the resonance parameters. The trend displayed by the curves of Figure 8 differs, however, in the two cases: for $E_r = 3$ keV, the efficiency of ^{3}He-burning is so enhanced to give a net destruction; conversely, a value of $E_r = 10$ keV is already adequate to produce a slight overabundance at the presolar epoch (for more details, see Galli et al. 1994).

An indirect evidence in favor of ^{3}He destruction in stars is

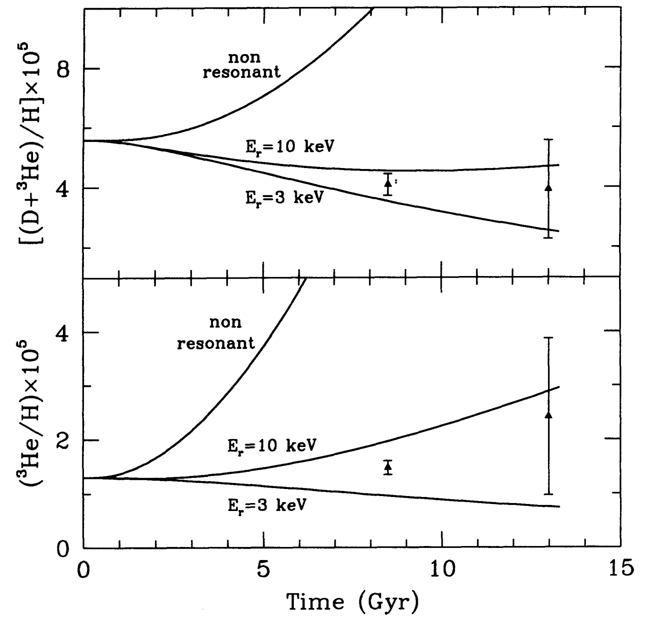

FIG. 8.—Effect of an hypothetical low-energy resonance on the evolution of ^{3}He (*bottom*) and (D + ^{3}He) (*top*). Two values of the resonance energy and strength are considered: $E_r = 3$ keV, $(\omega\gamma)_r = 5.0 \times 10^{-22}$ MeV, and $E_r = 10$ keV, $(\omega\gamma)_r = 1.2 \times 10^{-19}$ MeV. The evolution in the nonresonant case is also shown for comparison.

given by the observed galactic gradients of ^{3}He/^{4}He and ^{13}C/^{12}C. The production schemes for ^{3}He and ^{13}C as well as their distribution in stars, are similar (Rood et al. 1984); thus, one would expect that their galactic gradients show the same trend. The recent observations of Balser et al. (1994) indicate that the ratio ^{3}He/^{4}He increases in the outer Galaxy, unlike ^{13}C/^{12}C that is larger in the inner Galaxy where the star formation rate is higher than in the solar neighborhood. In conclusion, it appears that nuclear physics can provide an effective way to solve the problem of galactic ^{3}He. One should notice, however, that within the range of uncertainties the abundance of this isotope measured in the planetary nebula NGC 3242 is in agreement with the predictions of standard stellar models, even though a sensible reduction of the value given by Bania et al. (1993) cannot be excluded at the moment (T. Wilson, private communication).

6. CONCLUSIONS

In this paper, we have studied in detail the evolution of D and ^{3}He in our Galaxy. Several conclusions can be drawn from our analysis. In the case of D, since the processes affecting the abundance of this species are simple and rather well understood, models of chemical galactic evolution can be used to put stringent constraints on the amount of primordial D, and hence on the baryon-to-photon ratio η. The resulting allowed range for the baryon-to-photon ratio η results $4.2 \leq \eta_{10} \leq 4.8$, where the lower bound comes from the ISM value of D/H as determined by McCullough (1991), and the upper bound from the presolar D/H. This constraint corresponds to a baryon density parameter $\Omega_b h_{50}^2 = 0.066 \pm 0.004$ (or $\Omega_b h_{100}^2 = 0.016 \pm 0.001$) for a standard model with three neutrino flavors. Our lower bound to $\Omega_b h_{50}^2$ is $\sim$50% higher than the corresponding lower bound determined by Smith et al. (1993) and $\sim$20% higher than the one derived by Steigman & Tosi (1992) on similar grounds. This seems to reinforce the need for baryonic dark matter, since the amount of luminous matter in the universe is significantly lower, $\Omega_{\rm lum} \simeq 0.005$–$0.007$ (Persic & Salucci 1992; see, however, Bristow & Phillips 1994 for a different conclusion).

The conclusions above are based on our best model for the chemical evolution of the Milky Way, which predicts a moderate astration of D. If, indeed, the intergalactic absorption feature observed by Songaila et al. (1994) and Carswell et al. (1994) is produced by primordial D and its abundance is a factor of $\sim$15 larger than the present-day value, our model is not correct and its constitutive elements (SFR and/or IMF, occurrence of a galactic wind, etc.) must be radically revised.

Conversely, the problem of ^{3}He overproduction during normal galactic evolution, already stressed in several early works on the subject, is still with us. The results of current stellar evolutionary models, supported by the recent detection of ^{3}He in the ejecta of a planetary nebula, are irreconcilable with the abundance of this isotope inferred for the presolar material and observed in galactic H II regions. In our opinion, there are only two possible solutions to this dilemma:

1. Stellar models give correct values for the ^{3}He yield, *but* its abundance in the protosolar nebula and in H II regions has been reduced by about one order of magnitude by some unknown process;

or, conversely,

2. Observed abundances of ^{3}He are representative of the (uniform) composition of the ISM at the corresponding epochs, *but* the predicted stellar yields of ^{3}He need to be drastically reduced.

The first possibility cannot be completely dismissed: one should also keep in mind that the determination of the ^{3}He abundance in H II regions suffers various uncertainties the most serious being probably the conversion from $^3He^+/H^+$ to ^{3}He/H, which is virtually unknown (compare, e.g., the results of Rood et al. 1984 and Bania et al. 1987). Much more difficult is to accept that the ^{3}He abundance in the presolar material might be underestimated by a factor of 5–7, in face of the overall consistency of the available observational data (see Appendix A). The possibility that the Sun was born in a region locally depleted of ^{3}He by the same process that apparently reduces ^{3}He in H II regions appears more plausible (D. Schramm, private communication).

Possibility 2 is more appealing, but it seems in conflict with the recent detection of ^{3}He in the ejecta of a planetary nebula, where a value in excellent agreement with that predicted by standard stellar models has been measured. Should this value be substantially reduced by future observations, the possible existence of a mechanism able to destroy ^{3}He in stellar envelopes during the latest stages of stellar evolution would gain credit.

Of course, in this situation, ^{3}He cannot be used as a constraint on the value of η_{10}: in fact, if stellar production of ^{3}He is correctly predicted by standard stellar models, whatever initial abundance of ^{3}He in the range of a few 10^{-5} is rapidly offset already after the first 1 Gyr of galactic evolution.

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APPENDIX A

ABUNDANCES OF D, ^{3}He, AND ^{4}He AT THE TIME OF FORMATION OF THE SUN

A1. ABUNDANCE OF D

The abundance of D has been measured in the atmosphere of Jupiter and Saturn (see, e.g., Gautier 1983; Gautier & Owen 1985) where D is mainly in the form of CH_3D and HD. The more recent (after 1980) values for Jupiter are D/H = $(2.15 \pm 0.95) \times 10^{-5}$ from HD (Encrenaz & Combes 1982), D/H = $(3.6 \pm 1.4) \times 10^{-5}$ from CH_3D (Kunde et al. 1982), D/H = $(3.3 \pm 1.1) \times 10^{-5}$ from CH_3D (Courtin et al. 1984), and D/H = $(1.95 \pm 0.95) \times 10^{-5}$ from HD (Hayden Smith & Schempp 1987). The more recent accurate values for Saturn is D/H = $(1.6 \pm 1.3) \times 10^{-5}$ from CH_3D (Courtin et al. 1984).

According to Gautier & Owen (1985), the abundance of D in the atmospheres of Jupiter and Saturn should reflect the local interstellar value at the time these two planets formed, since, unlike the case of Uranus and Neptune, no substantial enrichment in D should have occurred (Hubbard & MacFarlane 1980). However, the D abundance inferred from HD lines measurments seems to be systematically lower than the value obtained from CH_3D measurements, probably owing to systematic errors in the model assumptions. In addition, there is apparently a depletion of D in Saturn with respect to Jupiter which, although small, may be a real effect reflecting an enhanced D fractionation in Jupiter (see discussion in Courtin et al. 1984).

In any case, we are confident that a weighted average of the above values should represent a sufficiently precise estimate of the D abundance in the ISM at the time of formation of the Sun, and we assume in this work $(D/H)_\odot = (2.52 \pm 0.49) \times 10^{-5}$.

A2. ABUNDANCE OF ^{3}He

The abundance of ^{3}He in the ISM at the time of formation of the Sun (4.6 Gyr ago) is usually inferred from measurements of the ^{3}He/^{4}He isotopic ratio in meteorites. It has long been known that the isotopic concentration of rare gases in meteorites falls in two significantly different ranges of values (Pepin & Signer 1965). These two kinds of elemental patterns, or components, are commonly denominated "planetary" (or "A") and "solar" (or "B"). Because of its similarity in composition with the present day solar wind (SW), the latter component, observed in carbonaceous chondrites (CC) and gas-rich meteorites (GRM) is attributed to ion implantation by the SW in the meteoritic material (Signer 1964). Component "A," which is characterized by a lower ^{3}He/^{4}He ratio and is not observed in GRM, is more puzzling. According to one hypothesis (Black 1972b), this component may result from a mixture of rare gas ions already present in the dust grains of the protosolar cloud and rare gas ions implanted by a very early protostellar SW on those grains which were not vaporized during the process of star formation. Black (1972a) attributed the increase in ^{3}He content from a SW of composition "A" to a SW of composition "B" to the intervening episode of D burning in the proto-Sun. It should be mentioned that this hypothesis raises some difficulties (see, e.g., the discussion following Black 1972c).

In an alternative scenario, the ^{3}He/^{4}He ratio in the SW may have been raised with time because of leakage from the interior of the Sun of ^{3}He brought to the surface by meridional circulation (Schatzman 1970, private communication to Eberhardt et al. 1970; Schatzman & Maeder 1981; see also Geiss & Bochsler 1981 for a discussion). But even in this case, however, the lower abundance of ^{3}He in CC may well be representative of the protosolar value.

If Black's (1972a, b) hypothesis is correct, then the ^{3}He abundance in the protosolar nebula can be obtained directly from the CC data; moreover, since a star of 1 $M_\odot$ converts all its initial D in ^{3}He during the protostellar and PMS phases (Palla & Stahler 1990, 1993), the protosolar D abundance can be determined by subtracting the value obtained in this way from the ^{3}He abundance in GRM or in the present-day SW *if* there is no leakage of internal ^{3}He. Let us consider this procedure quantitatively.

The "A" component of ^{3}He/^{4}He in CC has been measured with good accuracy several times, resulting $(1.25 \pm 0.76) \times 10^{-4}$ (Anders et al. 1970), $(1.76 \pm 0.40) \times 10^{-4}$ (from Fig. 1 of Black 1971), $(1.5 \pm 1.0) \times 10^{-4}$ (Black 1972a), $(1.40 \pm 0.30) \times 10^{-4}$ (from Fig. 6 of Black 1972b), $(1.43 \pm 0.40) \times 10^{-4}$ (Jeffery & Anders 1970), $(1.42 \pm 0.2) \times 10^{-4}$ (Reynolds et al. 1978), $(1.558 \pm 0.055) \times 10^{-4}$ (Frick & Moniot 1977), $(1.460 \pm 0.073) \times 10^{-4}$ (Eberhardt 1978). The weighted average of these values is ^{3}He/^{4}He = $(1.514 \pm 0.042) \times 10^{-4}$; with the value of $(^4He/H)_\odot$ given above, the presolar abundance of ^{3}He as estimated from the composition of CC results $(^3He/H)_\odot = (1.48 \pm 0.13) \times 10^{-5}$.

On the other hand, measured values for ^{3}He/^{4}He in GRM are ^{3}He/^{4}He = $(3.79 \pm 0.4) \times 10^{-4}$ (Jeffery & Anders 1970), $(3.9 \pm 0.3) \times 10^{-4}$ (Black 1972a). The abundance of ^{3}He in the SW has been recently reviewed by Bochsler, Geiss, & Maeder (1990). Geiss et al. (1972) have obtained in the SW an average ratio ^{3}He/^{4}He = $(4.25 \pm 0.21) \times 10^{-4}$ from the Apollo Foil experiment. The average ratio obtained from data collected by the *ISEE 3* spacecraft is ^{3}He/^{4}He = $(4.88 \pm 0.48) \times 10^{-4}$ (Coplan et al. 1984). Solar wind particles trapped in the lunar mineral ilmenite give values in the range ^{3}He/^{4}He = $(3.45 \pm 1.15) \times 10^{-4}$ (Benkert et al. 1988). The weighted average of the SW, lunar soil, and GRM data is $(^3He/^4He) = (4.16 \pm 0.16) \times 10^{-4}$, corresponding to $[(^3He + D)/H]_\odot = (4.08 \pm 0.36) \times 10^{-5}$.

A direct determination of the ^{3}He abundance in the photosphere of the Sun would provide an independent constraint on $[(^3He + D)/H]_\odot$; in practice, such a measurement is difficult and the result uncertain. Hua & Lingenfelter (1987) used the 2.223 MeV γ-line to obtain a photospheric value ^{3}He/H = $(2.3 \pm 1.2) \times 10^{-5}$. It seems that the photospheric abundance is significantly lower than the SW wind abundances, although the two values overlap within the uncertainties. Fractionation effects, or more likely, different Coulomb drag factors on ^{4}He and ^{3}He ions (see, e.g., Bochsler et al. 1990) may enhance the ^{3}He/^{4}He isotopic ratio in the SW relative to the solar photosphere. In any case, we rely here on the meteoritic value. It is interesting to notice that gas extracted from terrestrial diamonds is characterized by values of the isotopic ratio ^{3}He/^{4}He similar to the ones found in GRM and in the SW, $\sim 3 \times 10^{-4}$ (Ozima & Zashu 1983). This may indicate that the dust grains from which Earth and the GRM formed were subjected to irradiation by a SW of composition very similar to the present one, in contrast with dust incorporated in CC.

Following this discussion, the values of the abundance of the light elements at the time of formation of the Sun, with their 1 σ uncertainty, are given in equations (4)–(6). It is important to observe that, in support of Black's hypothesis, the value of $(D/H)_\odot$

obtained as a difference of the meteoritic determinations of $(^3He/H)_\odot$ and $[(D + {}^3He)/H]_\odot$ (eqs. [5] and [6]) is fully consistent with the value determined independently from planetary atmospheres (eq. [4]).

A3. ABUNDANCE OF ^{4}He

The ^{4}He content of the presolar nebula can be derived directly from measurements of the He abundance in the photosphere of the Sun or in prominences, or determined theoretically from recent standard solar models. Anders & Grevesse (1989) in their critical compilation of solar photospheric abundances prefer the latter method and determine a theoretically allowed range, $(^4He/H)_\odot = 0.098 \pm 0.008$, which we adopt here. In terms of mass fractions, the solar values adopted in this paper are $X_\odot = 0.707 \pm 0.017$, $Y_\odot = 0.274 \pm 0.016$ (Anders & Grevesse 1989).

APPENDIX B

EVOLUTION OF D AND ^{3}He: ANALYTICAL SOLUTIONS

In this Appendix we will derive simple analytical expressions for the evolution of D and ^{3}He. Similar results have appeared in various forms in the literature. We present them here in a consistent formalism for commodity of the reader to present a check on the numerical calculations, and to provide a simple, although approximate, way to evaluate the relative importance of the various physical quantities.

Consider a "closed box" model of galactic evolution for the solar neighborhood, with constant star formation rate ψ, total mass M_t and gas mass $M_g(t)$ in the form of gas, in the IRA approximation. Since the system is closed, initially $M_g(0) = M_t$. The evolution of the gas mass, of the abundance of D and ^{3}He is governed by equations analogous to the equation (11) with $f = 0$ and $i = 2, 3$. Written explicitly,

$$\frac{d}{dt}(M_g) = -(1 - R)\psi\ , \tag{22}$$

$$\frac{d}{dt}(M_g X_1) = -(1 - R_1)X_1 \psi\ , \tag{23}$$

$$\frac{d}{dt}(M_g X_2) = -X_2 \psi\ , \tag{24}$$

$$\frac{d}{dt}(M_g X_3) = -X_3 \psi + P_3 X_1 \psi + R_3\left(\frac{3}{2} X_2 + X_3\right)\psi\ , \tag{25}$$

$$\frac{d}{dt}(M_g Z) = -Z(1 - R)\psi + P_Z \psi\ , \tag{26}$$

where we have defined the returned fraction,

$$R = \int_{m_1}^{m_{\rm upp}} [1 - d(m)]\phi(m)dm\ , \tag{27}$$

and the upper limit is $m_{\rm upp} = 60\ M_\odot$; the ^{3}He returned fraction,

$$R_3 = \int_{m_1}^{m_{\rm upp}} [1 - q_3(m)]\phi(m)dm\ ; \tag{28}$$

the ^{3}He production factor,

$$P_3 = \int_{m_1}^{m_{\rm upp}} w_3(m)\phi(m)dm\ ; \tag{29}$$

and the metal production factor

$$P_Z = \int_{m_1}^{m_{\rm upp}} [q_C(m) - d(m)]\phi(m)dm\ ; \tag{30}$$

the lower limit of these integrals is the turn-off mass, taken as $m_1 = 1\ M_\odot$.

Expressed as function of the present-day gas mass fraction $\mu = M_g/M_t$, the solutions of equations (24)–(26) read

$$X_2(\mu) = X_{2p}\,\mu^{R/(1-R)}\ , \tag{31}$$

$$X_3(\mu) = X_{1p}\frac{P_3}{R - R_3}\,[1 - \mu^{(R-R_3)/(1-R)}] + \frac{3}{2}X_{2p}[\mu^{(R-R_3)/(1-R)} - \mu^{R/(1-R)}] + X_{3p}\,\mu^{(R-R_3)/(1-R)}\ , \tag{32}$$

$$Z(\mu) = -\frac{P_Z}{1 - R}\ln\mu\ , \tag{33}$$

TABLE 9

Comparison with the Simple Model at $Z = Z_\odot$

η_{10}	X_2 SM	X_2 Numerical	$X_3(w_3 = 0)$ SM	$X_3(w_3 = 0)$ Numerical	$X_3(w_3 \neq 0)$ SM	$X_3(w_3 \neq 0)$ Numerical
2......	1.2(−4)	1.3(−5)	1.5(−4)	7.3(−5)	2.8(−4)	1.4(−4)
4......	3.8(−5)	4.0(−5)	6.1(−5)	3.0(−5)	1.9(−4)	9.5(−5)
6......	1.9(−5)	2.0(−5)	4.0(−5)	1.9(−5)	1.7(−4)	8.2(−5)

where X_{2p} and X_{3p} are the primordial values and the initial metallicity of the gas is zero. The integrals (27)–(30) can be evaluated numerically from our prescriptions on the function $d(m)$, $q_3(m)$, q_C, $X_3^{\rm surf}(m)$, and the IMF. The resulting values are

$$R = 0.211 \,, \qquad R_3 = 0.180 \,, \qquad P_3 = 7.30 \times 10^{-5} \qquad P_Z = 7.94 \times 10^{-3} \,, \tag{34}$$

where we have adopted Schatzman's (1987) expression for $X_3^{\rm surf}$ (case with no mixing) to evaluate P_3. With a simpler notation, defining

$$f_2 = \mu^{R/(1-R)} \,, \qquad f_3 = \mu^{(R-R_3)/(1-R)} \,, \qquad K_3 = \frac{P_3}{R - R_3}\left[1 - \mu^{(R-R_3)/(1-R)}\right] , \tag{35}$$

equations (31) and (32) become

$$X_2 = f_2 X_{2p} \,, \tag{36}$$

$$X_3 = (3/2)X_{2p}(f_3 - f_2) + f_3 X_{3p} + K_3 X_{1p} \,, \tag{37}$$

already obtained on the basis of qualitative arguments by Black (1971) and derived rigorously by Tinsley (1974). From these two equations, with $K_3 = 0$, we can easily get

$$\frac{(3/2)X_2(\mu) + X_3(\mu)}{(3/2)X_{2p} + X_{3p}} = \left[\frac{X_2(\mu)}{X_{2p}}\right]^{1-g_3} , \tag{38}$$

where $g_3 = R_3/R$, adopted by Olive et al. (1990) to provide an upper limit on the primordial D + ^{3}He abundance and hence a lower bound on η.

It is convenient to eliminate the poorly constrained variable μ from equations (31) and (32), using the $\mu - Z$ relation, equation (33). Then the results of the analytical calculations above can be compared with the numerical results at any value of Z. This is shown in Table 9 for $Z = Z_\odot$ in the case of no stellar production of ^{3}He ($w_3[m] = 0$) and with stellar production of ^{3}He ("standard"). The analytical results approximate very well the abundance of D computed numerically (within 10%); however, the predicted abundances of ^{3}He results generally lower by a factor ~ 2 than the corresponding numerical values, clearly showing the inadequacy of the IRA approximation in this case.

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